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The effects of the star-disk interaction for the evolution of a
protoplanetary system

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Abstract

To understand the formation of planets and life itself, the evolution of protoplanetary disks, the cradle of young planets, has to be understood. A vast number of observations of protoplanetary disks and planetary systems raise important questions that can not be answered with current theoretical models. To gain additional insights into the evolution of young star-disk systems, different numerical methods are coupled in a self-consistent way allowing a combined treatment of the stellar and disk components. One particular model is used for each component of the star-disk system. The stellar component is modeled with a sophisticated long-term stellar evolution model. To simulate the inner disk regions and the results of the star-disk interaction, an implicit, one-dimensional disk model is used that includes important physical mechanisms such as the influence of the stellar magnetic field on the disk. Properties of the outer disk regions are based on a detailed, multi-dimensional model. The coupled star-disk model indicates that different stellar rotation periods or metallicities can considerably influence the disk's temperature structure, dynamic evolution, and lifetime. At the same time, the stellar spin evolution and distribution of rotation periods are influenced by the disk's accretion rate and evolution. The inner disk region plays a key role in this model. The magnetic star-disk interaction, which is strongest close to the star in the inner ~ 0.1 AU, controls how much mass and angular momentum is exchanged between the star and the disk. These results indicate that a coupled treatment of the stellar and disk evolution is important in understanding these complex star-disk systems. Further advancing the combined star-disk model, especially of the innermost disk region, will provide additional insights into previously unexplored processes.

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Part I.

Introduction

1. Motivation and context

The question of how our planet and life itself have formed is as old as mankind. Already in the 18th century, the basic idea that a star is formed in a collapsing cloud or nebula of gas and dust has been proposed by philosophers like Laplace or Kant. Based on this nebular hypothesis, planets were believed to be formed in the leftovers of the star-formation process. While this hypothesis is not able to explain all observational aspects, for example, the distribution of angular momentum in the solar system, it has been pursued until the 20th century. The advancement of observational capabilities drove the search for such nebular structures around young stars. The first indications of material that is located around a young star have been provided by spectroscopic and infrared studies [e.g., Joy, 1945, Herbig, 1960, Mendoza V., 1966]. A few years later, the idea of an accretion disk was pushed forward by the work of Safronov [1972]. The properties of gas and dust, as well as theories for forming planets, have been presented. The first direct observational indication of a disk surrounding a young star was provided in the following decade [e.g., Sargent and Beckwith, 1987, Strom et al., 1989]. With direct imaging by the *Hubble Space Telescope* (HST), the existence of protoplanetary disks has been indubitably proven [e.g., O'Dell and Wen, 1992, McCaughrean and O'Dell, 1996].

With the discovery of the first exoplanet [Mayor and Queloz, 1995] and exoplanets embedded in gaps of protoplanetary disks [e.g., Müller et al., 2018] the circle between accretion disks and the formation of planets closes. Today, thousands of exoplanets show a vast variety of planetary systems [e.g., He et al., 2019]. Among these systems are a growing number of rocky, Earth-like planets in the stellar habitable zone [e.g., Gillon et al., 2016]. Motivated by these discoveries, the search for the building blocks of life has been pushed forward [e.g., Seager, 2014]. The latest generation of space-based telescopes, like the *James Webb Space Telescope* (JWST), promises

to provide new and groundbreaking answers to the question of these signatures of life can be detected in exoplanets [e.g., Meadows et al., 2019].

To understand how these planetary systems can emerge, the evolution of the young star and its protoplanetary disk has to be understood. Theoretical stellar or disk evolution models are compared to observational data to gather new insights into the physical processes defining the young star disk system. This thesis highlights the importance of the interaction between the young star and its accretion disk by, self-consistently, combining stellar and disk evolution models. Such a combined model can provide valuable insights into explaining observations of young star-disk systems, in which new planets are formed.

1.1. Basic picture

The formation of stars and their disks starts in cold interstellar clouds. During the gravitational collapse of a rotating cloud, a rotationally supported disk containing gas and dust is formed around the new protostar due to the conservation of angular momentum (AM). The timescale on which the young star-disk system forms is relatively short, $\sim 10^4$ years [e.g., Yorke et al., 1993, Hueso and Guillot, 2005]. Once the PPD has formed, the infalling material of the parent cloud is assumed to fall on the disk [e.g., Vorobyov and Basu, 2005, Hueso and Guillot, 2005, Visser and Dullemond, 2010, Drążkowska and Dullemond, 2018, Schib et al., 2021] and the mass growth of the star depends on how fast and efficiently the disk material can be transported in the disk toward the star. With the number of detected exoplanets rising rapidly over nearly three decades (over 5300 exoplanets have been detected to date¹), it stands to reason that accretion disks are at least common around young stars. Indeed, ever-improving observation methods from the first direct images of PPDs with the *Hubble Space Telescope* [HST, O’Dell and Wen, 1992, McCaughrean and O’Dell, 1996] to recent high-resolution images of gas and dust structures by the *Atacama Large Millimeter/sub-millimeter Array* [ALMA; see the review of Andrews, 2020] provide growing number of the ubiquity of accretion disks around young stellar objects. Besides infall, other essential processes occur in PPDs from the disk formation until the end of their

¹see exoplanets.eu

lifetime. Fig. 1.1 summarizes some of these critical mechanisms. Note that this is an idealized cartoon and some of the processes shown can occur at different evolutionary stages.

In the schematic picture, the young star with a stellar radius $R_\star$ rotates with angular velocity $\Omega_\star$. In the protoplanetary disk (PPD), the material is transported from the outer disk regions toward the star (symbolized by dashed black lines). The accretion disk is connected to the young star via magnetic field lines that originate from the stellar surface and couple to ionized disk material. Different mechanisms can lead to ionization in PPDs [e.g., Vorobyov et al., 2020b], such as internal high energy UV/X-Ray radiation, external UV radiation, Cosmic rays (CR), or thermal ionization. While the innermost and outer disk regions, as well as the upper disk layers, are assumed to be partially ionized with ionization fractions of $\sim 10^{-10} - 10^{-7}$ [e.g., Vorobyov et al., 2020b], there is also a region that is shielded from high energy irradiation and too cold for effective thermal ionization. In this dead zone [DZ, e.g., Gammie, 1996a], the disk material is only weakly ionized [with ionization fractions $\lesssim 10^{-13}$, e.g., Vorobyov et al., 2020b]. In Fig. 1.1, the ionized regions and the dead zone are colored in green and grey, respectively.

Disk material from the innermost, ionized disk region is channeled along the field lines and accretes on the star in so-called 'funnel flows' [FF, colored in light green, e.g., Romanova et al., 2002, Bouvier et al., 2007, Bessolaz et al., 2008, Kurosawa and Romanova, 2013, Hartmann et al., 2016]. Accretion rates range from $\sim 10^{-5} M_\odot \text{ yr}^{-1}$ [shortly after the creation of the accretion disk or during phases of strong episodic accretion events, e.g., Kley and Lin, 1999, Vorobyov et al., 2017b] to $\lesssim 10^{-10} M_\odot \text{ yr}^{-1}$ at later evolutionary stages [e.g., Testi et al., 2022, and references therein]. Besides accretion flows onto the star, the disk material is also ejected in the form of jets, outflows, and winds (see Sec. 1.6). Magnetized disk winds (DW) and thermal winds due to photo-evaporation (PE) can transport material and angular momentum (in the case of DW) away from the disk. In addition, some outflows originate at the stellar surface. In this context, these stellar winds (SW) are either thermally driven or powered by the accretion process itself (see Sec. 1.9).

In the radial direction, PPDs extend over several orders of magnitude. The inner boundary of the accretion disk can be defined in different ways [e.g., Bessolaz et al., 2008]. Usually, the truncation radius, r_t , which is defined by the balance between the stellar magnetic pressure and

the ram pressure of the accreting disk material, is referred to as the inner disk boundary [e.g., Romanova et al., 2002, Bessolaz et al., 2008, Hartmann et al., 2016]. Typical estimates for the truncation radius are of the order of 10^{-2} AU [e.g., Hartmann et al., 2016]. In Sec. 1.8, a more detailed description of the physical processes shaping the innermost disk region is given. Another characteristic radius of the accretion disk is the co-rotation radius, r_{co} . At this radius, the angular velocity of the star corresponds to the angular velocity of the disk material. For typical stellar rotation periods of 1-10 days [e.g., Herbst et al., 2002, Rebull et al., 2016, 2020, Serna et al., 2021], the co-rotation radius is located at $\lesssim 0.1$ AU. The difference between the co-rotation radius and the truncation radius, for example, can influence the stellar spin evolution (see Sec. 1.9). The dead zone in the disk extends from $r_{\text{DZ,in}} \sim 0.5$ AU [e.g., Gammie, 1996a, Bai and Stone, 2013] to $r_{\text{DZ,out}} \lesssim 10$ AU [e.g., Bai and Stone, 2013, Vorobyov et al., 2020a, Kadam et al., 2021]. Outside the dead zone, the disk material extends to outer radii of $r_{\text{out}} \gtrsim 10^{1-2}$ AU [e.g., Tripathi et al., 2017, Andrews et al., 2018, Ansdell et al., 2018b, Andrews, 2020]. The outer disk radius marks the size of the PPD and can vary, for example, for different evolutionary stages of the star-disk system (see Sec. 1.3).

1.2. Classification of star-disk systems

Young star-disk systems are usually classified based on their observational characteristics. The spectral energy distributions (SED) of star-disk systems show an excess in the mid-infrared compared to the stellar black-body radiation [Lada and Wilking, 1984, Lada, 1987]. This excess has been used to classify PPDs into three classes (I, II, and III) based on the slope α_{SED} of the excess between wavelengths of ~ 2 and $25 \mu\text{m}$

$$\alpha_{\text{SED}} = \frac{\partial \log \nu F_\nu}{\partial \log \nu} = \frac{\partial \log \lambda F_\lambda}{\partial \log \lambda}, \quad (1.1)$$

with the measured flux per frequency ν , F_ν , and the flux per wavelength λ , F_λ , respectively [e.g., Lada, 1987, Williams and Cieza, 2011]. A Class I object is characterized by a slope of $\alpha_{\text{SED}} > 0.3$ and can be described as a deeply embedded star that is optically obscured [e.g., Williams and Cieza, 2011]. The mass of the disk is considered to be comparable to the envelope

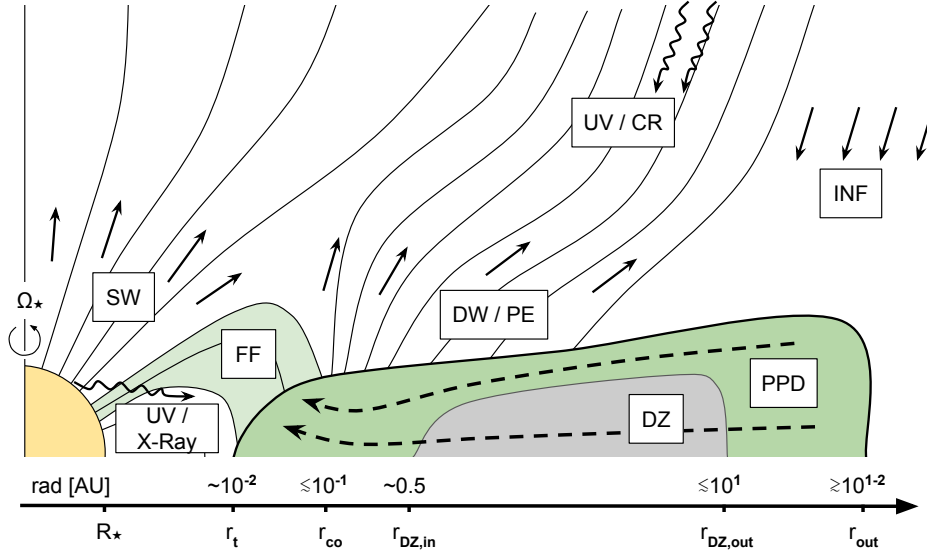


Figure 1.1.: Sketch of a young star-disk system in the radial direction. The star (rotating at $\Omega_\star$), accretion funnel flow (FF), proto-planetary disk (PPD), and dead zone (DZ) are shown in yellow, light green, green, and grey, respectively. Outflows (stellar winds, SW, disk winds, DW, photo-evaporation, PE) and infall from the parent cloud (INF) are marked as black arrows. Internal high-energy radiation (ultraviolet, UV, and X-rays), external UV, and cosmic rays (CR) are symbolized with serpent lines. Magnetic field lines are shown as black lines. In the radial direction, several characteristic radii are shown: the stellar radius, $R_\star$, the truncation radius, r_t , the co-rotation radius, r_{co} , and the beginning and end of the dead zone, $r_{\text{DZ,in}}$ and $r_{\text{DZ,out}}$, respectively, and the outer disk radius, r_{out} .

mass and smaller compared to the star. The SED slope of a Class II object ranges between $-1.6 < \alpha_{\text{SED}} < 0.3$ and is connected to an optically revealed accreting disk showing H_α and UV radiation [e.g., Williams and Cieza, 2011]. The envelope has been dissolved and the disk mass contains approximately 1 % of the stellar mass. Even further evolved systems are characterized as Class III with $\alpha_{\text{SED}} < -1.6$ and no or only weak accretion and disk SED features. The disk mass in this stage is very small and contains only $\ll 1$ % of the stellar mass. In addition to these three classes, due to better observational methods, Andre et al. [1993] introduced Class 0 objects. These systems are younger than Class I objects and show no optical or near/mid-infrared emission [e.g., Williams and Cieza, 2011]. Although the star has already formed, the envelope is still assumed to contain the majority of the mass of the Class 0 system. Finally, an intermediate class between Class I and Class II has been introduced by Greene et al. [1994]. The so-called flat-spectrum sources (FS) are characterized by a flat SED slope of $-0.3 < \alpha_{\text{SED}} < 0.3$. FS sources are already showing accretion features (H_α or UV emission) and still contain a weak envelope [e.g., Anderson et al., 2022]. A schematic overview of the different classes is given in Fig. 1.2.

In addition to the classification of a young star-disk system based on the SED profile, an accretion-based classification system is used for later, optically visible (Class II and older) sources [e.g., Kenyon and Hartmann, 1990, Williams and Cieza, 2011]. Classical T Tauri stars (CTTS) show X ray, UV, and H_α emissions [e.g., Edwards et al., 1994, Muzerolle et al., 1998, Williams and Cieza, 2011, Hartmann et al., 2016]. These emissions are linked to the disk material that accretes onto the stellar surface at near free-fall velocity resulting in an accretion shock near the stellar surface [e.g., Hartmann et al., 2016]. The shock heats the gas to $\sim 10^6$ K, causing the observed emission features. In contrast, weak-lined T Tauri stars (WTTS) show no or only weak emissions features [e.g., Kenyon and Hartmann, 1990, Hartmann et al., 2016]. This difference is either connected to different stellar ages and evolutionary stages of the star-disk system [e.g., Ingleby et al., 2014], or to the presence of an accretion disk assuming similar stellar ages [e.g., Hartmann et al., 2016].

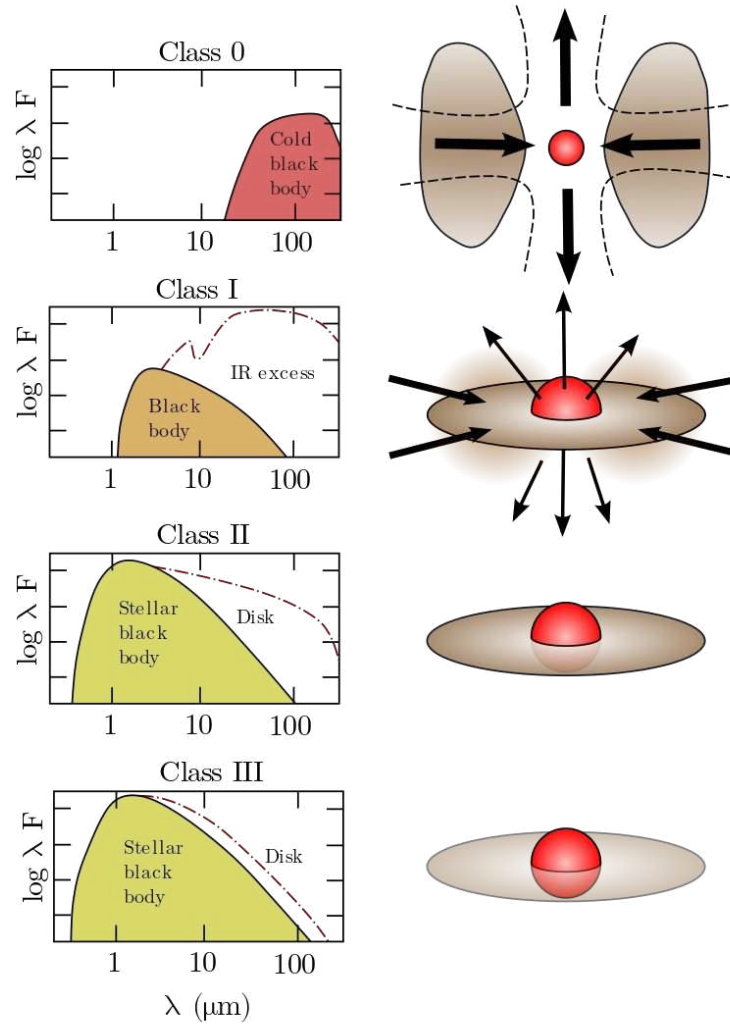


Figure 1.2.: Schematic overview of different evolutionary stages of protoplanetary star-disk systems (Classes 0-III). For each class, an exemplary spectral energy distribution (left) and a sketch (right) highlighting the star (red), the circumstellar material (brown), and the material in and outflows (black arrows) are given. This figure is adapted from Wikimedia Commons https://commons.wikimedia.org/wiki/File:Evolution_of_young_stellar_objects.svg produced by User:Vallastro and published under Creative Commons Attribution-Share Alike 4.0 International license (<https://creativecommons.org/licenses/by-sa/4.0/deed.en>).

1.3. Typical disk scales

Disk radii

An important observational benchmark for theories describing the evolution of the star-disk system is the radial extent of the accretion disk [e.g., Andrews, 2020]. The most abundant ingredient of a PPD (molecular hydrogen) is, unfortunately, almost invisible to observational techniques [e.g., Hoadley et al., 2017]. Thus, different tracers can be used to infer disk radii from measurements: stellar light scattered on small dust particles at micron-wavelengths [e.g., van Boekel et al., 2017], continuum emissions of solids in the disk at wavelengths of $\lambda \sim 1 \mu\text{m} - 1 \text{ cm}$, for example, pebble-sized emission continuum at $\lambda = 0.9 \text{ mm}$ [e.g., Andrews et al., 2016], the $\lambda = 1.33 \text{ mm}$ continuum [e.g., Ansdell et al., 2018b], or CO spectral line emissions for molecular gas [e.g., Huang et al., 2018, Ansdell et al., 2018b]. Based on the different tracers, the disk radius is usually inferred by an effective radius that encloses 90% to 95% of the respective flux [e.g., Tripathi et al., 2017, Andrews et al., 2018, Ansdell et al., 2018b, Andrews, 2020]. Over the last decade, hundreds of disk radii have been calculated based on the continuum emission of solids in the millimeter range, ranging from 10 to 500 AU [see the review of Andrews, 2020, and references therein]. There are fewer PPD radii inferred by the measurement of CO line emissions ranging between 100 to 500 AU [e.g., Ansdell et al., 2018b, Long et al., 2022]. The emissions of smaller disks are probably below the current detection limit [Andrews, 2020]. Disk radii derived from molecular tracers are systematically larger than those derived from dust emission [e.g., Testi et al., 2014, Andrews, 2015, Birnstiel et al., 2016]. One possible explanation for this difference is an efficient inward drift of dust particles within the disk.

Comparing observed disk radii from objects of different ages provides information on the development and evolution of the disk. During the early evolutionary stages of the star-disk system, the disk radii decrease from Class 0 to flat spectrum sources [e.g., Tobin et al., 2020, Sheehan et al., 2022]. One possible explanation for this observed trend is the existence of strong stellar outflows preventing high angular material from being added to the disk, which would lead to an increase in disk radius. Compared to Class II systems, disk radii of Class 0 to flat spectrum sources can be smaller, similar, or larger [e.g., Najita and Bergin, 2018, Tobin et al., 2020]. This

observation reveals information on how angular momentum might be transported in or removed from a Class II PPD [e.g., Ansdell et al., 2018b, Najita and Bergin, 2018, Andrews, 2020]. While the transport of angular momentum is dominated by gravitational torques during the early disk phases (Class 0 and I), angular momentum in disks is assumed to be primarily transported or removed by either turbulence-induced viscosity or magnetized winds during later evolutionary stages (Class II, see Sec. 1.6). In the case of viscosity, the outer disk region spreads out with time transporting angular momentum outward, while facilitating mass transport toward the central stellar object [e.g., Armitage, 2019]. This behavior is generally consistent with the trend of larger disk radii at later evolutionary stages. However, viscous spreading over-predicts the disk radii for older clusters [e.g., Trapman et al., 2020] and does not explain the smaller Class II objects. The radii of models including magnetized disk winds to remove angular momentum from the PPD remain approximately constant or decrease over time [e.g., Trapman et al., 2022].

The estimates and trends given above are based on the current accuracy of available observational data. The uncertainty on the radius based on the dust component ranges between 1% and 25% [e.g., Piétu et al., 2014, Tripathi et al., 2017]. In addition, the size of Class I disks could be systematically underestimated in the literature and additional models and observational data are required to draw further conclusions [e.g., Harsono et al., 2015, Najita and Bergin, 2018]. Furthermore, different techniques are used to estimate the disk radius in different clusters, impeding the comparability of different studies [e.g., Tobin et al., 2020].

Disk masses

A precise knowledge of the mass of PPD is necessary to constrain the reservoir of material that may form a planetary system around a young star. Calculations of the disk masses usually rely on the measured flux (similar to the disk radius). As continuum observations show the dusty disk component is less expensive compared to measurements of the gas component, most calculations of the mass of PPDs are based on the mass of the dust component. In the optically thin limit, the mass of the dust component scales as

$$M_d = \frac{d^2 F_\nu}{\kappa_\nu B_\nu(T_d)}, \quad (1.2)$$

with the distance to the object, d , the dust temperature, T_d , and the opacity and Planck function at wavelength ν , κ_ν and B_ν , respectively [e.g., Williams and Cieza, 2011, Tobin et al., 2020, van Terwisga et al., 2020, Anderson et al., 2022]. It is important to note that these mass estimates have certain caveats. As the disk is assumed to be optically thin, which is a valid assumption for disk radii $\gtrsim 10$ AU [e.g., Davis, 2005], the contributions of the inner 10 AU are not accounted for correctly. Furthermore, free-free emissions from jets, ionized winds, or external sources can contaminate the dust continuum [e.g., Anglada et al., 1998, Natta et al., 2004, Mann et al., 2015, van Terwisga et al., 2020, Tobin et al., 2020]. The dust mass does also depend on the assumed dust temperature, which usually ranges between $T_d = 20$ K and 40 K [e.g., Tobin et al., 2020, van Terwisga et al., 2020, Anderson et al., 2022] and is assumed to be constant throughout the disk to estimate the dust mass with Eq. 1.2. With the dust mass derived from the continuum flux, the disk mass is obtained by scaling up the dust mass by an assumed gas-to-dust fraction that is usually 100 [e.g., Ansdell et al., 2016]. This assumption, however, is idealized as numerical disk models show a radial temperature dependence over the whole disk radius [also beyond 10 AU, e.g., Vorobyov et al., 2020a], a dependence on the metallicity [e.g., Gehrig et al., 2023] and the irradiation by external sources [e.g., van Terwisga et al., 2020, Roquette et al., 2021]. High dust temperatures result in lower dust masses [e.g., van Terwisga et al., 2020]. In addition, short-term episodic accretion outbursts can change the temperature structure of the disk significantly, including the outer regions [e.g., Vorobyov et al., 2020a]. These uncertainties can lead to disk masses that are underestimated by up to one order of magnitude [Dunham et al., 2014].

With dust emissions being the most common observable to determine the mass of PPDs, there are also disk mass estimates based on gas observations of HD [hydrogen deuteride, e.g., McClure et al., 2016, Trapman et al., 2017] and CO lines [e.g., Ansdell et al., 2016, Long et al., 2017, Ansdell et al., 2018a]. Here, the gas mass of the disk is inferred by comparing the observed line luminosity with published disk models. The results of disk masses inferred by gas lines are usually significantly lower compared to the dust components (approximately an order of magnitude). Possible reasons for this distinction are that there are additional chemical reaction networks, which are not included in the disk models, that lower the abundance of the observed molecule [e.g., Andrews, 2020], optically thick gas lines that are hidden in the continuum [e.g.,

Trapman et al., 2017, Booth et al., 2019], or the inclusion of C and O in solids like grains, ices, or organics [e.g., Reboussin et al., 2015, Yu et al., 2016, Bosman et al., 2018].

During the evolution of PPDs from Class 0 to Class III, the disk mass decreases as several recent studies report [e.g., van Terwisga et al., 2020, Tobin et al., 2020, Lovell et al., 2021, Michel et al., 2021, Anderson et al., 2022]. The disk material in the most massive Class 0 objects is accreted onto the star, ejected as outflows or winds, or forms planets resulting in a decreasing disk mass with time. Reported median dust masses for Class 0, Class I, and flat-spectrum systems in Orion are $25.9 M_{\oplus}$, $14.9 M_{\oplus}$, and $11.6 M_{\oplus}$, respectively [Tobin et al., 2020]. Maximum values for the different classes are $1000 M_{\oplus}$, $320 M_{\oplus}$, and $280 M_{\oplus}$, respectively. A similar trend but higher dust masses by a factor of $\sim 2 - 3$ are reported by Anderson et al. [2022]. In the Serpens region, Class I, flat-spectrum systems, and Class II objects have median dust masses of $57 M_{\oplus}$, $57 M_{\oplus}$, and $48 M_{\oplus}$, respectively. For Class II objects, dust masses are reported for several young clusters at different ages. In general, the dust mass decreases with increasing age [e.g., van Terwisga et al., 2020, and references therein]. Proceeding further in the evolution of the star-disk system, Class III disks have significantly smaller dust masses compared to earlier stages. Dust masses range between $0.5 M_{\oplus}$ [Michel et al., 2021] to $0.03 M_{\oplus}$ [Lovell et al., 2021]. Assuming the usual gas-to-dust ratio of 100 [e.g., Ansdell et al., 2016], PPD masses range between $\gtrsim 0.3 M_{\odot}$ and $\sim 10^{-4} M_{\odot}$ from Class 0 to Class III.

Disk lifetimes

The material in protoplanetary disks forms the initial building blocks for planet formation and the period this material is maintained in the accretion disk marks an essential benchmark for planet evolution models. The lifetimes of PPDs are usually determined by comparing the fraction of disk-bearing stars to stars without disk signatures [the disk-fraction, e.g., Haisch et al., 2001, Ribas et al., 2014, Michel et al., 2021, Luhman, 2022]. The distinction between disk-bearing and disk-less stars is made by comparing the slope of the SED (see Sec. 1.2). While Class 0, I, FS, and II objects are considered to bear disks, the disks of Class III objects are assumed to be mostly dissolved [e.g., Michel et al., 2021]. An alternative method to determine the disk lifetimes is to determine the fraction of accreting stars [e.g., Fedele et al., 2010]. Disk fractions, f_{frac} , of clusters

with different ages are usually fitted by an exponential decay of the form

$$f_{\text{frac}} = f_0 \exp\left\{-\frac{t}{\tau_{\text{frac}}}\right\}, \quad (1.3)$$

with the age of the cluster, t , the initial disk fraction at an age of $t = 0$ Myr, f_0 , and the characteristic timescale, τ_{frac} [e.g., Mamajek, 2009].

The inferred disk lifetimes span over a large range from ~ 1 Myr to ~ 10 Myr [e.g., Pfalzner et al., 2022]. These differences can have different reasons. First, the density of young clusters can affect the disk lifetime. In dense clusters, external effects are assumed to disperse the disks faster, compared to sparse clusters [e.g., Stolte et al., 2015, Concha-Ramírez et al., 2021]. The value of τ_{frac} for dense clusters is 0.77 Myr and for sparse clusters 5.9 Myr [Pfalzner et al., 2022]. Another factor changing f_{frac} is the cluster’s distance [Pfalzner et al., 2022]. In clusters at large distances, smaller stars tend to fall below the detection limit and brighter stars dominate the sample. As these stars tend to lose their disks earlier compared to low-mass stars [e.g., Ribas et al., 2015], the resulting disk fraction is biased. For clusters with distances of $\gtrsim 1000$ pc, the characteristic timescale for disk dispersal is short, $\tau_{\text{frac}} = 1$ Myr [e.g., Richert et al., 2018]. Clusters closer to us, with distances of < 200 pc, have significantly larger characteristic timescales of $\tau_{\text{frac}} = 7 - 8$ Myr [Pfalzner et al., 2022]. In addition, the disk lifetime depends on the stellar evolution model that is used to determine the age of the stellar cluster. One stellar parameter that can significantly affect the stellar age is the magnetic field strength [e.g., Richert et al., 2018]. Including magnetic effects in stellar evolution models results in stars that are older compared to non-magnetic models. The corresponding disk lifetimes vary by a factor of ~ 3 between ~ 1 Myr for non-magnetic models and 3.5 Myr for models including magnetic effects.

Summarizing the different parameters, disk lifetimes of low-mass stars (5–10 Myr) are possibly longer than still recently assumed [Pfalzner et al., 2022]. These longer timescales give planets sufficient time to grow and reduce constraints on current planet evolution models [e.g., Mordasini et al., 2015]. The disk lifetimes for higher-mass stars ($> 1 M_{\odot}$) are shorter (4–5 Myr) and planets around these stars have to form considerably faster [Pfalzner et al., 2022]. A theoretical model, combining stellar and disk evolution, could provide further insights into the evolution of the disk

fraction with cluster age.

1.4. Structure of protoplanetary disks

With better observational capacities and a growing number of resolved PPDs, an average structure of the disk can be estimated by spatially resolved measurements of the disk's density and temperature [e.g., Andrews, 2020]. A resolved disk structure can provide insights and constraints on different aspects of the disk's evolution such as angular momentum transport [e.g., Hartmann et al., 1998], winds [e.g., Bai and Stone, 2013], and planet formation [e.g., Miguel et al., 2011, Baruteau et al., 2014].

Density structure

The density structure of PPDs, in the form of a radially dependent surface density $\Sigma(r)$, has been intensively studied in the outer, optically thin disk regions at radii larger than ~ 10 AU [e.g., Andrews and Williams, 2007, Isella et al., 2009, Piétu et al., 2014, Tazzari et al., 2017, Andrews, 2020]. Similar to estimating the masses and radii of PPDs, continuum observations (for the density of solids) and spectral line observations of CO (for gas surface densities) are used to infer the radial structure. Early estimates assume a single power-law relation $\Sigma(r) \propto r^{-p}$ with an exponent ranging between $p = 0$ and $p = 1.5$ [e.g., Mundy et al., 1996, Wilner et al., 2000, Andrews and Williams, 2007]. The value of $p = 1.5$ corresponds to the minimum mass solar nebular model [MMSN, Weidenschilling, 1977]. More recent profiles show a shallow inner, power-law profile with $p \sim 1$ inside a characteristic radius r_C [e.g., Isella et al., 2009, Andrews, 2020]. Estimates on the characteristic radius r_C range from $\sim 20 - 200$ AU [e.g., Hughes et al., 2008, Andrews et al., 2009]. Outside r_C , the profile steepens and can be estimated by a power-law factor of $p = 3$ [e.g., Andrews, 2020] or an exponential decay with $\Sigma(r) \propto \exp\{-(r/r_C)^q\}$ [e.g., Hughes et al., 2008, Williams and Cieza, 2011]. The exponent q ranges between 1.1 and 2.8 [e.g., Hughes et al., 2008, Isella et al., 2009]. Assuming a viscous accretion disk [e.g., Lynden-Bell and Pringle, 1974, Hartmann et al., 1998], an exponentially tapered power-law profile can be

predicted

$$\Sigma(r) = (2 - \gamma) \frac{M_d}{2\pi r_c^2} \left(\frac{r}{r_c} \right)^{-\gamma} \exp \left\{ - \left(\frac{r}{r_c} \right)^{2-\gamma} \right\}, \quad (1.4)$$

where γ describes the relation between the viscosity, ν , and radius, $\nu \propto r^\gamma$. A comparison between this theoretical relation and observations of Class II objects results in a good agreement with surface densities estimated by the MMSN of $\Sigma \sim 10 - 100 \text{ g/cm}^2$ at 20 AU [e.g., Andrews et al., 2009].

The vertical density structure can be estimated by assuming a disk in hydrostatic equilibrium, rotating at a Keplerian velocity, Ω_K , where the vertical component of the stellar gravity is balanced by the disk's pressure [e.g., Armitage, 2019],

$$\rho(z) = \frac{\Sigma}{\sqrt{2\pi} H_P} \exp \left\{ - \frac{z^2}{2H_P^2} \right\}, \quad (1.5)$$

with the pressure scale height, $H_P = c_S / \Omega_K$ and the isothermal speed of sound, c_S . To explain SED data of T Tauri stars, Kenyon and Hartmann [1987], suggest a flaring disk structure, meaning an increase of the scale height with radius, $H_P \propto r^b$, and $b > 1$. Direct indications of such a flaring disk are given by *Hubble* data [e.g., Burrows et al., 1996, Smith et al., 2005]. Analytic and numerical approaches to estimate the vertical disk structure find $b \approx 1.3 - 1.5$ [e.g., Chiang and Goldreich, 1997, D'Alessio et al., 1998, Dullemond et al., 2002].

Temperature structure

The vertical structure of the PPD depends directly on the disk's temperature, T_d ($H_P \propto c_S \propto T_d^{1/2}$). Thus, precise models for estimating the disk's temperature are required to analyze the disk's structure. Many factors can influence the disk's temperature structure such as stellar irradiation, the vertical disk structure [e.g., Dullemond et al., 2002, D'Alessio et al., 2006], viscous dissipation [e.g., D'Alessio et al., 1998], density waves [e.g., Rafikov, 2016], radioactive heating [e.g., Cleeves et al., 2013], external irradiation [e.g., Natta, 1993, Roquette et al., 2021], shadowing effects [e.g., Dullemond and Dominik, 2004], and chemical reactions [e.g., Woitke et al., 2016]. These mechanisms affect each other and are extremely difficult to combine in one self-consistent

model. Thus, based on different assumptions, there are simplified analytical radial temperature profiles for PPDs. Their derivation is explained in detail in, for example, Armitage [2019]. More complex systems are computed numerically. One example of a numerical disk model that includes several heating and cooling mechanisms is the ProDiMo (PROtoplanetary Disk MOdel) code [e.g., Woitke et al., 2009, 2016, 2022]. The temperature structure is calculated in the radial and vertical directions using various physical, chemical, and radiative processes. In addition, the feedback of the different processes on each other is taken into account. Other examples of a numerical calculation of the temperature of the disk can be found in Vorobyov and Pavlyuchenkov [2017] and Steiner et al. [2021]. Thin, vertically integrated disk models are improved by a reconstruction of the vertical disk structure. A vertical temperature structure can be derived by balancing the heating and cooling rates from the disk’s midplane to its surface. Depending on the strength of each process, the vertical temperature profile is influenced.

Observationally, radial and vertical temperature profiles can be inferred by highly resolved CO spectral lines [e.g., Fedele et al., 2016, Law et al., 2022] or the spatial distribution of optically thick emission lines [e.g., Weaver et al., 2018]. In addition, the location of ice lines can be used to roughly constrain the used method [e.g., Qi et al., 2019]. Recent results show radial temperature profiles that scale as $T_d \propto r^{-(0.4-0.6)}$ [Law et al., 2022]. In the vertical direction, the temperature profile is characterized by an increase from the midplane toward the disk surface due to irradiation by the star and external sources [e.g., Woitke et al., 2009, Weaver et al., 2018, Law et al., 2022].

The inner disk

Estimating the density and temperature structures of the inner 5 – 10 AU becomes increasingly difficult. The achievable resolution and the dominant stellar luminosity complicate the gathering of highly resolved data. In addition, the disk material becomes increasingly optically thick in the inner regions and emissions might be blocked [e.g., Weaver et al., 2018, Isella et al., 2018]. Estimates of the inner disk are extrapolated from the outer regions [e.g., Andrews, 2020]. Theoretical models, including the inner disk regions, have to take additional mechanisms into account. Close to the star, stellar magnetic torques change the rotational and radial velocity profile of the disk [e.g., Romanova et al., 2002, Bessolaz et al., 2008, Steiner et al., 2021] and,

thus, the disk's density distribution. In the inner 10 AU, outflows and winds also influence the structure of the disk [e.g., Bai and Stone, 2013]. Such outflows can block the irradiation of the outer regions, affecting the temperature structure of the whole disk [e.g., Takami et al., 2013]. A more detailed summary of the processes in the inner region of PPDs is given in Sec. 1.8.

1.5. The dusty disk component

The description of the dust component in a PPD and its evolution represents an important step in solving the question of planet formation. Variations in the local gas-to-dust fraction [that is usually assumed to be 100, e.g., Ansdell et al., 2016], different dust size distributions or grain growth processes can affect the density and temperature structure as well as the disk dynamics [e.g., Testi et al., 2014]. Over the past decades, several mechanisms have been identified that can affect the dust distribution in PPDs: vertical settling and mixing processes [e.g., Safronov, 1972, Goldreich and Ward, 1973, Dubrulle et al., 1995, Bai and Stone, 2010], inward radial drift induced by drag forces [e.g., Whipple, 1972, Weidenschilling, 1977], radial turbulent mixing that can oppose the radial inward drift [e.g., Youdin and Goodman, 2005, Squire and Hopkins, 2020], and the presence of dusty disk winds removing small dust grains from the disk [e.g., Takami et al., 2013, Ellerbroek et al., 2014, Miyake et al., 2016]. In addition, the grain growth of dust is explained by dust-to-dust collisions caused by turbulence. These collisions can have different outcomes, which depend on the dust properties, such as sticking, bouncing, fragmentation, erosion, or a proportionate mass transfer between dust particles [see the review by Testi et al., 2014]. Despite recent progress, it is still unclear how dust can grow from sub-micron sizes to planetesimals [e.g., Tu et al., 2022].

It is beyond question that a comprehensive understanding of the processes in PPDs includes the dusty component. The focus of this thesis, however, lies in combining different numerical models, in which the dust component is treated in a simplified manner (at least for now). The detailed description of dust dynamics would go beyond the scope of this thesis and has to be included in future work. For a detailed description of dust-related processes in PPDs, the reader is referred to the review of Testi et al. [2014] and Birnstiel et al. [2016] and the results of numerical models

including the evolution of a dusty component [e.g., Birnstiel et al., 2010, Vorobyov et al., 2022, Tu et al., 2022].

1.6. Transport of angular momentum in PPDs

In PPDs, the circumstellar material orbits the central object (in most regions) close to the Keplerian velocity, Ω_K , with a certain amount of angular momentum. Starting from the early star-disk evolution (Class 0 / I) and throughout the entire disk lifetime, disk material is accreted onto the young central star [there are hundreds of accretion rate measurements, e.g., Sicilia-Aguilar et al., 2010, Manara et al., 2012, Venuti et al., 2019, Alcalá et al., 2019, Frasca et al., 2020]. Thus, the disk material's angular momentum must be removed to allow mass transport toward the star. This inward movement of material is presumed to influence the formation of planets [e.g., Najita and Bergin, 2018] and, thus, the understanding of the mechanisms that drive angular momentum transport is of great importance. Unfortunately, our knowledge about angular momentum transport in PPDs is not yet sufficient, and different mechanisms are still under debate [e.g., Fromang and Lesur, 2017, Najita and Bergin, 2018, Lesur et al., 2022].

Depending on the evolutionary stage of the star-disk system and the location within the disk different mechanisms can result in torques acting on the disk, which are responsible for the transport of angular momentum within the disk. These torques can be calculated by multiplying the local, azimuthal forces that act on a mass element within the disk by the disk radius. Here, I want to review three different torque mechanisms. First, gravitational torques transport a large quantity of angular momentum within the disk [e.g., Larson, 1984, Vorobyov and Basu, 2007, Vorobyov and Basu, 2009]. Gravitational torques arise due to non-axisymmetric features in the disk resulting in an azimuthal gradient in the gravitation potential [e.g., Vorobyov and Basu, 2007]. The mass infall from the parent cloud during the early, embedded stages (Class 0 and Class I) leads to gravitational instabilities and non-axisymmetric fragmentation within the disk at radii $\gtrsim 20$ AU. Another possible source of angular momentum transport is viscosity. Viscosity leads to stress in the azimuthal direction transporting angular momentum outward. Usually, this viscous transport of angular momentum is assumed to be driven by turbulence in the disk [e.g.,

Shakura and Sunyaev, 1973, King et al., 2007, Najita and Bergin, 2018, Lesur et al., 2022] and is parameterized by the dimensionless quantity α_S [based on Shakura and Sunyaev, 1973] that controls how effectively a turbulent process can redistribute angular momentum radially and the kinetic viscosity, ν , can be written as

$$\nu = \alpha_S c_S H_P. \quad (1.6)$$

The quantity α_S ranges from $\sim 10^{-5}$ in the DZ [e.g., Barranco and Marcus, 2005] to $\sim 10^{-1}$ in turbulent disk regions [e.g., King et al., 2007, Kotko and Lasota, 2012]. During the early stages, viscous torques are small compared to gravitational torques [e.g., Vorobyov and Basu, 2009]. During later stages, however, viscosity can provide one possible explanation for the evolution of PPDs [e.g., Hartmann et al., 1998, Calvet et al., 2000]. In addition, angular momentum can be removed from the disk via outflows and winds [e.g., Wardle and Koenigl, 1993, Salmeron et al., 2011, Bai and Stone, 2013]. Under the presence of an inclined magnetic field in the disk, out-flowing disk material is accelerated centrifugally along the magnetic field lines anchored at the disk's surface [e.g., Bai and Stone, 2013]. The resulting torque due to these magnetized winds can be calculated by multiplying the mass-loss rate by the angular momentum of the wind [e.g., Weber and Davis, 1967, Mestel, 2012, Li and Cao, 2019]. In the following subsections, I highlight and summarize different physical approaches to explain torques, turbulence, and the angular momentum transport due to winds in PPDs.

Gravitational instabilities

Starting at the early stages of disk evolution, massive disks with disk masses $\gtrsim 0.1 M_\star$ [corresponding to Class 0/I systems, e.g., Kratter and Lodato, 2016] can be subjected to gravitational instabilities (GI) due to self-gravity of the disk. In the standard picture, the self-gravity of a disk can be expressed by the so-called Toomre parameter,

$$Q = \frac{c_S \Omega}{\pi G \Sigma}, \quad (1.7)$$

with the isothermal sound speed, c_s , the angular velocity of the disk (assuming a Keplerian disk), Ω , the gravitational constant, G , and the surface density, Σ [Toomre, 1964]. For $Q \lesssim 1.7$, the importance of self-gravity for the disk evolution increases, and disk material can start to form clumps, spiral arms, instabilities, or planets through fragmentation [e.g., Vorobyov and Basu, 2009, Vorobyov, 2010, Kratter and Lodato, 2016, Liu, 2017, Dong et al., 2018]. In this context, an effective measure of angular momentum transport compatible with the α prescription was first formulated by Lin and Pringle [1987], Lin and Pringle [1990]. The α value scales with Q^{-2} as

$$\alpha_{\text{S,GI}} = \eta \left(\frac{Q_{\text{crit}}^2}{Q^2} - 1 \right), \quad (1.8)$$

with a numerical constant, $\eta \lesssim 1$, and the critical Toomre value at which the transport of angular momentum sets in, Q_{crit} . For $Q > Q_{\text{crit}}$, $\alpha_{\text{S,GI}}$ is assumed to be zero. In numerical disk evolution models, $\alpha_{\text{S,GI}}$ ranges between $\lesssim 10^{-3}$ and $\sim 10^{-1}$ [assuming $Q_{\text{crit}} = 1.7$, e.g., Vorobyov, 2010]. In an alternative formulation, Schib et al. [2021] expressed the influence of self-gravity as

$$\alpha_{\text{S,GI}} = \exp\{-Q^4\}. \quad (1.9)$$

For $Q \rightarrow 1$, $\alpha_{\text{S,GI}}$ is also of the order of 10^{-1} . Furthermore, $\alpha_{\text{S,GI}}$ can be estimated by dividing the gravitational stress tensor by the gas pressure [e.g., Riols and Latter, 2018, Vorobyov et al., 2023]. The resulting values for $\alpha_{\text{S,GI}}$ can reach values up to 0.1 at radial distances of 10 – 100 AU.

Magneto-rotational instabilities

Over the past 30 years, outward transportation of angular momentum in a PPD has been predominantly modeled through the use of magneto-rotational instabilities (MRI) [e.g., Balbus and Hawley, 1991, 1998, Simon et al., 2015, Vorobyov et al., 2020a]. In the picture of MRIs, flows in Keplerian rotation around a central object are destabilized by the presence of a weak magnetic field, transporting angular momentum from fast-rotating inner disk regions to slow-rotating outer parts of the disk [Balbus and Hawley, 1991]. The formulation of MRIs in the context of accretion disks greatly advanced our understanding of how angular momentum is transported in PPDs. This is due to the low requirements of MRIs: inward increasing angular velocity and the presence of

a weak magnetic field. Both conditions are met in PPDs. Furthermore, MRIs grow quickly on dynamic timescales and can produce turbulence strong enough to transport angular momentum outward, matching observational limitations [e.g., Balbus and Hawley, 1998]. The viscous α value due to MRI usually ranges between $\sim 10^{-3}$ and 10^{-1} [e.g., Bell and Lin, 1994, Bai and Xue-Ning, 2014, Yang et al., 2018, Zhu et al., 2020, Vorobyov et al., 2020b].

During the last decade, however, the applicability of MRI in PPDs has been challenged [e.g., Bai and Stone, 2013, Turner et al., 2014, Bai et al., 2016, Najita et al., 2018, Lesur et al., 2022]. The disk material has to be sufficiently ionized [with ionization fractions of $\gtrsim 10^{-13}$, e.g., Vorobyov et al., 2020b, Lesur et al., 2022] so that the disk material can couple to the magnetic field lines and angular momentum can effectively be transported within the disk. Different processes can ionize the accretion disk [e.g., Vorobyov et al., 2020b]. In cool disk regions of the disk's midplane ($T_{\text{disk}} \lesssim 1000$ K), the disk material can be ionized by cosmic rays, collisions, and ionization by radionuclides. The layers close to the surface of the disk, in addition, can also be ionized by external high energy (extreme UV or X-ray) radiation. Above midplane temperatures of ~ 1000 K, the ionization of alkali metals (usually potassium is considered, due to the low ionization potential) can increase the ionization fraction to $\gtrsim 10^{-6}$ and enable the formation of MRIs [e.g., Dudorov and Khaibrakhmanov, 2014, Vorobyov et al., 2020b].

During the early Class 0/I evolutionary stages ($\lesssim 1$ Myr), high ionization fractions are possible in PPDs and MRIs can develop, leading to episodic events of high protostellar accretion [e.g., Bae et al., 2014, Vorobyov et al., 2020b]. In combination with GIs, these mechanisms can efficiently remove mass and angular momentum from embedded (Class 0 and I) disks. In disks around slightly older T Tauri stars, however, MRIs are expected to be suppressed as the disks are cooler, weaker ionized, and due to non-ideal MHD effects [e.g., Bai and Stone, 2013, Turner et al., 2014, Bai et al., 2016, Najita et al., 2018, Lesur et al., 2022]. The question of an alternative method to remove angular momentum from evolved disks past the GI evolutionary stage led to the development of alternative methods, such as disk winds, outflows, and hydro-dynamical instabilities (see subsections below). The latter instabilities do not require the presence of a magnetic field and can develop independently of disk ionization.

Thermal instabilities

Thermal instabilities (TIs) arise if the gas opacity steeply rises with temperature [e.g., Bell and Lin, 1994, Kadam et al., 2020]. An increase in temperature leads to a sharp increase in opacity and the temperature rises further, resulting in a runaway situation. When first adapted for PPDs [Bell and Lin, 1994], the steep opacity rise at the ionization temperature of hydrogen (about 7000 K) and the resulting thermal instability has been used to explain FU Orionis like outbursts. The disk temperatures in PPDs are in most regions well below the ionization temperature of hydrogen [e.g., Boss, 1998]. Very close to the central star or in combination with GIs or MRIs, TIs could develop in early, massive disks. More recently, Kadam et al. [2020] has verified the same mechanism at lower temperatures of approximately 1400 K and 3500 K. The reason for the TI in these temperature regions is the dependence of opacity due to molecular absorption by water vapor [Kadam et al., 2020]. Disk temperatures of ~ 1000 K are expected to be realistic in PPDs and thus, TIs can contribute to the transport of mass within the disk. It is important to note that the occurrence and strengths of TIs and the associated transport of mass highly depend on the opacity model used. Different opacity models can differ significantly in specific regions and depend on different parameter settings, for example, for different metallicity values [e.g., Pollack et al., 1985, Semenov et al., 2003, Caffau et al., 2011a, Woitke et al., 2016]. Thus, the occurrence of TIs should be carefully tested under different opacity models.

Hydro-dynamical instabilities

The previous instabilities occur predominately in hot, massive ($M_{\text{disk}} \gtrsim 0.1 M_{\star}$) disks in the early stages of the star-disk evolution (Class 0/I), or if the disk material is sufficiently ionized. In later evolutionary stages, Class II/III, the disk mass is expected to be considerably lower [$M_{\text{disk}} \lesssim 0.01 M_{\star}$, e.g., Ansdell et al., 2016]. Thus, the disk's self-gravity is not sufficient for the development of GIs, and the maximum disk temperature of < 1000 K is too low for TIs or MRIs within the dead zone (usually ranging between < 1 and $\lesssim 10$ AU, see Fig. 1.1). To explain the transport of angular momentum in this inner region, alternative methods have been developed. In a recent review, Lesur et al. [2022] summarized three hydro-dynamic instabilities

that can develop without the presence of a magnetic field in weakly-ionized disk regions. The development of these instabilities depends on the vertical disk structure and the local cooling rate, τ_{cool} [e.g., Lesur et al., 2022].

- *Vertical shear instability (VSI)*: The VSI can develop under the assumption of differential rotation in vertical direction, $\partial\Omega_z \neq 0$, and a rapid cooling rate, $\tau_{\text{cool}}\Omega_K \ll 1$ to overcome vertical stability in accretion disks [e.g., Arlt and Urpin, 2004, Nelson et al., 2013, Lin and Youdin, 2015]. Small and large-scale vertical disturbances can arise at disk radii $\gtrsim 5$ AU [e.g., Lin and Youdin, 2015, Manger et al., 2020, Flock et al., 2020], transporting angular momentum outward with an efficiency of $\alpha_{\text{S,VSI}} \sim 10^{-4}$.
- *Convective overstability (COS)*: The requirements for the COS are 1) a negative entropy gradient, which can be vividly described as a shallow or radially increasing surface density profile in combination with a decreasing radial temperature profile, and 2) a cooling time of the order of a dynamic timescale, $\tau_{\text{cool}}\Omega_K \sim 1$ [e.g., Klahr and Hubbard, 2014, Lyra, 2014, Latter, 2016, Lesur et al., 2022]. These requirements can be satisfied at the edges of the dead zone, opacity transitions, or gaps in the accretion disk [e.g., Latter, 2016, Delage et al., 2022]. The disturbance, caused by the COS, can redistribute angular momentum in a radial direction outward with an efficiency of $\alpha_{\text{S,COS}} \sim 10^{-3}$ [e.g., Raettig et al., 2013].
- *Zombie vortex instability (ZVI)*: In contrast to VSIs, ZVIs require a stable stratification in the vertical direction with sufficiently large cooling times, $\tau_{\text{cool}}\Omega_K \gtrsim 1$ [e.g., Barranco and Marcus, 2005, Marcus et al., 2013, 2015]. In case of too short cooling timescales, the stable vertical stratification would be diminished (see VSI). ZVIs require initial, small-scale perturbations that merge toward large-scale vortices under the influence of the disk's background angular rotation [see Fig. 1 in Marcus et al., 2015]. This vortex is the starting point for new small perturbations of an adjacent disk layer and the ZVI can "infect" its surroundings, leading to new ZVIs and so on [e.g., Marcus et al., 2015, Lesur et al., 2022]. This behavior and the location of ZVIs in the dead zone of PPDs give the instability its name [Barranco and Marcus, 2005]. How effectively the ZVI can transport angular momentum is unfortunately not well constrained [e.g., Lesur et al., 2022]. First estimates, based on the

calculations of Barranco and Marcus [2005] ranging from $\alpha_{S,ZVI} \sim 10^{-5} - 10^{-4}$, have to be confirmed in future, global simulations.

- *Streaming instability*: The streaming instability (SI) requires the presence of a dust component in the disk. It was first described by Youdin and Goodman [2005] based on the study of Goodman and Pindor [2000]. In its basic form, the SI can be summarized as follows [an intuitive visualization can be found in, e.g., Squire and Hopkins, 2020, Lesur et al., 2022]: 1) the inward motion of a dust clump produces a radial gas flow, 2) this gas flow is deflected in an azimuthal direction by the Coriolis force due to the near-Keplerian angular velocity of the disk, 3) the drag of the deflected gas flow causes super-Keplerian dust velocity, 4) dust is deflected outwards enhancing the dust clump. The growth rate of dust clumps depends on the dust-to-gas ratio and the stopping time of the dust with respect to the local azimuthal velocity. The effectiveness of radial angular momentum transport in the disk due to SIs is not well-constrained and needs further investigation.

Magnetized winds

Over the last few years, the drawbacks of turbulence-induced angular momentum transport (e.g., low α values in the DZ) have led to an alternative method for removing angular momentum from the disk, facilitating accretion. Even in weakly magnetized disk regions, winds can emerge from the ionized upper layers of PPDs [e.g., Bai and Stone, 2013]. These upper disk layers can be ionized, for example, through cosmic rays [e.g., Rab et al., 2017b] or high-energy stellar irradiation [e.g., Perez-Becker and Chiang, 2011]. Under the presence of a large-scale magnetic field, the ionized disk material can couple to the field lines and accelerate centrifugally along magnetic field lines [e.g., Bai and Stone, 2013]. This magnetized wind (or jet, if formed close to the inner edge of the disk) can carry away angular momentum from the upper disk layers (above aspect ratios of $H_p/r \approx 5$) and enable accretion rates of $\sim 10^{-8} M_\odot \text{ yr}^{-1}$, even for weakly magnetized disks [e.g., Blandford and Payne, 1982, Ouyed and Pudritz, 1997, Bai and Stone, 2013, Bai and Xue-Ning, 2014, Bai, 2017, Lesur, 2021], matching observed accretion rates of T Tauri stars [e.g., Hartmann et al., 2016]. Currently, different sources of energy for the magnetized winds are being discussed [e.g., Lesur et al., 2022]. First are magneto-hydrodynamic outflows, powered

by accretion and turbulence in the disk, and second, magneto-thermal winds, driven by internal or external photo-evaporation. Global simulations done by Bai [2017] revealed that in the DZ (which ranges from $\lesssim 1$ to ~ 10 AU), the energy from accretion and turbulence is not enough to drive the magnetized winds. In this region, the nature of the wind is magneto-thermal. Very close to the star and inside the DZ, MRIs can provide sufficient energy for magneto-hydrodynamic winds [e.g., Lesur et al., 2022].

1.7. Which processes dominate the dynamic disk evolution?

In the previous section, several mechanisms of transporting angular momentum within or away from the disk are summarized. The occurrence and strength of these mechanisms depend on a variety of factors, such as the evolutionary stage, disk/stellar mass, metallicity, stellar environment, disk structure, and so forth. During the early stages of the star-disk evolution (Class 0/I), the disk mass is expected to be high ($\gtrsim 0.1 M_\star$) and angular momentum can be effectively transported by GIs [e.g., Vorobyov and Basu, 2009, Kratter and Lodato, 2016]. For lower disk masses ($\sim 0.01 M_\star$), which correspond to later evolutionary stages (Class II/III), the influence of self-gravity decreases, and other mechanisms dominate the disk evolution and the transport of angular momentum. In their recent review, Lesur et al. [2022] have summarized the dominant mechanisms in different disk regions for low-mass, non-self-gravitating disks (see Fig. 1.3).

The innermost disk region close to the star (~ 0.1 AU) is assumed to have temperatures of $\gtrsim 1000$ K close to the midplane of the disk [e.g., Jankovic et al., 2021]. At these temperatures, thermal ionization and stellar X-rays can lead to ionization fractions high enough for the development of MRIs that transport angular momentum and facilitate accretion on the star [e.g., Dudorov and Khaibrakhmanov, 2014]. With decreasing temperatures and ionization fractions in the DZ at larger disk radii $\gtrsim 0.5$ AU [e.g., Gammie, 1996a], non-ideal MHD effects inhibit MRIs and angular momentum can be transported due to hydro-dynamic instabilities (VSI, COS, ZVI). In the dead zone, close to the disk midplane, where the dust-to-gas ratio is usually the largest, SIs can dominate the disk evolution [e.g., Squire and Hopkins, 2018, Lesur et al., 2022]. Outside the DZ ($\gtrsim 10$ AU), where the disk becomes optically thin and ionization increases

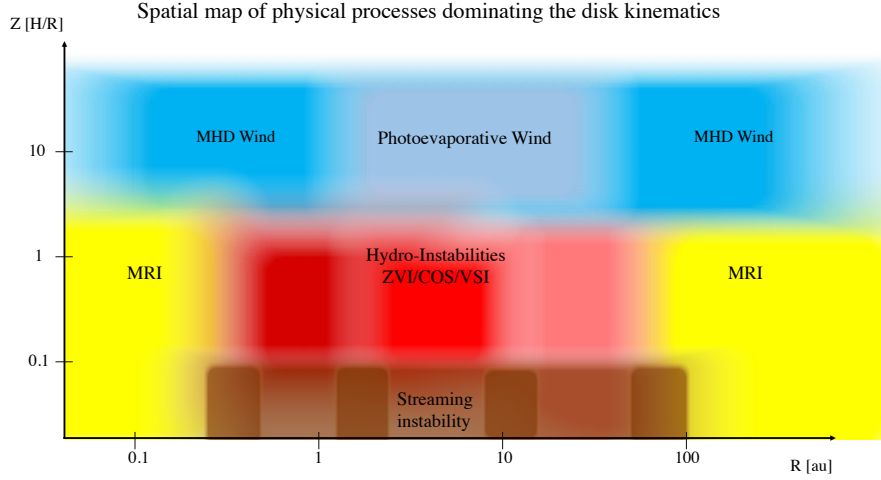


Figure 1.3.: The dominant physical process for the dynamic protoplanetary disk evolution. Depending on the structure of the disk, different mechanisms are dominant in different disk regions (see text). The figure has been taken from Lesur et al. [2022] (licensed under CC BY 4.0, <https://creativecommons.org/licenses/by/4.0/>)

due to external high-energy radiation or cosmic rays, the requirements for MRIs are met again. Magnetized outflows and winds can, in principle, emerge from all disk radii and dominate the disk evolution in the upper disk layers [above aspect ratios of $H_P/r \approx 5$, e.g., Lesur et al., 2022]. Purely magneto-hydrodynamic outflows need energy from accretion or turbulence in the disk to be launched [e.g., Ferreira and Pelletier, 1995, Lesur et al., 2022] and roughly overlap with the MRI active disk regions inside and outside the DZ. Within the weakly ionized regions, which range from $\lesssim 1$ to ~ 10 AU, the dominant mechanisms that remove angular momentum from the star are magneto-thermal winds [e.g., Bai, 2017].

1.8. The stellar influence on the inner disk region

Over the past years, the interest in a comprehensive understanding of the innermost disk region has increased significantly. The disk temperature structure and mass of the entire disk depend on the mass flow rate through the inner (sub-AU) disk region [Vorobyov et al., 2019]. Furthermore, the evolution of the accretion rate can be controlled by the processes in the inner disk region and the star-disk interaction [e.g., Vorobyov et al., 2019, Hennebelle et al., 2020]. A detailed

description of the processes in this region is thus necessary to explain the vast variety of observed disk features. In the current understanding of the innermost region of a PPD [e.g., Romanova et al., 2002, Bessolaz et al., 2008, Hartmann et al., 2016, Steiner et al., 2021], the disk material is transported from the outer regions inward. Moving closer to the star, the influence of a stellar magnetic field, which is usually assumed to be a dipole, increases rapidly. Close to the inner disk edge, the material is usually sufficiently ionized and can couple to the stellar magnetic field lines (see Sec. 1.6). At some point, the stellar magnetic field reduces the radially inward velocity of the disk material and enforces a movement in the vertical direction. The disk material is lifted on the stellar magnetic field lines and is accreted at near free-fall velocity on the star. Close to the stellar surface, the infalling material is abruptly slowed down and releases its energy in form of observable accretion luminosity from the optical to X-rays. In this section, several important aspects of the inner disk region are summarized and influences of different stellar and disk configurations on the inner disk are discussed.

Calculation of the truncation radius

One essential aspect of the inner disk region is the inner disk boundary. Usually, this inner disk boundary is defined via the magnetic truncation radius, r_t . To calculate this characteristic radius, the stellar magnetic pressure is equated with the ram pressure of the infalling material $B_\star^2/8\pi = \rho v^2$ [e.g., Romanova et al., 2002, Bessolaz et al., 2008, Hartmann et al., 2016]. Assuming a dipolar stellar magnetic field, the truncation radius can be expressed as

$$r_t = k \left(\frac{B_\star^4 R_\star^{12}}{2GM_\star \dot{M}_\star^2} \right)^{1/7}, \quad (1.10)$$

with the dipole field strength, $B_\star$, the stellar mass, $M_\star$, the gravitational constant, G , and the accretion rate onto the star, $\dot{M}_\star$ [e.g., Bessolaz et al., 2008]. The factor k usually ranges between 0 and 1 [e.g., Ghosh and Lamb, 1979a, Koenigl, 1991, Bessolaz et al., 2008, Hartmann et al., 2016]. For typical parameters of a Class II star ($M_\star = 0.5 M_\odot$, $R_\star = 2 R_\odot$, $\dot{M}_\star = 10^{-8} M_\odot/\text{yr}$, and $B_\star = 1 \text{ kG}$) the value of the truncation radius ranges between $4 - 6 R_\star$ or $\sim 0.03 \text{ AU}$ [e.g., Hartmann et al., 2016]. Alternative calculation methods of r_t include the disk's gas pressure,

P_{gas} , to counteract the stellar magnetic pressure, $B^2/8\pi = \rho v^2 + P_{\text{gas}}$ [e.g., Romanova et al., 2002, Steiner et al., 2021] or compare the magnetic torque acting on the inner disk with the viscous or turbulent torque [e.g., Bessolaz et al., 2008].

Influence of stellar magnetic torques

The stellar magnetic field does not only define the truncation radius. Stellar magnetic field lines can couple to the disk material over a region that is sufficiently ionized ($\lesssim 0.5$ AU). In this region, differential rotation between the stellar magnetic field lines, which are usually assumed to rotate with the same angular velocity as the star, and the disk material causes a torque transporting angular momentum between the disk and the star [e.g., Armitage and Clarke, 1996, Matt et al., 2012a, Gallet et al., 2019, Steiner et al., 2021]. The position at which the angular velocity of the star corresponds to the angular velocity of the disk, is called the co-rotation radius

$$r_{\text{co}} = \left(\frac{GM_{\star} P_{\star}^2}{4\pi^2} \right)^{1/3}, \quad (1.11)$$

with the stellar rotation period, $P_{\star}$. Inside r_{co} , the disk material rotates faster than the star and angular momentum is transferred from the disk to the star. As a consequence, the disk material is slowed down, which results in a sub-Keplerian rotation, and accelerates in the radial direction toward the star [e.g., Steiner et al., 2021]. Outside the co-rotation radius, the star rotates faster than the disk and stellar angular momentum is added to the disk material, resulting in a super-Keplerian velocity. The disk material outside r_{co} is accelerated and the radial velocity toward the star is slowed down. With disk material still moving inward from the outer regions, where the stellar influence is negligible, the material starts to pile up just outside the co-rotation radius, dependent on the stellar magnetic field strength and stellar rotation period [e.g., Steiner et al., 2021].

Episodic accretion outbursts

Other important mechanisms that are connected to the inner disk region are episodic accretion outbursts. During such outburst events, the accretion rate onto the star and the stellar luminosity

are abruptly enhanced by several orders of magnitude [e.g., Hartmann and Kenyon, 1996, Zhu et al., 2007, Eisner and Hillenbrand, 2011, Audard et al., 2014]. Early estimates suggest that these events are common and occur repeatedly [Hartmann and Kenyon, 1996]. During an accretion outburst, the disk is heated by the released energy of the accreted material. This additional heating affects the disk chemistry, the location of snow lines, and, in further consequence, the formation processes of planets [e.g., Cieza et al., 2016, Rab et al., 2017b, Wiebe et al., 2019, Vorobyov et al., 2020a,a]. Different possible mechanisms could induce such an outbursting behavior. During the early phases of the star-disk evolution (Class I and early Class II), when the accretion disk is still massive, gravitational instability can lead to the formation of gaseous clumps that move toward the star and ignite an outburst [Vorobyov and Basu, 2015]. Alternatively, gravitational instabilities can cause disturbances in the inner disk region that ignite MRIs and start the outburst [e.g., Bae et al., 2014]. Furthermore, thermal instabilities in the inner disk region can be responsible for an enhanced accretion rate onto the star [e.g., Bell and Lin, 1994, Kadam et al., 2020].

During later evolutionary stages of PPDs, the disk mass is usually too small for the development of gravitational or thermal instabilities. Yet, an accretion outburst has been observed in the late Class II object HBC 722 with a very low disk mass [e.g., Kóspál et al., 2016, Hillenbrand et al., 2018]. An additional explanation for episodic accretion outbursts without GIs or TIs is based on the layered accretion model [Gammie, 1996a]. Material that moves inward from the outer disk regions piles up in the weakly ionized dead zone. The accumulation of material increases the disk temperature due to viscous heating. Above a certain temperature (usually ~ 1000 K), the ionization fraction of the disk increases abruptly and the MRI can develop [e.g., Bae et al., 2014, Steiner et al., 2021]. In the very inner disk region (0.1 AU) high temperatures during MRIs and TIs can lead to the destruction of planets due to extreme evaporative events [e.g., Nayakshin et al., 2023]. Furthermore, a binary component or stellar fly-bys can ignite accretion outbursts during the entire star-disk evolution through tidal interactions [e.g., Reipurth and Aspin, 2004, Forgan and Rice, 2009, Vorobyov et al., 2021, Dong et al., 2022]. Except for tidal interactions and gaseous clumps, produced by GIs that fall toward the star, accretion outbursts are triggered in the inner disk regions. Thus, a detailed analysis of this region close to the star is essential for understanding accretion outbursts.

1.9. The disk's influence on the star

Protoplanetary disks are linked to their stars in the inner disk region by magnetic field lines. Via this connection, disk material and angular momentum can be transferred between the disk and the star. Recent studies on how this connection might affect stellar evolution suggest that the disk material can significantly alter the way a star evolves during the entire disk lifetime [e.g., Matt and Pudritz, 2005b, Matt et al., 2012a, Baraffe et al., 2012, Kunitomo et al., 2017, Steindl et al., 2022]. In this section, the effects of energy that is added to the stellar interior by the accreted material and the influence of the disk on the stellar spin evolution are reviewed.

Release of accretion energy

In the "classical" theory of stellar evolution [Hayashi, 1961], a (low-mass) protostar forms with an extended radius. During its further evolution the young star contracts and is heated by a large amount of gravitational energy. With a rising temperature in the stellar interior, the first nuclear reactions set in, contraction slows down and the star develops a radiative core. More recent studies, on the other hand, suggest that the evolution of stellar radii and the contraction phase depend on the amount of energy the accreted material adds to the stellar interior [e.g., Baraffe et al., 2009, 2012, Kunitomo et al., 2017, Vorobyov et al., 2017b, Steindl et al., 2022]. As a consequence, the stellar evolutionary tracks on the Hertzsprung-Russel Diagram (HRD) significantly vary, and inferred stellar properties change [e.g., Hosokawa et al., 2011].

In the updated picture, a small hydrostatic stellar seed forms with a mass of $\sim 0.01 M_{\odot}$ and a radius of $0.25 - 3.0 R_{\odot}$ [this choice corresponds to slightly evolved stars from the second Larson core, e.g., Kunitomo et al., 2017]. It is assumed that a disk forms around this stellar seed after a short period of $\lesssim 10^4$ years [e.g., Hueso and Guillot, 2005, Vorobyov et al., 2017b], which marks the start of the Class 0/I stage of the star-disk system. Material from the PPD accretes onto the star carrying a specific energy rate [e.g., Hartmann et al., 1997, Baraffe et al., 2009],

$$\dot{E} = \epsilon \frac{GM_{\star} \dot{M}_{\star}}{R_{\star}}. \quad (1.12)$$

The factor ϵ depends on the geometry of the accretion process. In the case of spherical accretion,

$\epsilon \leq 1$, and for accretion from a thin disk $\epsilon \leq 1/2$ can be assumed [e.g., Hartmann et al., 1997]. Dependent on the entropy of the accreted material, $\dot{E}$ can be divided into two parts [e.g., Hartmann et al., 1997, Baraffe et al., 2009]. If the entropy of the accreted material is smaller compared to the stellar object, the energy is mostly radiated away in the form of accretion luminosity, L_{acc} . If, on the other hand, the material carries a lot of entropy, a large portion of the energy is added to the stellar interior in the form of an additional energy term, $\dot{E}_{\text{add}}$. The ratio between L_{acc} and $\dot{E}_{\text{add}}$ is controlled by a free parameter β , leading to

$$L_{\text{acc}} = \epsilon(1 - \beta) \frac{GM_{\star} \dot{M}_{\star}}{R_{\star}} \quad (1.13)$$

and

$$\dot{E}_{\text{add}} = \epsilon\beta \frac{GM_{\star} \dot{M}_{\star}}{R_{\star}}. \quad (1.14)$$

For $\beta \ll 1$, the energy of the accreted material is mostly radiated away. Upper values for β and its dependence on stellar and disk properties are mostly free parameters in simulations. In its current understanding, β ranges between a lower and upper bound, β_{low} and β_{up} , respectively, dependent on the accretion rate [e.g., Baraffe et al., 2012, Vorobyov et al., 2017b, Elbakyan et al., 2019, Steindl et al., 2022]. During phases of high accretion rates, β approaches its upper boundary, which ranges between $\beta_{\text{up}} = 0.1 - 0.5$ [e.g., Baraffe et al., 2012, Elbakyan et al., 2019]. For low accretion rates, $\beta \sim \beta_{\text{low}} \sim 0.001$ [e.g., Vorobyov et al., 2017b, Steindl et al., 2022]. An additional free parameter is the location within the star at which the additional energy is released. Theoretical studies suggest that the energy is distributed between the surface of the star and a specific mass coordinate, M_{add} [e.g., Kunitomo et al., 2017]. The value of M_{add} ranges between 1.0 and 0.01 [e.g., Kunitomo et al., 2017, Steindl et al., 2022].

Stellar rotation

One fundamental challenge in astrophysics is the origin and evolution of stellar angular momentum. Recent observational campaigns have revealed thousands of stellar rotational periods in stellar clusters at different ages, stellar masses, or other properties [e.g., Herbst et al., 2002, Irwin et al., 2008, Irwin and Bouvier, 2009, Hartman et al., 2009, Meibom et al., 2011, Affer et al., 2013,

Gallet and Bouvier, 2013, Venuti et al., 2017, Serna et al., 2021]. As the ages of stars in such clusters are assumed to be (at least) similar, the rotational distributions obtained are important benchmarks to describe the evolution of the stellar spin and angular momentum. Following Gallet and Bouvier [2013], the stellar spin evolution starting from the accretion disk phase to an age of ~ 1 Gyr can be divided into three phases. For the first few million years, the star interacts with its accretion disk and the spin evolution depends on the star-disk interaction [e.g., Armitage and Clarke, 1996, Matt et al., 2012a, Gallet and Bouvier, 2013, Ireland et al., 2021]. Once the disk has been dissolved, the star is expected to spin up due to contraction during the pre-main sequence phase (pre-MS). With increasing pressure and temperature in the stellar interior, the first nuclear reactions stop the contraction and start the development of a radiative core. During the main sequence (MS), stellar winds spin down the star, and, after several gigayears, the stellar spin converges toward the solar value (in the case of a solar mass star).

Even though the stellar rotation period converges toward a similar value during the evolution on the MS for a given stellar mass [e.g., Gallet and Bouvier, 2013], the spread of observed rotation periods during earlier stages spans over at least one order of magnitude [e.g., Venuti et al., 2017, Rebull et al., 2020, Serna et al., 2021]. Already at an age of ~ 1 Myr, there are fast-rotating stars with rotation periods $\lesssim 1$ day and slow-rotating stars with periods $\gtrsim 10$ days [e.g., Rebull et al., 2020]. Forward-modeling of stellar rotation periods shows that a fast-rotating star at the end of the star-disk interaction phase continues to rotate faster compared to a slow-rotating star at young ages [e.g., Johnstone et al., 2021]. The rotation periods are linked to stellar high-energy radiation [e.g., Pizzolato et al., 2003] and, thus, a fast-rotating star produces more high-energy radiation compared to a slow-rotating star [e.g., Johnstone et al., 2021]. The amount of high-energy radiation, on the other hand, affects planetary atmospheres [e.g., Güdel et al., 1997, Lammer et al., 2003, Buccino et al., 2006, Cecchi-Pestellini et al., 2009, Sanz-Forcada et al., 2011, France et al., 2013, Nortmann et al., 2018]. Strong exposure to high-energy radiation can lead to a loss of the planetary atmosphere [e.g., Johnstone, 2021]. Consequently, understanding the formation and evolution of our planet requires an understanding of the stellar spin evolution, starting from the early star-disk interaction phase.

The rotational evolution after the disk phase has been studied in detail including different stellar

wind models that remove angular momentum from the star, leading to a stellar spin-down during the main sequence [e.g., Gossage et al., 2021]. During the disk phase, these models usually use the disk-locking hypothesis [e.g., Koenigl, 1991, Shu et al., 1994a]. Inferred from observations, angular momentum is assumed to be exchanged between the star and the disk resulting in a nearly constant rotation period when the accretion disk is present. More recent stellar spin models, however, include the effects of different torque contributions during the star-disk interaction phase [e.g., Matt and Pudritz, 2005b, Matt et al., 2012a, Gallet et al., 2019, Ireland et al., 2021]. These models indicate that a constant rotation period only occurs for specific conditions and stellar rotation periods can significantly vary during the lifetime of PPDs. To model the stellar spin evolution during the presence of an accretion disk, three torque contributions are usually considered [e.g., Gallet et al., 2019, Ireland et al., 2021].

The accretion process itself adds the angular momentum of the disk material at the inner disk boundary to the star. The resulting torque, Γ_{acc} , depends on the accretion rate, $\dot{M}_{\star}$, and the angular disk velocity at the truncation radius, $\Omega_{\text{d}}(r_{\text{t}})$, leading to

$$\Gamma_{\text{acc}} = \dot{M}_{\star} r_{\text{t}}^2 \Omega_{\text{d}}(r_{\text{t}}). \quad (1.15)$$

Winds originating at the stellar surface, on the other hand, can remove angular momentum from the star. Material ejected from the star couples to stellar magnetic field lines, accelerates, and causes the star to spin down. These winds can be thermally driven or powered by the accretion process, so-called accretion-powered stellar winds [APSW, e.g., Matt and Pudritz, 2005b, Gallet et al., 2019]. The mass-loss rates of APSW are assumed to scale with the accretion rate $\dot{M}_{\text{APSW}} = W \dot{M}_{\star}$, with the efficiency factor W . Theoretical estimates of W range between 1% and 2% [e.g., Cranmer, 2008, Pantolmos et al., 2020]. Thermally driven winds show mass-loss rates of $\lesssim 10^{-11} \text{ M}_{\odot} \text{ yr}^{-1}$ [e.g., Gallet et al., 2019]. Thus, during early disk phases with accretion rates $\gtrsim 10^{-9} \text{ M}_{\odot} \text{ yr}^{-1}$, the APSW dominates the stellar mass-loss $\dot{M}_{\text{wind}}$. Thermally driven winds become more important during later disk stages with smaller accretion rates and when the disk has been dissolved. The torque due to disk winds can be written as

$$\Gamma_{\text{wind}} = \dot{M}_{\text{wind}} r_{\text{A}}^2 \Omega_{\star}, \quad (1.16)$$

with the Alfvén radius, r_A , acting as the lever arm of the wind. The third contribution to the stellar spin model depends on the magnetic star-disk interaction [e.g., Armitage and Clarke, 1996, Matt and Pudritz, 2005a, Gallet et al., 2019, Ireland et al., 2021]. When coupled to the disk material, stellar magnetic field lines exert torque onto the stellar material due to a differential rotation between the star and the disk. The torque exerted on a small annulus of the disk, Δr , can be written as

$$\Delta\Gamma_{\text{mag}} = qr^2B_z^2\Delta r, \quad (1.17)$$

with the vertical magnetic field component, B_z , and the twist of the magnetic field lines, q . The total torque due to this magnetic star-disk interaction results from integrating over the whole region, at which the stellar magnetic field lines can couple to the disk material. Disk material inside the co-rotation radius rotates faster than the star and adds angular momentum to the star. Outside the co-rotation radius, disk material rotates more slowly compared to the star and the star adds angular momentum to the disk material. The total amount of angular momentum exchanged between the star and the disk due to the magnetic star-disk interaction depends on the relation between the truncation radius and the co-rotation radius and how much disk material can effectively couple to the stellar magnetic field lines. In summary, the stellar spin model during the accretion phase depends on the accretion rate onto the star, $\dot{M}$, and the inner disk region where the stellar magnetic field lines couple to the disk material. Thus, a detailed model of this region is vital to understanding the evolution of the stellar spin.

1.10. Modelling the star-disk system

In the previous subsection, PPDs and different processes during the star-disk interaction phase have been introduced. With lifetimes of millions of years, observations of the complete evolution of these objects are impossible. Each observation provides a snapshot of a certain system at a specific evolutionary stage. To understand the evolution of young star-disk systems that lead to the formation of planets and possibly life, numerical models have to be constructed. Observational data then provide important benchmarks for these models. In this section, the challenges in modeling a young star-disk system are highlighted and current star and disk models

are summarized.

A problem of scales

A young star-disk system is a three-dimensional object extending from the stellar radius outward to hundreds (or in some cases thousands) of astronomical units. The processes that transport angular momentum and material toward the star include the scales of small dust grains and magnetic field lines coupled to the disk material. In addition, timescales in PPDs can range between days (rotation of the inner disk region) and the lifetime of PPDs (millions of years). Ideally, three-dimensional models including various physical aspects such as instabilities, non-axisymmetric structures, mass loss, magnetic fields, different turbulence mechanisms, and the evolution of dust grains would be required to simulate planet formation during the evolution of PPDs. To account for small-scale processes, the model would require a computational domain with an extremely high sub-AU numerical resolution. Modeling the whole lifetime of PPDs with these high-resolution, multi-dimensional models is unfortunately impossible at the moment. The reason for this is a timestep limitation according to the Courant-Friedrichs-Lewy (CFL) condition [e.g., Courant et al., 1928]. In models utilizing an explicit numerical scheme, a stable simulation can only be achieved for timesteps Δt smaller than

$$\Delta t \leq \frac{\Delta r}{|v|}, \quad (1.18)$$

with the characteristic velocity of the system, v , and the size of a grid cell, Δr . Here, v corresponds to the maximum local velocity component. Physically interpreted, this criterion implies that a quantity cannot be transported over more than one grid cell during one timestep. Especially in the inner disk regions, velocities in the PPDs are high (e.g., the angular velocity increases with decreasing radius). At the same time, a good resolution (small Δr) is required to study small-scale processes. As a result, a maximum achievable timestep for the inner disk region is of the order of one second [e.g., Steiner et al., 2021]. To circumnavigate this restrictive criterion, the physical domain has to be divided. In this section, models to study the evolution of the disk component and the stellar component of the star-disk system are summarized.

Disk component

The model to study the evolution of PPDs depends on the characteristic timescale, δt_{char} , of the process that is about to be explored. In the case of short-term, small-scale processes with $\delta t_{\text{char}} \sim 1/\Omega_{\text{d}}$, detailed multi-dimensional models are available to model the star-disk system [e.g., Romanova et al., 2002, Fouchet et al., 2010, Zanni and Ferreira, 2013, Čemeljić, 2019, Flock et al., 2020, Ireland et al., 2021]. These models include important physical aspects such as the effects of magnetic fields, the influences of planets, or a detailed dust component. Due to the requirement for high resolution and explicit methods, the maximum simulation time (using reasonable computational sources) is restricted to $\sim 10^2$ to $\sim 10^3$ stellar rotation periods [e.g., Čemeljić, 2019, Ireland et al., 2021, Flock and Mignone, 2021]. While this period is sufficient to study short-term processes like different turbulence mechanisms [e.g., Flock et al., 2020], predictions of the entire evolution of the star-disk system cannot be made. In addition, the stellar component in these models is usually assumed to be constant during the simulation.

To utilize long-term models of PPDs, two different techniques are used. In the most commonly used approach, the hydrodynamic disk equations are simplified using the diffusion approach [e.g., Lynden-Bell and Pringle, 1974, Pringle, 1981, Armitage, 2015]. The number of dimensions is reduced to the radial direction assuming a vertically integrated, axisymmetric disk. Furthermore, the angular velocity of the disk scales with the Keplerian velocity. As a result, only one equation for the surface density of the disk has to be solved, reducing the computational requirements significantly. In its basic form [Lynden-Bell and Pringle, 1974], this diffusion equation only takes viscosity into account. In more recent studies, the diffusion equation has been extended for infall out of a parent cloud, fragmentation, magnetically driven disk winds, as well as internal and external photo-evaporation [e.g., Schib et al., 2021, Kunitomo et al., 2020]. Thanks to the reduced complexity of the numerical model, the diffusion approach allows long-term simulations to cover large parameter spaces from the truncation radius to the outer disk regions [e.g., Kunitomo et al., 2020]. These studies provide valuable insights into global disk properties like the lifetime and dispersal of PPDs [e.g., Kunitomo et al., 2020, Schib et al., 2021]. The assumptions on which the diffusion equation is based, on the other hand, exclude important physical aspects. The influences of stellar magnetic field lines or radial pressure gradients that change the angular disk velocity

locally cannot be included. Furthermore, non-axisymmetric features like disk spiral arms and fragmentation have to be approximated as an additional component in the diffusion equation [e.g., Schib et al., 2021]. The stellar component is usually included in the form of tabulated values from stellar evolution models [e.g., Schib et al., 2021]. These models are calculated assuming a certain accretion rate [e.g., Yorke and Bodenheimer, 2008, Sakurai et al., 2015] and, thus, the effects of an evolving or variable accretion rate [e.g., Elbakyan et al., 2019] cannot be included.

In an alternative method that allows long-term studies of PPDs to be made, the innermost disk region ($\lesssim 1$ AU) is considerably simplified while the outer disk is resolved in radial and azimuthal directions [e.g., Zhu et al., 2012, Vorobyov et al., 2019]. Outside a radius of ~ 1 AU, the characteristic disk velocities are large enough to use timesteps (see Eq. 1.18) that are suited for long-term simulations using the full set of hydrodynamic equations in the thin disk limit. Without the simplifications of the diffusion approach, local deviations from the Keplerian angular velocity and non-axisymmetric features like spiral arms and fragmentation can be resolved thanks to self-consistent computations of the disk's self-gravity [e.g., Vorobyov, 2010]. The accretion rate onto the central star is approximated based on the mass flux at the inner boundary of the computational domain [Vorobyov et al., 2019]. Depending on the chosen parameters, the accretion history varies significantly, affecting the stellar and disk evolution and highlighting the importance of the inner disk region. The stellar component is coupled to the disk evolution code [e.g., Vorobyov et al., 2019]. The star accretes disk material and heats the disk with its internal and accretion luminosity.

Implicit disk models of the inner disk region

To model the innermost disk region and the star-disk interaction, current long-term models are subjected to the simplified diffusion approach that ignores the effects of stellar magnetic torques. These torques, however, are important for describing the magnetic star-disk interaction and the evolution of the stellar rotation period (see Sec. 1.8 and Sec. 1.9). More detailed models, including these effects, are restricted to short simulation times ($\lesssim 100$ years) and consequently also not suited for long-term simulations of the inner disk region. For these detailed disk models, an explicit computational method is usually used that calculates the disk parameters at the new

timestep, X^{t+1} , based on the values at the old timestep, X^t ,

$$X^{t+1} = F(X^t), \quad (1.19)$$

where F denotes the system of governing equations. In contrast, an implicit method has to solve

$$X^{t+1} = F(X^{t+1}), \quad (1.20)$$

including the new timestep, X^{t+1} , in its calculation. This translates to a system of equations

$$G(X^{t+1}) = 0, \quad (1.21)$$

which has to be solved for obtaining X^{t+1} . An implicit disk model is not subjected to the CFL condition and, thus, larger timesteps (up to a factor of 10^6) are possible compared to an explicit integration scheme [e.g., Steiner et al., 2021]. In their recent work, Steiner et al. [2021] combines a detailed physical disk model and the ability to perform long-term disk modeling. This allows a more detailed calculation of the magnetic star-disk interaction and the accretion rate onto the star during the entire disk lifetime. Thus, the implicit model of the inner disk region promises to be an elegant link between the stellar component and the models describing the outer disk regions.

Despite the significant advantage of an implicit model regarding timestep limitations, the construction of the model requires a substantial amount of additional numerical work [e.g., Dorfi and Drury, 1987, Stoekl and Dorfi, 2014, Steiner et al., 2021]. First, for each time step, a Jacobian matrix of the system of equations, $\partial G/\partial X^t$, must be created, containing the derivatives of each variable at a particular grid point relative to every variable at every grid point within the computational domain. The matrix must be inverted until the system converges toward a new solution of the equations. Another complication lies in the construction of the initial model, from which the simulation is started. This initial model has to solve the hydrodynamic equations [e.g., Dorfi and Drury, 1987]. Especially for the inner disk region with complex physical mechanisms, such as magnetic torques, this model has to be constructed carefully to ensure convergence. A proven method is described in Steiner et al. [2021]. In the first step, a (relatively simple)

Keplerian disk with no radial velocity and no magnetic torques acting on the disk is assumed. The complexity of the model is then slowly increased in the form of a mass flux through the outer boundary condition and an increasing stellar magnetic field strength at the inner disk boundary. When the disk has adapted to these boundary conditions, the model becomes stationary and can be used as an initial model for the simulations.

Stellar component

During the past years, different stellar evolution codes have been developed and continuously improved. Most commonly, one dimensional radial codes are used [e.g., Siess et al., 2000, Yorke and Bodenheimer, 2008, Paxton et al., 2011, Baraffe et al., 2012, Feiden, 2016, Paxton et al., 2019a, Amard et al., 2019b]. These models include a variety of important physical aspects, such as convection, different metallicities, effects of magnetic fields, stellar rotation, and energy added to the star during the accretion process. The stellar evolution, starting from a small stellar seed to the end of the stellar life, can be accurately modeled. Amongst other things, the stellar radius and luminosity are calculated, which influence the evolution of the accretion disk. In the stellar evolution models, the mass accretion rate from the disk that feeds the star is derived previously by numeric models [e.g., Baraffe et al., 2012, Kunitomo et al., 2017, Steindl et al., 2022] and a back-reaction of the stellar evolution on the disk cannot be included. Similarly, models studying the stellar spin evolution assume a simplified disk accretion rate in the form of a power-law [e.g., Gallet et al., 2019] or an exponential decay [e.g., Matt et al., 2012a]. Different rotation periods and the resulting magnetic torques acting on the disk, however, might affect the accretion rate onto the star and thus, the stellar spin evolution itself. These effects cannot be included in these models in a self-consistent way. Thus, a combined stellar and disk evolution model will add new insights into the mechanisms that affect stellar evolution. In addition to the long-term radial codes that are capable of modeling the entire stellar life, there are multi-dimensional models [e.g., Viallet et al., 2011]. These models are suited for studying complex mechanisms operating on shorter time scales, for example, the excitation and damping of internal stellar gravity waves [Le Saux et al., 2022]. Similar to complex multi-dimensional disk models, numerical restrictions only allow for a limited simulation time of ~ 10 years.

1.11. Aims of this thesis

In pursuit of understanding the evolution of a star-disk system and the formation of planets, a well-founded theoretical knowledge of the basic properties and scales of the system and the underlying processes is vital. Even at a very early age (Class 0/I), the interaction between the star and its disk can shape the emerging planetary system. Due to the (relatively) long lifetime of PPDs, observations can provide snapshots of these systems at different evolutionary stages and can be used as benchmarks for numerical models. Modeling a complete star-disk system including all scales over its entire lifetime is currently not possible, because they extend over several spatial and temporal timescales and are subject to complex multi-dimensional physical processes. Multi-dimensional complex models for small-scale, short-term processes and simplified long-term codes have been developed. In these models, the different components (the star, the star-disk interaction, and the accretion disk) of the system are treated separately and missing components are often simplified or not considered at all. Thus, back-reactions that might affect the evolution of a star-disk system are not included.

This thesis aims to create the foundation for a coupled, long-term star-disk evolution model that combines the evolution of the stellar and disk components. Focus is laid on the effects of the star-disk interaction in the innermost disk region. By coupling the stellar and disk components, valuable insights into the back-reaction of different mechanisms can be obtained. For example, how does stellar spin influence the accretion disk and the accretion rate onto the star? And how does it affect the stellar evolution? The basis for this coupled star-disk model is an implicit model for the accretion disk including the innermost disk region and capable of modeling the star-disk interaction [e.g., Ragossnig et al., 2020, Steiner et al., 2021]. In its previous versions, this code did not include the evolution of the stellar component and was expanded accordingly in this work. The implicit disk model was augmented with a simplified model for the evolution of the stellar radius and a stellar spin evolution model (see Sec. 2). This allowed to study the effect of different stellar rotation periods on disk evolution is studied. The expanded implicit disk model can be self-consistently coupled to the stellar evolution code *Modules for Experiments in Stellar Astrophysics* [MESA, Paxton et al., 2011, 2019a] and allow for the exploration of the

effect of different metallicities on the disk and stellar spin evolution (see Sec. 3). Furthermore, the governing parameters that control the stellar spin evolution are studied during the early evolutionary stages (≤ 1 Myr) of the star-disk system (see Sec. 4) and during the classical T Tauri phase until the disk dispersion (see Sec. 5). Finally, an outlook is given on what is yet to be done to improve a coupled star-disk evolution model.

Part II.

Main publications

2. Paper I: The importance of stellar rotation in the evolution of protoplanetary disks

- **Title:** Time-dependent, long-term hydrodynamic simulations of the inner protoplanetary disk - II. The importance of stellar rotation
- **Authors:** L. Gehrig, D. Steiner, E. I. Vorobyov, and M. Güdel
- **Journal:** A&A 667, A46 (2022)
- **DOI:** 10.1051/0004-6361/202243549
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Contribution: In this study, I have added a stellar evolution and spin model to the inner disk code. All of the figures and the majority of the text were produced by me. The co-authors have written single paragraphs and have provided feedback in the form of comments.

Summary: The spin evolution of young protostars, surrounded by an accretion disk, still poses problems for observations and theoretical models. In recent studies, the importance of the magnetic star-disk interaction for stellar spin evolution has been elaborated. The accretion disk in these studies, however, is only represented by a simplified model, and important features are not considered. A more realistic representation of the accretion disk is indispensable for a better understanding of the star-disk interaction and the stellar spin evolution. The aim of this study is to investigate the influence of a hydrodynamic disk evolution on the stellar rotational period and vice versa during the accretion phase. We combined the implicit hydrodynamic TAPIR disk code with a stellar spin evolution model. The influence of stellar magnetic fields on the disk dynamics, the radial position of the inner disk radius, as well as the influence of stellar rotation on the disk, were calculated self-consistently. Within a defined parameter space, we can reproduce the majority of fast and slow rotating stars observed in young stellar clusters. Additionally, the back reaction of different stellar spin evolutionary tracks on the disk can be analyzed. Disks around fast-rotating stars are located closer to the star. Consequently, the disk midplane temperature in the innermost disk region increases significantly compared to slow-rotating stars. We can show the effects of stellar rotation on episodic accretion outbursts. The higher temperatures of disks around fast rotating stars result in more outbursts and a longer outbursting period over the disk lifetime. The combination of a long-term hydrodynamic disk and a stellar spin evolution model allows the inclusion of previously unconsidered effects such as the back-reaction of stellar rotation on the long-term disk evolution and the occurrence of accretion outbursts. However, a wider parameter range has to be studied to further investigate these effects. Additionally, a possible interaction between our model and a more realistic stellar evolution code (e.g., the MESA code) can improve our understanding of the stellar spin evolution and its effects on the pre-main sequence star.

2.1. Introduction

Observations of young pre-main sequence (PMS) stars show that the stellar rotation periods range between $\sim 1 - 10$ days, which is significantly lower than the breakup speed [e.g., Irwin and Bouvier, 2009, Gallet and Bouvier, 2013, Serna et al., 2021]. Usually, these stars are surrounded by a protoplanetary disk [e.g., Collier Cameron and Campbell, 1993, Hartmann et al., 1998] and are expected to spin up due to the accretion of disk material containing high angular momentum during the PMS phase, which usually lasts $\lesssim 10$ Myr [e.g., Mamajek, 2009]. Additionally, the contraction of the young star should reduce the rotational period during the first several mega-year. Throughout the PMS phase, observations of stellar rotation lead to the assumption that the stellar period remains approximately constant during the disk phase [e.g., Edwards et al., 1993, Bouvier et al., 1993, Irwin and Bouvier, 2009, Gallet and Bouvier, 2013], implying the interaction between the star and the disk results in efficient braking mechanisms that remove angular momentum from the star [e.g., Serna et al., 2021].

Early studies by Koenigl [1991], based on the model introduced by Ghosh and Lamb [1979b], describe angular momentum transfer via the star-disk interaction taking into account strong stellar magnetic fields. Such magnetic fields with a dipole strength of several kilo Gauß [e.g., Johns-Krull et al., 2009, Johnstone et al., 2014, and references therein] have been observed and the field lines can truncate the disk up to several stellar radii [e.g., Shu et al., 1994b, Romanova et al., 2009, Hartmann et al., 2016]. Along these field lines the star and the disk can exchange angular momentum and the stellar rotational period can be regulated by the accretion disk (disk-locking model). In more recent studies [Matt and Pudritz, 2005a, Zanni and Ferreira, 2013], the disk-locking model is shown not to be efficient enough to actually counteract stellar contraction and disk accretion, which in turn would spin up the star. The region, in which the stellar magnetic field lines connect to the disk is too narrow to extract sufficient angular momentum to match the observed stellar rotation periods.

Alternative braking mechanisms slowing down the star depend on the outflow of material from the star and (or) the disk. First, the so-called "X-wind" [Shu et al., 1994b] extracts material and angular momentum directly from the disk (close to the co-rotation radius) before it can be

accreted onto the star. The amount of angular momentum due to accretion and thus the stellar spin-up is reduced. To launch the "X-wind," no intrinsic, large-scale disk magnetic field is required [e.g., Ferreira et al., 2000] and the field lines are all anchored at the stellar surface. If a large-scale magnetic disk field is present, another kind of material outflow can carry away angular momentum from the star. The nature of this outflow depends on the orientation of the stellar magnetic moment compared to the disk's magnetic moment [Ferreira et al., 2000, Romanova et al., 2009, Zanni and Ferreira, 2013]. In the case, where the stellar and disk's magnetic moment are aligned, the so-called "ReX-wind" can be launched [Ferreira et al., 2000]. At the point inside the disk, the so-called "X-point," where the stellar poloidal field cancels out the disk field, open field lines can reconnect to the closed stellar magnetic field. Matter parameterized by a specific fraction of the disk's accretion rate in Ferreira et al. [2000] can be lifted onto and accelerated by stellar rotation along the open field lines at the "X-point" and thus removing angular momentum from the star. Magnetospheric ejections (MEs) can arise in case of a stellar and disk magnetic moment, which are anti-parallel [Romanova et al., 2009, Zanni and Ferreira, 2013]. A reoccurring process of inflation, opening, and re-connection of field lines leads to an out-bursting behavior, which can transfer angular momentum between the star and the disk. Another model to explain the removal of angular momentum from the star is the accretion-powered stellar wind (APSW) [Matt and Pudritz, 2005b,a, 2008a,b, Matt et al., 2012, Réville et al., 2015b, Finley and Matt, 2018]. The wind is powered by a specific fraction of the disk's accretion rate and accelerated along the open stellar field lines by stellar rotation removing angular momentum from the star.

An important parameter for the aforementioned theories is the disk's accretion rate [e.g., Matt and Pudritz, 2008b]. A precise calculation of the accretion rate including, for example, the influence of stellar magnetic torques and a back-reaction from the influence of stellar rotation on the disk appears rather difficult. Numerical time-step limitations, due to the Courant-Friedrichs-Lewy (CFL) condition [Courant et al., 1928], in the inner regions of the disk prevent the combined, self-consistent simulation of the stellar rotational evolution and hydrodynamic equations describing the protoplanetary disk. To circumvent this problem, previous studies have parameterized the disk's accretion rate by a simple power law or exponential decrease with respect to the disk lifetime [e.g., Matt et al., 2010a, Gallet et al., 2019]. An earlier attempt to combine disk and

stellar spin evolution by Armitage and Clarke [1996] has used the diffusion equation to simulate the disk. While the disk’s surface density can be evolved over time, assumptions have to be made, for example, concerning the angular velocity, which has to be made strictly proportional to the Keplerian angular velocity. Thus, a self-consistent calculation, especially of the inner disk parts is not possible as magnetic torques and local pressure gradient can alter the angular disk velocity [e.g., Romanova et al., 2002, Bessolaz et al., 2008, Steiner et al., 2021] and consequently the accretion process itself.

In this study, we want to continue the work of the first paper of this series. Steiner et al. [2021] have shown the importance of stellar magnetic torques for the long-term evolution of protoplanetary disks. We combine hydrodynamic disk and protostellar spin evolution to study the long-term effects of stellar rotation on the protoplanetary disk. The adaptive, implicit hydrodynamic (TAPIR) code for protoplanetary disks [Ragossnig et al., 2020, Steiner et al., 2021], which is based on Dorfi [1998] and Stoekl and Dorfi [2014], is extended with a stellar evolution model. Following Matt et al. [2010a], this stellar evolution model includes a simplified description of the protostellar contraction (Kelvin-Helmholtz contraction). Additionally, the stellar spin evolution includes contributions from disk matter accreted onto the star, an accretion-powered stellar wind (APSW), as well as magnetospheric ejections (MEs) [e.g., Zanni and Ferreira, 2013, Gallet et al., 2019].

By using a hydrodynamic rather than a diffusion approach, a self-consistent calculation of the innermost disk region is possible. This includes, for example, the radial position of the magnetic truncation radius, stellar magnetic torques, the influence of local pressure gradients, as well as episodic accretion outbursts. Those outbursts can be triggered by gravitational fragmentation in the disk’s outer parts followed by tidal destruction and accretion of inward-migrating clumps [e.g., Vorobyov and Basu, 2015], by thermal instabilities [TI, Bell and Lin, 1994], magneto-rotational instabilities (MRI) [Balbus and Hawley, 1991, Armitage et al., 2001] or a joint interaction of gravitational instability and MRI [Martin and Lubow, 2011, Martin et al., 2012, Martin and Lubow, 2013, Bae et al., 2014]. In this study, we focus on the occurrence of MRI-triggered outbursts. We refer to the first paper of the series [Steiner et al., 2021] for further details. Additionally, the effect of different stellar rotation periods on the disk’s structure and evolution

can be studied.

Due to the implicit nature of the TAPIR code, we can bypass the CFL condition and evolve our model for several mega-year [e.g., Dorfi, 1998]. Similar to Gallet et al. [2019] and Steiner et al. [2021], we focus on Class II star-disk systems, starting our simulations after the cloud infall has ended and the disk mass corresponds to $\lesssim 10\%$ of the stellar mass. Following Gallet et al. [2019], we choose an age of $t_0 = 1$ Myr as a start time for our simulations.

Finally, we note that "X-winds" and "ReX-winds" are not modeled in this study. A suitable expression for "X-winds" is currently not available. The formulation for "ReX-winds" presented in Ferreira et al. [2000] is very similar to the torque exerted by APSW and we follow Gallet et al. [2019] including APSW and omitting "ReX-winds" in our model.

2.2. Model description

In this chapter, the model used in our simulations is outlined. The focus of this study is the combination of a full hydrodynamic disk simulation and a stellar spin evolution model. The disk evolution is treated with the implicit hydrodynamics TAPIR code [Ragossnig et al., 2020, Steiner et al., 2021] and the stellar spin model is formulated in detail in Matt et al. [2010a], Matt et al. [2012] and Gallet et al. [2019]. Here, we want to give a recap of the respective model's key features. In Sec. 2.2.1, Sec. 2.2.2 and Sec. 2.2.3 some key features of the TAPIR code, the treatment of the inner disk boundary are shown and the boundary conditions are outlined, respectively. The stellar spin model is depicted in Sec. 2.2.4.

2.2.1. Hydrodynamic disk evolution with the TAPIR code

Contrary to long-term global disk simulations based on the diffusion evolution approach for the surface density Σ [e.g., Pringle, 1981, Armitage et al., 2001, Zhu et al., 2007, 2010], the TAPIR code allows the treatment of hydrodynamic equations. This includes a deviation from the Keplerian angular velocity $u_\phi \propto r^{-1/2}$. Compared to the model presented in Ragossnig et al. [2020], we added the influence of the stellar magnetic field, which is especially important in the inner part of the disk [e.g., Romanova et al., 2002, Bessolaz et al., 2008]. The disk is

assumed to be axisymmetric in the angular direction and to be in hydrostatic equilibrium in the vertical direction. Our model is formulated in cylindrical coordinates with the planar components (r, φ) , which are then discretized [for a comprehensive description regarding discretization and numerical properties, see Ragossnig et al., 2020]. The protoplanetary disk in our simulations is described in Eqs. (2.1) - (2.3) as a time-dependent, vertically integrated, viscous accretion disk [e.g., Shakura and Sunyaev, 1973, Armitage et al., 2001],

$$\frac{\partial}{\partial t} \Sigma + \nabla \cdot (\Sigma \vec{u}) = 0, \quad (2.1)$$

$$\begin{aligned} \frac{\partial}{\partial t} (\Sigma \vec{u}) + \nabla \cdot (\Sigma \vec{u} : \vec{u}) - \frac{B_z \vec{B}}{2\pi} \\ + \nabla P_{\text{gas}} + \nabla \cdot \mathcal{Q} + \Sigma \nabla \psi + H_P \nabla \left(\frac{B_z^2}{4\pi} \right) = 0, \end{aligned} \quad (2.2)$$

$$\begin{aligned} \frac{\partial}{\partial t} (\Sigma e) + \nabla \cdot (\Sigma \vec{u} e) + P_{\text{gas}} \nabla \cdot \vec{u} \\ + \mathcal{Q} : \nabla \vec{u} - 4\pi \Sigma \kappa_R (J - S) + \dot{E}_{\text{rad}} = 0, \end{aligned} \quad (2.3)$$

where Σ , $\vec{u}$, e , P_{gas} and H_P denote the gas column density, gas velocity in the planar components $\vec{u} = (u_r, u_\varphi)$, specific internal energy density, vertically integrated gas pressure and the vertical scale height of the gas disk, respectively. The gradient in planar cylindrical coordinates reads $\nabla = (\partial/\partial r, r^{-1}\partial/\partial\varphi)$ with $\partial/\partial\varphi = 0$ for our axisymmetric model. P_{gas} is utilized by the ideal equation of state $P_{\text{gas}} = \Sigma e(\gamma - 1)$, with the adiabatic coefficient $\gamma = 5/3$ [e.g., Bessolaz et al., 2008] and κ_R denotes the Rosseland-mean opacity [e.g., Mihalas and Mihalas, 1984], which is composed of a gaseous component $\kappa_{R,\text{gas}}$ [based on Caffau et al., 2011b] and a dust-dominated component $\kappa_{R,\text{dust}}$ [based on Pollack et al., 1985]; with $\kappa_R = \kappa_{R,\text{gas}} + f_{\text{dust}}\kappa_{R,\text{dust}}$, where $f_{\text{dust}} = 0.01$ is the dust-to-gas mass ratio. The gravitational potential of the star-disk system is denoted by ψ . Additionally, the effects of a stellar magnetic field on the disk are included. The vertical component of the magnetic field B_z is assumed to be a dipole and constant within the disk's vertical extent but can change radially. The planar magnetic field components $\vec{B} = (B_r, B_\varphi)$ are taken at the surface of the disk. We note that we neglect the radial component of the magnetic field in this study ($B_r = 0$, see Rappaport et al. 2004 or Kluźniak and Rappaport 2007). The last term on the left-hand side of Eq. (2.2) is the vertically integrated magnetic pressure gradient and

$B_z \vec{B}$ denotes the magnetic stress induced by a Lorentz force [as in Vorobyov et al., 2020b], which exerts a magnetic torque per unit area on the disk $\sim r B_z \vec{B}$ [e.g., Lubow et al., 1994]. The field is assumed to be in rigid rotation with the central star and is able to exchange angular momentum between the star and the disk (Sec. 2.2.4). Viscous angular momentum transport and viscous heating are represented by the viscous stress tensor Q in their corresponding terms in Eq. (2.2) and Eq. (2.3), respectively, which reads as,

$$Q = \frac{\mu}{2} \left[\nabla \vec{u} + (\nabla \vec{u})^T - \frac{2}{3} (\nabla \cdot \vec{u}) \mathbb{1} \right]. \quad (2.4)$$

The factor μ denotes the vertically integrated dynamical viscosity, which yields,

$$\nu = \alpha c_s H_p, \quad (2.5)$$

$$\mu = \nu \Sigma, \quad (2.6)$$

where α , c_s , H_p and ν depict the viscosity parameter [Shakura and Sunyaev, 1973], the speed of sound, the vertical pressure scale height, and the kinematic viscosity, respectively.

The disk is irradiated by the star and cools radiatively, which is described by the net stellar heating and cooling rate per unit surface area $\dot{E}_{\text{rad}}$. $\dot{E}_{\text{rad}}$ is obtained by balancing stellar irradiation, radiative cooling at the disk surface, and irradiation from an ambient medium [see Steiner et al., 2021]. The $(J - S)$ -term in Eq. (2.3) depicts radial radiative transport and is modeled in a radiative diffusion approximation with an Eddington factor of 1/3 [e.g., Ragossnig et al., 2020, Steiner et al., 2021]. We note that disk heating due to magnetic coupling to the star is not included in Eq. (2.3). Work done by the magnetic star-disk interaction is assumed to be comparable to $P\Delta V$ work only close to the inner disk edge as the stellar field strength decreases with r^{-3} (see Sec. 2.2.2). In this region close to the star, however, the disk temperature is dominated by stellar irradiation. If large-scale disk fields are taken into account, disk heating due to magnetic fields can affect the disk's structure and evolution and is added to Eq. (2.3) in future studies (Steiner et al. in prep). Furthermore, resistive dissipation in the disk is assumed to be negligible compared to $P\Delta V$ work [e.g., Miller and Stone, 1997]. Additionally, we utilized the adaptive grid formulation introduced by Dorfi and Drury [1987] to provide adequate radial grid point resolution

and allow the inner radius to move according to the disk evolution (see Sec. 2.2.2). The grid is solved together with Eqs. (2.1) - (2.3) to prevent artificial numerical oscillations [e.g., Dorfi and Drury, 1987, Ragossnig et al., 2020] from developing.

2.2.2. Inner disk boundary

In the innermost disk region, inside the co-rotation radius, the stellar magnetic field decelerates the disk material. As the disk's angular velocity u_ϕ becomes sub-Keplerian, the disk material accelerates radially toward the star [e.g., Romanova et al., 2002, Steiner et al., 2021]. This further leads to a drop in the surface density Σ [e.g., Bessolaz et al., 2008, Steiner et al., 2021] toward the star and strong localized pressure gradients dP_{gas}/dr . The stellar magnetic torque as well as pressure gradients imply a significant deviation from the commonly used 1D diffusion equation for describing the long-term evolution of Σ [e.g., Bell and Lin, 1994, Armitage et al., 2001, Zhu et al., 2009a], therefore it is necessary to fully solve the equations of hydrodynamics. The disk is assumed to eventually be disrupted by a strong stellar magnetic field at the magnetic truncation radius r_{trunc} . Inside r_{trunc} the gas does not behave as an accretion disk but flows along magnetic funnels toward the star [e.g., Romanova et al., 2002, Romanova and Kurosawa, 2014]. This magnetically dominated region cannot be described by a 1D model, which justifies the choice of our inner boundary r_{in} to be at $r_{\text{in}} = r_{\text{trunc}}$. The stellar magnetic field $\vec{B}$ is approximated as a dipole field being in corotation with the star. The vertical component B_z and angular component B_ϕ of $\vec{B}$ are modeled as [e.g., Rappaport et al., 2004, Kluźniak and Rappaport, 2007]

$$B_z(r) = B_\star \left(\frac{R_\star}{r} \right)^3 \quad \text{and} \quad (2.7)$$

$$B_\phi(r) \simeq -\alpha_{\text{cor}} B_z(r) \left[1 - \left(\frac{\Omega_\star}{\Omega(r)} \right)^{\alpha_{\text{cor}}} \right], \quad (2.8)$$

with

$$\alpha_{\text{cor}} \equiv \begin{cases} +1 & \text{if } r < r_{\text{cor}} \\ -1 & \text{if } r > r_{\text{cor}} , \end{cases} \quad (2.9)$$

$$(2.10)$$

as well as the co-rotation radius

$$r_{\text{cor}} = \left(\frac{G M_{\star} P_{\star}^2}{4 \pi^2} \right)^{1/3} . \quad (2.11)$$

Here, $B_{\star}$, $M_{\star}$, $\Omega_{\star}$, $P_{\star}$, and G denote the magnetic field strength at the stellar surface, the stellar mass, the stellar angular velocity, the stellar rotation period, and Newton's gravitational constant, respectively. The factor α_{cor} switches sign at the corotation radius r_{cor} and accounts for the braking region inside r_{cor} (field winds up as a trailing spiral) and the propelling region outside r_{cor} (field winds up as a leading spiral). Additionally, α_{cor} ensures that $|B_{\varphi}| \simeq B_z$ for all radii $r \ll r_{\text{cor}}$ and $r \gg r_{\text{cor}}$, which is a physically motivated model, since a tightly wound up field leads to open field lines due to magnetic buoyancy and therefore limits the maximum value of $|B_{\varphi}|$ [e.g., Lynden-Bell and Pringle, 1974, Rappaport et al., 2004].

The magnetic truncation radius r_{trunc} denotes the point where the magnetic pressure P_{mag} equals either the ram pressure P_{ram} of the in-falling material or the gas pressure P_{gas} , depending on which pressure contribution dominates [e.g., Koldoba et al., 2002, Romanova et al., 2002, Bessolaz et al., 2008],

$$P_{\text{magn}}(r_{\text{trunc}}) = \max \left[P_{\text{ram}}(r_{\text{trunc}}), P_{\text{gas}}(r_{\text{trunc}}) \right] . \quad (2.12)$$

This yields for $r_{\text{in}} = r_{\text{trunc}}$ the following description for the case $P_{\text{ram}} \geq P_{\text{gas}}$ [Hartmann et al.,

2016],

$$r_{\text{trunc}}(P_{\text{ram}} \geq P_{\text{gas}}) \approx 18 \xi R_{\odot} \left(\frac{B_{\star}}{10^3 \text{ G}} \right)^{4/7} \left(\frac{R_{\star}}{2 R_{\odot}} \right)^{12/7} \left(\frac{M_{\star}}{0.5 M_{\odot}} \right)^{-1/7} \left(\frac{\dot{M}_{\star}}{10^{-8} M_{\odot}/\text{yr}} \right)^{-2/7}, \quad (2.13)$$

with the accretion rate onto the star $\dot{M}_{\star}$ and a correction factor $\xi = 0.7$ [following Hartmann et al., 2016] accounting for non-spherical accretion and the oversimplified star-disk interaction. The case $P_{\text{ram}} < P_{\text{gas}}$ follows from equating the magnetic pressure of the vertical field component B_z (Eq. (2.7)) with the gas pressure,

$$r_{\text{trunc}}(P_{\text{ram}} < P_{\text{gas}}) = \frac{B_{\star}^{1/3} R_{\star}}{P_{\text{gas}}^{1/6} (8\pi)^{1/6}}. \quad (2.14)$$

2.2.3. Numerical boundary conditions

The numerical boundary conditions for the hydrodynamical quantities used in the TAPIR code (cp. Eqs. (2.1) - (2.3)) are chosen according to Romanova et al. [2002, 2004] as "free" (Neumann) boundary conditions (e.g., $\partial \Sigma / \partial r = 0$) at the inner boundary. At the outer boundary, the surface density and the internal energy are also treated with "free" boundary conditions, while the angular velocity u_{φ} is fixed to the Keplerian value and radial velocity u_r is set to zero representing no inflow of material from a parent cloud or outer envelope, as usual in late Class II systems. The position of the outer boundary is set to the position where $\Sigma = 0.5 \text{ g/cm}^2$ [similar to Vorobyov et al., 2020b] and is allowed to move during the simulation according to the disk evolution.

2.2.4. Stellar spin model

The spin evolution model in this work considers the torques acting on the young star from external contributors as well as the internal stellar evolution. The temporal derivative of angular momentum equals the sum of external torques $\dot{J}_{\star} = \Gamma_{\text{ext}}$. Following Matt et al. [2010a], we assume rigid body rotation with a moment of inertia $I_{\star} = k^2 M_{\star} R_{\star}^2$ with the normalized radius of gyration k [$k^2 = 0.2$, e.g., Armitage and Clarke, 1996, Matt et al., 2010a, Matt et al., 2012]. The

temporal derivative of the stellar angular momentum $J_\star = I_\star \Omega_\star$ reads

$$\dot{J}_\star = I_\star \dot{\Omega}_\star + \dot{I}_\star \Omega_\star = \Gamma_{\text{ext}}. \quad (2.15)$$

The evolution of the stellar spin results from rewriting Eq. (2.15) as follows

$$\dot{\Omega}_\star = \frac{\Gamma_{\text{ext}}}{I_\star} - \frac{\dot{I}_\star}{I_\star} \Omega_\star, \quad (2.16)$$

with

$$\frac{\dot{I}_\star}{I_\star} = \frac{\dot{M}_\star}{M_\star} + 2 \frac{\dot{R}_\star}{R_\star}. \quad (2.17)$$

Following the model of Gallet et al. [2019], the external torque Γ_{ext} acting on the star consists of three components; $\Gamma_{\text{ext}} = \Gamma_{\text{acc}} + \Gamma_{\text{w}} + \Gamma_{\text{ME}}$. The torques resulting from accreting disk material (gas and dust) onto the star Γ_{acc} , stellar winds and outflows Γ_{w} and magnetospheric ejections Γ_{ME} are described in this section. As we focus on young class II disk systems, we assume a fully convective star. Thus, internal torques resulting from the interaction between a radiative core and a convective envelope [as described in Gallet et al., 2019] are not considered in this work. The change of the stellar moment of inertia due to the contraction of the stellar radius is included in our model.

Stellar radius evolution

We assume a fully convective young star [e.g., Matt et al., 2010a, Pantolmos et al., 2020] contracting during the pre main-sequence until a radiative core develops [Siess et al., 2000]. During this phase, the stellar radius evolution can be approximated by a Kelvin-Helmholtz contraction [e.g., Collier Cameron and Campbell, 1993, Matt et al., 2010a, Matt et al., 2012] resulting in

$$\dot{R}_\star = 2 \frac{R_\star}{M_\star} \dot{M}_\star - \frac{28\pi\sigma R_\star^4 T_{\text{eff}}^4}{3GM_\star^2}, \quad (2.18)$$

with the stellar effective temperature T_{eff} and the Stefan-Boltzmann constant σ . The first term on the left-hand side of Eq. (2.18) accounts for accreted disk material. The second term represents

the contraction on the Kelvin-Helmholtz timescale.

We set the initial stellar mass of the star to $M_{\star} = 0.7 M_{\odot}$. A broader range of stellar masses is included in future studies. The respective $R_{\star, \text{init}}$ and $T_{\text{eff, init}}$ are taken from the isochrones of Baraffe et al. [2015]. The stellar radius then evolves according to Eq. (2.18) and the effective temperature updates according to the results of Baraffe et al. [2015] with time. We want to note that this stellar evolution model is a simple approximation as several important aspects of the pre-main-sequence evolution are ignored [e.g., Deuterium burning slows down the stellar contraction; Siess et al., 2000, Matt et al., 2010a]. The maximum difference between Eq. (2.18) and the results in Baraffe et al. [2015] is $< 2.5 \%$ for a $0.7 M_{\odot}$ star during the first 10 Myr of its lifetime, assuming no accretion.

Magnetospheric ejections

In star-disk systems, with a stellar magnetic field strong enough to connect to the disk material, angular momentum can be transferred between the star and the disk by magnetospheric ejection (ME) [Zanni and Ferreira, 2013], which is nicely summarized in Gallet and Bouvier [2013]. Differential rotation between the star and the disk generates a toroidal magnetic field. The resulting magnetic pressure inflates the magnetic field lines, followed by a re-connection and contraction [see Fig. 2 of Zanni and Ferreira, 2013]. The magnetic torque felt by an annulus (Δr) of the disk [e.g., Armitage and Clarke, 1996, Matt and Pudritz, 2005b, Gallet et al., 2019, Ireland et al., 2021] is given by

$$\Delta\Gamma_{\text{ME}} = qr^2 B_z^2 \Delta r, \quad (2.19)$$

with the twist of magnetic field lines

$$q \propto \frac{B_{\varphi}}{B_z} \sim \left[1 - \left(\frac{r_{\text{trunc}}}{r_{\text{cor}}} \right)^{3/2} \right]. \quad (2.20)$$

The additional assumption that the radius of the annulus of the disk connected to the MEs is of

the order of r_{trunc} [Gallet et al., 2019] yields the total torque felt by the star due to MEs,

$$\Gamma_{\text{ME}} = K_{\text{ME}} \frac{B_{\star}^2 R_{\star}^6}{r_{\text{trunc}}^3} \left[K_{\text{rot}} - \left(\frac{r_{\text{trunc}}}{r_{\text{cor}}} \right)^{3/2} \right], \quad (2.21)$$

with $K_{\text{ME}} = 0.21$ and $K_{\text{rot}} = 0.7$. This parameter choice represents the fact that the MEs rotate at a sub-Keplerian rate [see Gallet et al., 2019]. We want to note that Pantolmos et al. [2020] found smaller values for $K_{\text{ME}} = 0.13$ and $K_{\text{rot}} = 0.46$. The differences are explained by slightly different boundary definitions for the stellar wind and MEs. Furthermore, Ireland et al. [2021] include multi-dimensional and non-linear effects to the calculation of Γ_{ME} and also derive lower values for the constant of proportionality compared to Gallet et al. [2019]. In this study, we use the values given by Gallet et al. [2019], which can be seen as upper limits for Γ_{ME} .

The effect of the star-disk interaction depends on the position of the inner radius r_{trunc} in relation to the corotation radius r_{cor} . If $r_{\text{trunc}} < K_{\text{rot}}^{2/3} r_{\text{cor}}$, the stellar angular momentum increases and the star spins up, otherwise the star spins down. As argued in Zanni and Ferreira [2013], MEs reduce the effect of the accretion torque Γ_{acc} . In Gallet et al. [2019], a constant $K_{\text{acc}} < 1$ qualitatively represents this reduction of Γ_{acc} (see Sec. 2.2.4). K_{acc} serves as an approximation to quantify small-scale, time-dependent magnetically mediated mass transfer between the star and the disk. Following Gallet et al. [2019], we choose $K_{\text{acc}} = 0.4$ for all our simulations. As shown in Romanova et al. [2009], the process is very dynamic and thus cannot be adequately described in a 1D model. The following approximations should be seen as time-averaged values. We refer the reader to Romanova et al. [2009] and Zanni and Ferreira [2013] for a more elaborated view of the topic.

Accretion

During the accretion of material onto the young star, mass, and angular momentum are added to the central object. The accretion torque Γ_{acc} is given by the product of the specific angular momentum of disk material at the inner edge of the disk and the accretion rate;

$$\Gamma_{\text{acc}} = K_{\text{acc}} \dot{M}_{\star} r_{\text{trunc}}^2 \Omega_{\text{disk,in}} \quad (2.22)$$

with $r_{\text{trunc}}\Omega_{\text{disk,in}} = u_{\varphi,\text{in}}$.

Stellar outflow

In a magnetized stellar outflow, the material is transported along magnetic field lines. These field lines are rooted on the star and rotate with $\Omega_\star$. As the wind material is moving outwards, it magneto-centrifugally accelerates and thus removes angular momentum from the star.

In this study, we consider the APSW [Matt and Pudritz, 2005b], which takes its driving force from the accreting disk material. Consequently, the mass loss due to this wind is a fraction W of the disk accretion rate; $\dot{M}_w = W\dot{M}_{\text{disk}}$. The torque, caused by a stellar outflow [Weber and Davis, 1967], acting on the star, can be written as

$$\Gamma_w = \dot{M}_w r_A^2 \Omega_\star, \quad (2.23)$$

with the Alfvén radius

$$r_A = K_1 \left[\frac{B_\star^2 R_\star^2}{\dot{M}_{\text{wind}} \sqrt{K_2^2 v_{\text{esc}}^2 + \Omega_\star^2 R_\star^2}} \right]^m R_\star, \quad (2.24)$$

where $v_{\text{esc}} = \sqrt{2GM_\star/R_\star}$, $m = 0.2177$, $K_2 = 0.0506$ and $K_1 = 1.7$ arising as best fit values from magneto-hydrodynamic stellar wind simulations [see Matt et al., 2012b, Gallet et al., 2019]. We want to note that Γ_w acts as an angular momentum "sink" in our model. The exact driving mechanisms of the APSW and the order of W are still under debate (we refer to the discussion in Gallet et al. [2019] for further details).

Numerical integration

To include the spin evolution model in the TAPIR code, we updated the stellar parameters after each timestep Δt as follows. The current accretion rate resulting from solving the disk equation Eqs. (2.1) - (2.3) updates the stellar mass and radius. With the updated values the respective torque contributions acting during this timestep are calculated according to the formulation above. Finally, $\dot{\Omega}$ and the new stellar angular velocity $\Omega_{\text{new}} = \Omega_{\text{old}} + \dot{\Omega}\Delta t$ is derived and used in the next timestep. Ω_{old} and Ω_{new} denote the stellar angular velocity used for the current and the new

timestep, respectively.

2.3. Results

In this section, we want to define the parameter space and a reference model (Sec. 2.3.1), followed by the results of our simulations: the back-reaction of different stellar rotation rates on the disk (Sec. 2.3.4), the construction of a fast and a slow rotating protostar within our parameter space (Sec. 2.3.3), as well as the influence of stellar rotation on episodic accretion outburst events (Sec. 2.3.5).

2.3.1. Parameter space

We define a reasonable parameter space for stellar and disk properties covered in our simulations, which is based on observations and previous theoretical models. First, we define the stellar parameters appropriate for a 1 Myr-old T Tauri star. As described in Sec. 2.2, the initial stellar mass is $M_{\star} = 0.7M_{\odot}$ and the initial stellar radius, as well as the effective temperature, are taken from isochrones provided by Baraffe et al. [2015]. The initial stellar rotation period for a 1 Myr-old star can vary significantly. We choose a range for the initial period from 2 to 15 days. This choice is motivated by observations of the Orion Nebular Cluster (ONC) [Herbst et al., 2002]. The stellar dipole magnetic field strength has to be estimated. Measurements by Johnstone et al. [2014] suggest that the range of the dipole component of the magnetic field can range from several hundred Gauss to $\sim$ kG. Furthermore, M dwarfs can have very large magnetic field strengths of several thousand Gauss [e.g., Kochukhov and Reiners, 2020]. Based on these measurements we assume $B_{\star}$ to be in a range between 500 and 2000 G. The magnetic field strength remains constant during our simulations. We note that T Tauri stars are usually magnetically saturated [Rossby number $\lesssim 0.1$, e.g., Briggs et al., 2007, Johnstone et al., 2021] and the stellar magnetic field is roughly independent of the rotational period [e.g., Reiners et al., 2009, Lavail et al., 2019]. During later evolutionary stages, this relation is reverted and fast-rotating stars show more magnetic activity up to the state of magnetic saturation [e.g., Wright et al., 2011]. Finally, the efficiency W of the APSW has to be chosen. We choose three different values for W . Based

on the theoretical model by Cranmer [2008] and simulations by Pantolmos et al. [2020], we choose $W = 2\%$ as a reference value. $W = 0\%$ and $W = 5\%$ represent a weak and strong APSW, respectively.

The hydrodynamic disk model requires an initial disk accretion rate as well as a description of the viscous α parameter. Based on observations [e.g., Gullbring et al., 1998, Manara et al., 2016] and reproduced by numerical simulations [e.g., Vorobyov et al., 2017b], the accretion rate from the disk onto the star ranges between 10^{-10} and $10^{-7} \text{ M}_\odot \text{ yr}^{-1}$ for late-type star-disk systems. We therefore choose the initial accretion rate $\dot{M}_{\text{init}} = 1 \cdot 10^{-8} \text{ M}_\odot/\text{yr}$. The effects of a variation of $\dot{M}_{\text{init}}$ are covered in a further study. The viscous α -parameter ranges between $\lesssim 0.1$ and ~ 0.0001 [e.g., Zhu et al., 2007, 2009b, Vorobyov and Basu, 2009, Yang et al., 2018]. We adopt initially a constant value of $\alpha = 0.005$ used in Sec. 2.3.4 and Sec. 2.3.3 unless otherwise stated.

To generate episodic accretion outbursts (Sec. 2.3.5) due to magneto-rotational instabilities (MRI) [Balbus and Hawley, 1991], we adapt the layered disk model [Gammie, 1996b]. A detailed description of this model in the TAPIR code is presented in Steiner et al. [2021]. In the layered disk model, the disk is divided into two layers. On the one hand, we assume an MRI active layer, which is sufficiently ionized due to stellar irradiation, cosmic rays, or collisional ionization of, for example, potassium [Gammie, 1996b] with an effective viscosity α_{MRI} . On the other hand, we have an MRI inactive layer, the dead zone (DZ), with $\alpha_{\text{DZ}} \ll \alpha_{\text{MRI}}$, in which the mass transport rate is strongly diminished. This leads to mass piling up in the DZ up to a point where the increasing mass leads to increasing temperature and eventually to the onset of the MRI above a critical temperature T_{crit} . The turbulent viscosity and viscous heating increase sharply, which further causes ionization of the whole inner region and a highly increased mass transport rate $\dot{M}$. This ionized inner region is then accreted in a strong accretion event onto the star and the inner disk is replenished by material from the outer disk, which completes a burst cycle.

Depending on the choice of the previously mentioned parameters, the evolution of the accretion disk and therefore the spin evolution of the star can vary. One example is the viscosity parameter α that influences the viscous timescale of the accretion process; $\tau_v = r^2/\nu \sim 1/\alpha$. A large (small) α -value corresponds to short (long) disk lifetime, respectively, and the accretion rate changes accordingly [e.g., Vorobyov and Basu, 2009]. In our model, we can include these effects self-

consistently. In Fig. 2.1, the effect of different α -values on the disk evolution is shown. The red, black, and blue lines show the temporal evolution of the disk's accretion rate $\dot{M}_{\text{disk}}$ for a constant α -value of 0.01, 0.005, and 0.0025, respectively. The time for $\dot{M}_{\text{disk}}$ to drop differs by mega-years, which has an effect on the stellar spin evolution. Furthermore, our model enables the consideration of episodic accretion outbursts [see Steiner et al., 2021], based on the layered disk model [Gammie, 1996b] with $\alpha_{\text{MRI}} = 0.02$ and $\alpha_{\text{DZ}} = 0.0004$ [similar to Steiner et al., 2021]. During such an outburst, the accretion rate can rise for orders of magnitude (yellow line in Fig. 2.1), and the disk's structure and temperature change significantly. These effects are also calculated in a self-consistent manner. Previous stellar spin evolution studies tend to use a simplified description of the disk's accretion rate. Either an exponential decay [e.g., Matt et al., 2010a] or a power law description [Gallet et al., 2019] is used, which require free parameters to define for example the disk lifetime. Additionally, accretion outbursts as well as the back-reactions of such events on the disk cannot be included in these models.

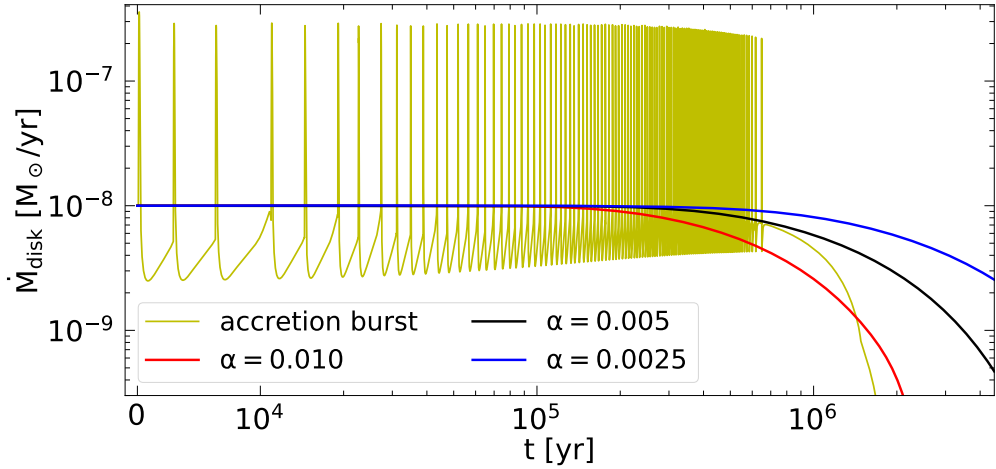


Figure 2.1.: Disk accretion rate $\dot{M}_{\text{disk}}$ with respect to the simulation time t in years starting from $t_0 = 1$ Myr. The red, black, and blue lines correspond to a constant α value of 0.01, 0.005, and 0.0025, respectively. The yellow line corresponds to a layered disk model [Gammie, 1996b] with $\alpha_{\text{MRI}} = 0.02$ and $\alpha_{\text{DZ}} = 0.0004$ [similar to Steiner et al., 2021].

2.3.2. Reference model

Based on the stellar and disk parameter ranges, defined in Sec. 2.3.1, we define a reference model for both a constant α -value and a layered disk model.

Constant α -value

The stellar and disk parameters for the reference model using a constant α -value (**Mref**) are summarized in Table 2.1. The stellar spin evolution for this model is shown in Fig. 2.2. During the simulation time of 5 Myr, the stellar rotation period $P_\star$ decreases slightly from the initial value of 6 toward 4 days (see Panel (a)). The individual external contributions to the total torque Γ_{tot} are shown in Panel (b) (black line). A positive (negative) torque value adds (removes) angular momentum to the star and the star spins up (spins down), respectively. The accretion torque Γ_{acc} (blue line) spins up the star during the whole simulation time, whereas the APSW (Γ_W , yellow line) spins down the star. If MEs (Γ_{ME} , red line) spin the star up or down, depends on the position of the inner disk radius r_{trunc} with respect to the co-rotation radius r_{cor} (see Eq. (2.21)), which is shown in Panel (c). For comparison, the value $K_{\text{rot}}^{2/3}$ is indicated by the dashed black line.

During the first 2×10^5 yr, r_{trunc} is close to r_{cor} and $r_{\text{trunc}}/r_{\text{cor}} > K_{\text{rot}}^{2/3}$. Γ_{ME} is negative and the MEs remove angular momentum from the star. The influence of the APSW and the ME, however, are not sufficient to counteract accretion and contraction and the star spins up slightly ($\Gamma_{\text{tot}} > 0$ after 10^4 years). While the star contracts and the stellar radius as well as the inner disk radius decreases, the co-rotation radius resides at approximately the same positions as the stellar rotational period and only changes slightly during the first couple of 10^5 years. Thus, $r_{\text{trunc}}/r_{\text{cor}}$ decreases until $r_{\text{trunc}}/r_{\text{cor}} < K_{\text{rot}}^{2/3}$ and the MEs also spin up the star. After 2 Myr of accretion, the disk is nearly depleted and the disk accretion rates drop (see Fig. 2.1). Consequently, r_{trunc} moves outwards and the MEs remove angular momentum from the star again.

Table 2.1.: Reference model parameters with a constant α -value

$B_\star$ [kG]	1.00	$R_{\star, \text{init}}$ [$R_\odot$]	2.096
P_{init} [days]	6.0	$M_{\star, \text{init}}$ [$M_\odot$]	0.7
$\dot{M}_{\text{init}}$ [$M_\odot/\text{yr}$]	10^{-8}	$T_{\text{eff, init}}$ [K]	4078
W [%]	2	α	0.005

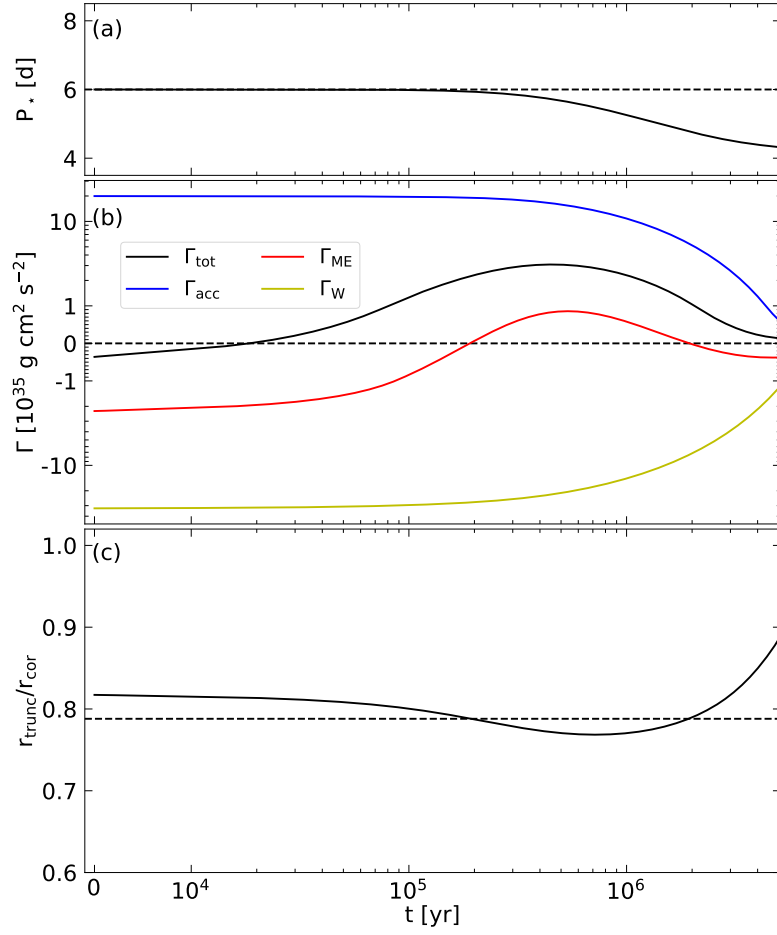


Figure 2.2.: Stellar spin evolution for **Mref** with respect to the simulation time t in years starting from $t_0 = 1$ Myr. Panel (a) shows the stellar rotation period $P_\star$ in days. The dashed black line indicates the initial period of 6 days. Panel (b) shows the external torque contributions (see Sec. 2.2.4) as well as the total torque acting on the star (black line). The blue, red, and yellow lines represent the accretion torque Γ_{acc} , torque due to MEs Γ_{ME} and the APSW torque Γ_{W} , respectively. The dashed line at $\Gamma = 0.0$ indicates the boundary between a spin-up and a spin-down contribution. Panel (c) shows the position of the inner disk radius r_{trunc} with respect to the position of the co-rotation radius r_{cor} . The dashed line indicate $K_{\text{rot}}^{2/3}$.

Layered disk model

The reference parameters (**MrefB**) used in combination with the layered disk model differ from **Mref** only by the viscous α value. Instead of a constant α -value trough out the disk, we use

$\alpha_{\text{MRI}} = 0.02$ and $\alpha_{\text{DZ}} = 0.0004$ [similar to Steiner et al., 2021]. The effect of different stellar spin evolutionary tracks based on this reference model is shown in Sec. 2.3.5. In Fig. 2.3, a detailed view of accretion bursts and the respective contributions to the total torque acting on the star is shown. In Panel (a) the disk's accretion rate is shown. Once the disk temperature reaches $T_{\text{crit}} = 1000$ K and the MRI is triggered, the disk's accretion rate $\dot{M}_{\text{disk}}$ increases sharply over an order of magnitude until the material piled up in the DZ is depleted and $\dot{M}_{\text{disk}}$ decreases again until the next outburst is triggered. In Panel (b) the position of the inner radius r_{trunc} with respect to the co-rotation radius r_{cor} is shown. During the time of increased $\dot{M}_{\text{disk}}$, r_{trunc} is pushed toward the star (see Eq. (2.13)). This again affects the total torque acting on the star (black line in Panel (c)). During the first several 10^1 yr, the total as well as the external torque contributions are similar to the model **Mref** and the star spins down, which is indicated by $\Gamma_{\text{tot}} < 0$. When the accretion outburst starts, Γ_{acc} (Γ_{W}) increases (decreases) due to the higher accretion rate, respectively, without changing their sign. The contribution of the MEs to Γ_{tot} , however, changes significantly. Before the outburst, r_{trunc} is close to r_{cor} and $\Gamma_{\text{ME}} < 0$ resulting in a stellar spin down effect. During the burst the distance between r_{trunc} and r_{cor} increases drastically and $r_{\text{trunc}}/r_{\text{cor}} < K_{\text{tot}}^{2/3}$. Now, the MEs spin up the star and the APSW alone cannot counteract the combined spin-up torque consisting of Γ_{acc} , Γ_{ME} , and the stellar contraction. As a result, during the accretion outburst, spins up the star. While the peak accretion rate of the outburst in Fig. 2.3 is similar to EXor outbursts, the duration is reminiscent of FUor outbursts [e.g., Audard et al., 2014]. We note that the peak and duration of these outbursts can be influenced by the choice of α_{MRI} , α_{DZ} and T_{crit} [e.g., Steiner et al., 2021] and the effects of different outburst types are investigated in further studies.

2.3.3. Fast and slow rotating stars

Within the predefined parameter space (Sec. 2.3.1), we define models for a fast and slow rotating star to test the limits of our approach and compare the results to observations of stellar rotation periods of young clusters. Based on our reference model **Mref**, we change the stellar magnetic field strength $B_{\star}$, the APSW value W , and the initial stellar period P_{init} . The fast (slow) rotation model **Mfast** (**Mslow**) has a stellar magnetic field strength of $B_{\star} = 500$ (2000) G, $W=0$ (5)% and

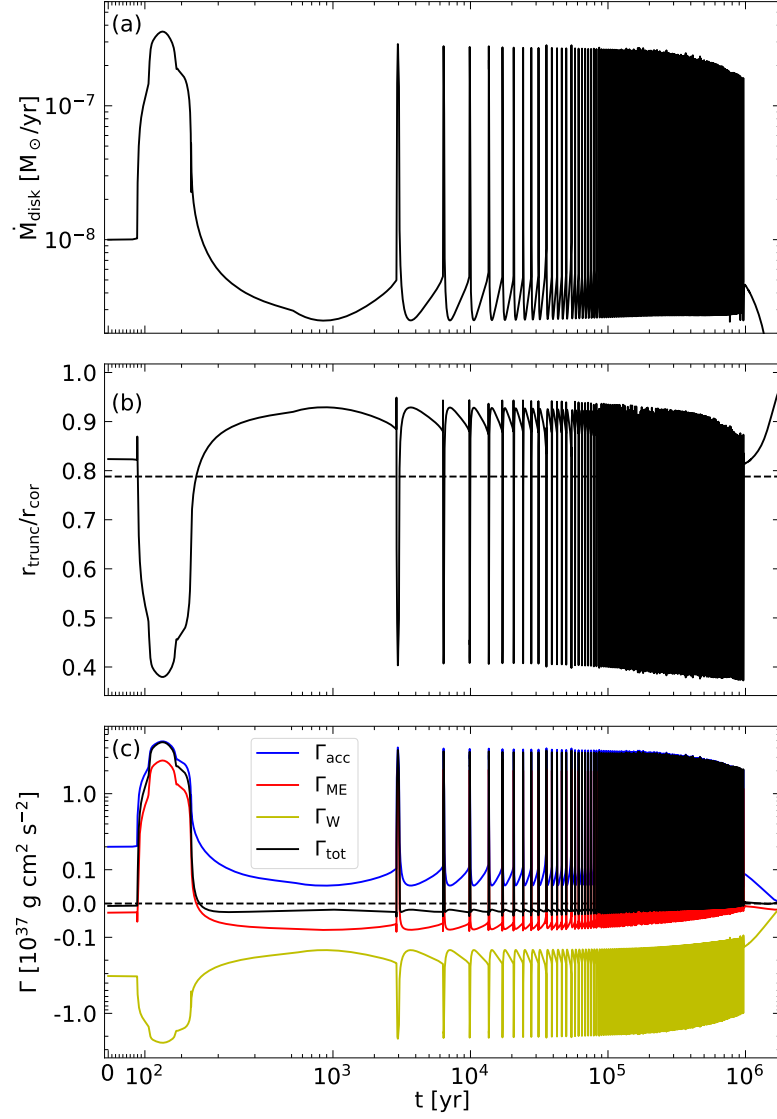


Figure 2.3.: Stellar spin evolution for **MrefB** with respect to the simulation time t in years starting from $t_0 = 1$ Myr. Panel (a) shows the disk’s accretion rate $\dot{M}_{\text{disk}}$. Panel (b) shows the position of the inner disk radius r_{trunc} with respect to the positions of the co-rotation radius r_{cor} . The dashed line indicate $K_{\text{rot}}^{2/3}$. Panel (c) show the total torque acting on the star Γ_{tot} as well as the external torque contributions, which are the same as in Fig. 2.2. Additionally, the dashed line at $\Gamma = 0.0$ indicate the boundary between a spin-up and a spin-down contribution.

$P_{\text{init}} = 2$ (15) days, respectively¹.

¹Following previous studies [Matt et al., 2012, Gallet et al., 2019], a weak stellar magnetic field results in a fast rotating star due to the weaker star-disk interaction and a less effective APSW.

In Fig. 2.4, the rotational periods of M_{ref} (solid line), M_{fast} (dashed line), and M_{slow} (dash-dotted line) are shown. The stellar rotation periods of four young clusters are also shown to compare the results to the observations. The Orion Nebular Cluster (ONC) with an age of 1 Myr [Herbst et al., 2002] was already used to define the initial parameters used in this study and is located at $t = 0$ yr. The following clusters from left to right are NGC 6530 [Henderson and Stassun, 2012] with an age of 1.65 Myr, NGC 2264 [Affer et al., 2013] with an age of 2 Myr and NGC 2362 [Irwin et al., 2008] with an age of 5 Myr. The solid black circles and the downward (upward) pointing triangles represent the 50th and 95th (5th) percentiles, respectively. The reference model M_{ref} remains close to the 50th percentile value (mean value) of all four clusters. Additionally, the majority of stellar rotation periods in NGC 6530, NGC 2264, as well as NGC 2362 are enclosed between M_{fast} and M_{slow} , which is visualized by the blue area between the respective curves. The spin evolution of M_{fast} and M_{slow} depends on the choice of $B_{\star}$ and W^2 . A weaker (stronger) magnetic field strength results in a shift toward shorter (longer) periods. In a similar way, a smaller (larger) APSW value W results in shorter (longer) periods. The order of magnitude of this dependence is symbolized by model $M_{\text{ref}W5}$ (solid blue line in Fig. 2.4). Reducing $B_{\star}$ by a factor of 2 spins up the star by 92% after 5 Myr (compare models M_{slow} and $M_{\text{ref}W5}$). An increase of W from 2% to 5% spins down the star by 60% after 5 Myr (compare models M_{ref} and $M_{\text{ref}W5}$).

2.3.4. Back-reaction on the disk

In this section, we show the effects stellar rotation has on the protoplanetary accretion disk. First, the effects of different spin evolutionary tracks are analyzed (Sec. 2.3.4), by comparing a star that spins up with a star that spins down. Second, we show the effects that different initial rotation periods have on the accretion disk (Sec. 2.3.4).

The effect of different stellar spin evolutionary tracks

To study the effects a star that spins up has on the accretion disk, compared with a star that spins down, we change the APSW efficiency value W . While the initial stellar and disk properties are

²The influence of the initial stellar period is studied in Sec. 2.3.4.

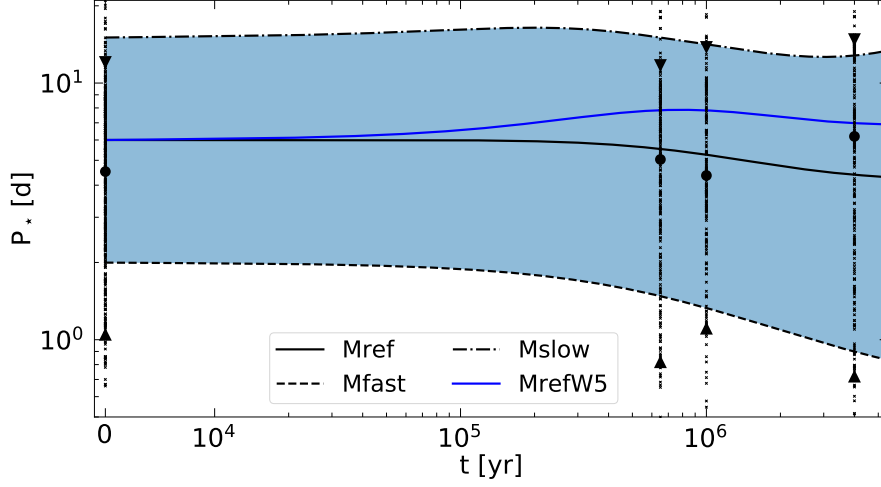


Figure 2.4.: Evolution of the stellar rotational period P_* with respect to the simulation time t in years starting from $t_0 = 1$ Myr for the models Mref (solid line), Mfast (dashed line) and Mslow (dash-dotted line). Model MrefW5 (solid blue line) highlights the dependence on B_* and the APSW value W (see text). The area between Mfast and Mslow is colored for better visualization. The observational data from four young stellar clusters are shown as "crosses" with the 50th percentile (big circle) and the 95th (5th) percentile (downward and upward pointing triangles). From left to right the open clusters are: Orion Nebular Cluster (ONC) [Herbst et al., 2002], NGC 6530 [Henderson and Stassun, 2012], NGC 2264 [Affer et al., 2013] and NGC 2362 [Irwin et al., 2008].

identical the APSW removes more (less) angular momentum from the star that spins down (up), respectively. Thus, the effect of different stellar spin evolutionary tracks on the accretion disk, starting from the same initial model, can be shown.

We choose, based on the reference value of $W = 2\%$, $W = 0\%$, as well as $W = 5\%$ representing a star that spins up and down, respectively. In Fig. 2.5, we show the resulting evolution of the stellar rotational period P_* (Panel (a)). The model Mref (red line) evolves as already shown in Fig. 2.2. As expected, a low APSW value $W = 0\%$ (black line) does not remove enough angular momentum from the star to counteract accretion and contraction and the star spins up during the whole simulation time. A large APSW value $W = 5\%$ (blue line) removes considerably more stellar angular momentum and the star spins down toward 8 days. The resulting effects of these different stellar spin evolutionary tracks on the accretion disk are shown in Panels (b) and (c). In Panel (b) the surface density distributions of the innermost disk part (< 0.5 AU) are shown for

different time snapshots: the initial model at $t = 0.0$ Myr (solid line), $t = 0.1$ Myr (dashed lines) and $t = 1.0$ Myr (dotted lines). Additionally, the co-rotation radii are shown for the respective time snapshots (vertical lines). Starting from the same initial model (the blue and black solid lines are identical), the surface density structure of the inner disk part differs clearly when changing the APSW value from 0 to 5%. Responsible for this difference is the position of the co-rotation radius r_{cor} . For a star that spins down (blue lines), the co-rotation radius moves outwards away from the star, and a larger part of the disk is located inside r_{cor} . Disk material inside r_{cor} rotates faster compared to the star and is slowed down due to magnetic torques. Thus, the disk's angular velocity (u_ϕ) decreases below the Keplerian velocity and the disk material inside r_{cor} accelerates in the radial direction toward the star [similar to Steiner et al., 2021]. For a star that spins up, r_{cor} moves inward toward the star. The disk material that is now located outside r_{cor} rotates slower compared to the star and is accelerated by magnetic torques to velocities faster than Keplerian. As a result, the radial inward velocity of the disk outside r_{cor} is decreased. Material transported from the outer parts toward the inner disk region, is piled up outside r_{cor} , pushing the inner disk radius r_{trunc} closer to the star.

With the position of r_{cor} also the position of the inner disk radius r_{trunc} changes. With r_{trunc} moving closer to or away from the star, the inner disk is heated differently due to stellar irradiation. This effect is shown in Panel (c). The maximum disk midplane temperature $T_{\text{disk, max}}$ shows a different behavior depending on the stellar spin evolutionary track. As expected, $T_{\text{disk, max}}$ decreases for models that show a stellar spin down (blue and red lines). For the model with $W = 0\%$ showing a stellar spin up, $T_{\text{disk, max}}$ increases for 1.0 Myr and differs by over 300 K compared to the model with $W = 5\%$. Such a difference can have a considerable effect on the disk's ionization fraction that is strongly temperature dependent in these regions of $T_{\text{disk, max}}$ [e.g., Fromang et al., 2002, Vorobyov et al., 2020b] as well as on the occurrence of episodic accretion outbursts.

The effect of different initial stellar periods

When changing the initial stellar rotation period P_{init} , we expect the period to converge toward a common value in the course of $\sim$ Myr as seen in Armitage and Clarke [1996] or Matt et al.

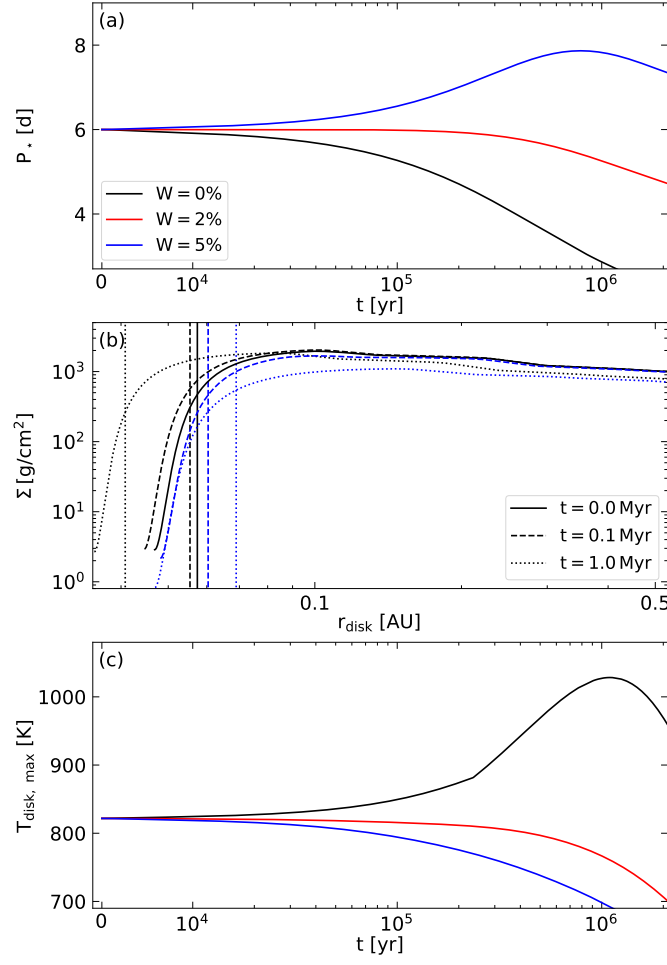


Figure 2.5.: Panel (a) shows the evolution of the stellar rotational period $P_\star$ with respect to the simulation time t in years starting from $t_0 = 1$ Myr. The reference model Mref with $W = 2\%$, the models with $W = 0\%$, and $W = 5\%$ are given by the red, black, and blue lines, respectively. Panel (b) shows the surface density of the innermost disk part (0.5 AU) for the models with $W = 0\%$ (black lines) and $W = 5\%$ (blue lines) at three different time snapshots: The initial model at $t = 0.0$ Myr (solid line), $t = 0.1$ Myr (dashed lines) and $t = 1.0$ Myr (dotted lines). The vertical lines represent the position of the co-rotation radius for the given APSW efficiency value and time snapshot. Note that the black and blue solid lines are identical as the same initial model ($t = 0.0$ Myr) is used for this comparison. Panel (c) shows the evolution of the maximum disk temperature $T_{\text{disk, max}}$ at the disk's midplane with respect to the simulation time t in years starting from $t_0 = 1$ Myr. The line colors represent the same models as in Panel (a).

[2010a]. In Fig. 2.6, we compare the stellar spin evolution for initial stellar periods ranging from 2 to 15 days. Panel (a) shows the rotation period for $P_{\text{init}} = 2$ days (red line), $P_{\text{init}} = 4$ days (magenta line), $P_{\text{init}} = 6$ days (black line), $P_{\text{init}} = 10$ days (blue line) and $P_{\text{init}} = 15$ days (yellow line). While the stellar periods approach each other, we still see a spread of 2 days after 2 Myr. The reason for this difference is the back-reaction of different stellar spin evolutionary tracks on the accretion disk, which is explained in Sec. 2.3.4.

The maximum disk midplane temperature $T_{\text{disk, max}}$ for each model is shown in Panel (b). Similar to Fig. 2.5, a fast-rotating star with an inner disk and co-rotation radius close to the star has a higher disk temperature. $T_{\text{disk, max}}$ varies by a factor of 2 for the models shown. Again, such a difference strongly affects the disk's ionization as well as influences the occurrence of accretion outbursts.

2.3.5. Effect on accretion outbursts

Finally, we want to show the effects of different stellar spin evolutionary tracks on the occurrence of accretion outbursts using the layered disk reference model **MrefB**. In Fig. 2.7, similar to Sec. 2.3.4, different stellar spin evolutionary tracks are shown. Starting from a common initial rotational period of 6 days, the stellar rotational period (Panel (a)) of the reference model **MrefB** (red line) does not change significantly over a simulation time of 1 Myr, whereas an APSW value of $W = 0\%$ results in a stellar spin up (black line) and $W = 5\%$ results in a slowly rotating star (blue line). In Panel (b), the respective disk accretion rates are shown. Similar to the models in Steiner et al. [2021], outbursts occur over a time span of $\lesssim 1$ Myr. The number of bursts as well as the time span on which outbursts are triggered depend on the stellar spin evolution. Fast-rotating stars tend to have more mass located closer to the star over longer periods of time (see Fig. 2.5). Thus, the disk temperature remains higher within these disks and the conditions that trigger an outburst are also met over longer periods of time. This results in a longer period as well as a larger number of outbursts for stars that spin up and rotate faster. In the model **MrefB** (red line in Fig. 2.7) 231 outbursts occur over a time span of 0.96 Myr. In the model with $W = 5\%$, representing a star that spins down, 85 outbursts occur over a time span of 0.53 Myr, whereas in the model with $W = 0\%$, representing a fast rotating star, 835 outbursts occur over a

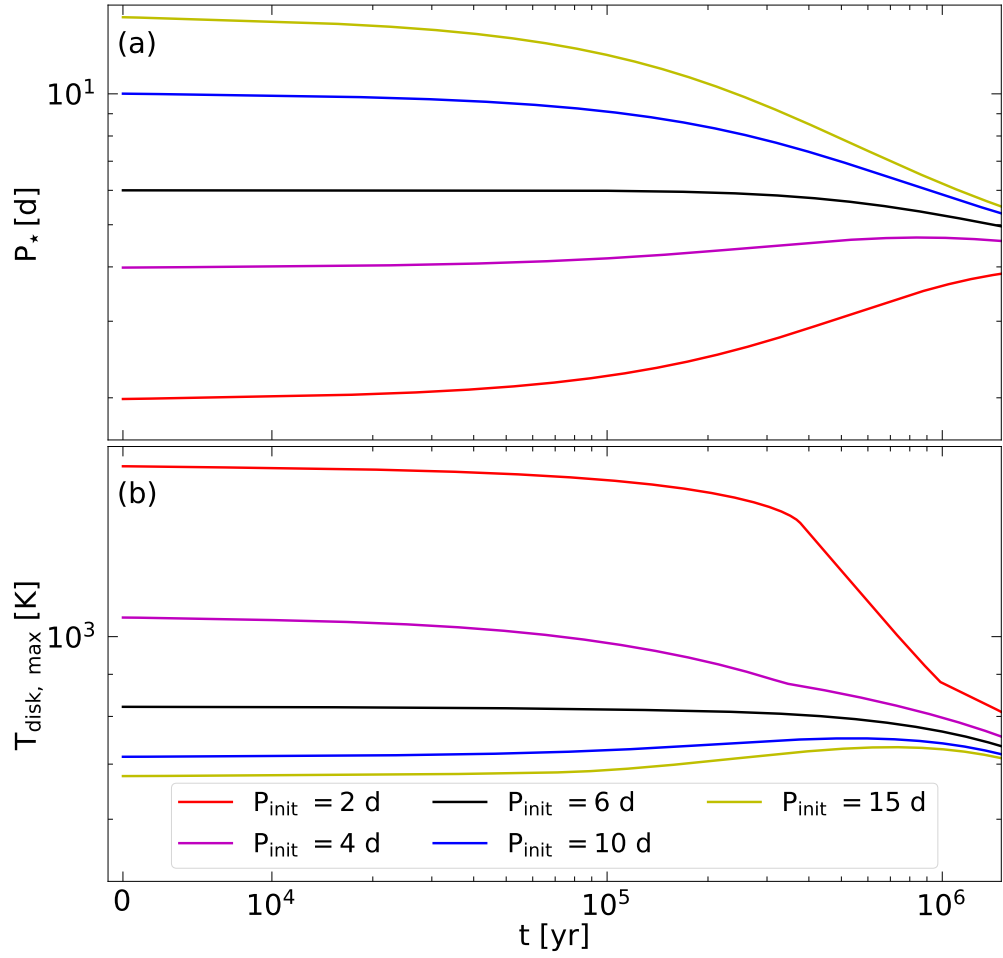


Figure 2.6.: Evolution of the stellar rotational period $P_\star$ (Panel (a)) and the maximum disk midplane temperature $T_{\text{disk, max}}$ (Panel (b)) with respect to the simulation time t in years starting from $t_0 = 1$ Myr. Different colors represent different initial stellar rotational periods: $P_{\text{init}} = 2$ days (red line), $P_{\text{init}} = 4$ days (magenta line), $P_{\text{init}} = 6$ days (black line), $P_{\text{init}} = 10$ days (blue line) and $P_{\text{init}} = 15$ days (yellow line).

time span of 1.11 Myr. Each outburst influences the temperature structure [e.g., Steiner et al., 2021] and leads to a change in the chemical composition of the disk [e.g., Rab et al., 2017a, and references therein]. This can change observables such as the snow line [e.g., Cieza et al., 2016], the abundance of HCO^+ [Jørgensen et al., 2013] or CO observations [Rab et al., 2017a].

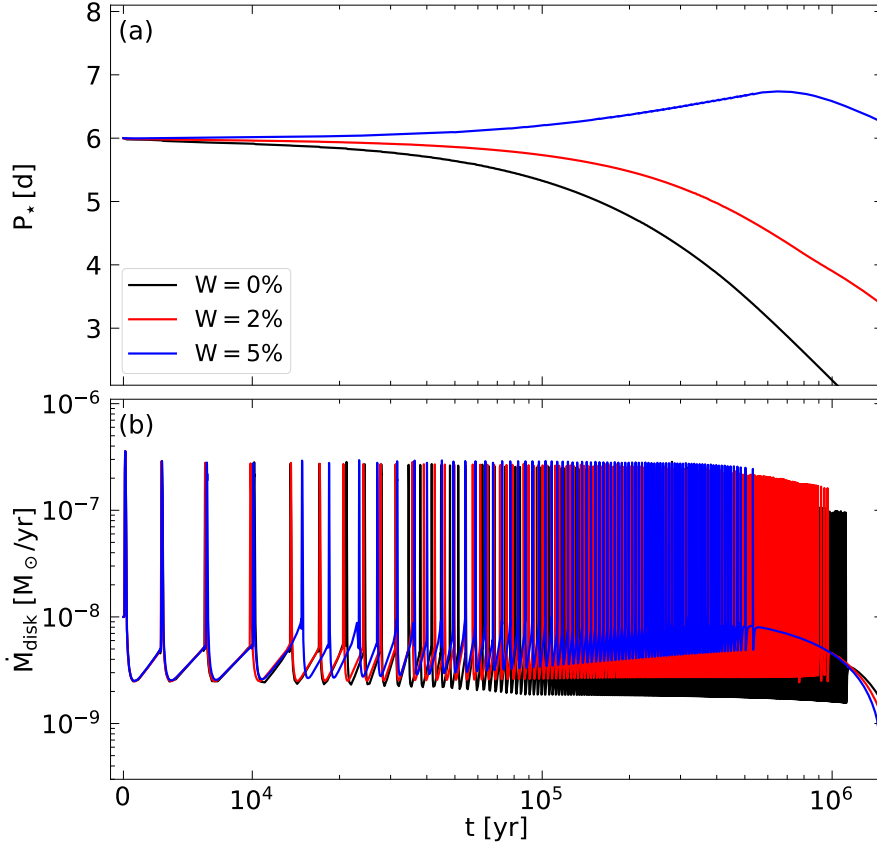


Figure 2.7.: Stellar spin evolution for different stellar spin evolutionary tracks of the layered disk reference model with respect to the simulation time t in years starting from $t_0 = 1$ Myr. Panel (a) shows the stellar rotational period P_* for the reference model $\mathbf{M}_{\text{refB}}$ with $W = 2\%$ (red line) as well as for models with $W = 0\%$ (black line) and $W = 5\%$ (blue line). Panel (b) shows the disk accretion rate $\dot{M}_{\text{disk}}$ for the respective models .

2.4. Discussion

To explain the rotational behavior of young stars, still surrounded by a protoplanetary disk, the star-disk interaction and its associated torques have to be considered (see Sec. 2.2.4). The disk component in previous studies, however, has been modeled in a simplistic way and important disk features and parameters have not been included in the stellar spin evolution [e.g., Matt et al., 2010a, Matt et al., 2012, Gallet et al., 2019]. To improve the treatment of the disk, we combine hydrodynamic disk and stellar spin evolution to simulate the disk's influence on the stellar rotation

and the stellar back-reaction on the disk. The influence of stellar magnetic torques on the disk, the position of the inner disk radius, disk radial density, temperature profiles, and the effects on episodic accretion outbursts can be calculated self-consistently during the whole simulation time.

A self-consistent treatment of both stellar and disk evolution introduces further complexity. As the stellar spin evolution depends on a precise formulation of the processes of the inner disk, this region has to be well resolved during the whole simulation ranging over several mega-year. Additionally, magnetospheric accretion is an intrinsically multi-dimensional problem [e.g., Romanova et al., 2002, Bessolaz et al., 2008, Romanova and Kurosawa, 2014]. In addition to timestep limitations imposed by the CFL condition [Courant et al., 1928], current explicit, multi-dimensional simulations either exclude the innermost disk regions or treat them in a simplified manner [e.g., Vorobyov et al., 2019] or converge toward a steady-state solution within several rotational periods of the inner disk [e.g., Romanova et al., 2004, Zhu et al., 2019, Pantolmos et al., 2020]. These scenarios are not ideal to study the dynamic stellar spin evolution over several mega-year. The TAPIR code is able to spatially resolve the inner disk region sufficiently well over the whole disk lifetime. Solving the hydrodynamic equations (cp. Eqs. (2.1) - (2.3)), the accretion disk can be modeled in a more self-consistent manner compared to previous stellar spin evolution studies [e.g., Matt et al., 2010a, Matt et al., 2012, Gallet et al., 2019]. We have to note that the one-dimensional TAPIR code requires a parameterization of the multi-dimensional processes occurring in the inner disk. Consequently, the results presented above are based on simplifications. Nevertheless, the general trends and dependencies shown in this study should be valid.

Young stars are assumed to have strong magnetic fields [e.g., Johns-Krull et al., 2009, Gregory et al., 2012, Johnstone et al., 2014]. The complex field topology and temporal variability of the field, however, are difficult to observe or predict and equally challenging to develop in a numerical model, which describes the field dynamics correctly and simultaneously maintains the ability to carry out long-term simulations. Based on the measurement technique, different values for the stellar magnetic field are observed for the same star. This issue can be explained on the basis of the young M dwarf AU Mic [Kochukhov and Reiners, 2020]. A mean magnetic field of $B_{\text{mean}} = 2.3 \text{ kG}$ results from Stokes I profile measurements. With a more detailed

Zeeman Doppler imaging technique based on the Stokes V parameter, a global field strength of $B_{\text{global}} = 0.09$ kG is observed. Finally, based on the Stokes Q and U parameters a dipole field strength of $B_{\text{dipole}} = 2.0$ kG is obtained. Thus, the observed magnetic field can change based on the used measurement technique by a factor of 20. Finley and Matt [2018], however, have shown in stellar wind calculations that usually the lowest order field component dominates Γ_w . Additionally, stellar magnetic fields can vary on short timescales. Johnstone et al. [2014] have shown that the dipole magnetic field strength of V2129 Oph varies by a factor of almost 4 over a four-year period.

Stellar winds, in this study represented by APSW, strongly affect the long-term stellar spin evolution. While an APSW influences not only the stellar spin evolution providing a possible solution to the occurrence of slow-spinning stars, it also affects the disk structure as well as its long term evolution (see Fig. 2.5 and Fig. 2.7). The exact nature and strength of these winds, however, are still debated [see the discussion in Gallet et al., 2019]. Previous authors need large APSW values of $W \gtrsim 10\%$ to explain slowly rotating stars [e.g., Matt et al., 2012, Gallet et al., 2019, Ireland et al., 2021], which contradicts the results of Cranmer [2008] and Pantolmos et al. [2020], limiting $W \sim 1\%$. We can reduce this value to $W \lesssim 5\%$ to reproduce the observed rotational period of most stars in young stellar clusters (see Fig. 2.4). Apart from APSW, other wind mechanisms can be important for stellar and disk evolution. Especially magnetic disk winds affect the disk evolution noticeably [e.g., Königl et al., 2011]. By removing matter and, in the presence of a large-scale magnetic field threading the disk, angular momentum from the disk before it can be accreted onto the star, the stellar spin evolution is affected. Similarly, photo-evaporation (internal or external) can influence the disk's lifetime and thus, the stellar spin evolution [e.g., Roquette et al., 2021]. These effects are added to our model in further studies.

The stellar spin evolution is strongly influenced by the initial stellar parameters. Besides the magnetic field strength, the initial radius as well as the temperature play an important role. The change of the stellar moment of inertia ($\dot{I}_\star$) can vary significantly based on the choice of initial parameters (Eq. 2.18).

We take initial stellar values from Baraffe et al. [2015] and evolve the star using a rather simplified model to describe $\dot{R}_\star$. As described in Matt et al. [2010a] such assumptions reproduce

more sophisticated stellar evolution models to a reasonable degree of accuracy. Siess et al. [2000], however, concludes that the variation of, for example, stellar metallicity can vary $T_\star$ by $\gtrsim 200$ K. Additionally, a spin-up or spin-down torque can alter the rotational period and thus stellar evolution [Zeipel, 1924, Endal and Sofia, 1976, Maeder and Meynet, 2000, Tassoul, 2000, Paxton et al., 2019a, Amard et al., 2019a]. Moreover, the amount of accreting energy that is radiated away from the star influences stellar evolution. The deviation in stellar radius and luminosity between models, in which most of the accreted energy is radiated away (cold accretion) and a certain amount of accreted energy is added to the star (hybrid or hot accretion), can reach up to 50% [e.g., Baraffe et al., 2009, Vorobyov et al., 2017b]. Especially during an accretion outburst, the stellar radius and effective temperature are influenced [e.g., Vorobyov et al., 2017b], which is not yet included in our model. Thus, the combination of our disk hydrodynamics model with a more elaborate stellar evolution code [e.g., the MESA code; Paxton et al., 2019a] could include these effects in a self-consistent way [similar to e.g., Vorobyov et al., 2017a] and would be a further improvement of the stellar spin evolution model.

2.5. Summary and conclusion

In this study, we have enhanced the hydrodynamic disk evolution TAPIR code [Ragossnig et al., 2020, Steiner et al., 2021] with a stellar spin evolution model (see Sec. 2.2.4). Several properties of the disk (e.g., the disk’s accretion rate or the disk lifetime) that have been approximated in previous studies [e.g., Matt et al., 2010a, Gallet et al., 2019] are now calculated in a self-consistent way. Additionally, the back reaction of stellar rotation on the disk can be included. Different stellar spin evolutionary tracks and the associated movement of the co-rotation radius influence the inner disk structure. The inner disk radius of a star that spins up is located closer to the star (see Fig. 2.5). More disk material is heated up due to the proximity to the star and the disk temperature increases. Similarly, the disk around a star with a large initial rotational period remains relatively cold, compared to an initially fast-rotating star (see Fig. 2.6). From our range of parameters presented in Sec. 2.3.1, we can produce fast and slow rotating models (see Fig. 2.4), which are able to enclose most of the stellar rotational periods observed in young stellar clusters [e.g.,

Henderson and Stassun, 2012, Gallet and Bouvier, 2013] with smaller APSW values compared to previous studies [e.g., Matt et al., 2012, Gallet et al., 2019, Ireland et al., 2021]. Furthermore, for the first time, the effects of episodic outburst events on the stellar rotation and vice versa can be studied over several mega-year. Disks around stars that spin up show a larger number of outbursts as well as a longer outbursting period, compared to slowly rotating stars. While we have shown the effects of different stellar spin evolutionary tracks on the occurrence of outbursts, further studies have to be conducted to analyze a wider range of parameters such as, for example, different stellar and disk masses, outburst strengths as well as including the early evolution of the star-disk system (setting t_0 to $\lesssim 10^4$ years). Our model that combines the stellar spin and hydrodynamic disk evolution enables us to include these effects in further studies and work toward a better understanding of the stellar spin evolution.

3. Paper II: The influence of metallicity on a combined stellar and disk evolution

- **Title:** The influence of metallicity on a combined stellar and disk evolution
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Contribution: In this work, I have coupled an implicit disk model to the stellar evolution code MESA. The co-author TS prepared the MESA code and I added the required changes to the disk model. With the initial stellar evolution model that I received from TS, I run the coupled star-disk evolution model and generated the figures. The majority of the text has been written by me, except Section 2.2 (written by TS) and Section 4.8 (written by RG). The co-authors provided feedback in the form of comments.

Summary: The effects of an accretion disk are crucial to understanding the evolution of young stars. During the combined evolution, stellar and disk parameters influence each other, motivating a combined stellar and disk model. This makes a combined numerical model, evolving the disk alongside the star, the next logical step in the progress of studying early stellar evolution. We aim to understand the effects of metallicity on the accretion disk and the stellar spin evolution during the T Tauri phase. We combine the numerical treatment of a hydrodynamic disk with stellar evolution, including a stellar spin model, allowing a self-consistent calculation of the back-reactions between the individual components. We present the self-consistent theoretical evolution of T-Tauri stars coupled to a stellar disk. We find that disks in low metallicity environments are heated differently and have shorter lifetimes, compared to their solar metallicity counterparts. Differences in stellar radii, the contraction rate of the stellar radius, and the shorter disk lifetimes result in faster rotation of low metallicity stars. We present an additional explanation for the observed short disk lifetimes in low metallicity clusters. A combination of our model with previous studies (e.g., a metallicity-based photo-evaporation) could help to understand disk evolution and dispersal at different metallicities. Furthermore, the stellar spin evolution model includes several important effects, previously ignored (e.g., the stellar magnetic field strength and a realistic calculation of the disk lifetime) and we motivate to include our results as initial or input parameters for further spin evolution models, covering the stellar evolution towards and during the main sequence.

3.1. Introduction

In the classical pre-main sequence (pre-MS) evolution, a star forms with a large radius and contracts isotropically along the Hayashi track [Hayashi, 1961]. During the contraction the star heats up due to the release of gravitational energy, nuclear reactions are triggered, a radiative core develops and the star evolves toward the main sequence (MS). Within this standard model, however, the way the star gets its mass is not taken into account.

Within a collapsing cloud, two hydrostatic cores form [e.g., Larson, 1969, Bhandare et al., 2018]. Starting from the second hydrostatic core (in the context of this work, the stellar seed), material from the parent cloud and the forming accretion disk accretes on the young star [e.g. Mercer-Smith et al., 1984, Palla and Stahler, 1992, Hartmann et al., 1997, Baraffe et al., 2009, Dunham and Vorobyov, 2012, Kunitomo et al., 2017, Vorobyov et al., 2017b, Steindl et al., 2021]. The nature of the accretion process depends on different parameters (e.g., parent cloud mass, temperature, metallicity, or rotational rate), resulting in a variety of possible accretion rates over time [e.g., Vorobyov and Basu, 2010, Vorobyov et al., 2017b, 2020b].

The pre-MS evolution starting from the stellar seed is strongly influenced by the amount of energy added to the stellar interior (L_{add}) during the accretion process [e.g., Kunitomo et al., 2017, Steindl et al., 2021]. If the accreting material loses its energy during the accretion process ($L_{\text{add}} \rightarrow 0$, cold accretion), the stellar radius is significantly (up to one order of magnitude) smaller compared to the classical pre-MS evolution [e.g., Kunitomo et al., 2017]. If, on the other hand, the accreting material can add its energy in the stellar interior (L_{add} large, hot accretion), the stellar radius inflates and can become comparable with classical evolutionary models. After accretion has finished and the energy added into the stellar interior ceases, the star can evolve along the Hayashi track [e.g., Kunitomo et al., 2017]. We note that the exact pre-MS evolutionary track, starting from a stellar seed, depends on how much energy can be added to the star and where this energy is added in the stellar interior [e.g., Kunitomo et al., 2017, Steindl et al., 2021]. Besides the cold and warm accretion models, L_{add} can also depend on the stellar accretion rate [hybrid accretion, e.g., Vorobyov et al., 2017b]. In periods of high accretion rates, L_{add} is large (similar to hot accretion) and during low accretion rates L_{add} stays small (similar to cold accretion).

During the early (Class 0 and Class I) phases of stellar evolution, the accretion rate on the star can vary significantly over time. During episodic outburst events (e.g., FU Orionis outbursts), the stellar accretion rate can rise over orders of magnitude during a short timescale [$\sim$ years, e.g., Vorobyov and Basu, 2010, Vorobyov et al., 2020b]. With L_{add} coupled to the accretion rate, episodic outburst events during the early disk phase (e.g., FU Orionis outbursts) can lead to radial oscillations [e.g., Bastien et al., 2011, Vorobyov et al., 2017b], affecting the stellar pre-MS evolution. In previous studies, the accretion history on the star is either assumed to be constant over time or pre-computed [e.g., Kunitomo et al., 2017, Vorobyov et al., 2017b, Jensen and Haugbølle, 2018, Elbakyan et al., 2019, Steindl et al., 2021]. In the latter case, the effect of episodic outbursts can be included. Up to now, however, such calculations only deal with the stellar structure and are not able to include any back-reaction of the evolving star on the accretion disk.

Star-disk interactions furthermore influence the stellar rotation period. Understanding the stellar spin evolution allows the determination of stellar ages [gyrochronology, Barnes, 2007], magnetic activity or high-energy stellar radiation [e.g., Pallavicini et al., 1981, Micela et al., 1985, Pizzolato et al., 2003, Wright et al., 2011, France et al., 2018]. During the disk phase, the stellar rotation period can also influence the evolution of the accretion rate. Fast-rotating stars tend to experience more outbursts compared to slow-rotating stars [e.g., Gehrig et al., 2022]. According to recent spin evolution models [e.g., Matt et al., 2010a, Matt et al., 2012, Gallet et al., 2019, Ireland et al., 2021], the stellar accretion rate is an important factor to determine the stellar rotation rate. A precise calculation of the accretion rate is crucial for understanding the rotational evolution of young stars. Furthermore, stellar metallicity has an effect on the stellar spin evolution [e.g., Amard et al., 2019a, Amard and Matt, 2020]. In their work, Amard et al. [2019a] present a grid of low-mass ($\leq 1.5 M_{\odot}$) isochrones including a spin evolution model. The disk phase, however, is treated in a simplified way. During the disk phase, the stellar period is assumed to be constant (disk-locking) and the disk lifetime is a free parameter and independent of stellar mass and metallicity. This assumption is in conflict with observations of disk fractions in low metallicity clusters [e.g., Yasui et al., 2010, Yasui et al., 2016, 2021, Guarcello et al., 2021] and theoretical, metallicity-dependent photo-evaporation models [e.g., Nakatani et al., 2018],

indicating short disk lifetimes in low metallicity environments.

In this study, we aim to combine "The Adaptive Implicit RHD" (TAPIR) code [Ragossnig et al., 2020, Steiner et al., 2021] with the software instrument Modules for experiments in stellar astrophysics [MESA, Paxton et al., 2011, 2013, 2015, 2018, 2019a] to provide a self-consistent numerical description of a star-disk system at a young age. This allows us to study, which parameters affect, e.g., the disk lifetime, the stellar spin, and the pre-MS evolution, including the back-reaction on the other component. In this first paper, introducing the new method, we want to study the effects of metallicity on the disk and stellar spin evolution of T Tauri stars (with ages $\gtrsim 1$ Myr). We seek to understand how stellar metallicity affects the accretion disk and stellar spin evolution of T Tauri stars.

This paper is structured as follows: Sec. 3.2 describes the disk and stellar evolution models used in this study. Moreover, the combination of the disk, spin, and stellar model is introduced. Our results are presented in Sec. 3.3 and discussed in Sec. 3.4. Finally, we draw our conclusions in Sec. 3.5.

3.2. Model description

3.2.1. Hydrodynamic disk evolution with the TAPIR code

Studying the long-term evolution of an accretion disk including the inner disk regions close to the star, we use the implicit TAPIR code, which is described in detail in Ragossnig et al. [2020] and Steiner et al. [2021]. The key advantage of our numerical method is the treatment of hydrodynamic equations, rather than a diffusive disk evolution approach [e.g. Pringle, 1981, Armitage et al., 2001, Zhu et al., 2007, Schib et al., 2021]. We can include the influence of pressure gradients and a stellar magnetic field, which is of particular importance in the inner disk regions [e.g. Romanova et al., 2002, Bessolaz et al., 2008]. The following summarizes some key features of the TAPIR code.

Basic equations: In our model, we use a time-dependent, vertically integrated, viscous accretion

disk formulation [e.g. Shakura and Sunyaev, 1973, Armitage et al., 2001, Steiner et al., 2021].

$$\frac{\partial}{\partial t} \Sigma + \nabla \cdot (\Sigma \vec{u}) = 0, \quad (3.1)$$

$$\begin{aligned} \frac{\partial}{\partial t} (\Sigma \vec{u}) + \nabla \cdot (\Sigma \vec{u} : \vec{u}) - \frac{B_z \vec{B}}{2\pi} \\ + \nabla P_{\text{gas}} + \nabla \cdot \mathcal{Q} + \Sigma \nabla \psi + H_P \nabla \left(\frac{B_z^2}{4\pi} \right) = 0, \end{aligned} \quad (3.2)$$

$$\begin{aligned} \frac{\partial}{\partial t} (\Sigma e) + \nabla \cdot (\Sigma \vec{u} e) + P_{\text{gas}} \nabla \cdot \vec{u} \\ + \mathcal{Q} : \nabla \vec{u} - 4\pi \Sigma \kappa_R (J - S) + \dot{E}_{\text{rad}} = 0, \end{aligned} \quad (3.3)$$

where Σ , $\vec{u}$, e , P_{gas} and H_P stand for the gas column density, gas velocity in the planar components $\vec{u} = (u_r, u_\varphi)$, specific internal energy density per surface area, vertically integrated gas pressure and the vertical scale height of the gas disk, respectively. The gradient in planar cylindrical coordinates reads $\nabla = (\partial/\partial r, r^{-1}\partial/\partial\varphi)$ with $\partial/\partial\varphi = 0$ for our axisymmetric model. P_{gas} is utilized by the ideal equation of state $P_{\text{gas}} = \Sigma e(\gamma - 1)$, with the adiabatic coefficient $\gamma = 5/3$ and κ_R denotes the Rosseland-mean opacity, which is composed of a gaseous component $\kappa_{R,\text{gas}}$, which are based on Ferguson et al. [2005] with the abundances in Caffau et al. [2011b] [see Fig. 1 in Steiner et al., 2021], and a dust-dominated component $\kappa_{R,\text{dust}}$ [based on Pollack et al., 1985]; with $\kappa_R = \kappa_{R,\text{gas}} + f_{\text{dust}}\kappa_{R,\text{dust}}$, where f_{dust} is the dust-to-gas mass ratio set to the respective stellar metallicity. In our model, f_{dust} equals the stellar metallicity, ranging from $Z = 0.0028$ to $Z = 0.014$. The gravitational potential of the star-disk system is denoted by ψ . The vertical component of the stellar magnetic field B_z is assumed to be a dipole and constant within the disk's vertical extent. The planar magnetic field components $\vec{B} = (B_r, B_\varphi)$ are taken at the surface of the disk. Similar to Rappaport et al. [2004] and Kluźniak and Rappaport [2007], we ignore the radial magnetic field component $B_r = 0$. Differential rotation between the stellar magnetic field lines, which are assumed to be in rigid rotation with the star, and the disk's material generates an angular magnetic field component B_φ [e.g. Rappaport et al., 2004, Kluźniak and Rappaport, 2007, Steiner et al., 2021, Gehrig et al., 2022]. The radial radiative transport (depicted by $(J - S)$) is modeled in a radiative diffusion approximation with an Eddington factor of $1/3$. We note that an Eddington factor of $1/3$ is usually valid within optically thick regions or in optically thin regions

in the presence of a uniform radiation field. We assume that the stellar radiation and the ambient temperature roughly meet this requirement. The method, however, has its weaknesses when transitioning between optically thick and thin regions and a variable method would produce better results. Here, J and S stand for the 0th moment of the radiation field and the source function, respectively [e.g. Ragossnig et al., 2020, Steiner et al., 2021]. Finally, $\dot{E}_{\text{rad}}$ and Q depict the net heating/cooling rate per unit surface area and the viscous stress tensor, respectively [e.g. Steiner et al., 2021, Gehrig et al., 2022]. The kinematic viscosity is formulated according to Shakura and Sunyaev [1973] $\nu = \alpha c_s H_p$; with the viscous α parameters, the isothermal sound speed c_s and the pressure scale height H_p . In addition to Eqs. (3.1) - (3.3), we utilize an adaptive grid [Dorfi and Drury, 1987] to allow a moving inner and outer disk boundary as well as adequate radial gridpoint resolution throughout the computational domain.

Computational domain: The innermost disk regions are assumed to be truncated and eventually disrupted by a strong ($\sim$ kG) stellar magnetic field inside the so-called truncation radius r_{trunc} [e.g. Romanova et al., 2002, Bessolaz et al., 2008, Romanova and Kurosawa, 2014]. With our one-dimensional code, the physical processes inside r_{trunc} can not be described. Thus, we choose r_{trunc} to be the inner boundary of our computational domain. The radial position of r_{trunc} can be calculated by balancing the stellar magnetic pressure with the maximum from the ram pressure of the accreting disk material and the gas pressure [e.g. Koldoba et al., 2002, Romanova et al., 2002, Bessolaz et al., 2008, Steiner et al., 2021, Gehrig et al., 2022]

$$P_{\text{magn}}(r_{\text{trunc}}) = \max \left[P_{\text{ram}}(r_{\text{trunc}}), P_{\text{gas}}(r_{\text{trunc}}) \right] . \quad (3.4)$$

In case $P_{\text{ram}} \geq P_{\text{gas}}$, the truncation radius can be calculated following Hartmann et al. [2016]

$$r_{\text{trunc}}(P_{\text{ram}} \geq P_{\text{gas}}) \approx 18 \xi R_{\odot} \left(\frac{B_{\star}}{10^3 \text{ G}} \right)^{4/7} \left(\frac{R_{\star}}{2 R_{\odot}} \right)^{12/7} \left(\frac{M_{\star}}{0.5 M_{\odot}} \right)^{-1/7} \left(\frac{\dot{M}_{\star}}{10^{-8} M_{\odot}/\text{yr}} \right)^{-2/7} , \quad (3.5)$$

with the correction factor $\xi = 0.7$, the stellar dipole magnetic field strength $B_{\star}$, the stellar radius $R_{\star}$, the stellar mass $M_{\star}$ and the accretion rate on the star $\dot{M}_{\star}$. In the case of $P_{\text{ram}} < P_{\text{gas}}$, the

position of the truncation radius is calculated by equating the magnetic pressure of the vertical field component and the gas pressure

$$r_{\text{trunc}}(P_{\text{ram}} < P_{\text{gas}}) = \frac{B_{\star}^{1/3} R_{\star}}{P_{\text{gas}}^{1/6} (8\pi)^{1/6}} . \quad (3.6)$$

r_{trunc} is allowed to move during our simulations according to stellar and disk evolution. Furthermore, the outer boundary of our domain is set to the radial position where the surface density has reached a certain value $\Sigma_{\text{out}} = 1.0 \text{ g/cm}^2$ [similar to, e.g., Vorobyov et al., 2020b].

Boundary conditions: At the inner boundary, we choose "free"/Neumann boundary conditions [e.g. $\partial\Sigma/\partial r = 0$, Romanova et al., 2002, 2004, Steiner et al., 2021]. At the outer boundary, we fix the radial (angular velocity) u_r (u_{φ}) to zero (the Keplerian value), respectively. This indicates no inflow of additional material from an outer envelope or parent cloud, as expected in late Class II star-disk systems. The surface density Σ and the specific internal energy e are again treated with a Neumann boundary condition.

3.2.2. Stellar evolution with MESA

The stellar evolution calculations in this work are performed with version v12778 of *Modules for Experiments in Stellar Astrophysics* [MESA Paxton et al., 2011, 2013, 2015, 2018, 2019a]. MESA is an open-source software instrument that solves the fully coupled structure and composition equations simultaneously for a spherically symmetric and hence one-dimensional stellar model [Paxton et al., 2011]. Details about the adopted micro-physics are given in Appendix 3.6.1. The treatment of rotation in MESA is summarized in Appendix 3.6.2.

The input physics for the stellar evolution calculations performed in this work is identical to Steindl et al. [2021]. The initial models used in this study are identical to the ones used in Steindl et al. [2021] with a composition of $X_{1\text{H}} = 1 - X_{2\text{H}} - X_{3\text{He}} - X_{4\text{He}} - Z$, where $X_{2\text{H}} = 20 \text{ ppm}$, $X_{3\text{He}} = 85 \text{ ppm}$, $X_{4\text{He}} = 0.276$ and $Z = 0.014$ for the models with solar metallicity or $Z = 0.0028$ for models with a fifth of solar metallicity. In the following, we briefly discuss implemented changes from the usual MESA calculations for the treatment of accretion.

We follow the description of Baraffe et al. [2009] for which the accretion onto the protostar

proceeds non-spherically, allowing the central object to freely radiate energy across most of the photosphere [Hartmann et al., 1997]. The energy budget of the accretion process is given by

$$\frac{dE_{\text{acc}}}{dt} = (\epsilon - 1) \frac{GM_{\star} \dot{M}_{\star}}{R_{\star}}, \quad (3.7)$$

which is a combination of the gravitational ($-GM_{\star}/R_{\star}$) and internal energy ($+\epsilon GM_{\star}/R_{\star}$) of the accreted material. The geometry of the accretion process is described by the free parameter ϵ . We follow the approach of previous works [Baraffe et al., 2009, Jensen and Haugbølle, 2018, Steindl et al., 2021] and set $\epsilon = 0.5$, corresponding to accretion from a thin disc around the equator [Hartmann et al., 1997, Baraffe et al., 2009]. Thus, the total energy rate accreted to the star from a thin disk is given as $\epsilon GM_{\star} \dot{M}_{\star}/R_{\star}$.

The energy of the accreted material is either added to the star as extra heat (L_{add}) or radiated away as accretion luminosity (L_{acc}). The accretion luminosity is added to the intrinsic stellar luminosity and the additional heating of the disk is considered in Eq. (3.3). A free parameter controls the corresponding amounts such that

$$L_{\text{add}} = \beta \epsilon \frac{GM_{\star} \dot{M}_{\star}}{R_{\star}} \quad (3.8)$$

and

$$L_{\text{acc}} = (1 - \beta) \epsilon \frac{GM_{\star} \dot{M}_{\star}}{R_{\star}}. \quad (3.9)$$

The value β and its dependence on star and disk properties are free parameters. From earlier studies it is evident that β varies with accretion rate [see e.g. Baraffe et al., 2012, Jensen and Haugbølle, 2018, Elbakyan et al., 2019]. In this work, we use a functional dependency first presented by Jensen and Haugbølle [2018] given as

$$\beta(\dot{M}_{\star}) = \frac{\beta_1 \exp\left(\frac{\dot{M}_{\text{m}}}{\Delta}\right) + \beta_2 \exp\left(\frac{\dot{M}_{\star}}{\Delta}\right)}{\exp\left(\frac{\dot{M}_{\text{m}}}{\Delta}\right) + \exp\left(\frac{\dot{M}_{\star}}{\Delta}\right)} \quad (3.10)$$

This describes a step function with a smooth transition between β_1 and β_2 . The functional dependency is controlled by four parameters: $\beta_1 = 0.005$, $\beta_2 = 0.2$, a midpoint accretion rate of

$\dot{M}_m = 6.2 \cdot 10^{-6} \text{ M}_\odot \text{ yr}^{-1}$ and a crossover width of $\Delta = 5.95 \cdot 10^{-6} \text{ M}_\odot \text{ yr}^{-1}$. For comparability, we use the same values as in Steindl et al. [2022], which were originally adapted from Jensen and Haugbølle [2018] for numerical stability. In our case, this translates to a constant value β at the beginning of the evolution which becomes gradually smaller as the accretion rate decreases (see Sec. 3.3.1). During the calculation of the initial models, the accretion rates and thus $\beta(\dot{M}_\star)$ vary significantly. Given the proportionately small accretion rates, β is basically constant during the joint evolution of MESA and TAPIR.

The dependence of β on the accretion rate can also be physically interpreted. During phases of low accretion rates, the stellar magnetic field can truncate the disk at several stellar radii, and the disk material is funneled on the star [e.g., Bessolaz et al., 2008, Hartmann et al., 2016]. The in-falling disk material generates a shock close to the stellar surface and most of the energy is radiated away [e.g., Koenigl, 1991]. As a consequence, only a small part of the disk material's energy is assumed to be added to the stellar interior. On the other hand, during phases of high accretion rates, the disk is pushed towards the star, and accretion along magnetic field lines ceases. Without the shocks at the end of the magnetic funnels, a larger amount of energy can be added to the star. The transition between the regimes of low and high β values is expected to occur at $\sim 10^{-6} - 10^{-5} \text{ M}_\odot \text{ yr}^{-1}$ [e.g., Vorobyov et al., 2017b, Jensen and Haugbølle, 2018], which motivates the choice of $\dot{M}_m$. We note, however, that the description of β in Eq. (3.10) does not take metallicity into account. We assume that, in case of low accretion rates, the inner part of the disk is sufficiently ionized due to stellar irradiation. Stellar magnetic field lines can couple to the disk material and the accretion geometry is unaffected by metallicity. During phases of high accretion rates, when the disk is pushed towards the stellar surface, differences in disk ionization due to metallicity are assumed to be negligible as well. The transition between the two accretion regimes, however, could be affected by the disk's ionization fraction, which depends on metallicity. Thus, the value Δ should include a metallicity-dependent factor, that would exceed the frame of this study.

The heat added to the stellar model is distributed according to

$$l_{\text{extra}} = \frac{L_{\text{add}}}{M_\star} \max \left\{ 0, \frac{2}{M_{\text{outer}}^2} \left(\frac{m_r}{M_\star} - 1 + M_{\text{outer}} \right) \right\}. \quad (3.11)$$

which follows the approach of Kunitomo et al. [2017]. In this picture, the heat is distributed only in an outer region of fractional mass M_{outer} and increases linearly with the mass coordinate m_r . In our study, we choose $M_{\text{outer}} = 0.01$, hence injecting the heat only in the outer 1% of the stellar structure.

3.2.3. Stellar spin model

The stellar spin model used in our simulations is based on Matt et al. [2010a] and Gallet et al. [2019] and summarized in detail in Gehrig et al. [2022]. In this study, we assume solid body rotation for the star. Angular momentum transport within the star is discussed in Sec. 3.4.3. The stellar angular momentum $J_\star = I_\star \Omega_\star$ is influenced by external torques Γ_{ext} and its temporal derivative reads

$$\dot{J}_\star = I_\star \dot{\Omega}_\star + \dot{I}_\star \Omega_\star = \Gamma_{\text{ext}}, \quad (3.12)$$

with the stellar angular momentum $\Omega_\star$ and moment of inertia $I_\star$. Rearranging Eq. (3.12) results in a time-dependent equation for the stellar spin

$$\dot{\Omega}_\star = \frac{\Gamma_{\text{ext}}}{I_\star} - \frac{\dot{I}_\star}{I_\star} \Omega_\star. \quad (3.13)$$

The stellar moment of inertia $I_\star$ is calculated by the MESA code. Within Γ_{ext} the effects of accretion Γ_{acc} , stellar winds Γ_{W} [in form of an accretion powered stellar wind APSW, Matt and Pudritz, 2005b] and the magnetic star-disk connection Γ_{ME} [in form of magnetospheric ejections MEs, Zanni and Ferreira, 2013] are combined. During the process of accretion, the star gains mass and angular momentum from the disk that increases $I_\star$ and spins up the star, respectively. An APSW, on the other hand, is assumed to eject a specific amount of the accreted material ($\dot{M}_{\text{W}} = W \dot{M}$) removing stellar mass and angular momentum. W is expected to be $\lesssim 2\%$ [e.g. Cranmer, 2008, Pantolmos et al., 2020] and thus, we choose $W = 2\%$ in this study. The effect of MEs on the stellar spin evolution depends on the position of r_{trunc} with respect to the co-rotation radius

$$r_{\text{cor}} = \left(\frac{G M_\star}{\Omega_\star^2} \right)^{1/3}, \quad (3.14)$$

with the gravitational constant G and the stellar mass M . MEs increase the stellar angular momentum and spin up the star if $r_{\text{trunc}} < K_{\text{rot}}^{2/3} r_{\text{cor}}$. Otherwise, the star spins down. Following Gallet et al. [2019], we use $K_{\text{rot}} = 0.7$ in this study¹ and the external torque contributions scale as:

$$\Gamma_{\text{acc}} = K_{\text{acc}} \dot{M}_{\star} r_{\text{trunc}}^2 \Omega_{\text{disk,in}} \quad (3.15)$$

$$\Gamma_{\text{W}} = \dot{M}_{\text{W}} r_{\text{A}}^2 \Omega_{\star} \text{ and} \quad (3.16)$$

$$\Gamma_{\text{ME}} = K_{\text{ME}} \frac{B_{\star}^2 R_{\star}^6}{r_{\text{trunc}}^3} \left[K_{\text{rot}} - \left(\frac{r_{\text{trunc}}}{r_{\text{cor}}} \right)^{3/2} \right], \quad (3.17)$$

with $K_{\text{acc}} = 0.4$, $K_{\text{ME}} = 0.21$, the disk rotation rate at the truncation radius $\Omega_{\text{disk,in}}$, and the stellar Alfvén radius

$$r_{\text{A}} = K_1 \left[\frac{B_{\star}^2 R_{\star}^2}{\dot{M}_{\text{wind}} \sqrt{K_2^2 v_{\text{esc}}^2 + \Omega_{\star}^2 R_{\star}^2}} \right]^m R_{\star}, \quad (3.18)$$

where $v_{\text{esc}} = \sqrt{2GM_{\star}/R_{\star}}$, $m = 0.2177$, $K_2 = 0.0506$, and $K_1 = 1.7$.

3.2.4. Numerical combination of the disk, spin and stellar model

The interaction between the disk (TAPIR, Sec. 3.2.1), spin (stellar spin model, Sec. 3.2.3), and stellar model (MESA, Sec. 3.2.2) is schematically shown in Fig. 3.1. Starting from the initial or old values at timestep t_0 , the new disk values, spin values, and stellar values at timestep $t_1 = t_0 + \delta t$ are calculated in 3 steps, which are represented by numbered arrows in Fig. 3.1. We note that we calculate the stellar spin evolution and the MESA evolution separately. During the MESA calculation, the star is assumed to rotate at a constant rate. In the first step, the TAPIR code calculates the disk evolution from timestep t_0 to t_1 using the old disk, spin, and stellar values (Arrow 1). Then, the spin values are updated based on the new disk values (Arrow 2). Finally, the MESA code is updated and evolves the stellar parameters from timestep t_0 to t_1 (Arrow 3).

The accuracy and reliability of our method depend on the timestep size δt . Choosing small timesteps results in accurate but long computational times. Large timesteps will reduce computa-

¹We note that in more recent studies of MEs [e.g. Pantolmos et al., 2020, Ireland et al., 2021] smaller values for K_{rot} are found. $K_{\text{rot}} = 0.7$ can be understood as an upper limit [Ireland et al., 2021].

tional time but accuracy suffers. We choose $\delta t = 100$ yr in this study. During 1 Myr the stellar values are updated 10^4 times resulting in an acceptable representation of the stellar evolution. We also tested a range of δt from 1 yr to 1000 yr and found no significant change in our results. We note that MESA and TAPIR use multiple (internal) timesteps to evolve the star and the disk from time t_0 to t_1 . For MESA, the internal timesteps range between $\sim 10^6$ s and $\sim 10^8$ s. For TAPIR, the timesteps range between $\sim 10^4$ s and $\sim 10^8$ s [see Fig. A.1 in Steiner et al., 2021]. We note that we do not model for example episodic outburst events in this study. During these outbursts, the accretion rate onto the star can change significantly within short time scales and influence the stellar evolution [e.g. Vorobyov et al., 2017b]. In further studies, such outbursts are included in our model and δt has to be adapted to lower values to assure a sufficiently accurate stellar evolution calculation.

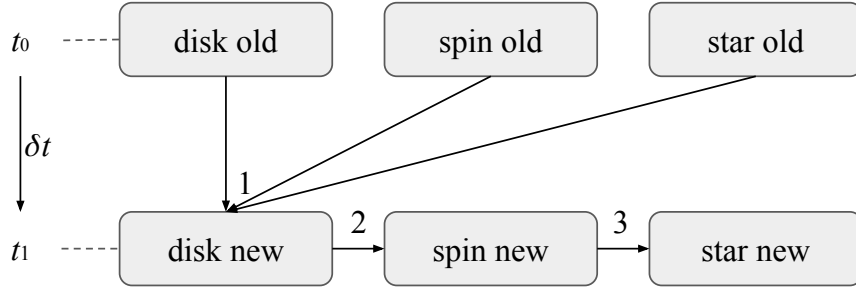


Figure 3.1.: Schematic representation of the interaction between TAPIR and MESA. Starting from an initial/old timestep t_0 , the disk, spin, and star values are evolved towards the new timestep $t_1 = t_0 + \delta t$. The new values at t_1 are obtained in three steps, represented by the numbered arrows (see text).

3.3. Results

To study the influence of metallicity on the combined evolution of a T Tauri star and its disk, several stellar and disk parameters have to be discussed. These parameters are defined based on observations, and previous theoretical or numerical studies and, within a reasonable range, initial models for a T Tauri star and its disk are constructed. We choose $t_0 = 2$ Myr as an initial time for our combined simulations. The effects of the early evolution and different values for t_0 are discussed in Sec. 3.4.2. For each model, the stellar metallicity is varied between $Z = 0.2 Z_\odot$

[motivated by observations of young, low metallicity clusters Yasui et al., 2010, Yasui et al., 2016, 2021] and $Z = 1 Z_{\odot}$.

Disk parameters: One important parameter that defines an accretion disk is the mass accretion rate. For T Tauri star-disk systems ($\lesssim 10$ Myr) accretion rates range within 10^{-10} and $10^{-7} M_{\odot} \text{ yr}^{-1}$ [e.g. Gullbring et al., 1998, Manara et al., 2016, Vorobyov et al., 2017b, Manara et al., 2022]. For $t_0 = 2$ Myr, Manara et al. [2012] and Testi et al. [2022] find disk accretion rates for a $\sim 1 M_{\odot}$ star between $10^{-8} M_{\odot} \text{ yr}^{-1}$ and $10^{-9} M_{\odot} \text{ yr}^{-1}$. Furthermore, the viscous α -parameter can have a wide range between 0.1 and ~ 0.0001 [e.g. Zhu et al., 2007, 2009b, Vorobyov and Basu, 2009, Mulders and Dominik, 2012, Pinte et al., 2016, Yang et al., 2018, Flaherty et al., 2020]. In our simulations, we choose the initial (starting) accretion rate $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_{\odot} \text{ yr}^{-1}$ and set the (constant) viscous α -value to 0.01 unless otherwise stated. This combination results in the expected disk lifetimes of $\lesssim 10$ Myr [consistent with e.g. Richert et al., 2018]. We stop our simulations when the disk accretion rate has decreased to $10^{-11} M_{\odot} \text{ yr}^{-1}$. This lower limit is motivated by the lowest detectable accretion rate for a $1 M_{\odot}$ star [e.g., Sicilia-Aguilar et al., 2016]. At such low accretion rates, the disk is expected to be dissolved rather quickly due to photo-evaporation, which is currently not included in our model. We note that an adaptive viscosity description will be added in further studies.

Stellar parameters: The inner disk region close to the star is influenced by a stellar magnetic field. The field strength of young T Tauri stars can vary between several hundred to several thousand Gauss [e.g. Johnstone et al., 2014, Lavail et al., 2017, 2019]. Unless otherwise stated, the magnetic field strength is chosen to be 2.0 kG. In addition, stellar parameters such as radius, luminosity and effective temperature at t_0 result from the initial stellar evolution ($t < t_0$). The initial stellar parameters are chosen as described in Kunitomo et al. [2017], which follow Stahler [1988] and Hosokawa et al. [2011]. We start from a fully convective stellar seed with a mass of $0.01 M_{\odot}$ (≈ 10 Jupiter masses) and a radius of $1.5 R_{\odot}$ [e.g. Kunitomo et al., 2017, Steindl et al., 2021]. As pointed out in Hosokawa et al. [2011], the initial entropy of the stellar seed influence the subsequent stellar evolution in the case of cold accretion. Cold accretion, however, is assumed to be unrealistic for most stars [e.g., Kunitomo et al., 2017, Vorobyov et al., 2017b]. In the case of hot or warm accretion, the differences due to a variation in the initial stellar seed

parameters are expected to be small [e.g., Stahler, 1988, Kunitomo et al., 2017]. Thus, we do not include a variation of the initial stellar seed parameter in this study and refer to Appendix D in Kunitomo et al. [2017] for further information. Each model accretes material until it has reached a stellar mass of $0.91 M_{\odot}$ and the desired accretion rate $\dot{M}_{\text{start}}$ at time $t_0 = 2$ Myr. The effect of the accretion history is described in Sec. 3.3.1. During the T Tauri phase, the rotational period can span over an order of magnitude from $\lesssim 1$ towards ~ 10 days [e.g. Serna et al., 2021]. To compare the effect of different metallicity values, we start all our models at the same initial stellar rotation period $P_{\text{start}} = 10$ days [comparable to the *slow* models in Amard et al., 2019a]. During the initial stellar evolution ($t < t_0$), the rotation period is fixed to 10 days. We note that the rotation period does change during the first 2 Myr. The respective effects, however, are beyond the scope of this work and will be treated in subsequent studies.

In order to successfully start a simulation combining stellar and disk evolution, initial models, which represent a T Tauri star and its disk at an age of t_0 have to be calculated. The initial models are based on the aforementioned stellar and disk parameters.

3.3.1. Initial stellar evolution: $t < t_0$

The early evolution of a protostar depends on the temporal evolution of the accretion rate [e.g. Kunitomo et al., 2017, Steindl et al., 2021]. Unfortunately, the accretion history of a young star is difficult to measure, hardly constrained, and, thus, treated as a free parameter in our model. Based on the numerical results by Vorobyov et al. [2017b], we start with a constant accretion rate $\dot{M}_{\text{init}}$ followed by a decrease towards our desired value of $\dot{M}_{\text{start}}$ according to $\dot{M}_{\star} \propto t^{-2}$ (see Fig. 3.2). We note that for the models with $\dot{M}_{\text{start}} = 1 \times 10^{-9} M_{\odot} \text{ yr}^{-1}$, we use $\dot{M}_{\star} \propto t^{-2.3}$ to avoid values for $\dot{M}_{\text{init}} > 2 \times 10^{-5} M_{\odot} \text{ yr}^{-1}$. The reason for this variation is convergence problems in the MESA initial model with higher initial accretion rates [e.g., Sec. 5.1.5 in Steindl et al., 2021]. The values of $\dot{M}_{\text{init}}$ and the time, at which the accretion rate begins to decrease, t_{dec} , depend on $\dot{M}_{\text{start}}$ as well as the desired stellar mass at t_0 . For a starting accretion rate of $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_{\odot} \text{ yr}^{-1}$, we find $\dot{M}_{\text{init}} = [1.87 \times 10^{-5}, 1.03 \times 10^{-5}, 5.29 \times 10^{-6}] M_{\odot} \text{ yr}^{-1}$ and $t_{\text{dec}} = [28, 45, 87]$ kyr, respectively. We note that the accretion history of the star is identical for both metallicities. We present our results

for the initial stellar evolution in three parts: from the stellar seed to the transition towards the Hayashi track corresponding to an evolution time τ_H , from τ_H along the Hayashi track towards t_0 and finally reaching t_0 , which marks the starting point for our combined stellar and disk model.

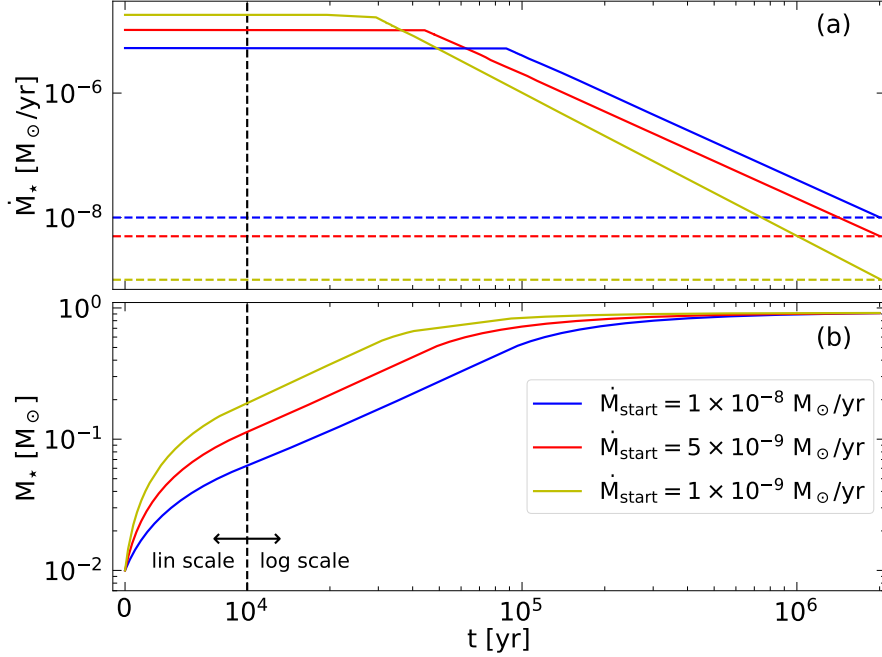


Figure 3.2.: Initial stellar evolution for $t < t_0$. Panel (a): Accretion history of the star. Initially, the star accretes with a constant accretion rate $\dot{M}_{\text{init}}$. At time t_{dec} , the accretion rate drops according to $\dot{M}_* \propto t^{-2}$ (see text). The colored dashed lines show the respective value of $\dot{M}_{\text{start}}$. Panel (b): The evolution of the stellar mass. Starting from the initial mass of $0.01 M_{\odot}$, all models reach their final mass of $0.91 M_{\odot}$ after 2 Myr. The yellow, red and blue lines show models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_{\odot} \text{ yr}^{-1}$, respectively. To show the entire initial stellar evolution, starting from $t = 0$ yr, we use the *symlog* representation [Hunter, 2007]. The vertical dashed line indicates the transition from the linear scale (for $t \leq 10^4$ yr) to the logarithmic scale (for $t > 10^4$ yr).

From the stellar seed to τ_H : Considering the evolution of the star on a Hertzsprung-Russell-Diagram (HRD) for $t < t_0$, it is noticeable that stellar metallicity and the accretion history have a distinct influence on the star (see Fig. 3.3). Stars with lower metallicity are usually more compact and hotter compared to higher metallicities and their position on the HRD is moved to the upper left. The accretion history influences the amount of accretion energy added to the stellar interior

(see Sec. 3.2.2 and Eq. 3.10). In our model, the fraction of added energy depends on the magnitude of the accretion rate. A higher accretion rate $\dot{M}_{\text{init}}$ results in a higher amount of energy added to the star.² We note that a higher $\dot{M}_{\text{init}}$ results in a lower $\dot{M}_{\text{start}}$ (see Fig. 3.2). The additional energy causes the star to expand and the stellar luminosity increases with an increasing amount of energy added to the star. On the HRD, the stars are shifted in a vertical direction dependent on their accretion history (see Fig. 3.3). In other words, for a given effective temperature and luminosity, a star is younger if less energy is added to the star during its formation [e.g. Kunitomo et al., 2017]. At a few kyr, Deuterium burning sets in (τ_{D} , marked with diamonds in Fig. 3.3) and additionally increases the stellar luminosity. The onset times range between 1.31 kyr and 2.72 kyr (see Table 3.1). The point of maximum luminosity (τ_{H} , marked with crosses in Fig. 3.3) indicates a transition. Before τ_{H} , the pre-MS star derives most of its energy from accretion and Deuterium burning and afterward from contraction on the Kelvin-Helmholtz timescale [e.g., Stahler and Palla, 2004]. The young star turns towards the Hayashi track.

Evolution on the Hayashi track: As the accretion rate decreases over time (see Fig. 3.2), the amount of energy added to the stellar interior decreases as well and the expansion of the radius stops. The star contracts subsequently on the Hayashi track, similar to the classical stellar evolution model. In Fig. 3.3, this transition towards the Hayashi track is marked with a cross ('x'). The respective stellar mass (M_{H}) and evolution time (τ_{H}) is summarized in Table 3.1. The stellar masses (M_{H}) range between 0.66 and 0.83 $M_{\odot}$ (with slightly lower values for $Z = 0.2 Z_{\odot}$). The time τ_{H} depends, as expected, on the accretion history. Shortly after the accretion rate decreases from $\dot{M}_{\text{init}}$ towards $\dot{M}_{\text{start}}$ the amount of energy added to the star is insufficient to expand the radius and counteract contraction. Our models are comparable with the results of Kunitomo et al. [2017] for warm/hot accretion models.

Reaching $t_0 = 2 \text{ Myr}$: When reaching an age of $t_0 = 2 \text{ Myr}$, stellar radii and temperatures (see Fig. 3.4) are used to calculate the initial disk model (see Sec. 3.3.2). While the stellar radius differs due to the accretion history, the effective temperature stays approximately constant. When connecting the colored triangles in Fig. 3.3 for $Z = 1 Z_{\odot}$, they are located on a vertical line

²The different behavior at the beginning of the evolution (the first ~ 100 years) can be explained by the outer boundary condition of the MESA code that reacts differently on different values of L_{add} , which depends on the accretion rate.

indicating evolution on the Hayashi track (stellar luminosity decreasing with age). For $Z = 0.2 Z_{\odot}$ however, the stars at t_0 already turn towards the Henyey track (stellar luminosity and effective temperature increase with age) indicating the development of a radiative core.

Table 3.1.: Onset of Deuterium burning (τ_D), stellar mass (M_H) and evolution time (τ_H) the star turns onto the Hayashi track

$\dot{M}_{\text{start}} [\text{M}_{\odot} \text{ yr}^{-1}]$	$Z [Z_{\odot}]$	$M_H [\text{M}_{\odot}]$	$\tau_H [\text{Myr}]$	$\tau_D [\text{kyr}]$
1×10^{-9}	1	0.82	0.09	1.45
5×10^{-9}	1	0.71	0.10	2.31
1×10^{-8}	1	0.69	0.16	2.72
1×10^{-9}	0.2	0.81	0.09	1.31
5×10^{-9}	0.2	0.66	0.08	1.84
1×10^{-8}	0.2	0.66	0.15	2.59

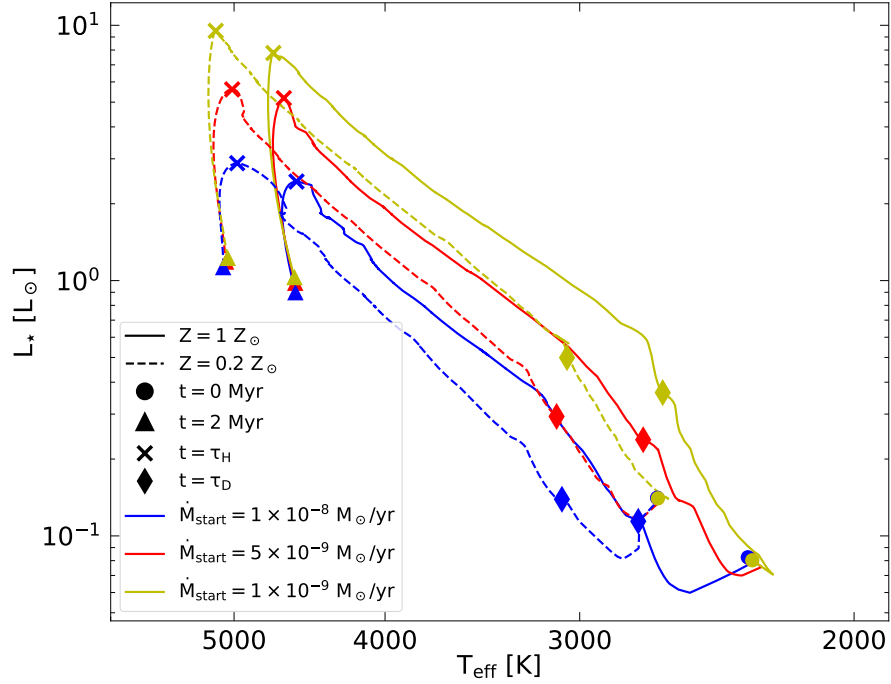


Figure 3.3.: Evolution on the HRD for $t < t_0$ for both metallicities $Z = 1 Z_{\odot}$ (solid lines) and $Z = 0.2 Z_{\odot}$ (dashed lines). The values at time $t = 0 \text{ Myr}$ ($t = 2 \text{ Myr}$) are marked with circles (triangles) and the yellow, red, and blue lines show models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] \text{ M}_{\odot} \text{ yr}^{-1}$, respectively. The onset of Deuterium burning ($t = \tau_D$) is marked with a diamond. The transition from expansion towards the Hayashi track ($t = \tau_H$) is marked with a cross.

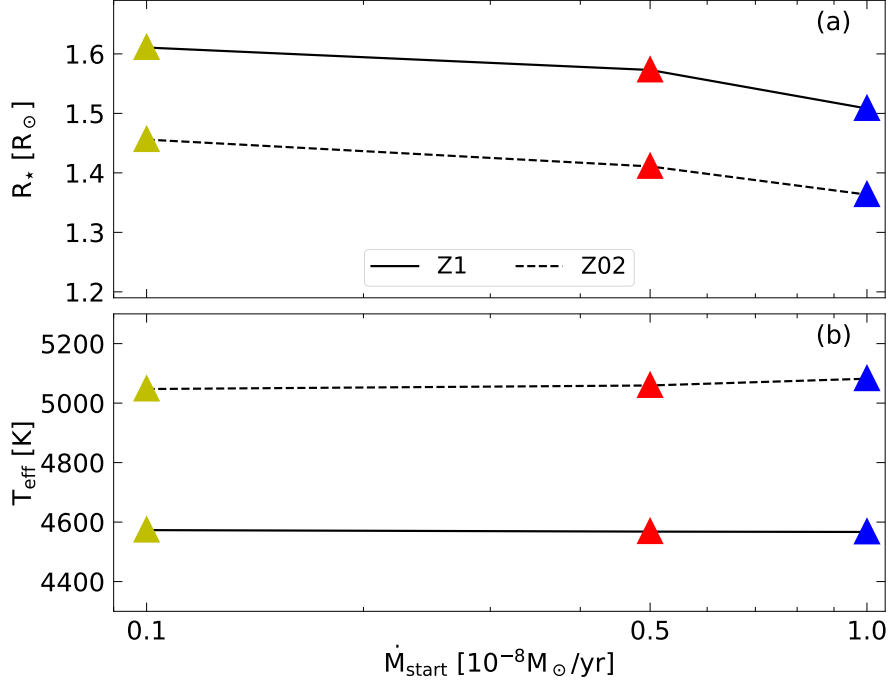


Figure 3.4.: Stellar radii and temperatures at $t_0 = 2$ Myr for $\dot{M}_{\text{start}} = 5 \times 10^{-9}$ (red triangles) and 1×10^{-8} (blue triangles). Solar (sub-solar) metallicities are connected by a solid (dashed) line, respectively.

3.3.2. Initial disk model

Starting a simulation with an implicit method requires an initial model that already solves Eqs. (3.1) - (3.3) [e.g. Dorfi et al., 2006]. Following Steiner et al. [2021], we construct a steady-state accretion disk that serves as the initial model in our simulation. The mass flow rate throughout the disk corresponds to the respective value of $\dot{M}_{\text{start}}$. In Fig. 3.5, the radial surface density (Panel a), disk midplane temperature (Panel b), and the optical depth τ (Panel c) profile of the initial disk models are shown. The stellar metallicities are represented by solid lines ($Z = 1 Z_{\odot}$) and dashed lines ($Z = 0.2 Z_{\odot}$) respectively. The yellow, red and blue lines symbolize models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_{\odot} \text{ yr}^{-1}$, respectively.

For larger initial accretion rates, the disks are more massive, have a higher temperature, and are pushed closer towards the star [e.g. Romanova et al., 2002, Bessolaz et al., 2008, Romanova and Kurosawa, 2014, Steiner et al., 2021]. Due to the smaller stellar radii at t_0 (see Fig. 3.4), the

absolute values of the inner disk edge r_{trunc} is smaller for $Z = 0.2 Z_{\odot}$. The radial temperature profile can be divided into three regions. Close to the star the disk is directly heated by stellar irradiation $r \lesssim 10^{-1}$. For low metallicities, stellar luminosity is larger and, thus, the disk temperature is higher. From $r \lesssim 10^{-1}$ AU to ~ 1 AU, the midplane of the disk is shielded from stellar irradiation and the disk is optically thick (Panel c in Fig. 3.5). This region is referred to as the optical thick zone, R_{thick} , from now on. In this region, the disk midplane temperature ranges from ~ 100 to 1000 K and is dominated by dust opacities [e.g., Semenov et al., 2003]. In our model, we assume that the disk gas-to-dust ratio corresponds to the stellar metallicity, $f_{\text{dust}} = 0.014$ and 0.0028 for $Z = 1 Z_{\odot}$ and $0.2 Z_{\odot}$, respectively. Consequently, larger stellar metallicity results in higher optical depths and temperatures in the disk midplane compared to lower stellar metallicities. We note that in the models with $\dot{M}_{\text{start}} = 1 \times 10^{-9} M_{\odot} \text{ yr}^{-1}$, these effects are almost negligible due to the low surface density values. Outside R_{thick} , the disk temperature is again dominated by stellar irradiation and smaller stellar metallicities result in higher disk temperatures. The disk temperature decreases with increasing disk radius until the ambient temperature of $T_{\text{amb}} = 20$ K is reached (see Panel b in Fig. 3.5).

To compare the initial disk models to observations and other theoretical models, we summarized the initial disk sizes, given by R_{c} , and masses in Table 3.2. The characteristic radius of the disk R_{c} includes 2/3 of the disk's mass. The values of R_{c} , ranging between 20 AU and 128 AU, and m_{disk} , ranging from $0.001 M_{\odot}$ to $0.055 M_{\odot}$ are within the limits of observed disk radii and masses in Ophiuchus II [Andrews et al., 2010] and Lupus star-forming regions [Ansdell et al., 2018a].

Table 3.2.: Characteristic radius R_{c} and disk mass m_{disk} of the initial disk model

$\dot{M}_{\text{start}} [M_{\odot} \text{ yr}^{-1}]$	$Z [Z_{\odot}]$	$R_{\text{c}} [\text{AU}]$	$m_{\text{disk}} [M_{\odot}]$
1×10^{-9}	1	23	0.003
1×10^{-9}	0.2	20	0.001
5×10^{-9}	1	78	0.019
5×10^{-9}	0.2	72	0.017
1×10^{-8}	1	128	0.055
1×10^{-8}	0.2	121	0.051

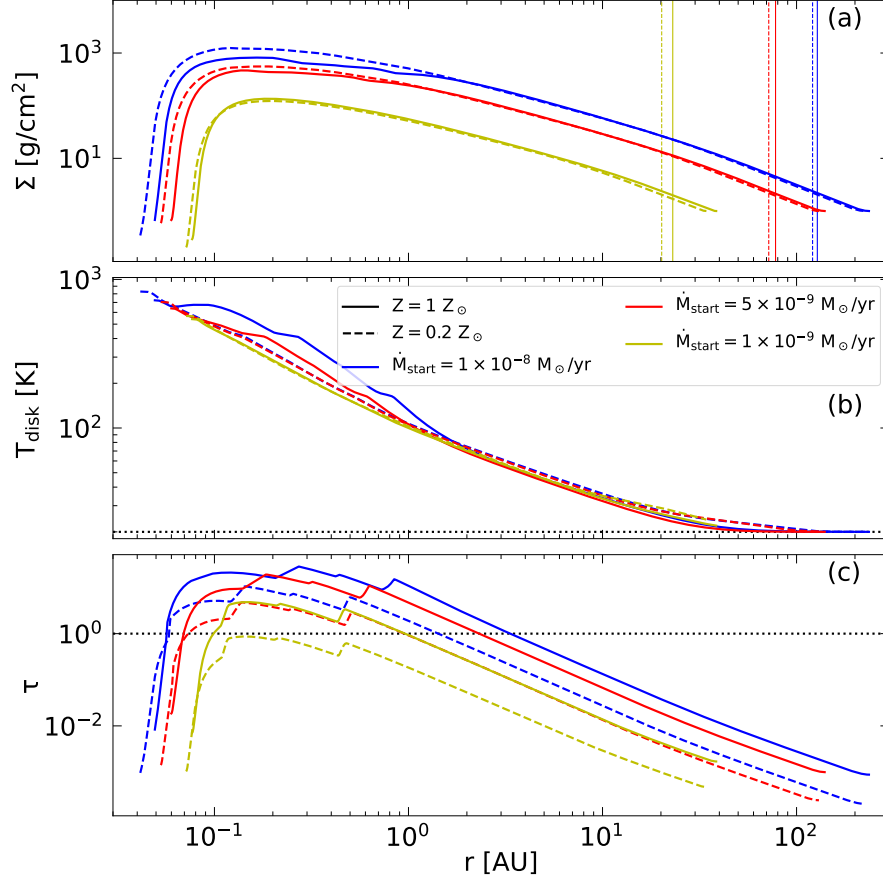


Figure 3.5.: The radial profile of the initial disk model for both metallicities $Z = 1 Z_{\odot}$ (solid lines) and $Z = 0.2 Z_{\odot}$ (dashed lines). The yellow, red and blue lines indicate models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_{\odot} \text{ yr}^{-1}$, respectively. Panel (a) shows the disk surface density Σ , Panel (b) the disk midplane temperature T_{disk} , and Panel (c) the optical depth τ , respectively. The thin vertical lines in Panel (a) indicate the respective position of R_c . The horizontal line in Panel (b) indicates the ambient temperature $T_{\text{amb}} = 20 \text{ K}$. The horizontal dashed line in Panel (c) symbolizes $\tau = 1$, dividing optical thin regions ($\tau < 1$) from optical thick regions ($\tau > 1$).

3.3.3. Disk evolution: $t > t_0$

Starting from the stellar and disk initial models, we evolve the star-disk system until the accretion rate onto the star decreases below $10^{-11} M_{\odot} \text{ yr}^{-1}$. In Fig. 3.6, the evolution of the disk accretion rate is shown for both metallicities $Z = 1 Z_{\odot}$ (solid lines) and $Z = 0.2 Z_{\odot}$ (dashed lines). The red and blue lines indicate models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_{\odot} \text{ yr}^{-1}$, respectively.

Independent of the accretion history of the star, a lower stellar metallicity results in a shorter disk lifetime³.

Observations of disk fractions in low metallicity clusters [Yasui et al., 2010] and statistical studies [Elsender and Bate, 2021] suggest that the disk lifetime is indeed shorter for lower stellar metallicities. Nakatani et al. [2018] explain this behavior with a metallicity-dependent photo-evaporation rate. As we did not include photo-evaporation in our model, we are dealing with a different mechanism affecting the disk lifetime.

Owing to the higher stellar luminosity at lower metallicities, the innermost disk region (inside R_{thick}) and the outer disk regions (outside R_{thick}) are hotter compared to higher metallicities (see Fig. 3.5). With higher disk temperatures, the disk's viscosity ν increases. Thus, the viscous timescale $\tau_\nu = r^2/\nu$ decreases. Since the viscous timescale is a measure of the time, in which the disk material is accreted onto the star, the disk material accretes faster onto the star for lower metallicities. As shown in Fig. 3.11, in which the evolution of the stellar parameters for $t \geq t_0$ are summarized, the differences in **intrinsic** luminosity between solar and sub-solar metallicity increase for older ages. As a consequence, the older a star-disk system is, the greater the difference in stellar irradiation and disk heating between solar and sub-solar metallicities. Usually, the (global) viscous timescale is measured in the outer (~ 10 AU) disk regions [e.g., Armitage, 2011]. During the past years, however, recent studies highlighted the importance of the inner disk, the sub-AU region [e.g., Vorobyov et al., 2019, Hennebelle et al., 2020]. The accretion rate onto the star is eventually regulated by the interaction between the star and the disk.

To test, which region (or perhaps even both) influences the disk lifetime, we artificially shield a certain part of the disk from 25 % of the stellar irradiation, resulting in two test models with a lower disk temperature in the inner disk region < 1 AU (model L_{inlow}) and the outer disk region > 1 AU (L_{outlow}), respectively (see Panel b of Fig. 3.7). We compare the disk evolution of the models L_{inlow} and L_{outlow} to a reference case (ref) without shielding certain regions from stellar irradiation. The evolution of the accretion rate (Panel a in Fig. 3.7) clearly shows that the different disk temperature in the outer region influences the disk lifetime. We note that in our current model, short-term effects such as episodic outbursts are not included. The number and strength

³In this context, disk lifetime is the time, at which the accretion rate has dropped to $10^{-11} \text{M}_\odot \text{yr}^{-1}$.

of these events depend strongly on the inner disk region [e.g., Vorobyov et al., 2020b, Steiner et al., 2021] and can also affect the long-term disk evolution. Including a more detailed physical model, could increase the importance of the inner disk region. In the simulation shown in Fig. 3.7, the stellar evolution has been switched off to preserve computational resources. The qualitative statement of the comparison should not be impaired by this choice.

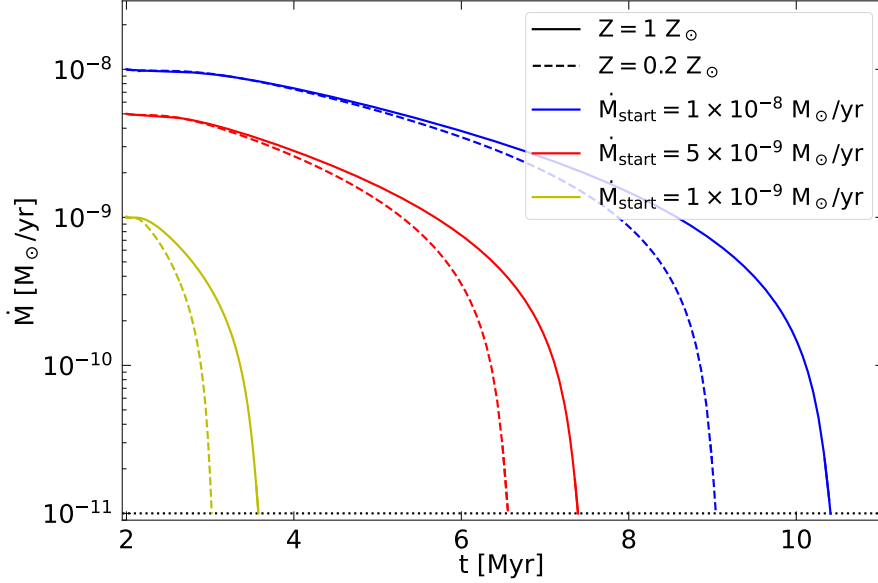


Figure 3.6.: Accretion rate onto the star over time for both metallicities $Z = 1 Z_{\odot}$ (solid lines) and $Z = 0.2 Z_{\odot}$ (dashed lines). The yellow, red and blue lines indicate models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_{\odot} \text{ yr}^{-1}$, respectively. The black dotted line shows the termination criterion used in our simulations of $\dot{M} = 10^{-11} M_{\odot} \text{ yr}^{-1}$.

3.3.4. Influence of metallicity on stellar rotation

The influence of metallicity on stellar rotation has been studied for older ($\gtrsim 1 \text{ Gyr}$) stellar populations based on the dependence of wind-braking formulations on stellar metallicity [Amard and Matt, 2020, Amard et al., 2020]. Low metallicity stars spin faster compared to their solar metallicity counterparts. During the disk phase, however, mechanisms besides wind-braking influence the stellar rotation rate (see Sec. 3.2.3), and these studies assume a constant stellar period (disk-locking).

In Fig. 3.8, the evolution of the stellar rotation period is shown over time for both metallicities

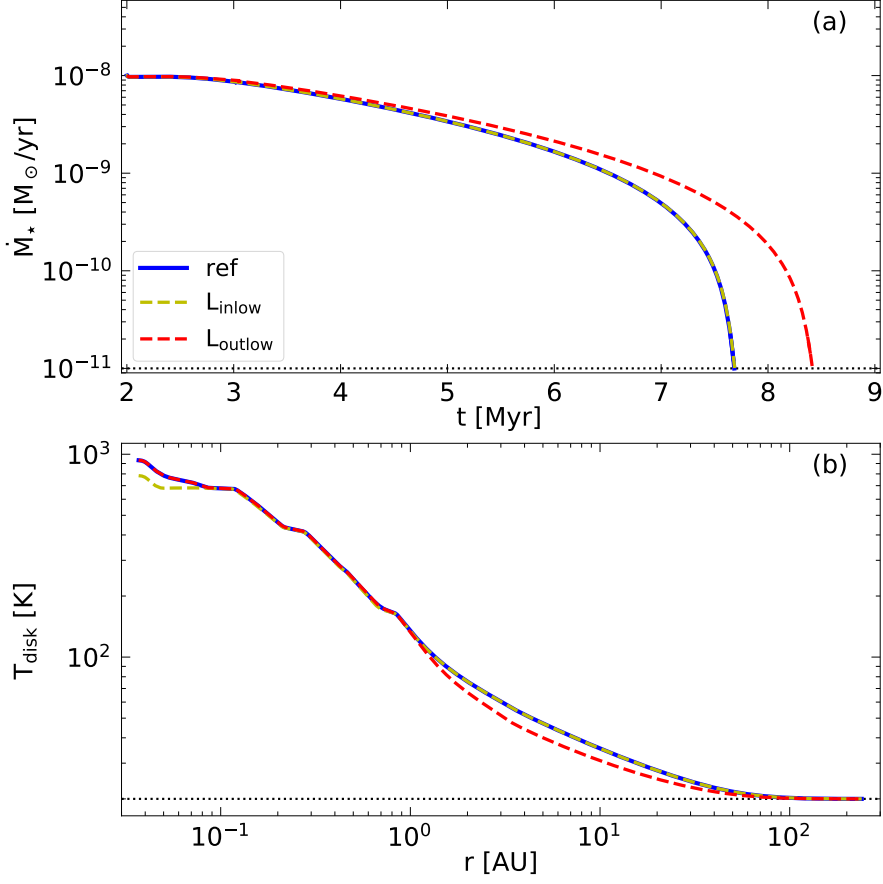


Figure 3.7.: Comparison between models L_{inlow} (yellow line) and L_{outlow} (red line). The reference model (ref, blue line) corresponds to a $Z = 1 Z_\odot$ star with $\dot{M}_{\text{start}} = 1 \times 10^{-8} M_\odot \text{yr}^{-1}$. The stellar luminosity is artificially reduced by 25 % in the inner 1 AU (L_{inlow}) and outside 1 AU (L_{outlow}). Panel (a) shows the evolution of the accretion rate. The black dotted line shows the termination criterion used in our simulations of $\dot{M} = 10^{-11} M_\odot \text{yr}^{-1}$. Panel (b) shows the temperature structure of the initial disk model. For model L_{inlow} , the inner disk region is cooler and for model L_{outlow} , the outer disk region is cooler compared to the reference model. The black dotted line shows the ambient temperature of 20 K.

$Z = 1 Z_\odot$ (solid lines) and $Z = 0.2 Z_\odot$ (dashed lines). The yellow, red and blue lines indicate models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_\odot \text{yr}^{-1}$, respectively. All simulations are started at a stellar period at t_0 of $P_0 = 10$ days. We expect that a variation in P_0 does not change the qualitative result shown in Fig. 3.8, but rather shifts the respective lines in the vertical direction. Independent of the stellar accretion history, stars with lower stellar metallicities rotate

faster compared to solar metallicity values. We can identify three reasons for this development during the disk phase.

Stellar radius: First of all, the metallicity-dependent stellar initial values have an effect on spin evolution. Low metallicities result in smaller stellar radii at t_0 . According to the stellar spin model described in Sec. 3.2.3, the two mechanisms that can remove angular momentum from the star (APSW and MEs) both scale with the stellar radius. Consequently, smaller stellar radius results in less effective stellar spin-down mechanisms, and low metallicity stars spin faster.

Stellar contraction: Contraction of the pre-MS star causes the star to spin up. Dependent on the stellar metallicity and the accretion history of the star, the evolutionary stage at a certain time and, thus, the stellar contraction can differ (see Fig. 3.3). Panel (a) in Fig. 3.11 shows the evolution of the stellar radii over time for both metallicities $Z = 1 Z_\odot$ (solid lines) and $Z = 0.2 Z_\odot$ (dashed lines). For low metallicities, the star contracts more slowly. A slower stellar contraction rate counteracts the effect of smaller stellar radii with respect to stellar rotation (see above).

Different disk lifetime: For low metallicities, the lifetime of an accretion disk is, within the scope of our models, ~ 1 Myr shorter (see Sec. 3.3.3). After the disk is dissolved, there are no mechanisms that can remove angular momentum and the star spins up due to contraction.⁴ Stellar spin-up due to contraction, without the interference of a disk, starts at an earlier age. As a result, the difference in rotation period between low and solar metallicity stars increases further.

3.3.5. Robustness of the results

The results presented in this section are based on certain, fixed stellar and disk parameters, for example, the stellar magnetic field strength $B_\star$ and the disk's viscous α -parameter. We want to verify the robustness of our results by changing these parameters. For our models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_\odot \text{ yr}^{-1}$, the stellar magnetic field is reduced to 0.5 kG (model MC1), the initial rotation period is set to 2 days (model MC2), and the viscous α -parameter is reduced to 0.005 (model MC3), respectively.

⁴We note that there is also a stellar wind mechanism acting after the disk phase. Compared to stellar contraction, however, its effect is small and can be neglected during a relatively short time period of ~ 1 Myr. After contraction has ended and the star evolves towards and on the MS, stellar winds slow down the star on longer timescales [e.g., Gallet and Bouvier, 2013].

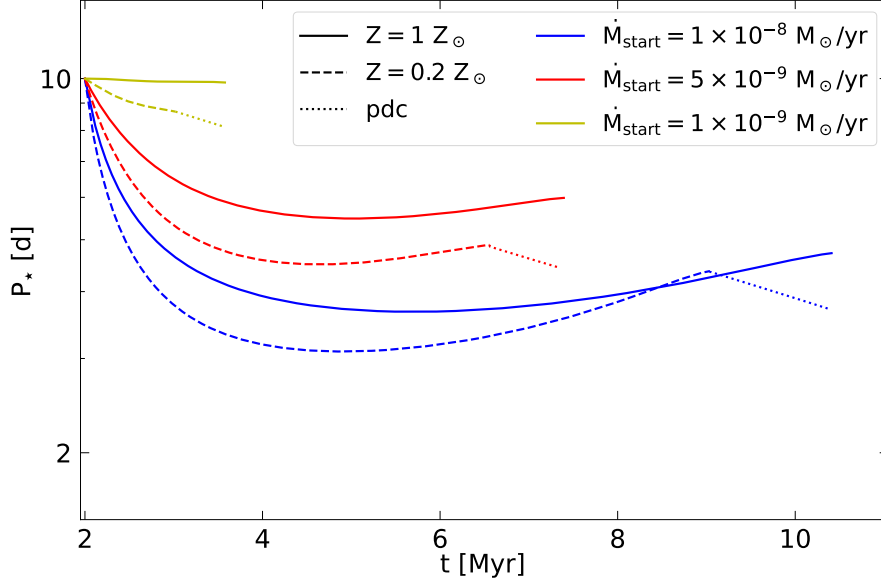


Figure 3.8.: Stellar rotational periods over time for both metallicities $Z = 1 Z_{\odot}$ (solid lines) and $Z = 0.2 Z_{\odot}$ (dashed lines). The yellow, red and blue lines indicate models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_{\odot} \text{ yr}^{-1}$, respectively. For low metallicities, stellar spin-up due to contraction (post disk contraction, pdc) until the disk lifetime of the solar metallicity is reached, is indicated by colored dotted lines.

Fig. 3.9 shows the evolution of the accretion disk (Panel a, see Fig. 3.6) and the stellar spin evolution (Panel b, see Fig. 3.8) for models MC1, MC2, and MC3. Similar to the results shown in Sec. 3.3.3 and Sec. 3.3.4, the trends regarding disk lifetime and stellar rotation period with respect to the metallicity are clearly visible. This indicates that our results are valid over a wider range of parameters.

3.4. Discussion

3.4.1. Comparing our results to observed rotation periods

At ages, $\gtrsim 10$ Myr, the rotation periods of young stars can directly affect the high energy radiation that interacts with planetary atmospheres [e.g., France et al., 2018]. Fast rotating stars usually have stronger high energy emissions compared to slow rotating stars [e.g., Johnstone et al., 2021]. Our results indicate that low-metallicity stars rotate faster than their solar-metallicity counterparts

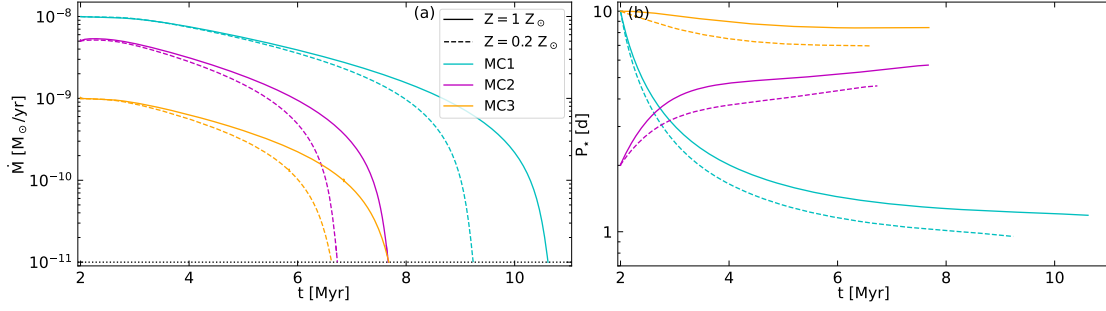


Figure 3.9.: Influence of the stellar magnetic field strength $B_\star$, initial rotation period, and viscous α -parameter on the results for both metallicities $Z = 1 Z_\odot$ (solid lines) and $Z = 0.2 Z_\odot$ (dashed lines). The cyan, magenta, and orange lines show models MC1, MC2, and MC3, respectively. Panel (a): Same as Fig. 3.6. Panel (b): Same as Fig. 3.8 without the post-disk contraction (pdc).

after the disk phase (with differences up to 2 days, see Sec. 3.3.4).

We compare our results to observed periods in young clusters to check the plausibility of the faster rotation periods for low-metallicity stars. Unfortunately, the age estimates of young clusters are subjected to large uncertainties, depending on the stellar evolutionary tracks used to fit the data. Additionally, within one cluster, the different stellar populations can co-exist expanding the age span within one cluster up to several Myr. With these limitations in mind, we choose 6 young clusters with ages between ~ 2 Myr and $\lesssim 5$ Myr. The clusters used in this comparison are NGC 6530, NGC 2362, NGC 2264, Orion star-forming cluster (OSFC), σ Ori and Taurus star-forming region with observed rotation periods taken from Henderson and Stassun [2012], Irwin et al. [2008], Affer et al. [2013], Serna et al. [2021], Cody and Hillenbrand [2010] and Rebull et al. [2020], respectively. The observed ages, mass ranges, and metallicities of these clusters are summarized in Table 3.3. We note that our models are not included in the mass range of observed periods in σ Ori and low-mass stars tend to rotate faster compared to higher stellar masses [e.g., Irwin et al., 2008]. Owing to the small number of available data, we choose to include σ Ori in this comparison.

The observed rotation periods are combined in two bins: the cluster metallicity is assumed to be solar or above (P_{Zhigh}) and the metallicity is assumed to be below the solar value (P_{Zlow}). In Fig. 3.10, the ranges of observed rotation periods from the 25th to the 90th percentile are shown

for P_{Zhigh} (light grey) and P_{Zlow} (light blue). The mean values ($\bar{P}$) are indicated by horizontal lines in the respective color. Consistent with our results, the range of observed periods is shifted to faster values for lower metallicities. The difference in rotation period due to lower metallicity values, however, is small, compared to other stellar parameters, e.g., the stellar magnetic field strength. We note that the lower stellar mass range of observed periods in σ Ori could further reduce the difference even further. Thus, the faster rotation periods due to lower metallicity values alone, will probably not affect the stellar high energy output and, in consequence, the planetary atmosphere significantly. If, on the other hand, the magnetic field strength (that is assumed to be constant during our simulations) can be related to metallicity, the importance of metallicity could increase. For such relations, however, additional theoretical and observational work is needed.

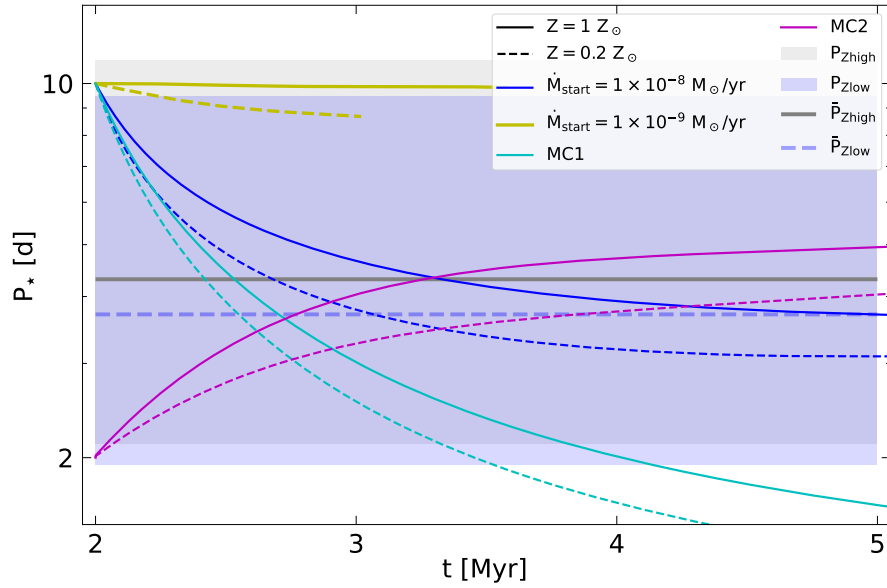


Figure 3.10.: Stellar rotational periods over time for both metallicities $Z = 1 Z_{\odot}$ (solid lines) and $Z = 0.2 Z_{\odot}$ (dashed lines). The colored lines represent the respective model. The ranges of the observed periods of young clusters with ages between ~ 3 Myr and $\lesssim 5$ Myr (see text) are indicated by the light grey (light blue) area for P_{Zhigh} (P_{Zlow}), respectively. The mean values ($\bar{P}$) are indicated by horizontal lines in the respective color.

3.4.2. Early evolution of the star-disk system: Episodic accretion during $t < t_0$

In this study, we start the combined star-disk simulation at $t_0 = 2$ Myr. For ages $t < t_0$ fixed parameters are used to define the accretion history on the star (see Fig. 3.2). The accretion rate for $t < t_0$, however, can vary significantly. Episodic accretion events (outbursts) triggered by gravitational instabilities and disk fragmentation, thermal and magneto-rotational instabilities, or a combination of those, can increase the accretion rate by orders of magnitude over short periods of time [e.g. Nayakshin and Lodato, 2012, Bae et al., 2014, Vorobyov and Basu, 2015, Vorobyov et al., 2017b].

The amount of energy added to the stellar interior depends on the accretion rate. Stronger and more numerous outbursts result in more energy added to the star. The number and intensity of outbursts also depend on the stellar mass [e.g., Vorobyov et al., 2017b, Kadam et al., 2020]. High stellar masses show more numerous and luminous outbursts compared to low-mass protostars.

Furthermore, the stellar rotation period is also affected by early evolution. The principle contributions summarized in Sec. 3.2.3 can influence the star starting from its formation. It is interesting to note that during accretion outbursts the additional energy can cause the star to inflate [e.g., Vorobyov et al., 2017b]. While the additional angular momentum carried by the accreted disk material spins up the star, the inflation of the stellar radius causes a spin-down effect. We plan, for future studies, a starting point t_0 reasonably close to the formation of the stellar seed, allowing the consideration of accretion outbursts and early stellar spin evolution.

3.4.3. Angular momentum transport within the star

In this study, we assume solid body rotation for the star. This simplification is justified as long as the star is fully convective without the presence of a radiative core [e.g. Amard et al., 2019a]. When a radiative core develops together with contraction on the pre-MS, a strong meridional circulation (flowing from North to South) transports angular momentum from the core towards the surface [e.g. Amard et al., 2016, 2019a].

In general, during its presence, a radiative core and the convective envelope above, which is still assumed to rotate as a solid body, have to be treated individually. The convective envelope

is still affected by external torques plus the angular momentum flux induced by the meridional circulation. Additionally, the core and envelope can exchange angular momentum due to shear-induced turbulence [e.g. Amard et al., 2019a]. The extent of this angular momentum transfer between the core and the surface of the star depends on the stellar rotation period and stellar mass. Based on the calculations of Amard et al. [2019a], stars rotate either close to solid-body rotation (valid for fast rotating stars during the first $\lesssim 10^8$ years) or the core rotates faster compared to the convective envelope.

Observational confirmation of such a model poses problems. Asteroseismic data reveal only a few estimates of the core angular velocity of low-mass, main sequence stars [Benomar et al., 2015]. In contrast to the model results presented in Amard et al. [2019a], these observations suggest slowly rotating stellar cores, motivating efforts to advance theoretical models and obtain more observational data.

On the other hand, helioseismic data suggest a fast-rotating solar core from g-mode observations [Fossat et al., 2017]. In addition, precise TESS observations of the massive star HD 192575 reveal differential rotation between the near-core region and envelope [Burssens et al., 2021]. Based on their results, the near-core region rotates at a higher frequency. To what extent these results can be applied to our model, however, is unclear due to the different structures and ages of the observed stars. Furthermore, recent theoretical works of Buldgen and Eggenberger [2022] and Eggenberger et al. [2022] present progress in combining the theory of internal stellar angular momentum transport and observational constraint, for example, the solar surface lithium abundance.

3.4.4. Evolution of $B_\star$

The stellar magnetic field strength is a highly variable parameter that can change its magnitude within a short period of time [several years, e.g. Johnstone et al., 2014]. The values used in this study for $B_\star$ must be understood as averaged over time. Furthermore, the field topology is assumed to depend on the evolutionary stage [e.g. Zaire et al., 2017, Emeriau-Viard and Brun, 2017]. With the presence of a radiative core, the magnetic field topology can become more complex, which can result in a weaker large-scale dipole component. In our model, the magnetic field strength $B_\star$ represents the large-scale, dipole component of the magnetic field, which usually

dominates the star-disk interaction region and the stellar spin evolution [Finley and Matt, 2018].

For low metallicities, the star evolves further and even begins its turn towards the Henyey track, indicating the development of a radiative core. During this time, the large-scale, dipole field strength can decrease, which affects the stellar spin evolution. Referring to Fig. 3.9, such a decrease of $B_\star$ for low metallicity stars could be visualized by a gradual transition from the blue-dashed line (where $B_\star = 2.0$ kG is assumed) towards the cyan-dashed line ($B_\star = 0.5$ kG in model MC1). As a result, the spread in rotational periods for different metallicity values increases further.

3.4.5. Disk lifetimes in low metallicity clusters

Our models predict a dependence of disk lifetimes on stellar metallicity. The disks around low metallicity stars dissolve faster compared to their solar metallicity counterparts (see Fig. 3.6). This correlation can be explained (within the scope of our model) by a hotter outer disk region enhancing viscosity, reducing the viscous timescale, and, thus, resulting in a shorter disk lifetime.

Over the last decade, several observational studies can confirm this trend between metallicity and disk lifetime [e.g. Yasui et al., 2010, Yasui et al., 2016, Guarcello et al., 2021, Yasui et al., 2021]. In general, the observations show a disk fraction in low metallicity disks that is lower compared to solar metallicity. For example, at an age of 1 Myr, the disk fraction in a $Z = 0.2 Z_\odot$ cluster is $\sim 10\%$ [e.g. Yasui et al., 2010] and considerably lower compared to solar metallicity values with disk fractions at 1 Myr of 60 – 80 %.

Different mechanisms affect the dispersal of the accretion disk in low-metallicity environments. The ionization fraction in low metallicity disks can be higher compared to solar metallicities. Smaller amounts of dust result in a slower recombination rate of free charges on dust grain surfaces. A higher ionization fraction in the disk increases the efficiency of magneto-rotational instabilities [Balbus and Hawley, 1991], which drive accretion [e.g., Hartman et al., 2009]. Alternatively, the efficiency of photo-evaporation also depends on metallicity. Lower metallicities and the resulting lower disk opacities allow deeper penetration of high energy radiation in the disk [e.g., Ercolano and Clarke, 2010]. As a result, a larger part of the disk is heated and the removal of gas and dust by photo-evaporation is facilitated. More recently, numerical simulations [e.g.,

Nakatani et al., 2018] confirm the metallicity dependence of photo-evaporation. For sub-solar metallicities, the photo-evaporative mass-loss rates are higher and the accretion disk disperses at a faster rate. It is worth noting that the observed disk lifetimes in low metallicity clusters cannot be conclusively explained by one mechanism alone [e.g., Nakatani et al., 2018] and results, more likely, from the combination of the aforementioned aspects. In a more recent study, Elsender and Bate [2021] proposes that increased stellar multiplicity in low-metallicity environments and resulting increased dynamical interactions might cause accretion disks to be smaller and have shorter lifetimes compared to higher metallicities.

3.4.6. Differences between our results and observations

The observations shown in Yasui et al. [2010] suggest disk lifetimes of ~ 3 Myr for low-metallicity clusters, matching our results for $\dot{M}_{\text{start}} = 1 \times 10^{-9} \text{ M}_{\odot} \text{ yr}^{-1}$. In this context, cluster ages can be used to estimate disk lifetimes. The disk lifetimes for models with $\dot{M}_{\text{start}} = 5 \times 10^{-9} \text{ M}_{\odot} \text{ yr}^{-1}$ and $1 \times 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$, on the other hand, are longer compared to observations by a factor of ~ 3 . We want to address three possible reasons for this discrepancy. *First*, our models do currently not take photo-evaporation into account. Including photo-evaporation in future models will reduce our model's lifetime towards the observed values. *Second*, the cluster ages presented by Yasui et al. [2010] are based on models that do not take magnetic fields into account. Cluster ages based on magnetic pre-MS stellar evolution models [e.g., Feiden et al., 2015] are up to 2 Myr older compared to previous models [e.g., Richert et al., 2018]. Disk lifetimes for low-metallicity stars could increase up to ~ 5 Myr. *Third*, the initial conditions in our models are independent of metallicity. It is, however, likely that the accretion rate, stellar rotation period, and stellar magnetic field strength depend on metallicity. Moving t_0 to earlier evolutionary stages in combination with additional observational constraints for low-metallicity clusters will improve the quality of our results in future work.

3.4.7. Previous spin evolution studies including metallicity

Many new formulations have emerged over the last decade to explain the stellar spin evolution [e.g., Matt et al., 2015, Finley and Matt, 2018, Gallet et al., 2019, Ireland et al., 2021]. These studies,

however, are often based on solar properties and composition. Thus, an explicit dependence on stellar metallicity is not given.

Amard et al. [2019a] presents, for the first time, a comprehensive set of stellar evolutionary tracks including rotation and different metallicities, using the *STAREVOL* code. They find that metallicity does influence the stellar spin evolution during the pre-MS and MS. Low metallicity stars are more compact and evolve faster compared to their solar metallicity counterparts. Thus, during the pre-MS, they spin up faster and retain faster rotation periods during the MS. Internally, low metallicity stars experience less differential rotation compared to solar metallicities. A more recent study Amard and Matt [2020] provides an in-depth analysis of the effects of metallicity on different wind-braking formulations. They also conclude that low-metallicity stars rotate faster than their solar metallicity counterparts. During the disk phase, a constant stellar period (disk-locking) is assumed for a fixed timescale independent of stellar mass and metallicity [Amard et al., 2019a, Amard and Matt, 2020]. This assumption is in conflict with observations of metallicity-dependent disk lifetimes [e.g., Yasui et al., 2010, Yasui et al., 2016, 2021, Guarcello et al., 2021], photo-evaporation models [e.g., Nakatani et al., 2018] and the results presented in Sec. 3.3.3.

Starting from the T-Tauri phase, our star-disk evolution model does include the effects of, e.g., metallicity, stellar magnetic field strength, stellar spin, and mass. Over a wide range of stellar and disk parameters, we can provide realistic, metallicity-dependent disk lifetimes and stellar periods after the disk phase. Those values can then be used as input parameters for the aforementioned spin evolution studies.

3.4.8. Impact on disk chemistry

Various studies confirm the influence of metallicity on the chemical evolution of disks. For example, by observing the chemical abundances of HII regions in 12 nearby dwarf galaxies, Berg et al. [2016] suggests that there are different chemistry cycles in non-solar metallicity environments. This evidence hints at the fact that disk evolution and planet formation in protoplanetary disks may proceed differently in diverse metallicity environments. Certain aspects of the physical conditions considered by our model could change the way metallicity influences chemistry.

One of the important features is the usage of passive or active disks. For passive disks, the only

heating source is stellar radiation and cosmic rays. Therefore, passive disks exhibit an increase in temperature with decreasing metallicity. In the case of active disks, the hydrodynamical processes provide additional heating via viscous and compressional heating, which operate predominantly in the disk midplane and depend on the assumed metallicity. A decrease in metallicity leads to a lower temperature in the midplane (see Fig. 3.5).

In a recent study, Guadarrama, R. et al. [2022] showed that snowlines of chemical species are pushed back as metallicity decreases. Assuming a passive disk, their model does not take viscous heating into account resulting in an increasing disk temperature throughout the disk with decreasing metallicity. This assumption might be appropriate for outer disk regions $\gtrsim 10$ AU, where the disk temperature is dominated by stellar irradiation. Within R_{thick} , however, the disk midplane temperature decreases with decreasing metallicity (see Sec. 3.3.2) and snowlines of species with sublimation temperatures $\gtrsim 50$ K (e.g., H_2O and CO_2) are shifted inward. As a consequence, dust grows and the radial chemical composition of the disk is affected. We note that our model starts at an already advanced evolutionary disk stage. To study the full impact of the aforementioned temperature structures for different metallicities on dust growth and planetesimal formation, the starting point of our model t_0 has to be shifted towards earlier stages.

3.4.9. Model limitations

Although several important aspects of stellar and disk evolution are combined in this study, the results are still preliminary due to certain limitations of our model:

One-dimensional model: The main limitation of our disk model is the one-dimensionality. Due to timestep limitations, long-term evolution models that include the innermost disk regions are currently restricted to one dimension [e.g. Steiner et al., 2021]. Unfortunately, several important stellar and disk properties require (at least) an extension in the vertical direction. Accretion via funnels, the interaction between stellar magnetic fields and the disk, vertical disk structure, and heating of the disk are intrinsic multi-dimensional processes. In the current version of our model, these aspects are included in a simplistic way. Nevertheless, the future goal can only be to expand our model by (at least) another dimension.

Viscous accretion disk: The transport of mass and angular momentum in our model is described

by a simple (constant α) viscosity prescription. α most certainly varies with respect to time and location [e.g. Gammie, 1996b, King et al., 2007, Vorobyov et al., 2020b], and other mechanisms such as disk magneto-centrifugal winds, photo-evaporation, and the combination of both also affect the disk evolution [e.g. Königl et al., 2011, Roquette et al., 2021, Kunitomo et al., 2020]. These are currently not included but will be added in some future versions.

Stellar multiplicity and close encounters: In the current model, only single stars are considered. Binary or even multiple companions can also influence the accretion disk [e.g. Messina et al., 2017]. Additionally, a fly-by or close encounter with a stellar object can impact the disk evolution [e.g. Vorobyov et al., 2017c]. These effects, however, are naturally non-axisymmetric events and are currently not feasible with our model.

Variable dust component: The dust-to-gas mass ratio f_{dust} in this study is kept constant at the respective stellar metallicity value through out the simulation and over the disk's radial extent. While this assumption might be reasonable in the first step, the dust fraction can differ within the disk over time due to a great number of processes, such as coagulation, drift, fragmentation, and vertical settling [e.g. Testi et al., 2014, Vorobyov et al., 2022]. Our results indicate a direct connection between the amount of dust in the disk and its evolution and a more sophisticated dust evolution model will improve our results.

3.5. Conclusion

In this study, we combine three important aspects of a young star: the stellar evolution with MESA, the disk evolution with TAPIR, and the stellar spin evolution based on Gallet et al. [2019]. With reference to the influence of stellar metallicity on the evolution of the accretion disk and stellar spin during the star-disk phase, the main conclusions are summarized as follows:

- The accretion history and metallicity can influence the evolutionary stage of a star. A low metallicity star is more compact and luminous. The higher the initial accretion rate $\dot{M}_{\text{init}}$, the more energy is added to the stellar interior during the early stellar evolution $t < t_0$ and the larger the stellar radius. On the HRD, the star appears less evolved compared to small values of $\dot{M}_{\text{init}}$ (see Fig. 3.2 and Fig. 3.3).

- Larger luminosities of low metallicity stars affect the radial temperature profile of the accretion disk. The disk midplane temperature of the innermost disk region $\lesssim 0.1$ AU as well as the disk beyond $\gtrsim 1$ AU are dominated by stellar irradiation and thus hotter compared to solar metallicity disks. Within R_{thick} , however, the disk midplane is shielded from stellar irradiation and the temperature is dominated by dust opacities. High stellar metallicities result in higher opacity values and the temperatures in the disk midplane are higher compared to lower stellar metallicities (see Fig. 3.5).
- The lifetime of an accretion disk is influenced by metallicity. The increased temperatures in the outer disk regions (see Fig. 3.5) result in a more effective (faster) accretion onto the star and the lifetimes are 0.6-1.4 Myr shorter for low metallicities, within the scope of our model (see Fig. 3.6). We present an additional explanation for the observed short disk lifetimes in young, low metallicity clusters [Yasui et al., 2010]. A metallicity-dependent photo-evaporation description [e.g. Nakatani et al., 2018] arrives at similar results. Their lifetimes, however, are still too long to match the observations. A combination of our star-disk evolution model, coupled with a metallicity-dependent photo-evaporation, could be a key step towards the understanding of disk lifetimes in different metallicity environments.
- Similar to their older counterparts [e.g. Amard and Matt, 2020, Amard et al., 2020], young, low metallicity stars rotate faster compared to solar metallicities (see Fig. 3.8). We can identify three mechanisms that affect stellar spin evolution. For low metallicities, stellar radii are smaller, and, thus, the mechanisms that can remove angular momentum from the star are less effective (see Sec. 3.2.3). Additionally, the stellar contraction rate depends on the stellar evolutionary stage. The stellar contraction rate is smaller if a star is further evolved. As a consequence, low metallicity stars contract more slowly (see Fig. 3.11) and the spin-up due to contraction is reduced. Finally, due to the shorter lifetime of low metallicity disks, stellar spin-up due to contraction, without the interference of a disk, starts at an earlier age. Thus, the difference in rotation period between solar and sub-solar metallicity stars increases further.
- Including the effects of metallicity, stellar magnetic field strength, and stellar mass in

the calculation of the disk lifetime and stellar rotation period during the disk phase, we motivate that our results can be used as input parameters for other recent spin evolution studies [e.g., Amard et al., 2019a, Amard and Matt, 2020, Gossage et al., 2021].

Our model clearly indicates the potential of the joint hydrodynamic evolution of the star and the disk. We motivate the step-by-step inclusion of additional physical aspects (e.g., photo-evaporation and magneto-centrifugal disk winds) and studies covering wider parameter spaces to improve our understanding of these young star-disk systems.

3.6. Appendix

3.6.1. MESA microphysics

The MESA EOS is a blend of the OPAL Rogers and Nayfonov [2002], SCVH Saumon et al. [1995], PTEH Pols et al. [1995], HELM Timmes and Swesty [2000], and PC Potekhin and Chabrier [2010] EOSes.

Radiative opacities are primarily from OPAL [Iglesias and Rogers, 1993, 1996], with low-temperature data from Ferguson et al. [2005] and the high-temperature, Compton-scattering-dominated regime by Buchler and Yueh [1976]. Electron conduction opacities are from Cassisi et al. [2007].

Nuclear reaction rates are a combination of rates from NACRE [Angulo et al., 1999], JINA REACLIB [Cyburt et al., 2010], plus additional tabulated weak reaction rates Fuller et al. [1985], Oda et al. [1994], Langanke and Martínez-Pinedo [2000]. Screening is included via the prescription of Chugunov et al. [2007]. Thermal neutrino loss rates are from Itoh et al. [1996].

3.6.2. Treatment of rotation in MESA

In MESA, rotation is treated as a modification of the stellar equations due to centrifugal acceleration and non-spherical symmetry, when the star is rotating [see Sec. 6 in Paxton et al., 2013]. Angular momentum transport due to rotation (rotational mixing) is implemented as a diffusion term, including five rotationally induced mixing processes: dynamical shear instabil-

ity, Solberg–Høiland instability, secular shear instability, Eddington–Sweet circulation, and the Goldreich–Schubert–Fricke instability [Paxton et al., 2013].

3.6.3. Stellar evolution: $t > t_0$

Fig. 3.11 shows the evolution of the stellar radius (Panel a), intrinsic luminosity (Panel b), and mass (Panel c) for $Z = 1 Z_\odot$ (solid lines) and $Z = 0.2 Z_\odot$ (dashed lines). The yellow, red, and blue lines indicate models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_\odot \text{ yr}^{-1}$, respectively.

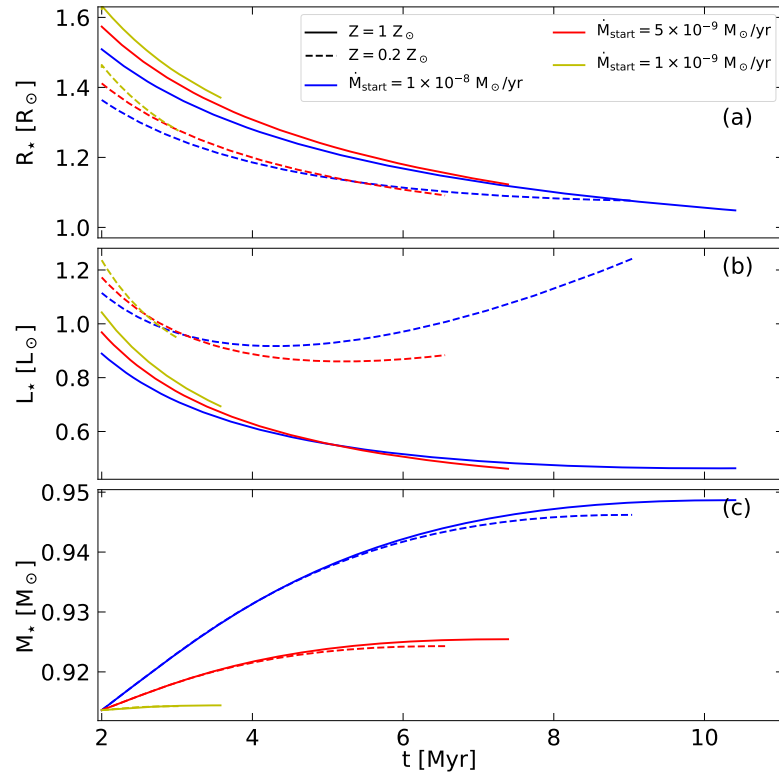


Figure 3.11.: Evolution of the stellar radius (Panel a), intrinsic luminosity (Panel b), and mass (Panel c) for $Z = 1 Z_\odot$ (solid lines) and $Z = 0.2 Z_\odot$ (dashed lines). The yellow, red, and blue lines indicate models with $\dot{M}_{\text{start}} = [1 \times 10^{-9}, 5 \times 10^{-9}, 1 \times 10^{-8}] M_\odot \text{ yr}^{-1}$, respectively.

3.6.4. Time evolution of radial disk structure

The radial profiles for the model with $\dot{M}_{\text{start}} = 1.0 \times 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$ at the starting point of the simulation $t_0 = 2 \text{ Myr}$, as well as at 5.0 Myr and 8.0 Myr are shown in Fig. 3.12. The solid (dashed) lines correspond to a metallicity of $Z = 1 \text{ Z}_{\odot}$ ($Z = 0.2 \text{ Z}_{\odot}$), respectively.

3.6.5. Age, stellar mass range and metallicity of young clusters

Table 3.3.: Age, stellar mass range, and metallicity of young ($\sim 2 - 5 \text{ Myr}$) clusters used in Sec. 3.4.1.

Name	Age [Myr]	Stellar mass range [$\text{M}_{\odot}$]	Metallicity
NGC 6530	$\lesssim 5(1)$	$0.2 - 1.9(2)$	$[\text{Fe}/\text{H}] = 0.11(3)$
NGC 2362	$\sim 5(4)$	$0.1 - 1.2(4)$	$[\text{Fe}/\text{H}] \approx 0(4)$
NGC 2264	$\sim 3(5)$	$0.2 - 2.5(5)$	$[\text{Fe}/\text{H}] < 0(6)$
OSFC	$4(7)$	$0.2 - 1.9(7)$	$[\text{Fe}/\text{H}] \approx 0(8)$
$\sigma \text{ Ori}$	$2 - 3(9)$	$0.2 - 0.5(10)$	$[\text{Fe}/\text{H}] = -0.27(9)$
Taurus	$\sim 3(11)$	$\lesssim 2(11)$	$[\text{Fe}/\text{H}] \approx 0(12)$

References: (1) Damiani et al. [2019]; (2) Henderson and Stassun [2012]; (3) Tadross [2003]; (4) Irwin et al. [2008]; (5) Affer et al. [2013]; (6) Cauley et al. [2012], we note that no exact metallicity value is given; (7) Serna et al. [2021]; (8) D’Orazi et al. [2009]; (9) Sherry et al. [2008]; (10) Cody and Hillenbrand [2010]; (11) Rebull et al. [2020]; (12) D’Orazi et al. [2011]

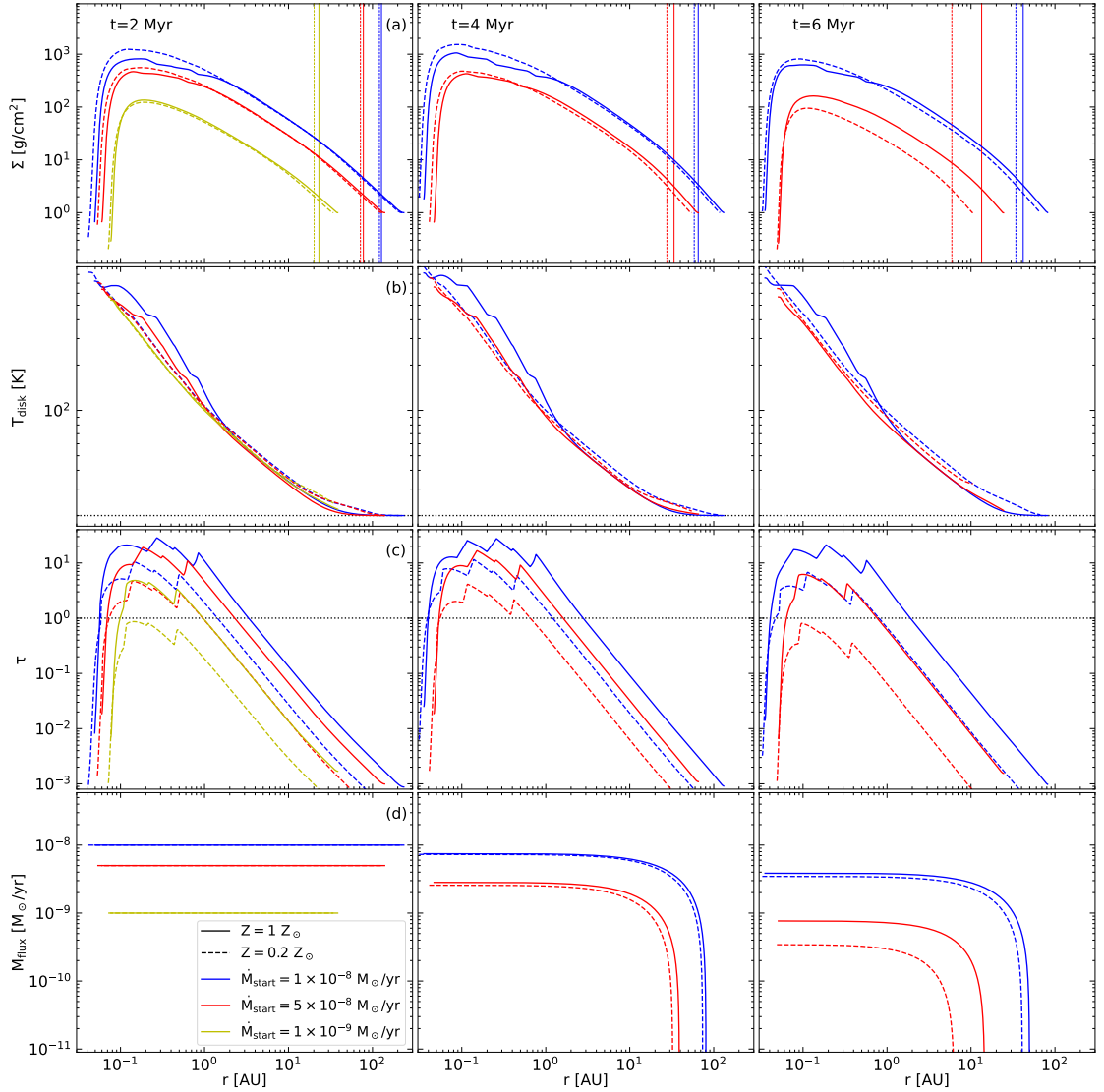


Figure 3.12.: Radial profiles for the model with $\dot{M}_{\text{start}} = 1.0 \times 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$ at the starting point of the simulation $t_0 = 2 \text{ Myr}$ (left column), as well as at 5.0 Myr (center column) and 8.0 Myr (right column). The solid (dashed) lines correspond to a metallicity of $Z = 1 Z_{\odot}$ ($Z = 0.2 Z_{\odot}$), respectively. Panels (a) shows the disk surface density Σ , panels (b) the disk midplane temperature T_{disk} , panels (c) the optical depth τ and panels (d) the mass flux $\dot{M}_{\text{flux}}$, respectively. The thin vertical lines in panels (a) indicate the respective position of r_c . The horizontal lines in panels (b) indicate the ambient temperature $T_{\text{amb}} = 20 \text{ K}$. The horizontal dashed lines in panels (c) symbolizes $\tau = 1$, dividing optical thin regions ($\tau < 1$) from optical thick regions ($\tau > 1$).

4. Paper III: What governs the spin distribution of very young < 1 Myr low mass stars

- **Title:** What governs the spin distribution of very young < 1 Myr low mass stars
- **Authors:** L. Gehrig, E. I. Vorobyov
- **Year:** 2023
- **Journal:** A&A 673, A54 (2023)
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Contribution: I have coupled a stellar spin model to the MESA stellar evolution code in this study. Setting up the parameter space, running the simulations, and writing the text have been performed by me. My co-author, EIV, has provided valuable comments throughout the process.

Summary: The origin of the stellar spin distribution at young ages is still unclear. Even in young clusters (~ 1 Myr), a significant spread is observed in rotational periods ranging from $\lesssim 1$ to ~ 10 days. We aim to study the parameters that might govern the spin distribution of low mass ($\lesssim 1.0 M_{\odot}$) stars during the first million years of their evolution. We compute the evolution and rotational periods of young stars, using the MESA code, starting from a stellar seed, and take protostellar accretion, stellar winds, and the magnetic star-disk interaction into account. Furthermore, we add a certain fraction of the energy of accreted material into the stellar interior as additional heat and combine the resulting effects on stellar evolution with the stellar spin model. For different combinations of parameters, stellar periods at an age of 1 Myr range between 0.6 days and 12.9 days. Thus, during the relatively short time period of 1 Myr, a significant amount of stellar angular momentum can already be removed by the interaction between the star and its accretion disk. The amount of additional heat added into the stellar interior, the accretion history, and the presence of disk and stellar winds have the strongest impact on the stellar spin evolution during the first million years. The slowest stellar rotations result from a combination of strong magnetic fields, a large amount of additional heat, and effective winds. The fastest rotators combine weak magnetic fields and ineffective winds or result from a small amount of additional heat added to the star. Scenarios that could lead to such configurations are discussed. Different initial rotation periods of the stellar seed, on the other hand, quickly converges and do not affect the stellar period at all. Our model matches up to 90% of the observed rotation periods in six young ($\lesssim 3$ Myr) clusters. Based on these intriguing results, we motivate to combine our model with a hydrodynamic disk evolution code to self-consistently include several important aspects such as episodic accretion events, magnetic disk winds, internal, and external photo-evaporation. Such a combined model could replace the widely-used disk-locking model during the lifetime of the accretion disk and provide valuable insights into the origin of the rotational period distribution of young clusters.

4.1. Introduction

The origin of the stellar spin and angular momentum (AM) distribution of young stars still poses problems for stellar evolution models. Observations of stellar rotation in young clusters over the last decade have provided valuable insights for the evolution of the stellar spin and AM [e.g., Herbst et al., 2002, Rebull et al., 2006, Irwin et al., 2008, Irwin and Bouvier, 2009, Hartman et al., 2009, Affer et al., 2013, Gallet and Bouvier, 2013, Rebull et al., 2020, Serna et al., 2021]. Especially the early distribution ($\lesssim 10$ Myr) of stellar spin and AM is an important benchmark for gyrochronology studies [e.g., Barnes, 2007] and can affect stellar high energy radiation of low mass stars [e.g., Pallavicini et al., 1981, Micela et al., 1985, Wright et al., 2011, France et al., 2018].

We want to focus on the first Myr of stellar evolution, in which the star is embedded in an envelope and surrounded by an accretion disk. During the accretion disk phase, it is assumed that the accretion disk plays an important role in regulating the stellar spin distribution. Mass and AM can be exchanged between the star and the disk via magnetic field lines, influencing the stellar rotation period [e.g., Armitage and Clarke, 1996, Matt and Pudritz, 2005a, Matt et al., 2010b]. Early models propose that the star rotates at the orbital period of the inner disk boundary [disk-locking, e.g., Ghosh and Lamb, 1979b, Koenigl, 1991]. However, because of a limited region, in which the magnetic field lines can connect to the accretion disk, the disk-locking mechanisms cannot remove enough AM from the star to counteract accretion and contraction and does not explain the existence of slowly rotating stars [e.g., Matt and Pudritz, 2005b,a, Zanni and Ferreira, 2013]. Since then additional mechanisms have been proposed that can remove AM from the star during the accretion disk phase: the process of accretion itself can drive accretion-powered stellar winds [APSW, Matt and Pudritz, 2005b] and disk material can be ejection along magnetic field lines [magnetospheric ejections, MEs, e.g., Zanni and Ferreira, 2013] removing AM from the star.

Even at a very young age of ~ 1 Myr, young clusters already show a significant spread in stellar rotation periods. Observations in the Orion Nebular Cluster (ONC) [e.g., Herbst et al., 2002, Serna et al., 2021], NGC 2264 [e.g., Venuti et al., 2017], Taurus [Rebull et al., 2020],

NGC 6530 [Henderson and Stassun, 2012], and other young clusters [YSOVAR, Rebull et al., 2014] show rotation periods ranging from < 1 to ~ 10 days. We note that there is a small number of stars rotating even faster or slower than this range at a young age [e.g., Rebull et al., 2014]. To date, the parameters responsible for this large spread are not constrained. Bouvier et al. [2014b] argues that processes during the cloud collapse or the embedded phase might govern the rotational distribution at ~ 1 Myr. Later studies have not solved for the very early stellar spin evolution and adopted the observed spread from ~ 1 day to ~ 10 days as initial conditions [e.g., Gallet et al., 2019, Roquette et al., 2021].

To explore the parameter and processes responsible for the rotational distribution of low-mass stars ($\lesssim 1 M_{\odot}$), we combine *Modules for Experiments in Stellar Astrophysics* [MESA Paxton et al., 2011, 2013, 2015, 2018, 2019b] with a recent stellar spin evolution model presented by Ireland et al. [2021]. Similar to current stellar evolution models, we start with a stellar seed ($M_{\star} = 0.05 M_{\odot}$) that accretes mass from the surrounding disk [e.g., Baraffe et al., 2012, Vorobyov et al., 2017b, Kunitomo et al., 2017, Steindl et al., 2021]. A fraction of the gravitational energy of the accreted material is deposited into the stellar interior as additional heat, which can lead to an inflation of the stellar radius and affect the stellar evolution significantly [e.g., Vorobyov et al., 2017b, Kunitomo et al., 2017]. The amount of additional heat depends on the fraction of energy deposited in the stellar interior, the region, in which the energy is deposited, and the accretion rate onto the star. The effect of those parameters on the stellar spin evolution, however, is not clear and will be tested in this study. Our model considers the influence of stellar inflation on the spin evolution of stars. It also takes into account different rotation rates of the stellar seed for the spin evolution of these stars. Furthermore, we can test a wide parameter space including the stellar magnetic field strength, the initial angular velocity distribution of the stellar seed, winds originating from the star and the disk, and the initial radii of the stellar seed.

Our results suggest that certain mechanisms and parameters have a major impact on the early stellar spin evolution, while others are virtually unimportant. Distinct slow and fast rotating models differ in the rotation period at an age of 1 Myr by up to two orders of magnitude. A comparison with observed rotational periods of six young clusters allows us to test these model results. This study is structured as follows: Sec. 4.2 summarizes the model highlights and defines

the parameter space used in our simulations. The results are presented in Sec. 4.3, discussed, and compared to observations in Sec. 4.4. Finally, we summarize our conclusions in Sec. 4.5.

4.2. Model and parameter space

In this study, we combine version v12778 of MESA with the stellar spin evolution model of Ireland et al. [2021]. In the following, we highlight important aspects of the respective models and provide the respective references for additional information. The adopted micro-physics and the treatment of rotation in MESA are summarized in Appendix 4.6.1 and Appendix 4.6.2, respectively. The MESA input files for re-creating our results are publicly available¹.

4.2.1. Stellar evolution with MESA

Similar to recent stellar evolution studies [e.g., Kunitomo et al., 2017, Steindl et al., 2021], we start our simulations with a stellar seed with an initial stellar mass of $M_{\text{init}} = 0.05 M_{\odot}$ ², solar metallicity, $Z_{\star} = 0.014$, and a Deuterium abundance of 20 ppm. During the initial stages of stellar evolution, disk material is accreted onto the star. The associated energy of the accreted material is partly radiated away in the form of an accretion luminosity

$$L_{\text{acc}} = (1 - \beta)\epsilon \frac{GM_{\star}\dot{M}_{\star}}{R_{\star}} \quad (4.1)$$

and partly added to the stellar interior as additional heat

$$\dot{E}_{\text{add}} = \beta\epsilon \frac{GM_{\star}\dot{M}_{\star}}{R_{\star}}, \quad (4.2)$$

with the stellar mass $M_{\star}$, the stellar radius $R_{\star}$, the gravitational constant G and the accretion rate onto the star $\dot{M}_{\star}$. The geometry of the accretion process is described by ϵ . Accretion from a thin accretion disk is described with $\epsilon \leq 0.5$ [Hartmann et al., 1997] and we choose $\epsilon = 0.5$ in this study [e.g., Baraffe et al., 2009, Steindl et al., 2021]. The amount of additional energy added

¹MESA input files to reproduce the results can be found at <https://zenodo.org/record/7763138>

²We note that Kunitomo et al. [2017] and Steindl et al. [2021] start with $M_{\text{init}} = 0.01 M_{\odot}$. The combination of high accretion rates and fast stellar rotation leads to convergence issues with MESA for stellar masses $< 0.05 M_{\odot}$

to the star is controlled by β and is assumed to depend on the accretion rate [e.g., Baraffe et al., 2012, Vorobyov et al., 2017b, Kunitomo et al., 2017, Elbakyan et al., 2019, Steindl et al., 2021]. During phases of low accretion rates, the disk is truncated at several stellar radii by magnetic field lines anchored in the star [e.g., Bessolaz et al., 2008, Hartmann et al., 2016]. Disk material is channeled through so-called funnels and is accreted onto the star at nearly free-fall speeds. Close to the stellar surface, a shock front develops, and most of the material’s energy is radiated away [e.g., Koenigl, 1991]. As a result, only a small fraction of the accreted energy is added to the stellar interior and β is small. During phases of high accretion rates, the disk material is pushed toward the star and the stellar magnetic field is no longer able to channel the disk material into the funnel flows. Without the radiative loss of energy at the shocks close to the stellar surface a larger amount of energy can be added to the star. Based on previous studies, accretion rates of $\dot{M}_\star \gtrsim 10^{-5} \text{ M}_\odot \text{ yr}^{-1}$ result in high values of $\beta = \beta_{\text{up}}$ and accretion rates of $\dot{M}_\star \lesssim 10^{-7} \text{ M}_\odot \text{ yr}^{-1}$ result in low values of $\beta = \beta_{\text{low}}$ [e.g., Baraffe et al., 2012, Vorobyov et al., 2017b, Jensen and Haugbølle, 2018]. In between the value of β transitions smoothly between β_{up} and β_{low} according to

$$\beta(\dot{M}_\star) = \frac{\beta_{\text{up}} + \beta_{\text{low}}}{2} + \frac{\beta_{\text{up}} - \beta_{\text{low}}}{2} \tanh \frac{\log \dot{M}_\star + 6}{0.5}, \quad (4.3)$$

with the accretion rates in units of $\text{M}_\odot \text{ yr}^{-1}$. In this study, we choose $\beta_{\text{low}} = 0.005$ and vary β_{up} between 0.005 (corresponding to the case of cold accretion, $\beta = \text{const.} = 0.005$) and 0.5 (corresponding to the case of warm or hybrid accretion). $\dot{E}_{\text{add}}$ is distributed linearly starting from the stellar surface, down to a relative mass coordinate M_{add} [e.g., Kunitomo et al., 2017, Steindl et al., 2021]. We choose M_{add} to be in the range of 0.01 – 0.1, meaning that $\dot{E}_{\text{add}}$ is distributed within the outer 1 – 10% of the stellar mass. $\dot{E}_{\text{add}}$ can cause the star to inflate, depending on the value of β and the position in the stellar interior, up to which this energy is added M_{add} [e.g., Baraffe et al., 2009, 2012, Vorobyov et al., 2017b, Kunitomo et al., 2017, Steindl et al., 2021].

The outer, atmospheric boundary condition is chosen according to Kunitomo et al. [2017]. For stellar radii larger than $0.7 \text{ R}_\odot$, we choose the “photospheric” table with an optical depth of $\tau_s = 2/3$, a surface temperature $T_s = T_{\text{eff}}$ and the surface pressure P_s obtained from PHOENIX [Hauschildt et al., 1999a,b] and ATLAS9 [Castelli and Kurucz, 2003] model atmospheres. For

smaller stellar radii, we adopt the “ $\tau_s = 100$ ” model, in which P_s and T_s are calculated at $\tau_s = 100$ from ATLAS9 and COND [Allard et al., 2001] model atmospheres.

4.2.2. Stellar spin evolution

In the scope of our model, the star rotates as a solid body with the stellar AM, $J_\star = I_\star \Omega_\star$; $I_\star$ and $\Omega_\star$ are the stellar moments of inertia and angular velocity $\Omega_\star$, respectively. External torques Γ_{ext} change $J_\star$ according to $\dot{J}_\star = I_\star \dot{\Omega}_\star + \dot{I}_\star \Omega_\star = \Gamma_{\text{ext}}$. This results in a stellar spin evolution according to

$$\dot{\Omega}_\star = \frac{\Gamma_{\text{ext}}}{I_\star} - \frac{\dot{I}_\star}{I_\star} \Omega_\star. \quad (4.4)$$

The moment of inertia $I_\star$ is calculated by the MESA code and the external torque contributions Γ_{ext} are calculated according to the description provided by Ireland et al. [2021]. The contributions to Γ_{ext} is composed of the influence of accretion Γ_{acc} , magnetospheric ejections Γ_{me} and stellar winds Γ_w . The internal contributions due to accretion, mass loss, and changes in the stellar radius are combined in $\Gamma_{\text{int}} = -\dot{I}_\star \Omega_\star$. An increase in stellar mass or the stellar radius results in a positive $\dot{I}_\star$ and the star spins down, for example, during phases of stellar radius inflation. On the other hand, the star spins up if the star contracts. According to Ireland et al. [2021], the external torque contributions depend on different stellar parameters and disk parameters. We want to highlight the general dependencies and specific trends of the individual torque values. The applicability of the spin model presented in Ireland et al. [2021] for this study is discussed in Sec. 4.4.1. For additional information, we refer to the original article.

Accretion torque Γ_{acc} : Material that accretes onto the star usually adds the AM of the disk material at the inner disk boundary to the star [e.g., Matt et al., 2010a, Gallet et al., 2019, Ireland et al., 2021], increasing the stellar AM, and causing a stellar spin-up according to

$$\Gamma_{\text{acc}} = |\dot{M}_\star|(\Omega(R_t) - \Omega_\star)R_t^2, \quad (4.5)$$

with the disk rotation rate at the truncation radius R_t , $\Omega(R_t)^3$. The truncation radius in units of the

³We note that the formulation of Γ_{acc} in Ireland et al. [2021] do not include the $-\Omega_\star$ term. They take into account stellar rotation periods much smaller compared to the rotation rates at the inner disk boundaries. At young ages, however, fast rotation rates close to break-up should be considered and the additional term should be included.

stellar radius can be written as

$$\frac{R_t}{R_\star} = K_t \left(\frac{B_\star^2 R_\star^2}{4\pi |\dot{M}_\star| v_{\text{esc}}} \right)^{m_t}, \quad (4.6)$$

with $K_t = 0.756$, $m_t = 0.34$, the stellar dipole field strength $B_\star$, and the escape velocity from the stellar surface $v_{\text{esc}} = (2GM_\star/R_\star)^{1/2}$. The disk at the truncation radius is assumed to rotate slower compared to the Keplerian velocity [e.g., Bessolaz et al., 2008, Zanni and Ferreira, 2013] due to the removal of AM from the disk due to the magnetic star-disk interaction [e.g., Zanni and Ferreira, 2013] or disk winds [e.g., Zhang et al., 2017]. Thus, $\Omega(R_t)$ can be written as

$$\Omega(R_t) = K_{\text{acc}} \left(\frac{R_t}{R_\star} \right)^{m_{\text{acc}}} \Omega_K(R_t), \quad (4.7)$$

with the correction factor for the non-Keplerian rotation $K_{\text{acc}} = 0.775$, $m_{\text{acc}} = -0.147$, and the Keplerian rotation rate at the truncation radius $\Omega_K(R_t)$.

Magnetic star-disk interaction: The influence of the magnetic star-disk interaction in Ireland et al. [2021] is modeled as magnetospheric ejections [MEs, Zanni and Ferreira, 2013]. MEs can transfer angular momentum between the star and the disk due to the twist of magnetic field lines that are rooted at the stellar surface and coupled to the disk material [e.g., Armitage and Clarke, 1996, Zanni and Ferreira, 2013, Ireland et al., 2021]. The amount of angular momentum transferred to or away from the star depends on the relation between the truncation radius R_t and the co-rotation radius

$$R_{\text{co}} = \left[K_{\text{acc}} \left(\frac{R_t}{R_\star} \right)^{m_{\text{acc}}} \right]^{2/3} \left(\frac{GM_\star}{\Omega^2} \right)^{1/3}, \quad (4.8)$$

which includes the effect of non-Keplerian rotation. Disk material, which is coupled to the stellar magnetic field lines inside (outside) R_{co} rotates faster (slower) compared to the star and increases (decreases) the stellar angular momentum. This process is highly dynamic and complex. Thus, the following formulation should be treated with caution and interpreted as a time-averaged approximation. Following Ireland et al. [2021], the total torque due to MEs can be written as

$$\Gamma_{\text{ME}} = K_{\text{ME}} B_\star^2 R_\star^3 \left(\frac{R_t}{R_\star} \right)^{m_{\text{ME}}} \left[1 - \left(\frac{R_t}{R_{\text{co}}} \right)^{3/2} \right], \quad (4.9)$$

with numerical fit parameters $K_{\text{ME}} = 0.00772$ and $m_{\text{ME}} = -2.54$. If $R_t < R_{\text{co}}$ ($R_t > R_{\text{co}}$), the total torque due to MEs is positive (negative) and the star spins up (spins down).

Accretion-powered stellar wind (APSW): In the picture of an accretion-powered stellar wind, a certain fraction W of the accretion rate is ejected by the star; $\dot{M}_w = W\dot{M}_\star$. According to Weber and Davis [1967], the torque due to such a stellar outflow can be written as

$$\Gamma_w = -|\dot{M}_w|\Omega_\star r_A^2, \quad (4.10)$$

with the Alfvén radius r_A . The Alfvén radius can be related to the open stellar magnetic flux [Réville et al., 2015a] and based on numerical best-fit parameters [Ireland et al., 2021], the torque due to an APSW can be written as

$$\Gamma_w = -K_{w1}|\dot{M}_\star|(GM_\star R_\star)^{1/2} W^{m_{w1}} f^{m_{w2}} \left[1 + \left(\frac{f}{K_{w2}} \right)^2 \right]^{-m_{w3}} \left(\frac{R_t}{R_\star} \right)^{m_{w4}}, \quad (4.11)$$

with best-fit parameters $K_{w1} = 57.92$, $m_{w1} = 0.254$, $m_{w2} = 1.092$, $K_{w2} = 0.0356$, $m_{w3} = 0.373$, $m_{w4} = 0.195$, and the fraction of the break-up rotation rate $f = \Omega_\star/\Omega_{\text{crit}}$. The torque due to stellar winds Γ_w always removes AM from the star.

4.2.3. Effects of different stellar rotation rates

Based on this spin evolution model presented above, we can predict the behavior of a star that rotates at a certain rate Ω_{old} and spins up to $\Omega_{\text{new}} > \Omega_{\text{old}}$, with all other parameters unchanged. First, the accretion torque Γ_{acc} is reduced (see Eq. 4.5). Second, the torque due to the magnetic star-disk interaction Γ_{ME} is smaller as the co-rotation radius moves closer to the truncation radius for fast rotation rates (see Eq. 4.9). Third, the rate of angular momentum that is removed by the APSW, Γ_w , increases (see Eq. 4.10). An increasing stellar rotation rate ($\Omega_{\text{new}} > \Omega_{\text{old}}$) would thus result in a spin-down torque and the star would return to Ω_{old} . Based on the same principles, a decreasing stellar rotation rate ($\Omega_{\text{new}} < \Omega_{\text{old}}$) would result in a spin-up torque. Hence, Ω_{new} evolves back to Ω_{old} , and the timescale, on which this change occurs is studied in Sec. 4.3.2.

4.2.4. Accretion rate onto the star

In this study, we start from an initial mass of $M_{\text{init}} = 0.05$ and consider three different stellar masses at an age of 1 Myr; $M_{\text{final}} = 0.3, 0.7$, and $1.0 M_{\odot}$. The accretion rate is defined in two regimes. Initially, we assume a high, constant accretion rate of $\dot{M}_{\text{init}} = 5 \times 10^{-6} M_{\odot} \text{ yr}^{-1}$, which is equivalent to the mass infall rate $\dot{M}_{\text{infall}} \propto c_s^3/G$ with the isothermal sound speed c_s and an external temperature of $\approx 21 \text{ K}$ [e.g., Vorobyov et al., 2017b]. Thus, an initial stellar mass of $M_{\text{init}} = 0.05$ corresponds to an age of $\sim 10^4$ years, assuming a roughly constant accretion rate during this time. After a specific time t_{acc} , the accretion rate decreases according to a power law with an exponent of -1.5, $\dot{M}_{\star}(t) \propto t^{-1.5}$. A stellar mass of 0.3, 0.7, and $1.0 M_{\odot}$ can be obtained at 1 Myr for $t_{\text{acc}} = 21.4, 54.5$ and 81.5 kyr , respectively. Fig. 4.1 shows the evolution of the stellar accretion rate (Panel a) and the stellar mass (Panel b) for a final stellar mass of 0.3, 0.7 and $1.0 M_{\odot}$ (solid colored lines). To study the influence of a different accretion history, we increase the initial accretion rate to $\dot{M}_{\text{init}} = 2 \times 10^{-5} M_{\odot} \text{ yr}^{-1}$ and decrease the power-law exponent to -2 ($\dot{M}_{\text{high}}$ dashed line). The final stellar mass at 1 Myr is also $0.7 M_{\odot}$ in case of using $\dot{M}_{\text{high}}$.

4.2.5. Parameter space

Finding reasonable ranges for stellar and accretion parameters for very young, embedded objects is challenging due to theoretical and observational uncertainties and a limited number of observations. The ranges for the free parameters presented below are based on previous studies, theoretical models, and observational results.

Magnetic field strengths: One important stellar parameter that influences the spin evolution model presented above is the magnetic field strength. It is assumed that the large-scale dipole field components spin evolution of young stars [e.g., Finley and Matt, 2018]. Small-scale field components, on the other hand, play only a minor role. We note that the results presented in Finley and Matt [2018] do not cover very young, embedded objects with high accretion rates (see Sec. 4.4.1). Due to observational limitations, there is only a limited number of magnetic field estimates for Class I (0) objects. Based on the observation of magnetically-sensitive line broadening (Zeeman effect), a total magnetic field strength of $2.9 \pm 0.4 \text{ kG}$ for WL 17 [Johns-Krull

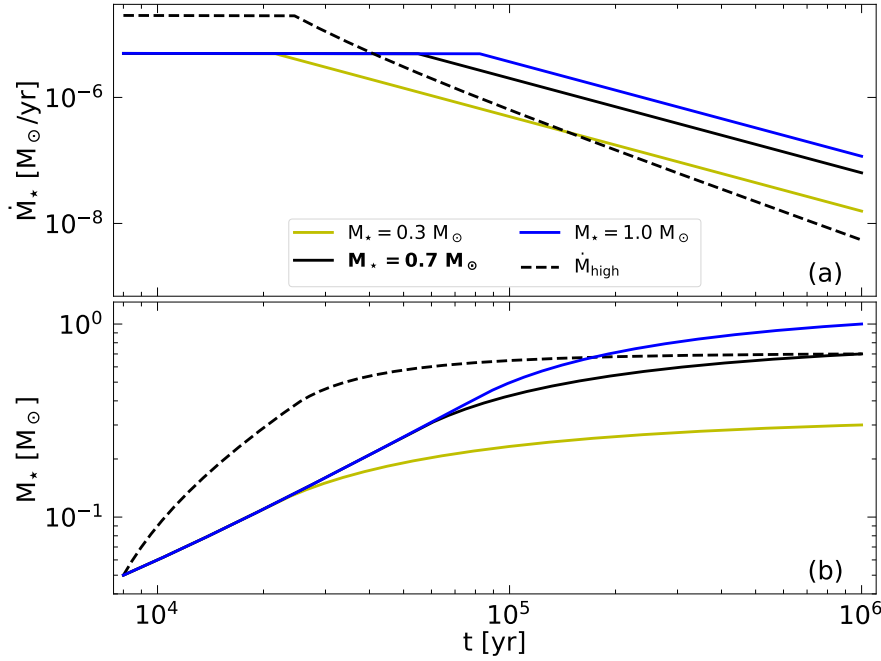


Figure 4.1.: Panel (a) shows the accretion rate and Panel (b) shows the stellar mass over time. The yellow, black, and blue lines represent a final stellar mass of 0.3, 0.7, and $1.0 M_\odot$, respectively. The black dashed line ($\dot{M}_{\text{high}}$) represents a different accretion history, which also results in a final stellar mass of $0.7 M_\odot$ at 1 Myr.

et al., 2009] and 1.36 ± 0.06 kG for V347 Aur [Flores et al., 2019] can be inferred. Unfortunately, the total field strength combines the small and large-scale features of the magnetic field and a distinction is only possible with additional (more precise) measurement [e.g., Zeeman-Doppler imaging, ZDI, Johnstone et al., 2014]. Such observations are currently not available for embedded Class I (0) objects. In addition, the small number of available measurements does not have to be representative. There are, however, estimates on the magnetic field strength of Class I (0) objects ranging from ~ 0.1 kG to ~ 1 kG based on flux conservation during the core collapse [Tsukamoto et al., 2022], near-infrared K-band spectra [Laos et al., 2021], and laboratory plasma experiments [Burdonov et al., 2021]. These estimates match the inferred field strengths for WL 17 and V347 Aur. Furthermore, Laos et al. [2021] argue that Class I (0) objects should have similar accretion mechanisms and magnetic field strengths to Class II objects. For these slightly further evolved objects, the large-scale field component ranges from ~ 0.1 kG to $\gtrsim 1$ kG [e.g., Johnstone

et al., 2014]. Thus, we choose a range for the large-scale magnetic field strength, $B_\star$, from 0.1 kG to 2.0 kG. These values are kept constant throughout the simulations in Sec. 4.3. In addition, we explore the possibility of mass-dependent magnetic field strengths as presented in Browning et al. [2016] (see Sec. 4.4.5).

Initial rotation rates: While there are rotational studies of young (~ 1 Myr) T Tauri stars in NGC 2264 [Venuti et al., 2017], ONC [Rebull et al., 2006], σ Ori [Cody and Hillenbrand, 2010], and Taurus [Rebull et al., 2020], there are no large-scale surveys of stellar rotation periods of Class I (0) objects. Similar to magnetic fields, there are only singular estimates for individual stars. The Class I object Elias 29 has a rotation period of ~ 2 days [Pillitteri et al., 2019]. Furthermore, White et al. [2007] suggest that Class I objects rotate slightly faster than their Class II counterparts. This would result in rotation periods of ~ 1 day during the Class I phase. The initial rotation periods at the beginning of the Class I (0) phase, however, are not well constrained. Thus, we explore a wide range of initial rotation rates from very slow-rotating stars with $\Omega_{\text{init}} = 0.01\Omega_{\text{crit}}$ to fast-rotating stars with $\Omega_{\text{init}} = 0.9\Omega_{\text{crit}}$, with Ω_{crit} being the critical or brake-up rotation rate, which is usually estimated as [e.g., Maeder, 2009]

$$\Omega_{\text{crit}} = \sqrt{\frac{GM_\star}{(1.5R_\star)^3}}. \quad (4.12)$$

Initial stellar radius: Studies of early stellar evolutionary stages that start shortly after the second collapse and the formation of the second hydrostatic core (with an initial stellar mass of $0.01 M_\odot$) assume stellar radii of $0.25 R_\odot$ to $3.0 R_\odot$ [e.g., Hosokawa et al., 2011, Kunitomo et al., 2017, Steindl et al., 2021]. According to Kunitomo et al. [2017], initial stellar radii of $\gtrsim 1.5 R_\odot$ are required to explain the most luminous stars and small initial radii can explain the low luminosities, and the observed luminosity spread in young clusters [Hillenbrand, 2009, Soderblom et al., 2014]. Although we start with slightly further evolved initial stellar models (with an initial mass of $0.05 M_\odot$), we assume similar initial stellar radii ranging from $0.6 R_\odot$ to $2.5 R_\odot$.

Accretion powered stellar wind efficiency: In this study, winds originating from the star are represented by accretion-powered stellar winds (APSW). Such winds extract a certain amount

of angular momentum from the star and cause a spin-down (see Eq. 4.10). The exact driving mechanisms of these winds are still under debate [see, e.g., the discussion in Gallet et al., 2019]. We adapt the driving mechanisms proposed by Decampli [1981] and Matt and Pudritz [2005b]. Angular momentum can be removed from the star by Alfvén waves, excited by the accreting material at the stellar surface and traveling outwards. The mass-loss rates of such winds are assumed to be in the order of one percent of the accretion rate onto the star [$W \sim 1\%$, e.g., Cranmer, 2008, Pantolmos et al., 2020]. Thus, the APSW efficiency value W ranges from 0% to 5% in this study. We note that there are also other wind-driving mechanisms affecting the star-disk system (e.g., jets launched at the inner disk). We discuss the possible impact of such winds or jets on the rotational evolution of the young star in Sec. 4.3.4.

4.2.6. Reference model

We define a reference model, to qualitatively show how each parameter affects the rotational evolution of young stars. In the following sections, the reference values are marked in boldface. The reference parameters are summarized in Table 4.1.

Table 4.1.: Reference model parameters used in this study. We note that the stellar mass corresponds to an age of 1 Myr.

$M_\star$	$0.7 M_\odot$	Ω_{init}	$0.1 \Omega_{\text{crit}}$
$B_\star$	1.0 kG	R_{init}	$1.1 R_\odot$
W	2%	β_{up}	0.2
M_{add}	0.1		

4.3. Results

4.3.1. Evolution of the reference value

Before going into detail on the individual parameters, we want to show the evolution of the reference model (see Fig. 4.2). The choice of initial conditions leads to an initial rotation period of 6 days. During the first few thousand years, the star spins up to ≈ 4 days. This initial spin-up can be explained by adjusting the rotation period to the accretion rate when starting the simulation.

Starting from the initial stellar radius of $R_{\text{init}} = 1.1 R_{\odot}$, the star inflates due to the energy of the accreting material added into the stellar interior and the onset of Deuterium burning. Up to an age of ≈ 18 kyr, the inflation of the star (combined with the APSW) can counteract the spin-up influence of accretion and the star-disk interaction and the star spins down towards 6 days ($\Gamma_{\text{tot}} < 0$, in Panel c of Fig. 4.2). With the increasing stellar radius, the spin-up influence of the MEs increases as well, and the stellar period decreases toward 4 days up to an age of 80 kyr. At ~ 80 kyr, the stellar inflation increases for a short period of time, resulting in a brief spin-down. Afterward, the star contracts for the remaining simulation, reaching a final stellar radius of $1.84 R_{\odot}$ after 1 Myr. During contraction, Γ_{int} increases the stellar spin up and the star reaches a period of 1.6 days after 1 Myr.

4.3.2. Effects of individual parameters

Now we examine how the stellar rotational evolution is affected by the initial rotation rate Ω_{init} , the stellar magnetic field strength $B_{\star}$, the APSW efficiency W , the initial stellar radius R_{init} , the final stellar mass at 1 Myr $M_{\star}$, and the amount of energy of the infalling material added into the star, β_{up} (see Fig. 4.3). For a clear description of different effects, we show, in addition to the stellar rotation period, the evolution of the stellar radius in Fig. 4.4, the individual torque contributions in Appendix 4.6.3, and the HRD for each model in Fig. 4.15.

The initial rotation rate: First, we want to show the effects of different initial rotation rates, Ω_{init} (Panel a in Fig. 4.3 and Fig. 4.4). We vary Ω_{init} from 1% to 90% of the critical (or break-up) velocity. The black, blue, and yellow lines symbolize the reference, fast, and slow rotating models. Within the first ~ 5 kyr, the initial difference in the stellar periods of ≈ 59 days converges towards a common value (see Sec. 4.2.3). As a consequence of the fast initial rotation and strong centrifugal forces, the initial stellar radius for a star with $\Omega_{\text{init}} = 0.9 \Omega_{\text{crit}}$ is slightly larger compared to the other models. However, this difference disappears after a few thousand years and the stellar evolution toward 1 Myr is completely unaffected by the initial rotation rate Ω_{init} .

Stellar magnetic field strength: The stellar magnetic field strength does affect all external torque contributions, such as Γ_{acc} , Γ_{ME} , and Γ_{W} . The torque due to MEs, however, is affected most strongly (see Eq. 4.9 in Sec. 4.2.2). As shown in Panel (b) of Fig. 4.3, a strong magnetic

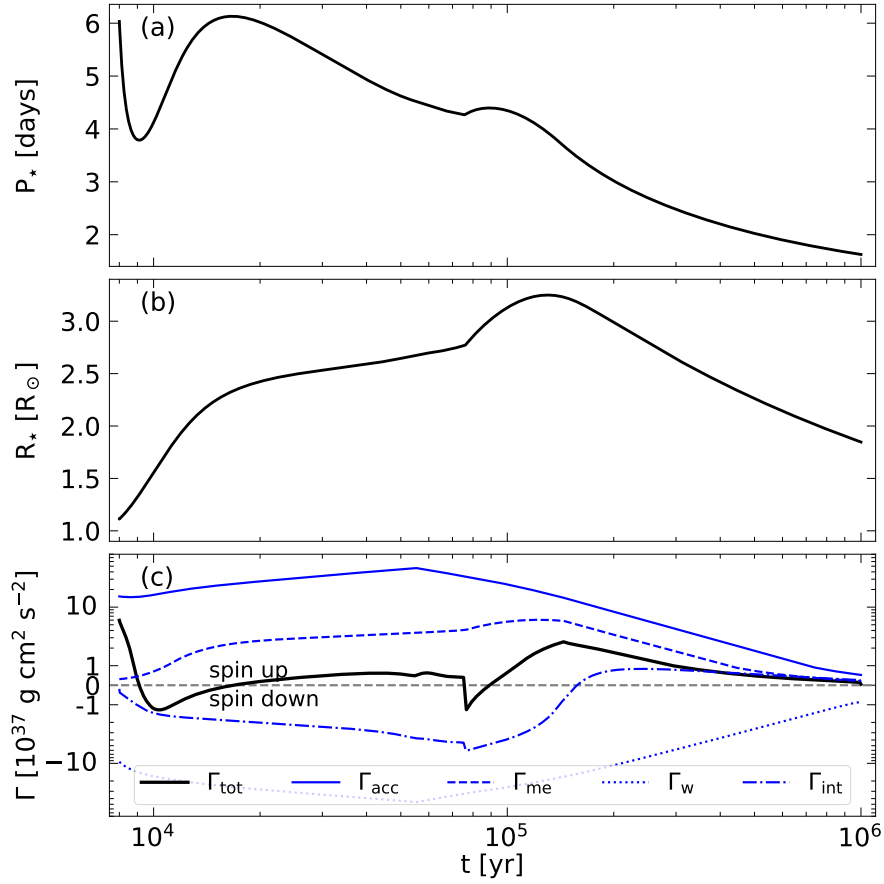


Figure 4.2.: Evolution of the reference model. The stellar rotation period and radius are shown in Panels (a) and (b), respectively. The different torque contributions are shown in Panel (c). The accretion torque, Γ_{acc} , the influence of the magnetic star-disk interaction, Γ_{ME} , the APSW torque, Γ_{W} , and the internal contributions due to variation of the stellar moment of inertia, Γ_{int} , are shown as a solid blue line, dashed blue line, dotted blue line, and dash-dotted blue line, respectively. The total torque, Γ_{tot} , defined as the sum of the other torque contributions is shown as a solid black line. A grey dashed line marks the change between the spin-up and spin-down regime.

field causes the star to spin up during the first ~ 100 kyr. After ~ 100 kyr, a strong magnetic field causes a spin down. The strong and weak magnetic field strengths are symbolized by the blue and yellow lines, respectively. This behavior can be explained with Eq. (4.9). During the initial ~ 100 kyr, the star is still heavily embedded and the accretion rates can be assumed to be high (see Fig. 4.1). The truncation radius, R_t , is pushed inwards towards the star (regardless of the magnetic field strengths used in this study), and the ratio between R_t and the co-rotation radius,

R_{co} , is small. Thus, the torque due to the magnetic star-disk interaction, Γ_{ME} , is positive, and a stronger magnetic field strength results in a stronger spin-up torque. After ~ 100 kyr, the accretion rate decreases, and a stronger magnetic field pushes the truncation radius outwards closer to the co-rotation radius. The disk's rotation at the truncation radius, $\Omega(R_t)$, and the difference between $\Omega(R_t) - \Omega_\star$ decreases. Consequently, stronger magnetic fields result in a weaker accretion spin-up torque (see Eq. 4.5), a weaker torque due to the magnetic star-disk interaction, Γ_{ME} (see Eq. 4.9), and a slower rotation period compared to weak magnetic field strengths. The differences in the rotation periods are relatively small (~ 1 day) and the stellar radius evolution is almost unaffected (see Panel b of Fig. 4.4).

APSW efficiency W : In the current spin evolution model, a stellar wind removes angular momentum from the star and causes a spin-down. The amount of AM that is removed scales with the mass loss, which is proportional to the APSW efficiency W (see Eq. 4.10). Thus, a higher efficiency value results in a stronger spin-down torque, and the star slows down compared to a weaker or no APSW (see Panel d in Fig. 4.3). The difference, especially during the early evolution ($\lesssim 10^5$ years) can be significant. With no angular momentum removed by an APSW ($W = 0\%$), the star spins up to 81.7% of the critical rotation rate (see the dotted lines in Fig. 4.3). The stellar rotation period differs by more than 7 days, equivalent to a factor of ≈ 4 . The stellar radius is also affected by the different rotation periods. A fast-rotating star shows a larger radius (Panel c in Fig. 4.4) and a lower effective temperature (see Panel c in Fig. 4.15). Thus, the first nuclear reactions ignite at a later age (see the shifted radius evolution in Panel c of Fig. 4.4). While, toward an age of 1 Myr, the rotation periods still show a difference of a factor of ~ 2 , the stellar radii converge to a similar value.

Maximum amount of additional heat $\dot{E}_{\text{add}}$: The amount of additional heat can be controlled by the free parameter β_{up} ⁴. The value of β_{up} does have a significant impact on stellar evolution. While the evolution of most stars can be explained by a substantial amount of additional heat added to the star [corresponding to $\beta_{\text{up}} \gtrsim 0.1$, e.g., Vorobyov and Pavlyuchenkov, 2017, Kunitomo et al., 2017, Steindl et al., 2021], objects with very low luminosity require $\beta \sim \beta_{\text{low}}$ [Kunitomo et al.,

⁴We expect that the stellar evolution is largely independent of β_{low} , as long the value is chosen to be small enough, $\beta_{\text{low}} \ll 0.1$.

2017, Vorobyov et al., 2017a]. The amount of additional heat affects the evolution of the stellar radius. The more energy added to the stellar interior, the stronger the star inflates (see Panel d in Fig. 4.4). In the cold-accretion scenario ($\beta = \beta_{\text{low}}$), the stellar radius does not inflate and remains approximately constant at a value of $\sim 1 R_{\odot}$. These different evolution tracks influence stellar rotation (see Panel d in Fig. 4.3). The strong initial inflation to $R_{\star} > 3 R_{\odot}$ for $\beta_{\text{up}} = 0.5$ results in a spin-down to 9.2 days. With larger radii, MEs become more effective and start to spin up the star slowly. After 1 Myr, the stellar period has reached ~ 2 days. Without the initial inflation, the star continuously spins up during its evolution (excluding the short period of inflation due to the first nuclear reactions). After 1 Myr, the stellar period has reached ~ 0.57 days. Within the parameter ranges studied so far, a variation in β_{up} marks the largest impact on the stellar rotation period.

Stellar mass at 1 Myr: To reach different stellar masses at 1 Myr, we vary the age, at which the initially constant accretion rate decreases (see Fig. 4.1). During the first $\sim 10^5$ years, smaller final stellar masses result in faster-rotating stars (see Panel e in Fig. 4.3). To explain this behavior, the amount of additional heat has to be considered. For a decreasing accretion rate, β_{up} and consequently $\dot{E}_{\text{add}}$ decrease as well. Thus, the stellar inflation stops and the star starts contracting, resulting in a stellar spin-up. After 10^5 years, the influence of the additional heat diminishes as accretion rates decrease further. Now, a lower accretion rate results in less accretion torque, Γ_{acc} , and a smaller stellar radius results in less effective MEs. As a consequence, a smaller accretion rate causes a slower rotation compared to high accretion rates.

Initial stellar radii: Finally, we show the impact of different initial stellar radii on the spin evolution. For a given stellar mass, a larger radius results in a cooler temperature throughout the star. Starting at a large initial stellar radius, gravity prevails over the inflation due to the additional heat, and the star contracts slightly (see Panel f in Fig. 4.4). After ~ 20 kyr, the inflation of the stellar radius has balanced the initial differences. The temperature deep in the stellar interior, however, is still cooler in case of an initially larger radius and the first nuclear reactions start slightly later (maximum radii are reached later). Reaching 1 Myr, the stellar radii converge toward a common value and the initial differences are erased. A similar picture emerges for the stellar rotation periods. Initial differences converge fast and throughout the evolution, only

small differences arise due to different stellar radii. After 1 Myr, the rotation period evolved independently of the initial stellar radius. In our reference case, we choose $\beta_{\text{up}} = 0.2$, which corresponds to a warm accretion scenario. Kunitomo et al. [2017], however, pointed out that the importance of the initial stellar radius increases, when decreasing the amount of additional heat. Hence, additional simulations have been carried out with different initial radii and $\beta_{\text{up}} = 0.005$ (see Appendix 4.6.4). Without the additional heat and the resulting stellar inflation, the stellar radii maintain a significant difference throughout their evolution and after 1 Myr, the stellar radii differ by a factor of ~ 2 (see Panel b in Fig. 4.14). As a result, the stellar rotation period does not converge and shows a dependence on the initial stellar radius. Smaller stellar radii result in faster-rotating stars and after 1 Myr, the rotation periods differ by a factor of 2.6 (see Panel a in Fig. 4.14).

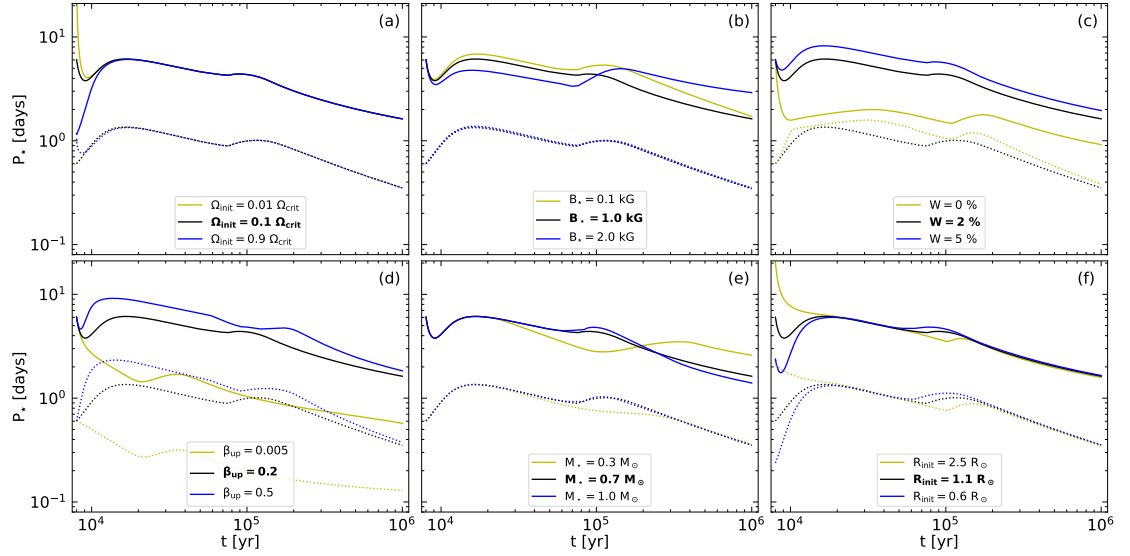


Figure 4.3.: Evolution of the stellar rotation period, P_* , for each parameter variation: the initial rotation rate Ω_{init} (Panel a), the stellar magnetic field strength B_* (Panel b), the APSW efficiency W (Panel c), the amount of energy of the infalling material added into the star, β_{up} (Panel d), the final stellar mass at 1 Myr, M_* (Panel e), and the initial stellar radius, R_{init} (Panel f). The reference values are marked in boldface. For comparison, the critical rotation period is shown for each evolution as dotted lines.

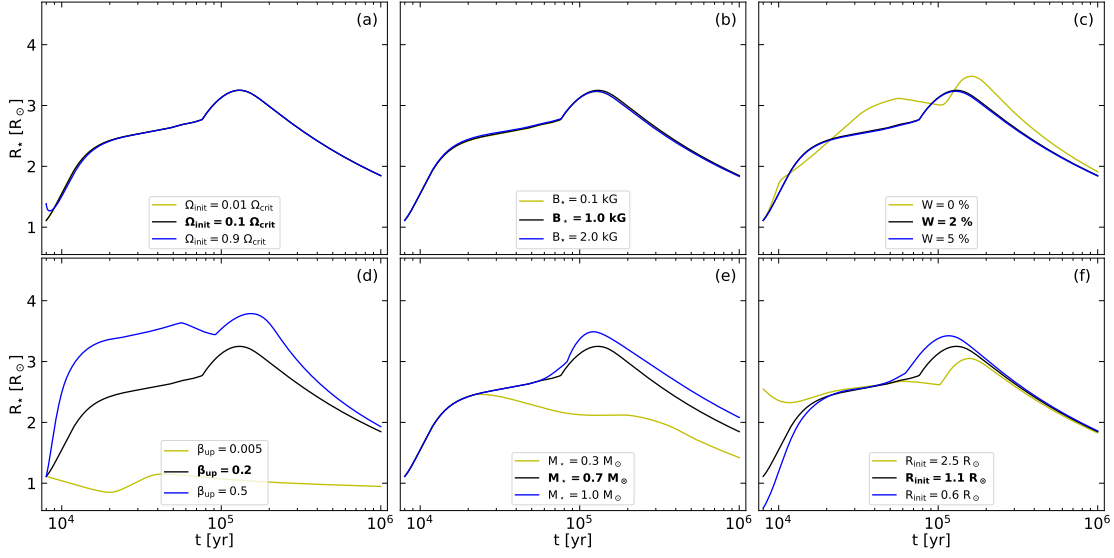


Figure 4.4.: Evolution of the stellar radius, R_* , for each parameter variation (same as in Fig. 4.3). The reference values are marked in boldface.

4.3.3. Other factors affecting $\dot{E}_{\text{add}}$

The additional heat, $\dot{E}_{\text{add}}$, is not only controlled by β . The region, in which $\dot{E}_{\text{add}}$ is added, and the accretion history can influence the stellar evolution. We reduce the region, in which the additional heat, $\dot{E}_{\text{add}}$, is distributed to $M_{\text{add}} = 0.01$. With more additional heat injected close to the stellar surface, the initial stellar inflation increases (see the sharp increase of the stellar radius in Column a of Fig. 4.5). During the whole evolution, the stellar radius remains slightly larger compared to the reference model. Reaching 1 Myr, the stellar radius converges toward the reference value as the influence of $\dot{E}_{\text{add}}$ becomes less important as the accretion rate decreases. As discussed before, larger inflated stars tend to rotate more slowly. During the strong initial inflation, the difference to the reference model is the largest (~ 3 days). Similar to the stellar radius, reaching an age of 1 Myr, the stellar period converges toward the reference value.

Next, we change the accretion history. Starting from an initial accretion rate of $2 \times 10^5 \text{ M}_\odot \text{ yr}^{-1}$, which is factor 4 larger compared to the reference value, the accretion rate drops earlier and faster to reach the same stellar mass ($0.7 \text{ M}_\odot$) compared to the reference model (see Fig. 4.1). The strong accretion rate injects a large amount of additional heat into the stellar interior and the

star inflates rapidly to $> 4 R_{\odot}$ (see Column a of Fig. 4.5). Due to this immense expansion, the star slows down to over 10 days. After the accretion rate drops, the stellar expansion stops and the star contracts toward the reference value, reaching the end of the simulation. In contrast to the previous model with $M_{\text{add}} = 0.01$, the smaller accretion rate at ages $> 10^5$ years reduces the accretion torque and the torque due to MEs and the star spins down. At the age of 1 Myr, the stellar rotation period differs by a factor of 1.9.

While the region, in which the additional heat is injected, plays only a minor role in the stellar spin evolution, the accretion history significantly affects the stellar period. In our model, the accretion rate is assumed to be smooth, without short-term variations. This idealistic assumption, however, should merely be considered as an average over time. In hydro-dynamic disk simulations, especially the embedded phase ($< 10^5$ years), show episodic outbursts with peak accretion rates of $> 10^5 M_{\odot} \text{ yr}^{-1}$ [e.g., Vorobyov and Basu, 2015, Vorobyov et al., 2017b, Hsieh et al., 2019, Kadam et al., 2021]. During such an outburst, a huge amount of additional heat is injected into the stellar interior, the star inflates, and the rotation period increases. The nature of these outbursts depends, amongst other things, on the stellar mass. More massive stars undergo more and stronger outbursts over a larger period of their early life. If each outburst spins down the star slightly, more massive stars should tend to rotate more slowly compared to low-mass stars. This trend is, in fact, reported in Henderson and Stassun [2012] for young ($\lesssim 5$ Myr) clusters.

4.3.4. Disk winds

In the presence of a large-scale magnetic field in the accretion disk, disk winds, launched by magneto-rotational processes in the inner disk regions, can remove a significant amount of angular momentum from the disk [e.g., Zhang et al., 2017]. The exact regions, in which the wind is launched, and the total amount of mass and angular momentum transported away, depend on a variety of factors [e.g., Guilet and Ogilvie, 2014, Steiner et al. in prep]. A detailed description of this process would go beyond the scope of this work. However, we can approximate the effect of a strong disk wind on the stellar rotation. The loss of angular momentum in the inner disk region causes the disk material to slow down and reduces the accretion torque. For our simulation, we extend Eq. 4.7 by a factor of $K_{\text{DW}} = 0.5$. We note that this factor is most likely not constant

and depend on different parameters, but for showing the qualitative effect of a disk wind, this assumption should be viable. The rotational evolution including a disk wind is shown in Column (b) of Fig. 4.5 (yellow line). As expected, the reduced accretion torque leads to a rapid spin-down during the radius inflation. Reaching 1 Myr, the rotational period is still slower by a factor of 2.0 compared to the reference model, which is similar to the effect of the model with $\dot{M}_{\text{high}}$.

4.3.5. Stellar metallicity

Finally, stellar metallicity can affect the rotational evolution of the star during the pre-main sequence and the main sequence. Low-metallicity stars are expected to rotate faster compared to their solar-metallicity counterparts [e.g., Amard et al., 2019a, Amard and Matt, 2020, Amard et al., 2020, Gehrig et al., 2023]. As low-metallicity stars are more compact compared to solar metallicities, mechanisms that remove angular momentum from the star and scale with the stellar radius are less effective in a low-metallicity environment during the Class II phase and beyond the main sequence [e.g., Amard and Matt, 2020, Gehrig et al., 2023]. In addition, the disk lifetime of low-metallicity stars is shorter compared to solar metallicities, and the pre-main sequence spin-up due to contraction starts at an earlier age [Yasui et al., 2016, 2021, Guarcello et al., 2021, Gehrig et al., 2023]. We find similar results for the evolution during the first million years for a metallicity of $Z_{\star} = 0.1 Z_{\odot}$ (see Column b in Fig. 4.5, blue line). The initial inflation is weaker and the star is more compact during the whole simulation time compared to the reference model. Thus, the star spins up during the embedded phase. Reaching the age of 1 Myr, the APSW is less effective in removing angular momentum from the star due to the smaller stellar radius and the star continues to spin up compared to the reference model.

4.3.6. Summarizing the effects on $P_{\star}$

In this work, we have studied the influence of various stellar and disk parameters on the stellar spin evolution during the first million years of stellar evolution. We want to summarize our results in an illustrative way in Table 4.2. We divide the stellar evolution into two phases: the embedded phase $\lesssim 10^5$ year and the later ($\gtrsim 10^5$ years) phase, which corresponds to the Class II or T Tauri phase. For each parameter variation, we indicate if the star rotates faster (red), slower (blue), or

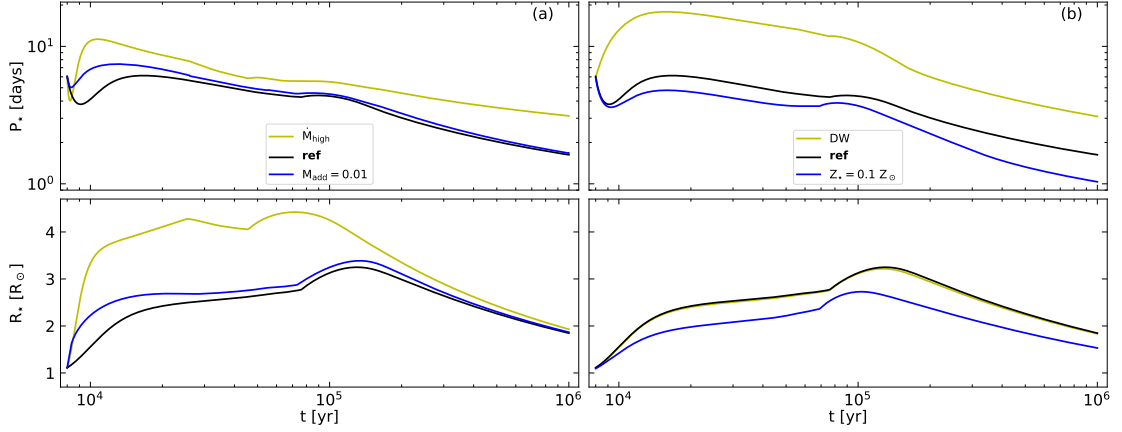


Figure 4.5.: Evolution of the stellar rotation period, P_* , and the stellar radius, R_* , for different variations from the reference model. Column (a): region, in which $\dot{E}_{\text{add}}$ is distributed, $M_{\text{add}} = 0.01$ (blue line) and using a different accretion history, $\dot{M}_{\text{high}}$ (see Fig. 4.1, yellow line). Column (b): low metallicity, $Z_* = 0.1 Z_{\odot}$ (blue line), and including the influence of disk winds (model DW, yellow line). The reference model is marked in boldface.

approximately unchanged (grey) compared to the reference model.

4.4. Discussion

4.4.1. Applicability of the spin model for embedded stars

In this study, we use the stellar spin evolution model presented in Ireland et al. [2021] that depends strongly on best-fit parameters and proportionality constants (see Eqs. 4.5 - 4.11). These parameters are based on models with smaller accretion rates $\dot{M}_{\text{acc}} \lesssim 10^{-8} M_{\odot} \text{ yr}^{-1}$. Such accretion rates are typically found in older, further evolved systems with ages of ~ 1 Myr, which marks the end of our simulations. In their work, Ireland et al. [2021] recommend the use of their formulation for $R_* \leq R_t \leq R_{\text{co}}$, which is the case for all our models. However, the large accretion rates can result in effects not considered by Ireland et al. [2021].

Disk truncation radius: The disk truncation radius, R_t , depends on the balance between stellar magnetic pressure and ram pressure of the infalling material [e.g., Hartmann et al., 2016]. For accretion rates of $\dot{M}_{\text{acc}} \lesssim 10^{-8} M_{\odot} \text{ yr}^{-1}$, the truncation radius is usually located at several stellar

Table 4.2.: Summary of the influence of different stellar and disk parameters on the stellar spin evolution. For each parameter, we distinguish between 3 different effects: the star rotates faster (red), slower (blue), or approximately unchanged (grey) compared to the reference model.

	Evolution phase	
	Embedded	Class II
fast Ω_{init}	unchanged	unchanged
strong $B_{\star}$	fast	slow
high W	slow	slow
high β_{up}	slow	slow
low $M_{\star}$ at 1 Myr	fast	slow
high R_{init} ($\beta_{\text{up}} = 0.2$)	unchanged	unchanged
high R_{init} ($\beta_{\text{up}} = 0.005$)	slow	slow
high $\dot{M}_{\star}$	slow	fast
$M_{\text{add}} = 0.01$	slow	unchanged
disk winds (DW)	slow	slow
low $Z_{\star}$	fast	fast

radii and the dipolar magnetic field component dominates the magnetic field strength at this position [e.g., Finley and Matt, 2018, Ireland et al., 2021]. During phases of higher accretion rates, however, the truncation radius is pushed toward the star and can reach $R_t = R_{\star}$. Close to the stellar surface, the importance of high-order magnetic field components increases. Quadro- or octopole components can have field strengths of several kilogauss [e.g., Kochukhov et al., 2020], which could also affect the position of the truncation radius close to the star. On the other hand, trends in the evolution of the stellar magnetic field suggest that young, fully-convective stars have axisymmetric fields dominated by a dipole component [e.g., Folsom et al., 2016]. Thus, we assume that the calculation of the truncation radius as presented in Ireland et al. [2021] provides a good approximation for young, fully-convective stars.

Accretion torque: The accretion torque, Γ_{acc} , depends on the ‘Keplerianity’ of the disk material at the truncation radius. Angular momentum removed by MEs or stellar magnetic torques results in sub-Keplerian rotation of the disk material [e.g., Bessolaz et al., 2008, Matt et al., 2012, Ireland et al., 2021]. In the current formulation, these effects depend on the proportionality constant $K_{\text{acc}} = 0.775$ and the position of the truncation radius (see Eq. 4.7). While the formulation should provide reasonable estimates for the accretion torque for $R_t > R_{\star}$, Γ_{acc} could be underestimated

in case the truncation radius would be located inside the stellar radius, due to strong accretion rates or, for example, weak magnetic field strengths. In such cases, we set $R_t = R_\star$. As a result, the ram pressure of the accreting material exceeds the stellar magnetic pressure, and the accreting material could be close to Keplerian rotation.

Magnetospheric ejections: The torque due to magnetospheric ejections, Γ_{ME} , depends, similarly to the truncation radius, on the stellar magnetic field strength and the location of the truncation radius (see Eq. 4.9). For truncation radii close to the stellar surface, small-scale magnetic field components could affect the magnitude of Γ_{ME} . Assuming primarily axisymmetric fields dominated by a dipole component [e.g., Folsom et al., 2016], the current formulation should provide reasonable estimates for Γ_{ME} .

APSW torque: In the current model, accretion-powered stellar winds (APSW) remove angular momentum from the star [see Eq. 4.11 and, e.g., Matt and Pudritz, 2005b, Gallet et al., 2019, Ireland et al., 2021]. These studies, however, do not consider the very early, embedded stages in their formulations. It is unclear how winds emerging from the stellar surface interact with an envelope falling onto the disk and how much angular momentum can be transported away from the star under these circumstances. In case the envelope around the young star has little or no influence on the APSW, the current formulation with efficiency rates of $W \sim 2\%$ provides a realistic estimate for the torque due to APSW. If, on the other hand, the infalling envelope interferes with the APSW, the efficiency rate could be reduced significantly, which is demonstrated in this study by $W = 0\%$.

In summary, the model presented in Ireland et al. [2021] should provide reasonable estimates for the torques acting on the young protostar during the embedded phase with high accretion rates. The individual relations between stellar and disk parameters are assumed to be valid. We note, however, that the proportionality constants used in Eqs. 4.5 - 4.11 could change when repeating the calculations presented in Ireland et al. [2021] with higher accretion rates.

4.4.2. Magnitude and distribution of β_{up}

The evolution of young stars strongly depends on the amount of additional heat, $\dot{E}_{\text{add}}$, added into the stellar interior [e.g., Baraffe et al., 2012, Vorobyov et al., 2017a,b, Kunitomo et al., 2017].

Usually, $\dot{E}_{\text{add}}$ is assumed to be a certain fraction, β , of the energy of the accreted material. In the hybrid accretion scenario [e.g., Vorobyov et al., 2017b], β depends on the accretion rate on the star, ranging from small values, β_{low} , during phases of low accretion rates to high values, β_{up} , during phases of high accretion rates [e.g., Jensen and Haugbølle, 2018, Elbakyan et al., 2019, Steindl et al., 2021]. According to Kunitomo et al. [2017], relatively large values of $\beta \gtrsim 0.1$ are necessary to explain the evolution of a majority of stars in young clusters⁵. Low values of β (e.g., $\beta = \beta_{\text{low}}$), on the other hand, result in stellar luminosities of $\lesssim 0.1 L_{\odot}$ and can explain very low-luminosity objects [VeLLOs, e.g., Vorobyov et al., 2017a]. To explain the luminosity spread in a young cluster, a distribution of β_{up} has to be assumed. The exact processes that constrain the magnitude of β_{up} are still unknown and more detailed modeling of the accretion process and the distribution of the additional heat is required to formulate reasonable limits. Our results show that the amount of additional heat affects the stellar spin evolution significantly. Low values of β_{up} result in very fast rotating stars (see Panel d in Fig. 4.3). It would be interesting to test if low-luminosity stars tend to rotate faster compared to objects with higher luminosities. Unfortunately, the available observational data do not allow any conclusion to be drawn so far.

4.4.3. External influences

The previous results are valid for a single star without any external or environmental influences. Recently, Roquette et al. [2021] studied the influence of external photo-evaporation caused by massive stars in young regions. The presence of a high-mass star can remove mass from the disk due to high-energy radiation and reduce the disk's lifetime. The timescale, on which the star and the disk can exchange angular momentum is reduced and the stars start to spin up due to contraction at an early age. Thus, such stars show faster rotation periods after several million years. During the initial stellar evolution, however, Roquette et al. [2021] assume disk-locking, meaning that the stellar rotation period is kept constant as long as the accretion disk is present. The effect of external photo-evaporation on the stellar spin evolution during the first million years can be summarized as follows: During the embedded phase the disk material is effectively

⁵We note that for a higher Deuterium abundance lower values of β suffice to explain the evolution of a majority of stars [Kunitomo et al., 2017]. Since a variation in the Deuterium abundance has similar effects on the stellar evolution compared to a variation in β_{up} , we forego the simulation of specific models.

shielded by the envelope and external photo-evaporation is assumed to have little effect on the disk evolution [e.g., Winter et al., 2020]. Without the envelope, on the other hand, the high-energy radiation removes material from the disk, and the disk accretion rate decreases for $\gtrsim 10^5$ years. As shown in the previous models with reduced accretion rates at $\gtrsim 10^5$ years (e.g., model $\dot{M}_{\text{high}}$, see Sec. 4.4.2), the star spins down. While assuming disk-locking, this effect is not seen in Roquette et al. [2021]. To model the extent of the spin-down due to lower accretion rates and the spin-up due to shorter disk lifetimes in a high-energy radiation environment, a combination of our early spin evolution model and a hydrodynamic disk model including the effects of external photo-evaporation will provide valuable insights in the early spin evolution.

Another environmental factor is the tidal interaction between different stars in close encounters, binary, or multiple systems. Tidal interactions in close encounters can remove disk material and reduce the disk accretion rate and lifetime [e.g., Winter et al., 2018]. In contrast to external photo-evaporation, these close encounters can act on the star-disk system during all phases of its evolution. The amount of mass removed, depends on the mass of the 'intruding object' [e.g., Vorobyov et al., 2020c] and the number of encounters [e.g., Winter et al., 2018]. After the embedded phase, a reduced accretion rate results in a stellar spin-down (similar to the effect of external photo-evaporation). During the embedded phase, a reduced accretion rate will reduce $\dot{E}_{\text{add}}$ and the stellar inflation. As a consequence, the star rotates faster (see our model with $M_{\star} = 0.3 M_{\odot}$, Panel e in Fig. 4.3). On the other hand, close encounters can ignite outburst events [e.g., Vorobyov et al., 2021] that result in a stellar spin-down. Which of the two mechanisms predominates is unclear and requires a combined model.

Binaries or multiple stellar companions can have different effects on the disk's lifetime and thus, on the stellar spin evolution. Accretion disks, truncated by a wide binary companion, are assumed to have shorter lifetimes as mass is removed from the accretion disk [e.g., Kraus et al., 2012]. As a consequence, the star spins up during the embedded phase and spins down afterward compared to a single star. On the other hand, circumbinary disks have longer lifetimes compared to single stars [e.g., Alexander, 2012]. In multiple stellar systems, certain configurations can also result in long disk lifetimes [e.g., Ronco et al., 2021].

4.4.4. Rotational distributions in young clusters

Our results show the impact of different parameters on stellar spin evolution. Now, we want to overview the observed rotational distributions of young clusters and estimate how fast (slow) a ~ 1 Myr old star probably spins. In the next step, we infer the parameters necessary to reproduce the observed distribution and discuss the physical consequences of these choices (see Sec. 4.4.5). We have selected the rotational period distribution of six young ($\lesssim 3$ Myr) clusters: NGC 6530 [Henderson and Stassun, 2012], Orion SFR [Serna et al., 2021], Taurus [Rebull et al., 2020], NGC 2264 [Venuti et al., 2017], σ Ori [Cody and Hillenbrand, 2010], and ONC [Rebull et al., 2006]. In Fig. 4.6, we show the 5. (left-facing triangle), 50. (diamond), and 95. (right-facing triangle) percentile for each cluster. For all six clusters, the 95. percentiles indicating the slow-rotating stars are located at $\gtrsim 10$ days with a mean value of 13.0 days (vertical dashed line). The fast-rotating boundary is at $\lesssim 1$ day with the exception of Taurus, which contains faster-rotating stars with the 5. percentile at 0.15 days. The mean value of the 5. percentile is 0.76 days (vertical dotted line). One explanation for the rapidly rotating stars in Taurus could be its age and the evolutionary stage of the star-disk systems. With an age of ~ 3 Myr, the sample of Taurus consists of disk-bearing stars and stars without an accretion disk [Rebull et al., 2020]. In addition, the ages of these young clusters are subjected to large scatters in the order of up to ~ 1 Myr [e.g., Testi et al., 2022]. Some stars might be significantly older than the final age of our models (1 Myr). During this additional time, the accretion disk could be dissolved due to accretion, winds, or photo-evaporative effects and the star starts to spin up due to contraction. This theory is supported by a comparison of stars with and without an accretion disk in Rebull et al. [2020]. Stars, which are still surrounded by an accretion disk, tend to rotate slower than stars without a disk. The fastest, disk-bearing stars have rotation periods slightly below 1 day, which matches the 5. percentiles of the other five clusters in our selection.

4.4.5. Slow and fast rotating configurations

Based on the rotational distribution of young star-forming regions (see Sec. 4.4.4), 90% of the stars in young clusters show rotation periods between 0.76 days and 13.0 days. While the rotation

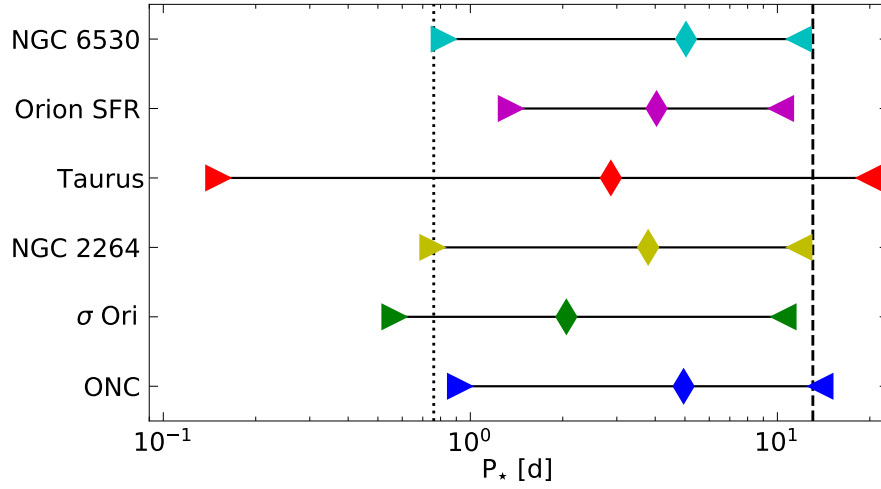


Figure 4.6.: Comparison between the rotational distribution in young clusters and rotation period range at 1 Myr of our fast and slow-rotating models (blue area). The diamonds, left-pointing triangles, and right-pointing triangles are the 50., 5., and 95. percentile of the rotational distribution for the respective cluster. The mean values of the 95. (5.) percentiles are indicated by the vertical dashed (dotted) line. See text for references.

periods of our models match the fast-rotating stars in observed clusters (especially in the case of $\beta = 0.005$ and $W = 0\%$, see Fig. 4.3), observed slow-rotating stars have rotation periods that are higher compared to our model results by a factor of $\gtrsim 4$. We want to explore possible scenarios that can lead to these rotation periods.

A rising number of observed rotational distributions of young clusters allows for deriving empirical relations between the stellar rotation period and other parameters. For accreting T Tauri stars, a positive correlation between the rotation period and the stellar magnetic field strength has been reported [e.g., Vidotto et al., 2014]. For Class II stars, a strong magnetic field result in a stellar spin down. This relation agrees with our results (see Panel b in Fig. 4.3) and other studies [e.g., Matt et al., 2010a, Ireland et al., 2021, Gehrig et al., 2022]. Since the magnetic field strength alone does not spin down the star sufficiently to match the observed slow-rotating stellar periods (see Fig. 4.3), the question arises if strong magnetic fields (in the star or the disk) might affect other parameters. Bu and Mosallanezhad [2018] report that disk winds in a strongly magnetized environment are more powerful compared to a weakly magnetized outflow. Similarly, Bai et al.

[2016] explored the relation between angular momentum loss in protoplanetary disks due to the disk winds and the magnetic field strength. The higher the magnetic field strength, the more effective is the wind in removing angular momentum from the disk and the higher the accretion rate induced by the wind. A stronger accretion rate increases the amount of additional heat, $\dot{E}_{\text{add}}$, which is added to the star. In addition, a strong magnetic field affects the accretion-powered stellar wind. The Alfvén radius, r_A , which acts as the lever arm of the APSW grows with an increasing stellar magnetic field strength [e.g., Gallet et al., 2019, Ireland et al., 2021]. The effect of the magnetic field strength on the mass-loss rate of the APSW, however, is poorly constrained by both, theory and observations. To summarize, a strong magnetic field strength in the star and the disk can result in an effective disk wind, APSW, and can increase the accretion rate onto the star. On the other hand, a weak magnetic field strength would be accompanied by weak winds (APSW and disk winds).

To test the effects of these configurations on the stellar spin evolution, two models `slow` and `fast` are constructed. Model `slow` is characterized by a magnetic field strength of $B_\star = 2.0$ kG, the inclusion of disk winds removing angular momentum from the inner disk (see Sec. 4.3.4), and $\dot{M}_{\text{high}}$ (see Sec. 4.3.3). All other values match the reference model. Model `fast` combines a weak magnetic field strength, $B_\star = 0.1$ kG, with an ineffective wind (no disk winds and $W = 0\%$). In Fig. 4.7, the evolution of the rotation periods for the model `slow` (blue line) and `fast` (yellow line) are shown in comparison to the reference model (black line). The slow-rotating model spins down during the initial inflation to 22 days. After spinning up to ~ 10 days during the embedded phase, the strong magnetic field strength spins down the star, reaching 1 Myr with a period of 12.9 days. The fast-rotating model spins up initially to a rotation period of 1.7 days due to the weak APSW wind. This period is kept roughly constant during the embedded phase. During the Class II phase, the weak magnetic field strength results in a further spin-up to a rotation period of 0.68 days at an age of 1 Myr. Thus, the effect of the magnetic fields in the young star-disk system on other parameters like the effectiveness of winds or the accretion rate can result in rotation periods that agree with the observed range in young clusters (blue shaded are in Fig. 4.7). We note that the influence of magnetic fields on the amount of additional energy controlled by β is unclear. Additionally, the magnetic field strength itself is a free parameter in our model. The

generation of the magnetic field and its evolution is still unclear. One possible evolution scenario for the magnetic field strength is given by Browning et al. [2016]. The magnetic energy is scaled with the convective energy in the stellar interior and a positive correlation between the stellar mass and the magnetic field strength is proposed. Further investigation is needed to constrain the evolution of the free parameters used in our model.

Another empirical relation connects the stellar rotation period and the stellar mass. In Herbst et al. [2007], a higher stellar mass is associated with longer rotation periods in ~ 1 Myr old clusters. We note that in the context of this work, a high-mass star corresponds to a solar-mass star. On the other hand, a weak and even negative correlation between the stellar rotation period and the stellar mass is reported in Henderson and Stassun [2012] for cluster ages of ~ 1 Myr. Our results presented in Fig. 4.3 agree with the negative correlation between stellar period and mass. However, the disks surrounding high-mass stars, which are more massive compared to low-mass stars [e.g., Vorobyov, 2011, Andrews et al., 2013], could launch stronger disk winds as these winds scale with the density in the disk [e.g., Bai et al., 2016]. In addition, the stellar mass scales with the accretion rate [e.g., Testi et al., 2022, Somigliana et al., 2022] and a high-mass star could be injected with a larger amount of additional heat, $\dot{E}_{\text{add}}$. Thus, a combination of high stellar masses, strong accretion rates, and an effective disk wind can also result in a slow-rotating star. To test the relation between the stellar mass and the rotation period in more detail and constrain the mentioned parameters, a combined stellar and disk evolution model, including disk winds, and further observations are required.

4.5. Conclusion

In this work, we present the evolution of stellar rotation periods for young (≤ 1 Myr), low-mass ($\leq 1 M_{\odot}$) stars. We start our simulation from a stellar seed ($0.05 M_{\odot}$) and take protostellar accretion, stellar winds, the magnetic star-disk interaction as well as the internal stellar evolution into account. Our model allows us to test the influence of a variety of parameters on the stellar rotation period. Based on our results, we can draw the following conclusions:

- The stellar rotation period is significantly influenced by the amount of additional heat, $\dot{E}_{\text{add}}$,

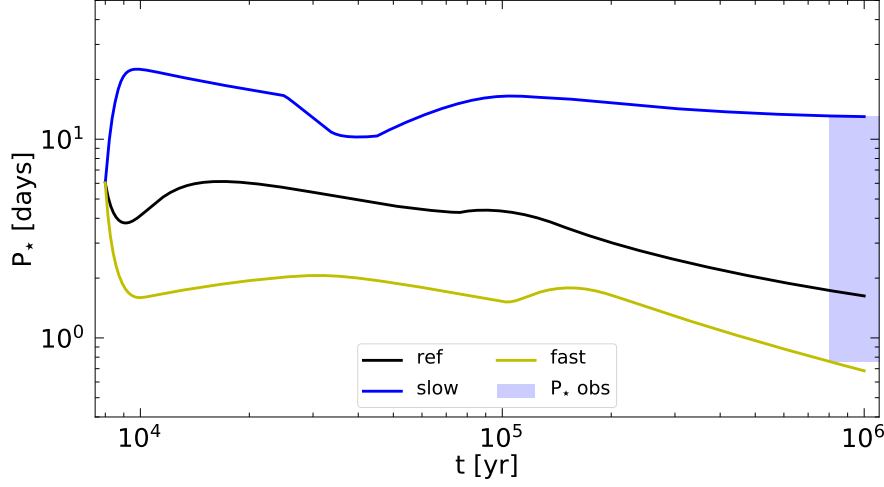


Figure 4.7.: Evolution of the stellar rotation period, P_* for the slow (blue line), reference (black line), and fast rotating models (yellow line). For comparison, the range of the mean values of the 5. and 95. percentile is highlighted as a blue area, $P_{* \text{ obs}}$ (see Sec. 4.4.4).

(controlled by β_{up}) that is injected into the stellar interior, the accretion history, and the presence of strong winds (APSW and disk winds). The magnetic field strength, which is dominant during later evolutionary stages, plays a noticeable but smaller role during the first million years.

- Differences in the initial rotational periods of the stellar seed, Ω_{init} , are quickly forgotten ($\lesssim 10^4$ years) and do not impact the rotational distribution at later stellar ages. Similarly, the initial stellar radius, R_{init} , (assuming $\beta_{\text{up}} \sim 0.2$) and the extent of the region, in which the additional heat is distributed, M_{add} , do not affect the rotation period noticeably.
- We explore the possibility of fast and slow-rotating models. The magnetic field strength in the star-disk system can affect the amount of additional heat, $\dot{E}_{\text{add}}$, and the effectiveness of both, stellar and disk winds. A strong magnetic field can lead to an enhanced accretion rate and effective winds (APSW and disk winds), resulting in slow-rotating stars. Weak magnetic fields or a small amount of additional heat ($\beta = 0.005$) lead to the fastest rotating stars. The stellar rotation periods in our models cover a range from 0.6 – 12.9 days.
- Our models match the rotational distribution of six young ($\lesssim 3$ Myr) clusters. The rotation

periods of up to 90% of all observed stars are located within our results.

- Motivated by these intriguing results, our model, extended over the whole disk lifetime, can be used as an alternative to the widely used disk-locking model during the presence of an accretion disk. In combination with a hydrodynamic disk model, including effects such as external photo-evaporation and episodic accretion outbursts, valuable insights into the origin of the rotational distribution of young clusters can be obtained.

4.6. Appendix

4.6.1. MESA microphysics

The MESA EOS is a blend of the OPAL Rogers and Nayfonov [2002], SCVH Saumon et al. [1995], PTEH Pols et al. [1995], HELM Timmes and Swesty [2000], and PC Potekhin and Chabrier [2010] EOSes.

Radiative opacities are primarily from OPAL [Iglesias and Rogers, 1993, 1996], with low-temperature data from Ferguson et al. [2005] and the high-temperature, Compton-scattering-dominated regime by Buchler and Yueh [1976]. Electron conduction opacities are from Cassisi et al. [2007].

Nuclear reaction rates are a combination of rates from NACRE [Angulo et al., 1999], JINA REACLIB [Cyburt et al., 2010], plus additional tabulated weak reaction rates Fuller et al. [1985], Oda et al. [1994], Langanke and Martínez-Pinedo [2000]. Screening is included via the prescription of Chugunov et al. [2007]. Thermal neutrino loss rates are from Itoh et al. [1996].

4.6.2. Treatment of rotation in MESA

In MESA, rotation is treated as a modification of the stellar equations due to centrifugal acceleration and non-spherical symmetry, when the star is rotating [see Sec. 6 in Paxton et al., 2013]. Angular momentum transport due to rotation (rotational mixing) is implemented as a diffusion term, including five rotational-induced mixing processes: dynamical shear instabil-

ity, Solberg–Høiland instability, secular shear instability, Eddington–Sweet circulation, and the Goldreich–Schubert–Fricke instability [Paxton et al., 2013].

4.6.3. Detailed evolution of individual torque components

We show each component affecting the stellar spin evolution: the external torques due to the accretion process Γ_{acc} , APSW Γ_{W} and the magnetic star-disk interaction Γ_{ME} , the internal contribution Γ_{int} and the resulting total torque Γ_{tot} .

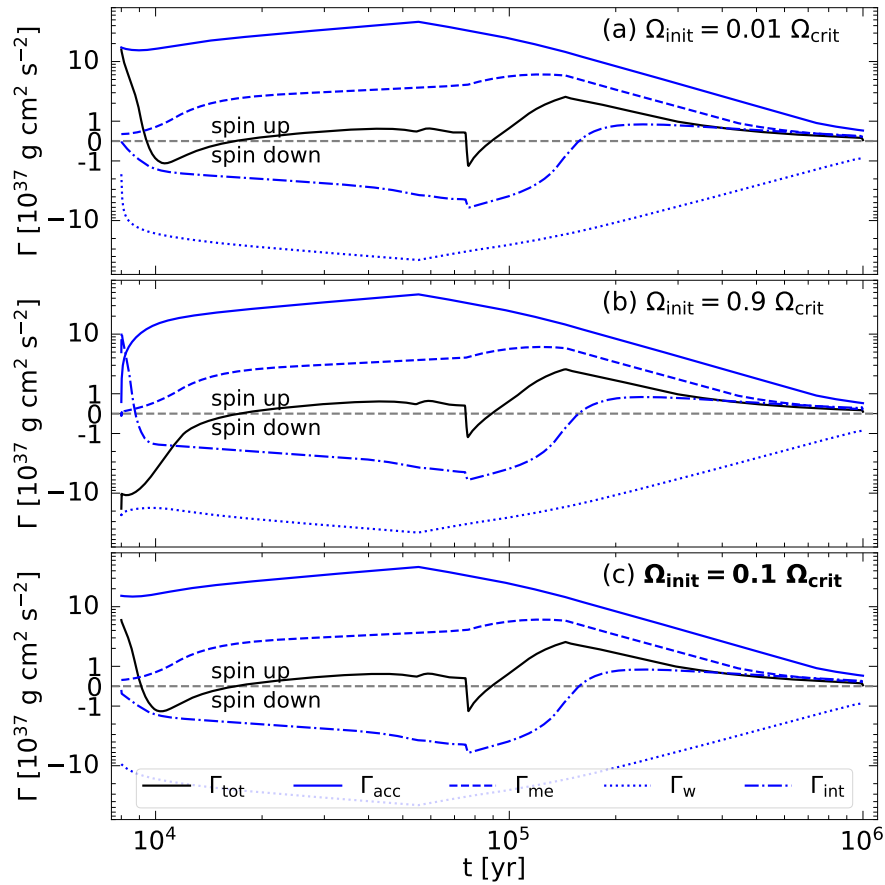


Figure 4.8.: Evolution of the different external torque components (Γ_{acc} , Γ_{me} and Γ_{W} ; solid blue, dashed blue, and dotted blue lines, respectively), the internal contribution to the stellar spin evolution Γ_{int} (blue dash-dotted line) and the combination of these mechanisms Γ_{tot} (black solid line) for different initial rotation rates, Ω_{init} . The value of the reference model is marked in boldface. The horizontal dashed black line divides the regions, in which the star spins up and down.

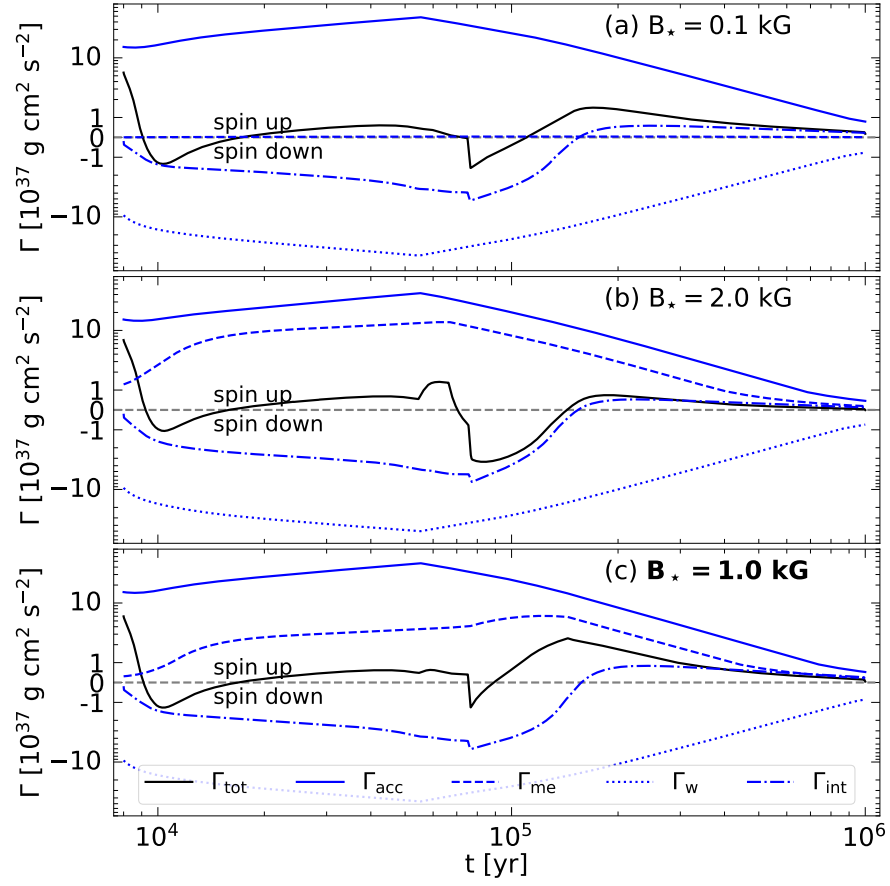


Figure 4.9.: Same as Fig. 4.8 for different magnetic field strengths, B_* .

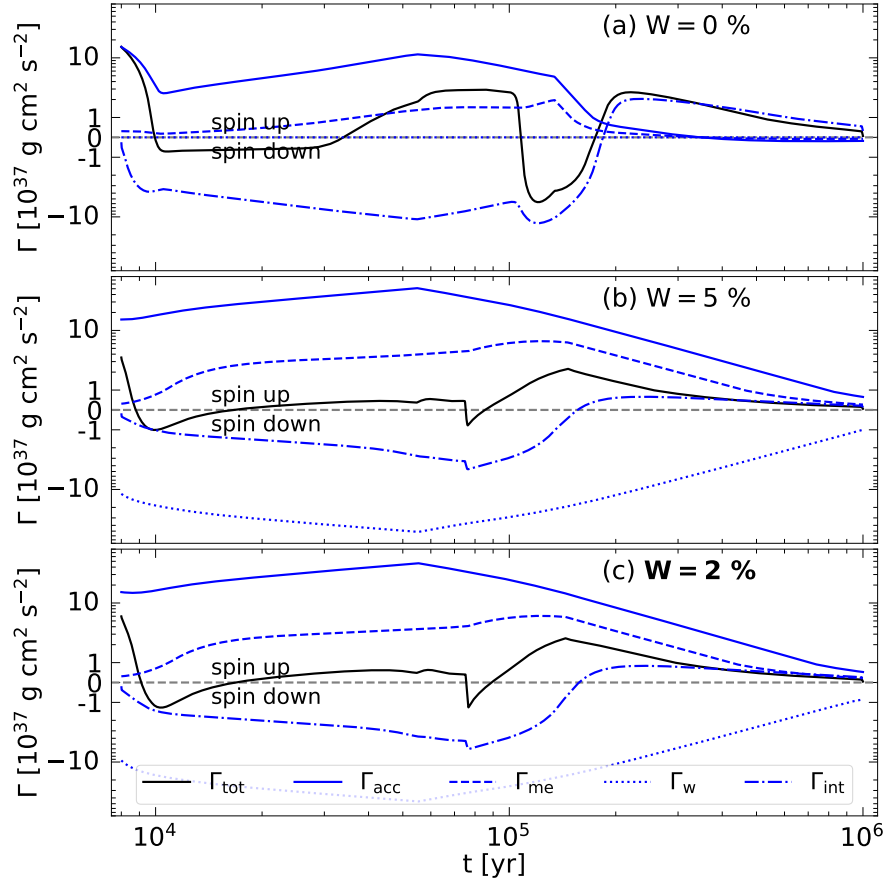


Figure 4.10.: Same as Fig. 4.8 for different APSW efficiencies, W .

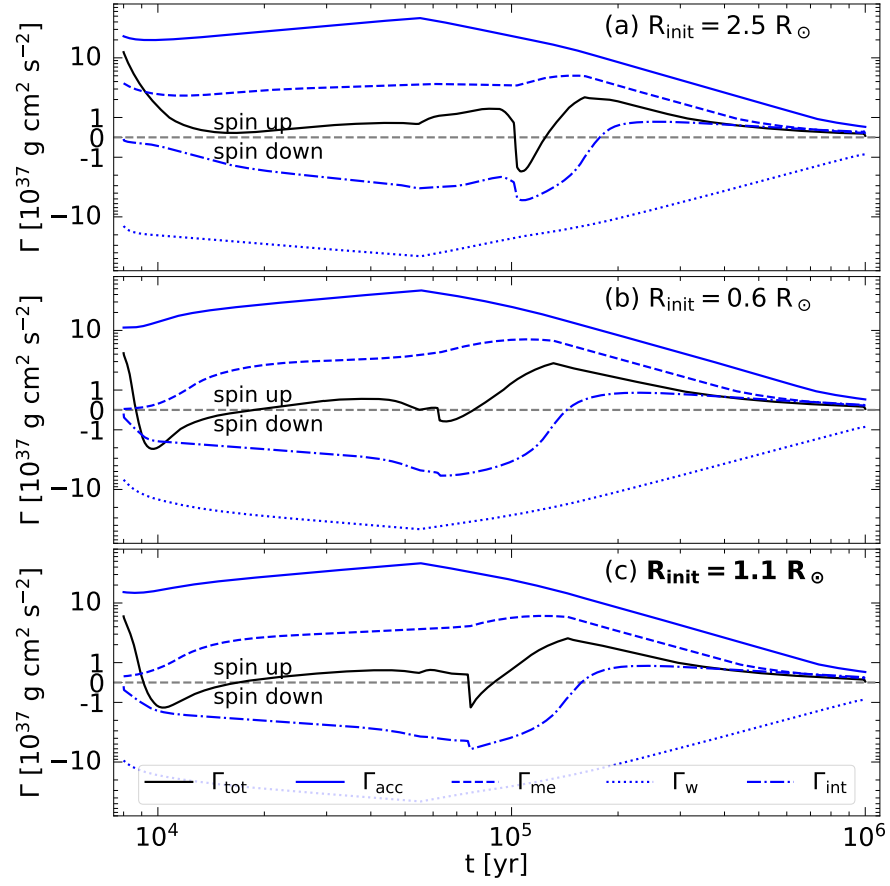


Figure 4.11.: Same as Fig. 4.8 for different initial stellar radii, R_{init} .

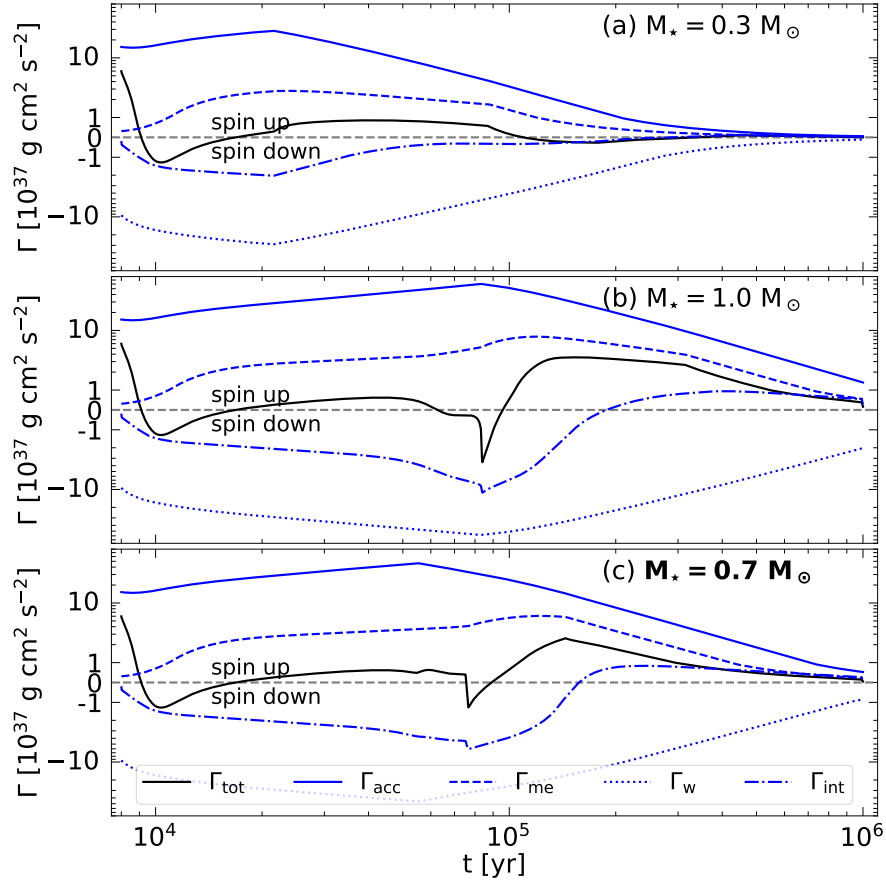


Figure 4.12.: Same as Fig. 4.8 for different stellar masses, M_* , at an age of 1 Myr.

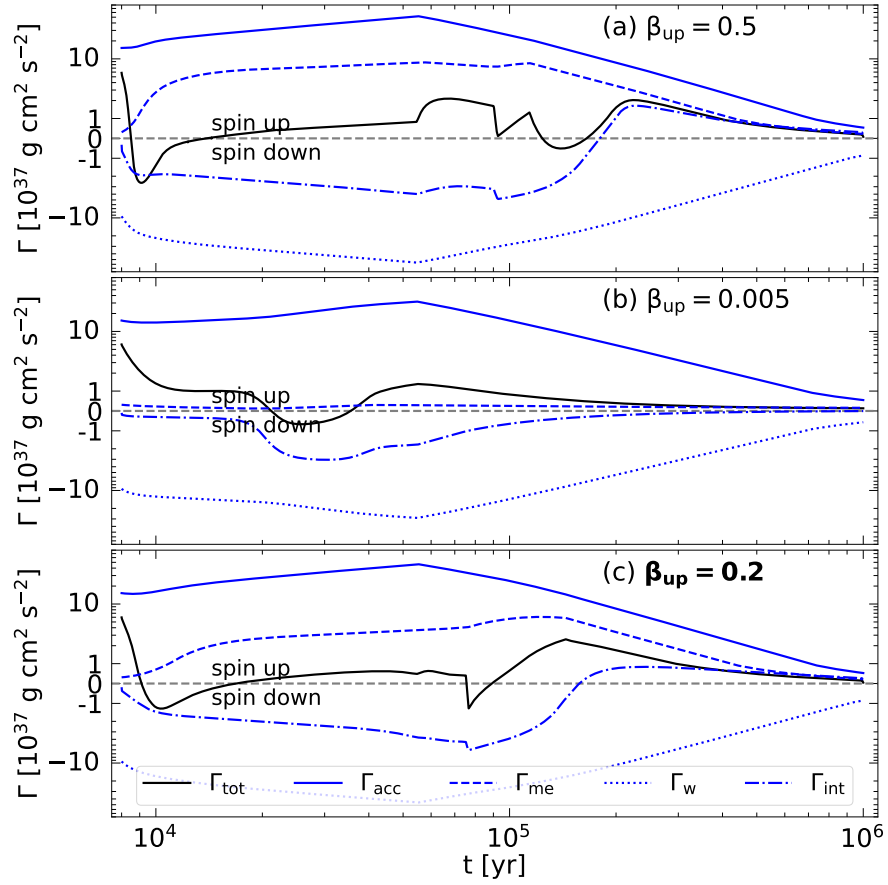


Figure 4.13.: Same as Fig. 4.8 for different maximum energy fractions that can be added to the stellar interior, β_{up} .

4.6.4. Influence of the initial stellar radius for $\beta_{\text{up}} = 0.005$

Evolution of the stellar rotation period and radius for $\beta_{\text{up}} = 0.005$ and different initial stellar radii.

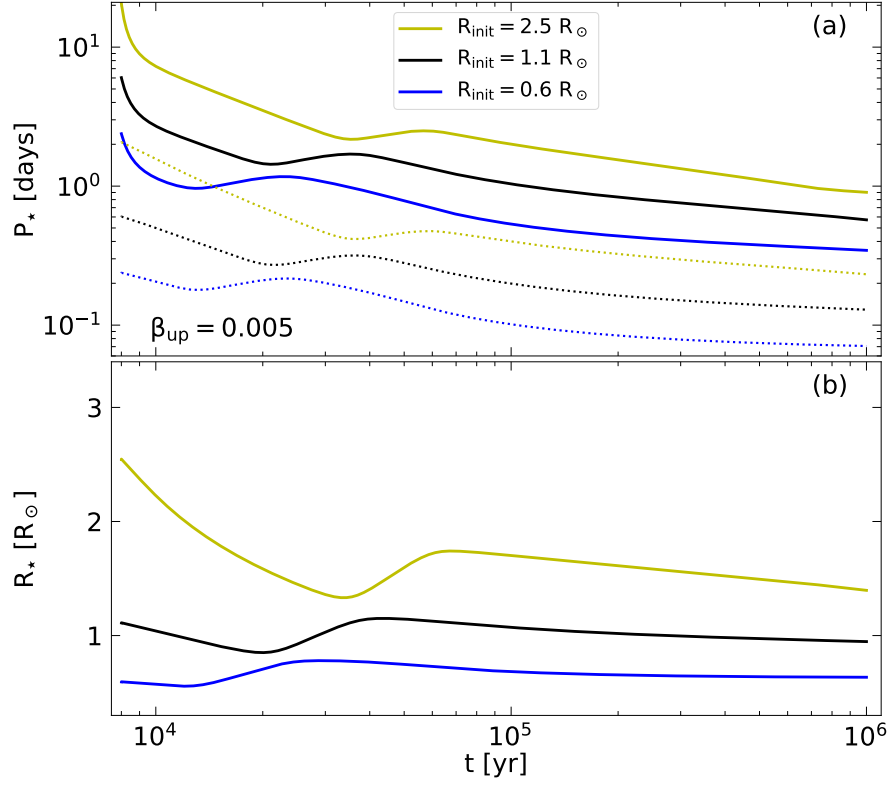


Figure 4.14.: Evolution of the stellar rotation periods, P_* (Panel a), and the stellar radius (Panel b) for different initial stellar radii, R_{init} . For these simulations we choose $\beta_{\text{up}} = 0.005$.

4.6.5. HRD for individual simulations

Here, we show the influence of each parameter on the HRD.

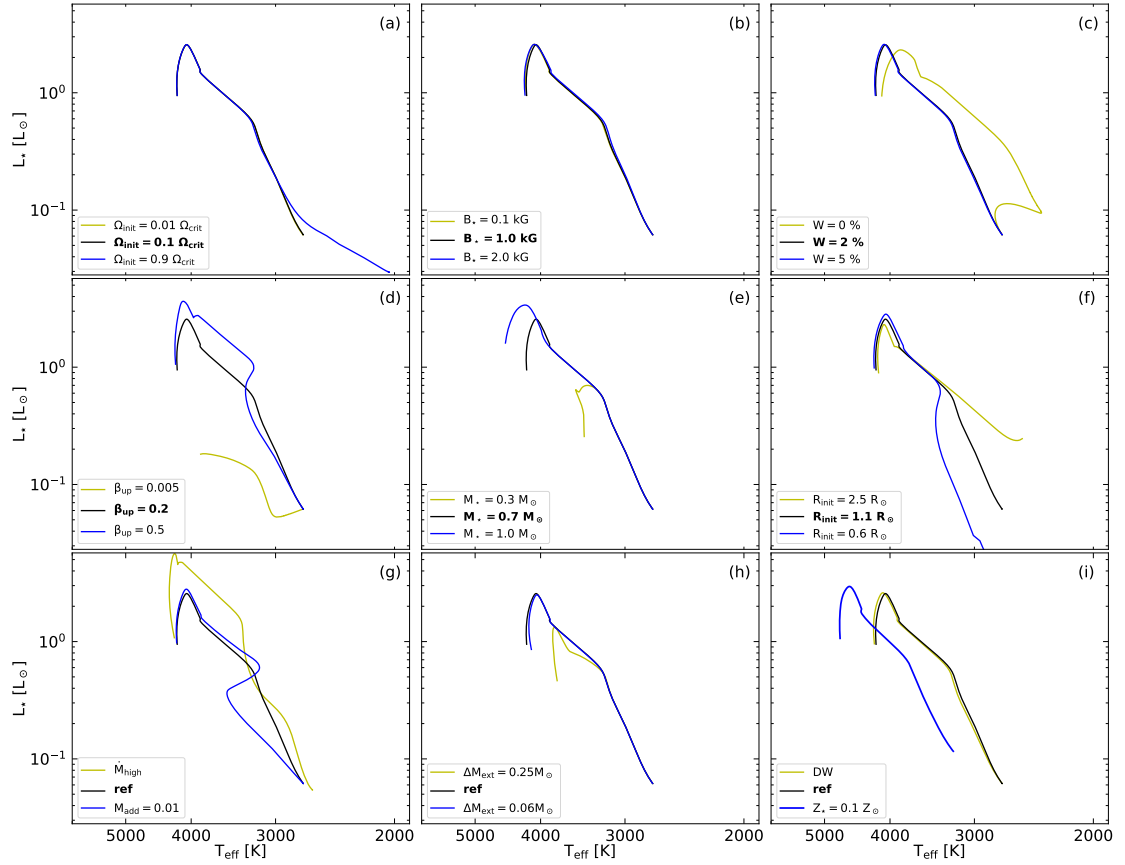


Figure 4.15.: Intrinsic stellar luminosity, $L_{\star}$, over effective temperature, T_{eff} , for each model presented in Sec. 4.3 and Sec. 4.4. The reference model is marked in boldface.

5. Paper IV: The post-disk (or primordial) spin distribution of M dwarf stars

- **Title:** The post-disk (or primordial) spin distribution of M dwarf stars
- **Authors:** L. Gehrig, E. Gaidos, and M. Güdel
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Contribution: In this work, the combined stellar and disk evolution model has been used to run over 500 long-term simulations. I have set up the parameter space, conducted the simulations, and performed the data analysis. My co-authors provided valuable feedback and comments during these processes and during the process of writing the paper.

Summary: The rotation periods of young low-mass stars after disks have dissipated ($\lesssim 10$ Myr) but before magnetized winds have removed significant angular momentum is an important launch point for gyrochronology and models of stellar rotational evolution; the rotation of these stars also regulates the magnetic activity and the intensity of high-energy emission that affects any close-in planets. A recent analysis of young M dwarf stars suggests a distribution of specific angular momentum (SAM) that is mass-independent, but the physical basis of this observation is unclear. We investigate the influence of an accretion disk on the angular momentum (AM) evolution of young M dwarfs, whose parameters govern the AM distribution after the disk phase, and whether this leads to a mass-independent distribution of SAM. We used a combination of protostellar spin and implicit hydrodynamic disk evolution models to model the innermost disk (~ 0.01 AU), including a self-consistent calculation of the accretion rate onto the star, non-Keplerian disk rotation, and the influence of stellar magnetic torques over the entire disk lifetime. We executed and analyzed over 500 long-term simulations of the combined stellar and disk evolution. We find that above an initial rate of $\dot{M}_{\text{crit}} \sim 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$, accretion "erases" the initial SAM of M dwarfs during the disk lifetime, and stellar rotation converges to values of SAM that are largely independent of initial conditions. For stellar masses $> 0.3 \text{ M}_{\odot}$, we find that observed initial accretion rates $\dot{M}_{\text{init}}$ are comparable to or exceed $\dot{M}_{\text{crit}}$. Furthermore, stellar SAM after the disk phase scales with the stellar magnetic field strength as a power law with an exponent of -1.1 . For lower stellar masses, $\dot{M}_{\text{init}}$ is predicted to be smaller than $\dot{M}_{\text{crit}}$ and the initial conditions are imprinted in the stellar SAM after the disk phase. To explain the observed mass-independent distribution of SAM, the stellar magnetic field strength has to range between 20 G and 500 G (700 G and 1500 G) for a $0.1 \text{ M}_{\odot}$ ($0.6 \text{ M}_{\odot}$) star. These values match observed large-scale magnetic field measurements of young M dwarfs and the positive relation between stellar mass and magnetic field strength agrees with a theoretically motivated scaling relation. The scaling law between stellar SAM, mass, and the magnetic field strength is consistent for young stars, where these parameters are constrained by observations. Due to the very limited number of available data, we advocate for efforts to obtain more such measurements. Our results provide new constraints on the relation between stellar mass and magnetic field strength and they can be used as initial conditions for future stellar spin models, starting after the disk phase.

5.1. Introduction

One of the major challenges in stellar astrophysics is the origin and evolution of stellar angular momentum (AM) and spin. Over the last decade, observations of young clusters have yielded new insight into the evolution of premain sequence (PMS) and young ($\lesssim 1$ Gyr) main sequence dwarf stars [e.g., Herbst et al., 2002, Irwin et al., 2008, Irwin and Bouvier, 2009, Hartman et al., 2009, Meibom et al., 2011, Affer et al., 2013, Gallet and Bouvier, 2013]. Such coeval groups of stars with established ages and a range of masses provide benchmarks for understanding how that rotational evolution varies with stellar mass and age. The spin evolution of low-mass stars (starting from the Class II phase) can be divided into three phases [Gallet and Bouvier, 2013]: (1) exchange of angular momentum between a young stellar object (YSO) and its primordial disk; (2) simple contraction and spin-up of the PMS star after the disk has dissipated; and (3) gradual spin-down on the main sequence by loss of AM through a magnetized wind.

Understanding the initial rotation distribution and subsequent evolution of young stars is important to relating stellar rotation to ages ["gyrochronology"; Barnes, 2007]. Furthermore, rotation drives a magneto dynamo operation in stars with convective envelopes, and magnetic activity is responsible for heating the upper stellar atmosphere, which is the primary source of X-rays and far- and extreme-ultraviolet (XUV) emission from low-mass stars [e.g., Pallavicini et al., 1981, Micela et al., 1985, Pizzolato et al., 2003, Wright et al., 2011, France et al., 2018]. XUV irradiation by the host star heats and inflates the upper atmospheres of close-in planets and can lead to evaporation and even complete loss of the atmosphere [e.g., Güdel et al., 1997, Lammer et al., 2003, Cecchi-Pestellini et al., 2009, Lecavelier Des Etangs et al., 2010, Sanz-Forcada et al., 2011, France et al., 2013, Nortmann et al., 2018]. XUV-driven loss of water as atomic H can affect the habitability of rocky planets [e.g., Buccino et al., 2006].

These issues come to the fore for studies of M dwarf stars ($M_{\star} \approx 0.1 - 0.6 M_{\odot}$) as hosts of planets, particularly rocky and potentially habitable planets. M dwarfs are the most common type of star [e.g., Chabrier, 2003], and the small size and low mass of M dwarfs favor the detection of planets by the transit and radial velocity methods on close-in orbits. In addition, the lower luminosities of M dwarfs mean that their "habitable zones," where the equilibrium temperatures

of planets would permit surface liquid water, are much closer to the stars [e.g., Zechmeister et al., 2009]. However, these stars are intrinsically faint and more difficult to study at the distances of benchmark clusters and as a result, studies and modeling of their rotational evolution and establishment of gyrochronological benchmarks are at an early stage.

Robust measurements of rotation periods for M dwarfs in a few star-forming regions and young clusters have been provided by the precise photometry of the Kepler space telescope, repurposed to observe multiple ecliptic fields in the two-wheel K2 mission [Howell et al., 2014]. These groups include the Taurus [1-5 Myr; Rebull et al., 2020], ρ Ophiuchus (3-5 Myr), and Upper Scorpius [$\sim$ 10 Myr, Rebull et al., 2018] star-forming regions, and the Pleiades [$\approx$ 125 Myr Rebull et al., 2016], Hyades [$\approx$ 650 Myr Douglas et al., 2019], and Praesepe [$\approx$ 800 Myr Rebull et al., 2017, Douglas et al., 2017] clusters. Somers et al. [2017] showed that the distributions of specific (per unit mass) angular momentum (SAM) of M dwarfs in Upper Scorpius and the Pleiades were similar, approximately independent of stellar mass, but with a significant scatter. Since by the age of Upper Scorpius circumstellar disks have largely (but not completely) dissipated [Mamajek, 2009], the similarity of the two distributions means that angular momentum loss through winds during the first $\leq$ 125 Myr is $\lesssim$ 40%.¹ It also follows that the origin of the mass-independent SAM distribution lies in the earlier YSO phase and could be a product of the interaction between stars and their disks.

Our study models the YSO phase [$\lesssim$ 10 Myr, e.g., Mamajek, 2009] of single M dwarfs to determine if star-disk interactions can explain the observed distribution of SAM and rotation at post-disk times. While the rotational evolution of young, disk-free M dwarfs on the PMS has been previously studied [e.g., Bouvier et al., 2014a, Matt et al., 2015], the influence of disks on their rotation is less well explored and most of the work has consisted of long-term spin evolution calculations [e.g., Armitage and Clarke, 1996, Matt et al., 2010a, Matt et al., 2012, Gallet et al., 2019]. During this phase, mass is accreted from the disk onto the star, altering the star’s angular momentum and subsequent rotational evolution. It has been proposed that this exchange leads to the star rotating at the orbital period of the disk inner edge [“disk-locking”, e.g., Ghosh and Lamb, 1979b, Koenigl, 1991]. Calculations by Armitage and Clarke [1996] suggest that massive

¹We note that contracting PMS stars spin up so the distributions of rotation periods differ.

disks coupled to the stellar magnetic field can regulate a star's spin evolution while low mass disks fail to slow down the star as it contracts and spins up. While Armitage and Clarke [1996] include simplified time-dependent disk evolution in their model, they assume the entire disk couples to the stellar magnetic field lines. However, due to the limited strength of the coupling between the disk and the stellar magnetic field lines the disk may not be able to remove the AM required to balance spin-up due to stellar accretion and stellar contraction and keep the star's rotation "locked" to the disk [e.g., Matt and Pudritz, 2005b, Zanni and Ferreira, 2013]. Moreover, disk-locking does not explain the existence of very slowly rotating stars [e.g., Matt and Pudritz, 2005a].

Other phenomena may be in play during the disk phase: Accretion-powered outflows or winds could also remove angular momentum from the star [APSW, e.g., Matt and Pudritz, 2005a, 2008a,b, Matt et al., 2012, Finley and Matt, 2018, Gallet et al., 2019]. Magnetized disk winds can remove mass and angular momentum from the disk before these are accreted to the star [e.g., Shu et al., 1994b, Ferreira et al., 2000, Romanova et al., 2009, Zanni and Ferreira, 2013]. Magnetospheric ejections, in which the stellar and/or disk magnetic field is loaded with mass from the disk [Zanni and Ferreira, 2013], can also transport angular momentum between the star and disk.

Self-consistent modeling of the accretion rate in the disk $\dot{M}_{\text{disk}}$ and onto the star is important to robustly describe mass and angular momentum exchange. Previous studies fixed the accretion rate to a power law or an exponential decay [e.g., Matt et al., 2010a, Gallet et al., 2019] and did not account for feedback between the star and disk. Modeling the disk's inner edge ($\lesssim 0.1$ au) is numerically challenging because of the short time steps that are required, but this can be achieved with implicit hydrodynamic codes which circumvent the time-step limitations of explicit codes, making long-term simulations feasible [Courant et al., 1928, Ragossnig et al., 2020, Steiner et al., 2021].

Here we combine a model of PMS stellar evolution [Matt et al., 2010a] with an implicit hydrodynamic model of disk evolution [TAPIR; Ragossnig et al., 2020, Steiner et al., 2021] that can more realistically and self-consistently calculate disk accretion, the effects of stellar magnetic torques and local pressure gradients, the position of the inner disk boundary, and the effect of the

star on the disk. We seek to understand (1) how the parameters describing the star-disk interaction govern the final distribution of SAM when disks have dissipated; and (2) whether the observed, mass-independent distribution of SAM is a natural outcome of our model.

5.2. Model description

We combine a model of stellar spin evolution [e.g., Matt et al., 2010a, Matt et al., 2012, Gallet et al., 2019] with a hydrodynamic description of a protostellar accretion disk. The model used in this study is already described in full detail in Ragossnig et al. [2020] and Steiner et al. [2021] and we refer the reader to these papers for further information. The most important aspects, as well as differences from the version used in Steiner et al. [2021], are highlighted below. Figure 5.1 schematically illustrates the major features and important parameters of the model.

5.2.1. Hydrodynamic disk model

Our simulations model the protoplanetary disk as a time-dependent, viscous accretion disk [e.g., Shakura and Sunyaev, 1973, Armitage et al., 2001]. Accretion rates are computed self-consistently and are not fixed. We use the one-dimensional implicit TAPIR code [Ragossnig et al., 2020, Steiner et al., 2021], which solves the equations for the surface density Σ , internal energy e and the radial velocity u_r as well as the azimuthal velocity u_ϕ [Ragossnig et al., 2020, Steiner et al., 2021]. Additionally, the adaptive grid equation introduced by Dorfi and Drury [1987] is solved to ensure sufficient grid resolution and enable moving inner and outer disk boundaries. The accretion disk is assumed to be axisymmetric and in hydrostatic equilibrium in the vertical direction, and thus quasi 2-d. The model allows the velocity to deviate from the Keplerian azimuthal velocity $u_\phi \sim r^{-1/2}$, and effects such as the influence of a stellar magnetic field can be included, being of particular importance in the inner part of the disk [e.g., Romanova et al., 2002, Bessolaz et al., 2008]. Models based on the diffusion approximation [e.g., Pringle, 1981, Armitage et al., 2001, Zhu et al., 2007, 2010] are not able to include these effects over the whole disk lifetime of $\lesssim 10$ Myr. The viscosity model used in this study is based on Shakura and Sunyaev [1973] with the kinematic viscosity $\nu = \alpha c_S H_P$. Here, α , c_S and H_P denote the dimensionless viscosity

parameter [assumed to range between 0.005 and 0.01, e.g., Zhu et al., 2007, Vorobyov and Basu, 2009, Yang et al., 2018], the speed of sound, and the vertical pressure scale height, respectively.

In the inner disk, the stellar magnetic field acts on the partially ionized disk, and the motion of the disk alters the ambient magnetic field. The magnetic field of the star is assumed to be corotating with the star and described as a dipole aligned with the rotation axis and perpendicular to the disk plane. However, the Keplerian motion of the disk tends to wind up locally vertical field lines and generates an azimuthal component of the magnetic field B_ϕ [e.g., Rappaport et al., 2004, Kluźniak and Rappaport, 2007, Steiner et al., 2021]. The motion of ionized disk gas with respect to the local magnetic field induces a torque in the azimuthal direction and thus leads to an acceleration in the radial direction [e.g., Romanova et al., 2002]. The resulting drop in disk surface density toward the star [e.g., Bessolaz et al., 2008, Steiner et al., 2021] creates a pressure gradient. The torque, induced by the stellar magnetic field, changes sign at the corotation radius

$$r_{\text{cor}} = \left(\frac{G M_\star P_\star^2}{4 \pi^2} \right)^{1/3}, \quad (5.1)$$

with G being the gravitational constant, $M_\star$ the stellar mass and $P_\star$ the stellar rotational period. Interior to r_{cor} , the disk gas is decelerated in the azimuthal and accelerated in the radial direction toward the star. Exterior to r_{cor} , the disk material is accelerated in the azimuthal direction, pushing it away from the star. Toward the innermost part of the disk, the magnetic pressure ($\propto B_\star^2 (R_\star/r)^6$) overwhelms the gas pressure as well as the ram pressure of the disk material moving toward the star. At a so-called truncation radius r_{trunc} , the disk material is forced along magnetic field lines and is “funneled” toward the star [e.g., Romanova et al., 2002, Romanova and Kurosawa, 2014]. We choose r_{trunc} to be the inner radius of the accretion disk r_{in} ; this radius is calculated by equating the stellar magnetic pressure to the disk’s ram pressure or gas pressure at the inner boundary, whichever is higher [see Steiner et al., 2021]. Similar to Romanova et al. [2002] and Steiner et al. [2021], the inner boundary is treated as a “free” Neumann boundary (i.e., $\partial_r \Sigma = 0$). At the outer boundary, the azimuthal velocity u_ϕ equals the Keplerian velocity, the radial velocity u_r is set to zero, and the surface density Σ as well as the internal energy e are again treated with Neumann boundary conditions. The outer disk radius is set to the position where the disk surface

density has declined to $\Sigma_{\text{out}} = 1 \text{ g/cm}^2$ [similar to Vorobyov et al., 2020b].

5.2.2. Stellar spin model

The stellar spin evolution model used in this study is based on the work of Matt et al. [2010a] and Gallet et al. [2019]. The simulations are started from a stellar age of $t_0 = 1 \text{ Myr}$ [similar to Gallet et al., 2019], an age corresponding approximately to the beginning of the Class II YSO phase when the only significant accretion is through a disk. The star is considered to be fully convective and to rotate as a solid body. The stellar angular momentum can be written as $J_\star = I_\star \Omega_\star$, with a moment of inertia $I_\star = k^2 M_\star R_\star^2$ including the stellar radius $R_\star$ as well as the dimensionless radius of gyration k and the stellar angular velocity $\Omega_\star$. We assume $k^2 = 0.2$ to be constant in time, which is a reasonable assumption over the first 10 Myr [e.g., Armitage and Clarke, 1996, Matt et al., 2010a, Matt et al., 2012]. $J_\star$ is altered by external torques Γ_{ext} of which, following the formulation in Gallet et al. [2019], we model three torque components:

First, the accretion torque, Γ_{acc} : The accretion of disk material adds mass and angular momentum to the star. Γ_{acc} is the product of the accretion rate onto the star and the specific angular momentum of the disk at the inner disk boundary [e.g., Gallet et al., 2019]. This torque contribution is positive, which means that the star spins up due to the accretion of disk material.

Second, accretion-powered stellar winds (APSW), Γ_{W} : In the picture of a magnetized stellar outflow, the ejected material is magneto-centrifugally accelerated along stellar magnetic field lines. As the wind propagates outward, it removes angular momentum from the star and results in stellar spin-down. In this study, the driving force of this wind is considered to be accretion [accretion-powered stellar wind, APSW, e.g., Matt and Pudritz, 2005b, Gallet et al., 2019] and the mass-loss rate scales as $\dot{M}_{\text{W}} = W \dot{M}_\star$, with the APSW efficiency parameter, W . The Alfvén radius r_{A} acts as a lever arm for this stellar outflow and controls how much angular momentum is carried away [e.g., Gallet et al., 2019, Pantolmos et al., 2020, Ireland et al., 2021]. We note that the order of W as well as the driving forces of the APSW are still under debate [see Sec. 5.4.3 or, e.g., Gallet et al., 2019].

Third, Magnetospheric ejections (ME), Γ_{ME} : Angular momentum can be transferred between the star and the disk by magnetospheric ejections if the stellar magnetic field is strong enough to

couple to the disk material [e.g., Zanni and Ferreira, 2013, Gallet and Bouvier, 2013]. Starting from vertical stellar magnetic field lines that corotate with the star, differential rotation between the star and the disk generates a toroidal field component. The toroidal field results in a pressure component in the vertical direction that induces a periodic inflation and reconnection process of the magnetic field lines [Zanni and Ferreira, 2013]. The angular momentum transferred by this process, Γ_{ME} , depends on the twist of the field lines, the stellar magnetic field strength, and the radial disk region that couples to the magnetic field lines [e.g., Zanni and Ferreira, 2013, Gallet et al., 2019]. The sign of Γ_{ME} results from the relation between the inner disk boundary, r_{trunc} , and the corotation radius, r_{cor} . If the ratio between the truncation radius and the corotation radius is smaller than a given value K_{ME} , $r_{\text{trunc}}/r_{\text{cor}} < K_{\text{ME}}$, the MEs spin up the star. For larger truncation radii, $r_{\text{trunc}}/r_{\text{cor}} \geq K_{\text{ME}}$, the MEs spin down the star. The dimensionless parameter, K_{ME} , results from a best-fit value based on multidimensional models of the star-disk interaction [e.g., Gallet et al., 2019, Pantolmos et al., 2020]. The value of K_{ME} varies between different studies and ranges between 0.60 [Pantolmos et al., 2020] and 0.79 [Gallet et al., 2019]. In this study, we choose the latter.

Internal redistribution of angular momentum resulting from the development of a radiative core [as described in Gallet et al., 2019] is neglected because very low-mass stars will remain fully convective during the disk lifetime and rotate as solid bodies. The evolution of the stellar rotation is governed by changes in AM and changes in the moment of inertia due to mass accretion and PMS contraction [e.g., Matt et al., 2010a, Gallet et al., 2019]. The rate of stellar mass change is calculated by the hydrodynamic disk model as the disk accretion rate at the inner boundary and we describe the stellar radius evolution by Kelvin-Helmholtz contraction with an index 3/2 polytrope [e.g., Collier Cameron and Campbell, 1993, Matt et al., 2010a, Matt et al., 2012], resulting in

$$\dot{R}_{\star} = 2 \frac{R_{\star}}{M_{\star}} \dot{M}_{\star} - \frac{28\pi\sigma R_{\star}^4 T_{\text{eff}}^4}{3GM_{\star}^2}, \quad (5.2)$$

with the stellar effective temperature T_{eff} and the Stefan-Boltzmann constant σ . The deviation of Eq. (5.2) from radii predicted by the PMS model of Baraffe et al. [2015] is only $\sim 1\%$ over the first 10 Myr.

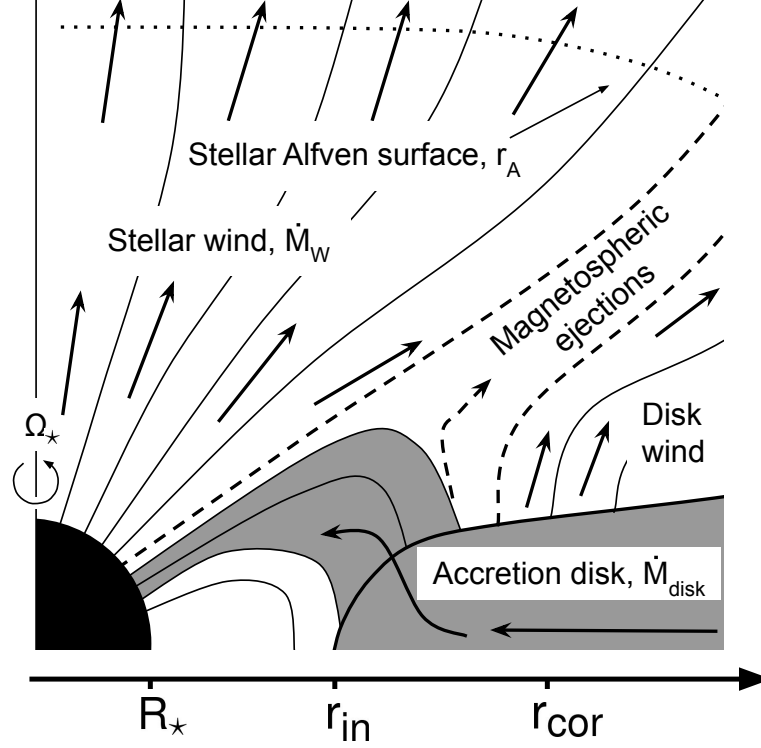


Figure 5.1.: Schematic cartoon of inner disk-star connection. The star (black quarter circle on the left) has a radius of R_* and rotates with Ω_* . Disk material (gray area) flows from the outer disk toward the inner disk radius r_{in} , where it is forced into funnel flows and is accreted onto the star [e.g., Romanova et al., 2004, Bessolaz et al., 2008, Steiner et al., 2021] with an accretion rate of $\dot{M}_{\text{disk}}$. The accreted disk material increases the stellar angular momentum by Γ_{acc} . Based on the model, a specific fraction of the accreted mass is driven into a stellar wind, which removes mass and angular momentum from the star ($\dot{M}_W$ and Γ_W , respectively). The Alfven radius r_A acts as a lever arm for this stellar outflow. We note that this cartoon is not to scale and we can usually assume $r_A \gg R_*$. Angular momentum can be transferred toward or away from the star (Γ_{ME}), depending on the position of r_{in} with respect to the corotation radius r_{cor} [e.g., Zanni and Ferreira, 2013]. Finally, magneto-centrifugally driven disk winds remove angular momentum and mass from the disk before it can be accreted on the star [e.g., Königl et al., 2011, Bai and Stone, 2013]. We note that disk winds are not included in our model.

5.2.3. Parameter range

We define a parameter range for our simulation that agrees reasonably with observations and theoretical estimates of stellar and disk parameters.

Stellar rotation period of T Tauri stars

Over the last two decades, the rotation periods of a large number of young ($\lesssim 10$ Myr) stars have been measured [e.g., Herbst et al., 2002, Rebull et al., 2006, Cody and Hillenbrand, 2010, Rebull et al., 2010, 2015, Venuti et al., 2017, Rebull et al., 2020, Serna et al., 2021, as well as other publications by the YSOVAR consortium]. In Fig. 5.2, the distributions of rotation periods of five of these studies are shown. The majority of periods are located between ~ 1 and 10 days. The distribution decreases strongly toward ~ 20 days and there are very few stars with rotation periods > 20 days. This distribution of values, however, has to be treated with caution. Studies based on TESS observations [e.g., Rebull et al., 2015, Serna et al., 2021] are biased against periods longer than TESS' lunar-synchronous orbit of 13.7 days [e.g., Howard et al., 2021, Claytor et al., 2022] and longer undetected periods are possible². Furthermore, ground-based studies are subjected to the known effect of decreased sensitivity to longer periodic signals that are near the threshold of detection (only a few periods can be coadded for detection). We adopt a conservative range of 1-10 days, recognizing that there are some stars with initial rotation periods outside this range.

Stellar magnetic dipole field strength

The dipole magnetic field strength is an important parameter for the spin model used in this study. Obtaining reliable estimates of $B_\star$, however, is challenging due to the limited number of observations and complex physical processes that generate the magnetic field [e.g., Kochukhov, 2021]. There are different techniques for measuring magnetic fields [e.g., the review of Kochukhov, 2021, and references therein]. In the most common method, the broadening of magnetically sensitive lines due to the Zeeman effect is measured and the strength of the magnetic field is described in terms of the Stokes I parameter. We refer to this B-field measurement as B_I . B_I is a

²We refer to the sharp drop in rotation periods at ~ 15 days in Fig. 19 of Rebull et al. [2015].

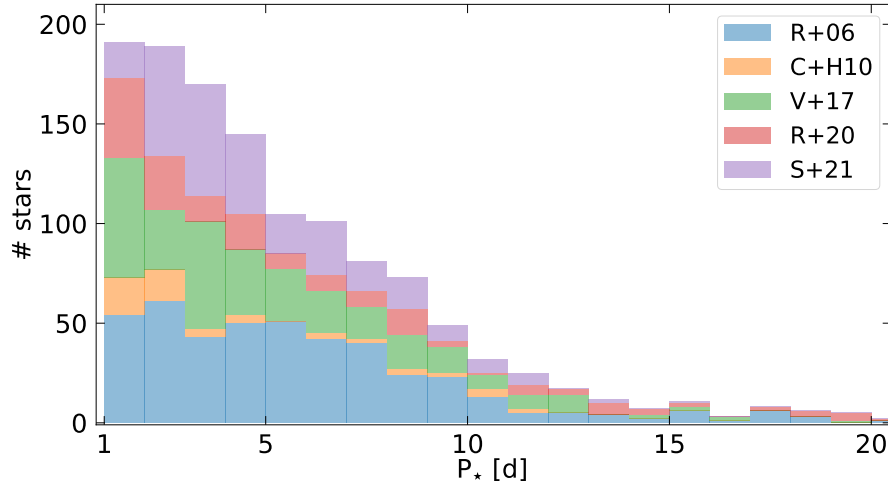


Figure 5.2.: Distribution of stellar rotation periods in young clusters: ONC [R+06, Rebull et al., 2006], σ Ori [C+H10, Cody and Hillenbrand, 2010], NGC 2264 [V+17, Venuti et al., 2017], Taurus [R+20, Rebull et al., 2020] and the Orion star forming region [S+21, Serna et al., 2021].

measure of the total intensity of the surface magnetic field and it is not possible to differentiate between small and large-scale components. A more detailed technique, Zeeman-Doppler imaging (ZDI), resolves the distribution of line broadening on the stellar surface and obtains the Stokes V parameter [e.g., Johnstone et al., 2014]. A large-scale magnetic field dipole component (which we refer to as B_V) can be calculated from this. We note that there is the possibility that B_V can underestimate the strength of the magnetic field as it can be partially canceled out by higher-order components [see the case of AU Mic, Kochukhov and Reiners, 2020]. If, in addition to the Stokes I and V parameters, the Stokes Q and U parameters are available, the overall structure of the stellar magnetic field and a more precise value for the large-scale field can be calculated [e.g., Kochukhov and Reiners, 2020]. We refer to this large-scale field measurement as B_{QU} . Unfortunately, this requires observationally expensive high-resolution spectropolarimetry, and these have been performed on very few stars [Kochukhov and Reiners, 2020]. The large-scale dipole field component (inferred indirectly based on Stokes V or Stokes Q and U parameters) usually dominates the magnetic star-disk interaction and governs the loss of AM due to the APSW [e.g., Finley and Matt, 2018]; it is this component that is parameterized by the surface field parameter $B_\star$ in this study.

Aggravatingly, the stellar magnetic field evolves with time [e.g., Folsom et al., 2016]. Young, fully convective stars have a strong, axisymmetric, and predominantly dipolar field. As the stars evolve and develop a radiative core (stars with masses $\lesssim 0.3 M_{\odot}$ stay fully convective during their entire lifetime), the field strength decreases, and the field complexity increases, reducing the dipole component [e.g., Folsom et al., 2016]. Based on the data for 8 T Tauri stars, the large-scale field comprises up to $b_{\text{lsf}} \approx 42\%$ of the total field intensity [e.g., Lavail et al., 2019] ranging from ~ 0.1 kG up to ~ 2 kG in their sample. For M dwarfs, the total field strength can range up to ~ 8 kG [Reiners et al., 2022]. Additionally, we want to highlight the case of AU Mic [Kochukhov and Reiners, 2020]. AU Mic is a 22 Myr-old, partially radiative M dwarf with a total field intensity of 2.3 kG and a dipole component of 2.0 kG, corresponding to $b_{\text{lsf}} \approx 87\%$ of the total intensity. As AU Mic has already developed a radiative core, the previously mentioned evolutionary trends [Folsom et al., 2016] allow the possibility of an even larger proportion of the large-scale field component b_{lsf} at younger ages.

Based on the aforementioned limitations and the small number of available observations, we note that reliable estimates of magnetic field strengths are difficult to obtain. The range of dipole field strengths from 0.1 kG to 2.0 kG used in this study covers most of the dipole field strengths inferred from observations. However, we do not exclude the possibility of higher and lower dipole field strengths.

Initial disk accretion rates

As shown in previous studies, the spin evolution of young stars depends on the initial accretion rate [e.g., Gallet et al., 2019]. Therefore, it is important to define realistic initial accretion rates over the M dwarf mass range at the starting point of our simulations, which is the beginning of the Class II phase. Age estimates of young Class II objects start at $\gtrsim 1$ Myr but show a significant spread of ~ 1 Myr, which complicates reliable estimates for the accretion rate at the beginning of the Class II phase [e.g., Manara et al., 2012, Fiorellino et al., 2021, Testi et al., 2022]. As a consequence, we motivate the choice of our initial accretion rates in this study not based on age estimates, but on the comparison of accretion rates of different evolutionary stages of the star-disk system.

In Fig. 5.3, observed accretion rates of Class I [assumed ages < 1 Myr, e.g., Fiorellino et al., 2022], Class II, and “flat spectrum” objects are summarized (see caption for references). Flat spectrum (or transition) objects are characterized by a flat spectral energy distribution [SED, e.g., Tobin et al., 2020], indicating the transition between Class I and Class II. Based on the observed sources, accretion rates of flat spectrum sources (blue squares) are located between Class I (gray lines) and Class II (red lines) objects. Thus, we choose the flat spectrum sources as initial values for our simulations. The solid blue line represents the fit and is taken as a reference value ($\dot{M}_{\text{ref}}$) for the initial accretion rates in the further course of this study and scales as

$$\dot{M}_{\text{init}}(M_{\star}) = \dot{M}_{\text{ref}}(M_{\star}) = 3.7 \times 10^{-8} \left(\frac{M_{\star}}{0.3 M_{\odot}} \right)^{1.60} M_{\odot} \text{ yr}^{-1}. \quad (5.3)$$

Due to large uncertainties in the observed accretion rate [e.g., Caratti o Garatti et al., 2012], we also consider smaller (by a factor of 5) initial accretion rates $\dot{M}_{\text{low}}$; these values are located at the upper edge of the typical values for Class II objects for all values of $M_{\star}$ (dashed blue line in Fig. 5.3).

5.3. Results

We describe two sets of results: first, the influence of an accretion disk on stellar SAM and how the spin evolution of M dwarfs varies with different initial stellar and disk parameters (Sec. 5.3.1); and second, how the post-disk specific angular momentum (SAM) of M dwarfs depends on stellar and disk parameters (Sec. 5.3.2). In all simulations, the disk is considered to be dissipated and simulations are stopped when the accretion rate falls below $10^{-13} M_{\odot} \text{ yr}^{-1}$. This is the minimum disk accretion rate that can be detected on M dwarfs with current observational methods [Sicilia-Aguilar et al., 2016]. At this accretion rate, the residual disk typically has a mass of $\sim 10^{-6} M_{\odot}$ and ceases to influence the stellar rotation.

Our simulations yield both disk accretion rates and lifetimes that are consistent with observations and empirical descriptions of disk accretion based on observations. Figure 5.4 compares disk accretion rate vs. time from our simulations to an empirical power-law relation $\dot{M} \propto t^{-\chi}$ [Eq. 3 in Gallet et al., 2019], where χ is a fitting parameter. Two cases are shown: The blue (red)

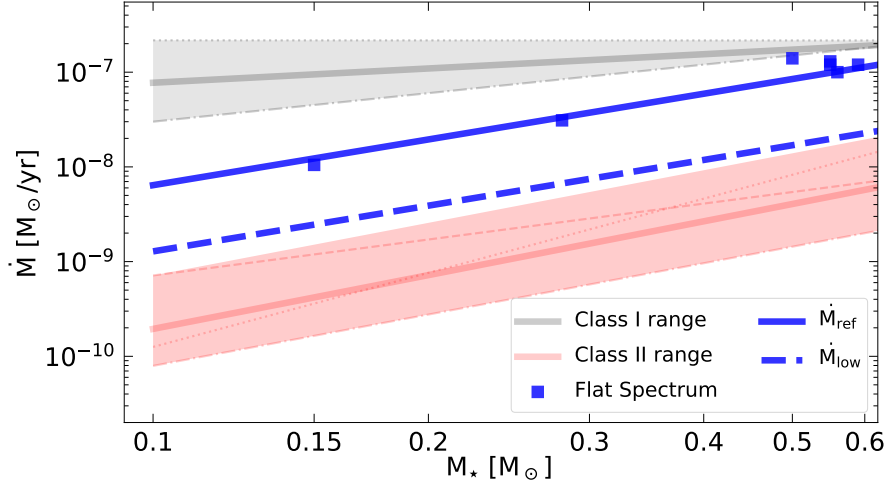


Figure 5.3.: Observed accretion rates of Class I (gray), Class II (red), and flat spectrum (blue) sources. The gray dash-dotted line is taken from Fiorellino et al. [2021] and the gray dotted line from Caratti o Garatti et al. [2012]. The gray-shaded area is limited by the two lines. The red dashed, dotted, and dash-dotted lines are taken from Manara et al. [2012], Fiorellino et al. [2021], and Testi et al. [2022], respectively. The red-shaded area has a minimum but constant logarithmic height so that all red lines are comprised. The thick gray and red lines indicate the mean values of the Class I and Class II objects. The blue squares represent accretion rates of flat spectrum sources (summarized in Table 5.6.2). The blue line is the best fit for the flat spectrum objects and the dashed line is five times less than that, just above the Class II regime.

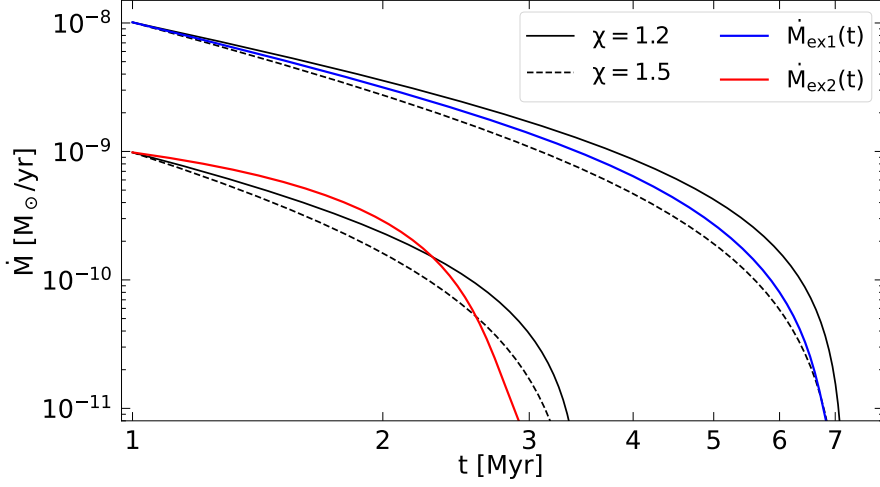


Figure 5.4.: Model disk accretion rates vs. time for two cases: The $\dot{M}_{\text{ex1}}(t)$ (blue line) and $\dot{M}_{\text{ex2}}(t)$ (red line) cases are simulations of stars with $M_\star = 0.3$ (0.1) $M_\odot$ and a disk with a viscosity parameter $\alpha = 0.005$ (0.010), respectively. The stellar magnetic field strength is the same ($B_\star = 1.0$ kG) in both cases. We compare these to empirical relations based on the power-law relation $\dot{M} \propto t^{-\chi}$ of Gallet et al. [2019]) with disk lifetimes of 3.5 and 7.2 Myr and exponents $\chi = 1.2$ (solid black line) and $\chi = 1.5$ (dashed black line) [Vasconcelos and Bouvier, 2017, Gallet et al., 2019].

line $\dot{M}_{\text{ex1}}(t)$ ($\dot{M}_{\text{ex2}}(t)$) represents our simulations with a star with $M_\star = 0.3$ (0.1) $M_\odot$ and a disk with a viscous parameter $\alpha = 0.005$ (0.010), respectively. The stellar magnetic field strength is the same ($B_\star = 1$ kG) in both cases. The power-law exponent takes values of $\chi = 1.2$ [solid black line, Gallet et al., 2019] and $\chi = 1.5$ [dashed black line, Vasconcelos and Bouvier, 2017]. For disk lifetimes of 7.2 and 3.5 Myr, our simulations are in reasonable agreement with the empirical power law. We note that the lifetimes among our simulations range between ~ 2 and 11 Myr. This range is comparable to disk lifetimes, inferred from observations [e.g., Richert et al., 2018, Pfalzner et al., 2022, and references therein]. Accretion rates that are initially $\gtrsim 10^{-8} M_\odot \text{ yr}^{-1}$ are also predicted to fall below $10^{-9} M_\odot \text{ yr}^{-1}$ within a few Myr (Fig. 5.4), consistent with the range of accretion rates observed in clusters at such ages [Manara et al., 2022].

5.3.1. The influence of a disk on stellar AM

To study the influence of an accretion disk on the central star, we consider different initial values for the stellar and disk AM and compare the final stellar AM to its initial value once the disk has

dissipated. For a given initial stellar radius and mass, the stellar AM in our model can be changed by varying the initial rotational period P_{init} . Since the disk is in quasi-Keplerian rotation, its AM can be adjusted by changing the disk mass, which we do by varying its initial accretion rate $\dot{M}_{\text{init}}$ for a fixed α [$\dot{M} \propto \alpha \Sigma$ for a viscous accretion disk; see Armitage, 2013].

To illustrate the varying effect of disks with different initial accretion rates on the final stellar AM and rotation period, we consider, as a fiducial case, a $0.3 M_{\odot}$ star with a photospheric magnetic field of $B_{\star} = 0.5$ kG. The disk's viscous α parameter is set to 1×10^{-2} and we set the efficiency of a stellar wind to $W = 2$ %. P_{init} is varied from 1 to 10 days and we vary $\dot{M}_{\text{init}}$ from 10^{-12} to $3 \times 10^{-8} M_{\odot} \text{ yr}^{-1}$ (see Sec. 5.2.3). Figure 5.5 plots the final stellar rotation period at disk dissipation as a function of P_{init} and $\dot{M}_{\text{init}}$. For a more intuitive interpretation of this plot, we show the initial and final rotation periods instead of the SAM values. The same conclusions can be drawn for SAM because the duration of these simulations is short compared to the Kelvin-Helmholtz contraction time ($\tau_{\text{KH}} \sim GM_{\star}^2/(R_{\star}L_{\star})$) of the star and thus the stellar moment of inertia does not change significantly.

In our stellar spin model (see Sec. 5.2.2), P_{final} in Fig. 5.5 is influenced by accretion, the magnetic star-disk connection, and the APSW. These quantities are calculated self-consistently and depend on time and on stellar and disk parameters [see Steiner et al., 2021]. Contrary to the effects of accretion and the APSW, which (in our model) always spin up and spin down the star, respectively, MEs can result in a spin-up or spin-down torque that acts on the star. As described in Sec. 5.2.2 and, for example, in Gallet et al. [2019], the position between the truncation radius and the corotation radius determines the sign of this torque contribution. MEs spin up the star if the truncation radius is located well within the corotation radius. In case the truncation radius moves outward and toward the corotation radius, MEs can spin down the star [e.g., Gallet et al., 2019].

For a given truncation radius ($r_{\text{trunc}} < r_{\text{cor}}$), the corotation radius of a star that spins up moves inward, and the distance between r_{trunc} and r_{cor} decreases. Thus, the spin-up torque of the MEs is reduced or even turned into a spin-down torque in the case of a faster-rotating star. Following the same logic, the MEs tend to increase the spin-up torque onto the star, if the star spins down. For a given stellar rotation and corotation radius, an increased accretion rate can also alter the effects of the MEs. A high accretion rate pushes the truncation radius inward and the distance between

r_{trunc} and r_{cor} increases. Thus, the MEs tend to spin up the star during phases of high accretion rates. Analogously, the distance between the truncation and corotation radius decreases during phases of low accretion rates, and the MEs spin down the star.

Furthermore, the strength and sign of the MEs depend on the magnetic field strength. A strong stellar magnetic field strength results in larger truncation radii. As a consequence, the MEs tend to show a spin-down tendency for strong stellar magnetic field strength. For weak stellar magnetic fields, on the other hand, the distance between the truncation and corotation radius increases, and the MEs show a spin-up tendency. Furthermore, the magnitude of the torque due to MEs is proportional to $B_{\star}^2$ [e.g., Gallet et al., 2019] and the overall effect of MEs on the stellar spin evolution increases in case of strong magnetic field strengths.

We note that the sign and magnitude of the torque due to MEs, Γ_{ME} , are strongly coupled to different (in turn coupled) parameters such as the accretion rate, stellar rotation period, and magnetic field strength. We show the evolution of Γ_{ME} in Fig. 5.11 for different configurations. As expected, the MEs result in a spin-up torque for high accretion rates during the beginning of our simulations, weak magnetic field strengths, or slow rotation periods. For fast-rotating stars and strong magnetic field strengths during phases of low accretion rates (especially toward the end of the disk lifetime), MEs spin down the star. Compared to the spin-down torque of APSW (see Fig. 5.11), the effect of MEs is comparable for very fast-rotating stars. MEs are even the dominant source of stellar AM loss during the last $\gtrsim 1$ Myr of the disk lifetime.

We can divide the parameter space in Fig. 5.5 into three regions (demarcated by red lines). In Region 1 (2), the star-disk interaction causes the stellar period to increase (decrease), respectively. The influence of the disk, however, is not strong enough to erase the imprint of the initial stellar spin values. A rapidly rotating star remains rapidly rotating compared to a slow rotator and the dispersion in the final states is comparable to the spread of initial values. In this regime, at the bottom of Fig. 5.5, the contours are more vertical. In Region 3, the accretion disk is the dominant factor that controls the stellar spin evolution during the disk phase, the initial conditions are forgotten, and P_{final} converges toward a narrower range of values, nearly independent of P_{init} . This is the regime at the top of Fig. 5.5 where the contours are nearly horizontal. The star's initial rotation state is "erased" by the interaction with the disk. This phenomenon has been seen in

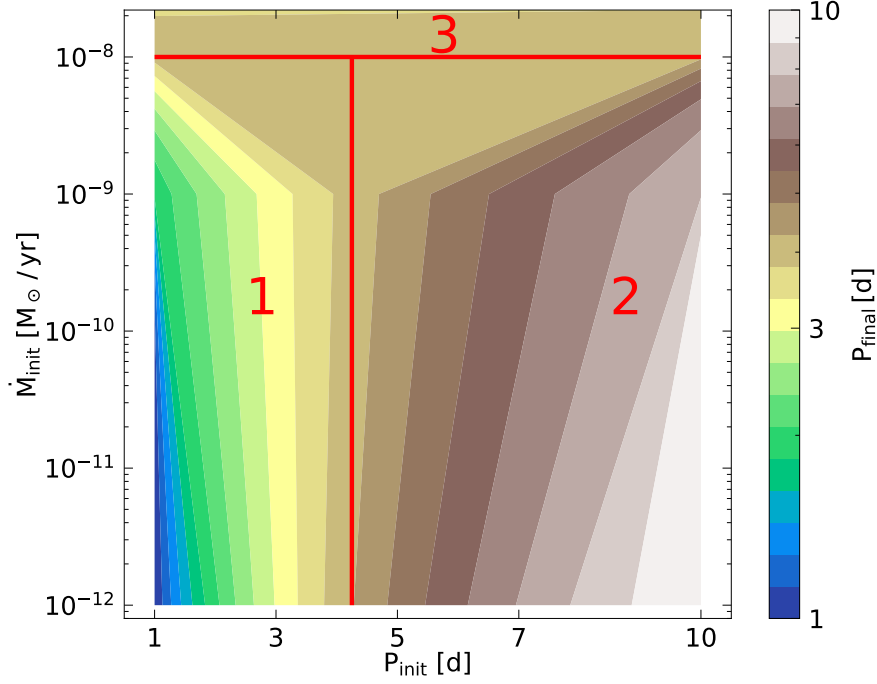


Figure 5.5.: Final stellar rotation period P_{final} for different initial periods P_{init} and disk accretion rates $\dot{M}_{\text{init}}$. The stellar and disk parameters are: $B_{\star} = 1000G$, $M_{\star} = 0.3M_{\odot}$, $W = 2\%$ and $\alpha = 0.01$. We can divide the parameter space into three different regions, separated by red lines (see text).

previous studies [e.g., Armitage and Clarke, 1996, Matt et al., 2012] but the dependence of this behavior on star and disk parameters has not been investigated.

To investigate how this behavior depends on stellar and disk properties other than accretion rate, we vary these parameters: the strength of the star's magnetic dipole field at its surface $B_{\star} \in [0.5, 1 \text{ kG}]$ (see Sec. 5.2.3), the initial stellar mass $M_{\star} \in [0.3, 0.6 M_{\odot}]$, and the disk viscous parameter $\alpha \in [0.005, 0.01]$. The resulting effect of the accretion disk on the stellar spin evolution is visualized by comparing the ratios between maximum and minimum values of the stellar rotational period and the specific angular momentum after the disk has been dispersed relative to the ratios of the initial values. A parameter C is introduced to quantify the change of these ratios by the following relations

$$\left(\frac{P_{\text{max}}}{P_{\text{min}}}\right)_{\text{final}} = \left(\frac{P_{\text{max}}}{P_{\text{min}}}\right)_{\text{init}}^{C_p} \quad (5.4)$$

and

$$\left(\frac{j_{\max}}{j_{\min}}\right)_{\text{final}} = \left(\frac{j_{\max}}{j_{\min}}\right)_{\text{init}}^C. \quad (5.5)$$

The left-hand sides of Eqs. (5.4) - (5.5) denote the stellar rotational periods and specific angular momentum values after the disk has been dissolved (final values) and the right-hand sides the respective initial values. C can be understood as the slope of the contours in a log-log plot (Fig. 5.5). A value of C close to 1 corresponds to a weak influence of the accretion disk. A strong influence on the stellar spin evolution on the other hand is indicated by $C \ll 1$. In this case, the initial rotational values are "forgotten" and the star's final state is independent of the initial rotation. We consider $C = 0.01$ (corresponding to a final difference of $\approx 2\%$) sufficiently low for final conditions to be convergent and define the initial accretion rate at which this condition is reached as a critical accretion rate $\dot{M}_{\text{crit}}$.

We quantify the effect of variation of the parameters disk viscosity (α), stellar wind (W), stellar magnetic field ($B_{\star}$), and stellar mass ($M_{\star}$) on $\dot{M}_{\text{crit}}$ in Fig. 5.6. These parameters, α in Panel a, W in Panel b, $B_{\star}$ in Panel c, and $M_{\star}$ in Panel d, determine the spin-down behavior. Therefore, for each set of parameters, we can identify a value of $\dot{M}_{\text{crit}}$ that is $\sim 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$. Above this value, the stellar rotational period, as well as the SAM, do not depend on their initial values.

5.3.2. Relations for the post-disk specific angular momentum

We examine the dependencies of the SAM after the disk phase on stellar and disk parameters. We compare these to the distribution of SAM among M dwarfs in the Upper Scorpius star-forming region reported by Somers et al. [2017]. This sample is especially suitable for studying the AM distribution after the disk phase because the cluster has an age of $\sim 10 \text{ Myr}$ [Feiden, 2016] and most accretion disks dissipate by that time [Mamajek, 2009]; also there's been no significant loss of AM from a thermal stellar wind up to this age. Somers et al. [2017] provided a mean and standard deviation for SAM in mass bins of $0.10 - 0.25 \text{ M}_{\odot}$, $0.25 - 0.40 \text{ M}_{\odot}$ and $0.40 - 0.60 \text{ M}_{\odot}$, respectively.

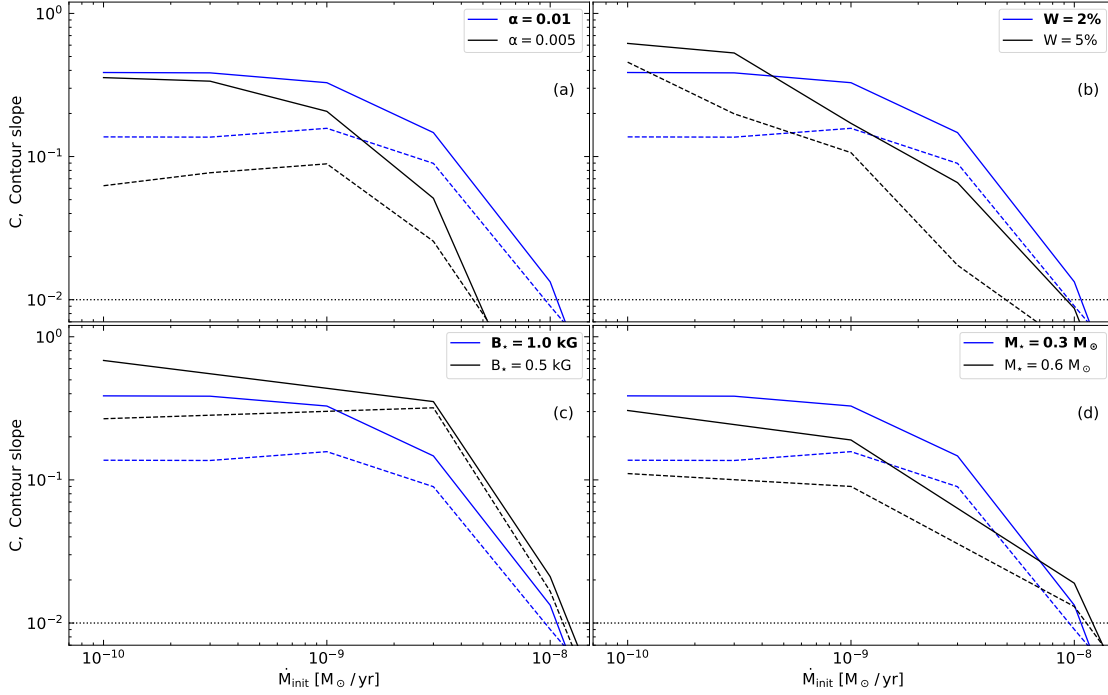


Figure 5.6.: C -parameter (the gradient of the contours in Fig. 5.5) with respect to initial disk accretion rates $\dot{M}_{\text{init}}$. The solid lines represent C_P and the dashed lines C_J values, respectively. Each subplot shows variations in (a) viscous α parameter: 0.01 (blue lines) and 0.005 (black lines). The stellar mass is $0.3 M_{\odot}$ with a magnetic dipole field strength of $B_{\star} = 1 \text{ kG}$. (b) W values: 2% (blue lines) and 5% (black lines). The stellar mass is $0.3 M_{\odot}$ with a magnetic dipole field strength of $B_{\star} = 1 \text{ kG}$ and a viscosity parameter of $\alpha = 0.01$. (c) $B_{\star}$ values: 1.0 kG (blue lines) and 0.5 kG (black lines). The stellar mass is $0.3 M_{\odot}$ with a viscosity parameter of $\alpha = 0.01$. (d) $M_{\star}$ values: $0.3 M_{\odot}$ (blue lines) and $0.6 M_{\odot}$ (black lines). The viscosity parameter is $\alpha = 0.01$. The threshold of $C = 0.01$ for the star to be in the "locked"-state is indicated by the horizontal dotted line. For better visualization, the highlighted parameters show the reference value in the other subplots.

Influence of stellar parameters on the post-disk SAM

We perform a set of simulations to examine the effect of varying stellar mass and magnetic field on the final (post-disk) SAM of the star, j_{final} . The initial stellar mass is set to 0.1, 0.3, $0.5 M_{\odot}$, reflecting the three mass bins of Somers et al. [2017]. We note that during the simulations the stellar mass is increased by $\lesssim 5 \%$ due to the accretion of disk material. Three different initial stellar rotation periods are assumed: 1, 3, and 10 days. We adopt initial accretion rates

that correspond to an evolutionary stage just preceding the classical T Tau stage, namely those of flat-spectrum transitional Class I to Class II sources [Tobin et al., 2020]; these empirically determined accretion rates are mass dependent (see Eq. 5.3). In addition, the magnetic dipole field strength is set to values of 0.1, 0.2, 0.4, 0.7, 1.0, and 2.0 kG. The viscous parameter is set to $\alpha = 0.01$. The SAM after the disk phase j_{final} for each possible combination is shown in Fig. 5.7. The colored circles, crosses, and squares represent initial rotation periods of 1, 3, and 10 days, respectively, and the solid, dash-dotted, and dashed lines show the respective fits. Mean and standard deviation values for the *observed* SAM taken from Somers et al. [2017] for each mass bin are represented by gray solid and dashed lines, respectively.

For stellar masses, $\gtrsim 0.3 M_{\odot}$ the initial accretion rate is $\gtrsim \dot{M}_{\text{crit}}$ (see Eq. (5.3)) and rotation period at the end of the disk phase is largely independent of the initial period (see Panel c). For stellar masses, $< 0.3 M_{\odot}$, the initial accretion rates are below the critical value and the initial rotation period influences the final value for $B_{\star} < 1$ kG (see panel a). The ensemble of outcomes is well represented by power laws of the form $j_{\text{final}} \sim B_{\star}^{-b}$. The magnetic field strengths of the intersection between the fitted data and the mean values from Somers et al. [2017] ($\bar{B}_{\star}$) and the exponents b are summarized in Table 5.1.

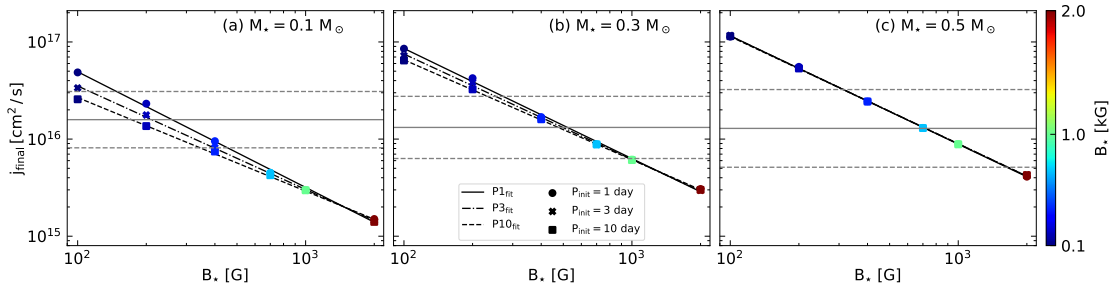


Figure 5.7.: The SAM after the disk phase j_{final} for different stellar magnetic field values and stellar masses of $0.1 M_{\odot}$ (Panel a), $0.3 M_{\odot}$ (Panel b), and $0.5 M_{\odot}$ (Panel c). The colored circles, crosses, and squares represent initial rotation periods of 1, 3, and 10 days, respectively, and the solid, dash-dotted, and dashed lines show the respective fits. Mean and standard deviation values for the SAM taken from Somers et al. [2017] are represented by gray solid and dashed lines, respectively.

Influence of disk parameters on j_{final}

To study how the j_{final} relations derived in Sec. 5.3.2 depend on the assumed values for disk parameters, we vary the initial accretion rate, the viscosity value α , as well as the APSW efficiency W and compare the results to the standard cases described in the previous subsection. In Fig. 5.8, the results are shown and in Table 5.2, the magnetic field strengths of the intersection between the fitted data and the mean values from Somers et al. [2017] ($\bar{B}_\star$) and the exponents b are reported. In the following cases, the initial stellar mass is set to $0.3 M_\odot$. For comparison, the yellow lines show the fits from Fig. 5.7 for a stellar mass of $0.3 M_\odot$

We reduce the initial accretion rate by a factor of 5 compared to the values generated by Eq. (5.3). At the lower initial accretion rates $\dot{M}_{\text{low}}$ (Panel a), the star-disk interaction is weaker than in the reference case, and j_{final} is more strongly influenced by the initial conditions. As a result, the difference between $\bar{B}_\star$ and b for fast and slow initial rotation periods is larger. More intuitively, if the predicted values of j_{final} are to agree with the observations, a larger variation in $B_\star$ and the exponent b is needed to compensate for variations in the initial rotation period.

The amount of AM that is removed by the APSW is controlled by the efficiency parameter W . As a consequence, the values for j_{final} are shifted in the vertical direction toward higher (lower) values for smaller (larger) values of W , respectively. The values for the magnetic field strengths of the intersection between the fitted data and the mean values from Somers et al. [2017], $\bar{B}_\star$, are nearly equally increased (reduced) by 530 G (137 G) in case of $W = 0\%$ ($W = 5\%$) (see Panels b and c in Fig. 5.8 and Table 5.2). At a given accretion rate, a lower value of α increases the disk mass and lifetime, allowing star-disk interaction to influence the stellar AM over a longer time period and reducing the critical accretion rate (see Panel a in Fig. 5.6). The values in $\bar{B}_\star$ and b for different P_{init} values converge and the initial conditions are now (nearly) completely forgotten (Panel d in Fig. 5.8).

The stellar SAM after the disk phase can be described by a power law such as

$$j_{\text{final}} = j_{\text{ref}} \left(\frac{B_\star}{0.5 \text{ kG}} \right)^{-b}, \quad (5.6)$$

with the proportionality factor j_{ref} , depending on the stellar mass and, if $\dot{M}_{\text{init}} < \dot{M}_{\text{crit}}$, also on the

initial stellar rotation period P_{init} . The values for j_{ref} for our reference case are summarized in Table 5.1.

Table 5.1.: Values of $\bar{B}_\star$, b , and j_{ref} for models with $\dot{M}_{\text{ref}}$. The units for $M_\star$, P_{init} , $\bar{B}_\star$, and j_{ref} are $M_\odot$, days, G, and $10^{16} \text{ cm}^2/\text{s}$ respectively.

$M_\star$	0.1	0.3	0.5
P_{init}	1,3,10	1,3,10	1,3,10
$\bar{B}_\star$	259,211,172	521,497,475	720,718,715
b	1.19,1.06,0.96	1.13,1.08,1.02	1.11,1.11,1.10
j_{ref}	0.73,0.63,0.57	1.38,1.31,1.25	1.93,1.92,1.92

Table 5.2.: Values of $\bar{B}_\star$ and b for models with $\dot{M}_{\text{low}}$, $W = 5 \%$, and $\alpha = 0.005$. The values $\Delta_{\bar{B}_\star}$ and Δ_b show the difference with respect to the models with $\dot{M}_{\text{ref}}$. The initial stellar mass is $0.3 M_\odot$ for all models. The units for P_{init} , $\bar{B}_\star$, and j_{ref} are days, G, and $10^{16} \text{ cm}^2/\text{s}$, respectively.

P_{init}	$\dot{M}_{\text{low}}$	$W = 0 \%$	$W = 5 \%$	$\alpha = 0.005$
	1,3,10	1,3,10	1,3,10	1,3,10
$\bar{B}_\star$	536, 426, 298	1058, 1021, 1005	381, 362, 338	461, 458, 456
$\Delta_{\bar{B}_\star}$	+15, -71, -177	+537, +524, +530	-140, -135, -137	-60, -39, -19
b	1.23, 1.08, 0.90	1.12, 1.08, 1.03	1.13, 1.08, 1.01	1.10, 1.09, 1.09
Δ_b	+0.10, <0.01, -0.12	-0.01, <0.01, +0.01	<0.01, <0.01, -0.01	-0.03, +0.01, +0.07

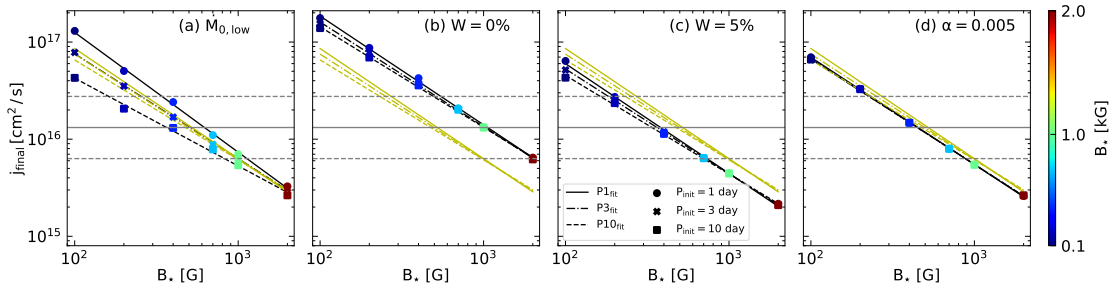


Figure 5.8.: The SAM after the disk phase j_{final} as a function of stellar magnetic field values (x-axis and colors) and for different initial rotation periods (different sets of point styles and curves). Each panel shows the variation of one parameter with respect to the reference case in Fig. 5.7. In Panel (a), $\dot{M}_{\text{low}}$ is used as the initial accretion rate. In Panel (b) and (c), the APSW efficiency is set to 0 % and 5 %, respectively. In Panel (d), the viscous α parameter is set to 0.005. In all cases, the stellar mass is $0.3 M_\odot$. The symbols and lines are identical to Fig. 5.7. For comparison, the yellow lines show the fits from Fig. 5.7 for a stellar mass of $0.3 M_\odot$.

5.4. Discussion

We now wish to discuss the consistency of our results with observations, resulting implications, limitations, and uncertainties due to initial conditions and the adopted ranges of our principal parameters.

5.4.1. Comparison of $\bar{B}_\star$ with inferred dipole field strengths

We compare the values of $\bar{B}_\star$ that are required to explain the observed SAM to the dipole field strengths observed or inferred on young, low-mass stars. In Fig. 5.9, dipole magnetic field strengths are shown with respect to the stellar mass. Solid black circles show dipole field strengths inferred by the Stokes V parameter (B_V), open circles indicate an estimate based on the total field intensity (B_I), and a dipole field strength fraction of $b_{\text{lsf}} = 22\%$ [Lavail et al., 2019], and the black diamond indicates the dipole field strength of AU Mic, the only young star in our sample with direct measurement of B_{QU} . The data is taken from Yang and Johns-Krull [2011], Johnstone et al. [2014], and Table 5.4. It is important to note that, for this comparison, young accreting and nonaccreting stars are combined to increase the number of available data points. To minimize the effect of an evolving magnetic field, the age of the nonaccreting stars is restricted to the disk dispersal age (see Table 5.4).

The range of the required magnetic field strengths to match the observed median value of the SAM distribution [Somers et al., 2017] is marked in blue ($\bar{B}_\star$) and the respective standard deviations given in Somers et al. [2017] are indicated by the blue dashed lines (B_σ). The changes of $\bar{B}_\star$ due to different disk and stellar parameters ($\Delta_{\bar{B}_\star}$, see Table 5.2) are extrapolated from $0.3 M_\odot$ over the M dwarf mass range. To reproduce the observed mass-independent distribution of SAM, our model requires a positive relation between stellar mass and magnetic field strength. The observed values can explain the disk-star coupling required to explain the post-disk SAM distribution [Somers et al., 2017]. The data, however, are not sufficient and the uncertainties are too large to test the required mass dependence.

The theoretical justification for a mass-dependent magnetic field strength is described by Browning et al. [2016]. By equating the magnetic energy density $B_\star^2$ to the kinetic energy

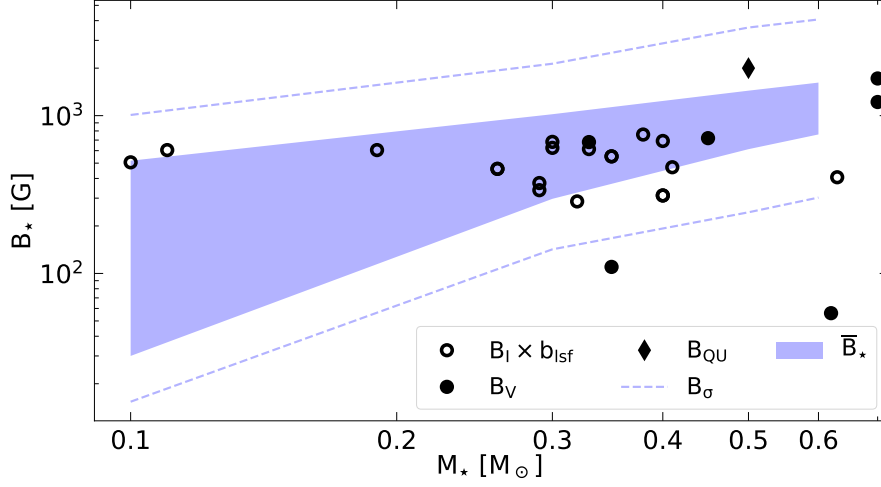


Figure 5.9.: Observed dipole field strengths of young stars with respect to stellar mass. Solid black circles show dipole field strengths inferred by the Stokes V parameter (B_V), open circles indicate an estimate based on the total field intensity (B_I) and a dipole field strength fraction of $b_{lsf} = 22\%$ [Lavail et al., 2019], and the black diamond indicates the dipole field strength of AU Mic, the only young star in our sample with direct measurement of B_{QU} . The errors are expected to reach up to one order of magnitude. The range of the required magnetic field strengths to match the observed median value of the SAM distribution [Somers et al., 2017] is marked in blue ($\bar{B}_\star$) and the respective standard deviations given in Somers et al. [2017] are indicated by the blue dashed lines (B_σ).

density in the turbulent flow of the star's convective envelope ρv_c^2 , with v_c the convective velocity, Browning et al. [2016] derive a theoretical kinetic limit to the magnetic field strength on rapidly rotating stars. Since turbulent convection is responsible for energy transport from the stellar interior, $B_\star$ can then be related to the luminosity of the star and thus to T_{eff} via the Stefan-Boltzmann relation: $B_\star \propto (M_\star T_{\text{eff}}^8 / R_\star^3)^{1/6}$. Power-law fits of T_{eff} vs. $M_\star$ and $R_\star$ vs. $M_\star$ for Baraffe et al. [2015] ≤ 10 Myr isochrones return power-law indices of ≈ 0.15 and 0.33 , respectively, depending on age. This yields $B_\star \propto M_\star^{0.2}$. If Ohmic dissipation and magnetic buoyancy are considered in addition [Browning et al., 2016], the magnetic field strength scales like $B_\star \propto M_\star^{0.3}$. Compared to our results, this relation shows a slightly weaker trend in the same direction, but can at least partially explain the observed distribution of SAM.

5.4.2. Comparisons with individual stars after the disk phase

We compare our results for j_{final} to available measurements of stellar rotation, mass, and magnetic field strength for stars that are old enough for disks to have dissipated, but not significantly older than ~ 100 Myr so that the AM lost in a magnetized wind is small. Unfortunately, the set of published observations that are relevant to our case is very limited. Due to the nature of the different measurement techniques, we use B_I as an upper limit and B_V as a lower limit of the surface field associated with the large-scale dipole component. While there are ~ 50 stars that have estimates of B_I and B_V [Kochukhov, 2021], only five of these fall within the required mass ($< 0.6 M_\odot$) and age ranges ($\lesssim 100$ Myr) and B_{QU} is only available for AU Mic [Kochukhov and Reiners, 2020]. In total, 12 stars are chosen to be shown in this comparison (see Table 5.4). For some stars, there is no large-scale magnetic field measurement available. Finally, large uncertainties in the observed quantities impede the comparison. Our analysis should be considered as a coarse interpretation of possible trends and an order of magnitude consideration.

The observations are summarized in Table 5.4 and we compare these to our predicted values of $j_{\text{ref}}(M_\star)$ scaled by $B_\star^{1.1}$ (see Sec. 5.3.2) in Fig. 5.10. The upper and lower limits of $\xi = j_{\text{obs}} \times (B_\star/0.5 \text{ kG})^{1.1}/j_{\text{ref}}$ due to B_I and B_V are indicated by left and right facing triangles, respectively. The more precise estimate for AU Mic based on B_{QU} (black diamond) is plotted with an error bar. A value of $\xi = 1$ (vertical dashed line) indicates a perfect agreement with our reference case. Due to the variations in stellar and disk parameters, j_{ref} can vary (see Fig. 5.8), which is indicated by the gray area. This gray area falls within the range bracketed by the observed values of B_I and B_V . Note that we assume $P_{\text{init}} = 3$ days for all stellar masses for this comparison. The agreement is encouraging but the limited number of observations, all but one of which are lower or upper limits on the magnetic field, does not allow us to draw firm conclusions. Many more detailed measurements of the magnetic fields of young M dwarfs are needed for a more robust test of our models.

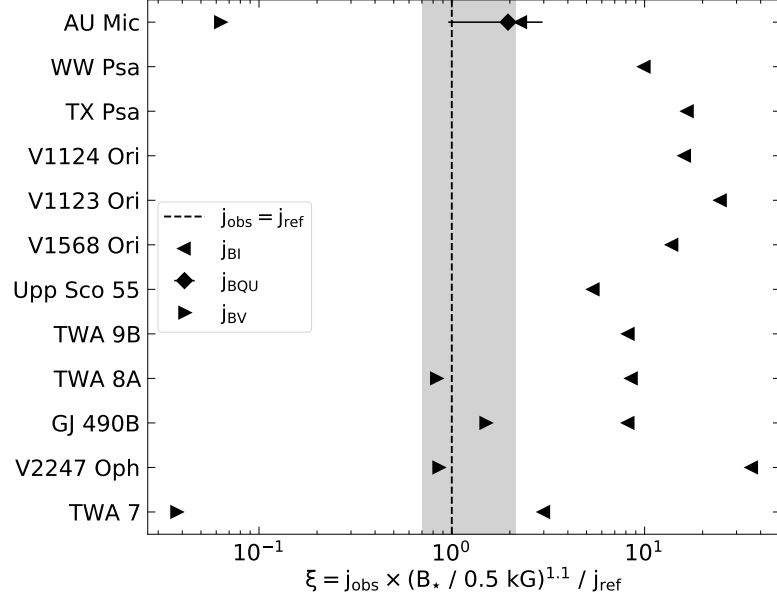


Figure 5.10.: The ratio of the observed to predicted SAM $j_{\text{obs}}/j_{\text{final}}$. To facilitate the comparison, j_{obs} is multiplied by a factor of $(B_{\star}/0.5 \text{ kG})^{1.1}$ (see text). The vertical dashed line indicates a perfect agreement between the observed value and our model $j_{\text{ref}}(M_{\star})$. The gray area accounts for different stellar and disk parameters. The SAM values j_{BI} and j_{BV} are based on B_{I} and B_{V} functions as the upper and lower boundary for the stellar SAM. They are indicated by left and right-facing triangles, respectively. The single measured value (and uncertainties) of j_{BQU} (diamond marker) is based on B_{QU} for the star AU Mic. The measurements are tabulated in Table 5.4.

5.4.3. Accretion-powered stellar wind (APSW) efficiency W

Winds, in this study represented by APSW, can strongly affect the distribution of stellar SAM as well as the structure and long-term evolution of its disk [e.g., Matt and Pudritz, 2005b, Gallet et al., 2019]. APSW is a potential explanation for the existence of slowly spinning stars, but their exact mechanisms and strength are the subject of active research [see the discussion in Gallet et al., 2019]. Observations of Doppler-shifted line absorption features [Dupree et al., 2005, Edwards et al., 2006] as well as statistical evaluations of the observed ratio of outflow mass loss to disk accretion rate [e.g., Cabrit, 2002, Ferreira et al., 2006] indicate that there are indeed strong outflows from young star-disk systems. These outflows are expected to be too cool to originate from a thermal pressure gradient at the stellar corona where temperatures are $\sim 10^6$ K [e.g., Gallet et al., 2019] and their exact launch point and driving mechanism are not known [e.g., Ireland et al., 2021] but they are supposed to be disk winds launched in the inner disk, possibly by the magneto-centrifugal process [e.g., Königl et al., 2011]; some of these winds can even form jets. A possible alternative driving mechanism for winds is proposed in Decampli [1981] and Matt and Pudritz [2005b]. The accreting disk material can excite Alfvén waves originating at the stellar surface, traveling outward and removing angular momentum from the star. The mass loss rate through these winds is expected to be $W \sim 2\%$ [Cranmer, 2008, Pantolmos et al., 2020].

While previous authors need large APSW values of $\sim 10\%$ to explain slowly spinning stars, which contradicts the results of Cranmer [2008] and Pantolmos et al. [2020], we can reduce this value considerably to reproduce the observed rotational period of slowly rotating stars. In this work, we can limit W to $\lesssim 5\%$, which is in better agreement with observational constraints and consistent with the theoretical predictions of Cranmer [2008] and Pantolmos et al. [2020]. The reduction of W might be explained by the combined stellar and disk model that includes the back-reactions of the rotational evolution on the accretion disk and the spin-down effect of MEs, especially during phases of fast stellar rotation and toward the end of the disk lifetime. Apart from APSW and MEs, other wind mechanisms can be important for star and disk evolution. Especially, disk winds noticeably affect the disk evolution [e.g., Königl et al., 2011]. By removing mass and (if magnetized) angular momentum from the disk before it can be accreted onto the star, the evolution of stellar spin evolution is affected. Likewise, photo-evaporation (internal or external)

can influence the disk’s lifetime and thus, the stellar spin evolution [e.g., Roquette et al., 2021]. These effects will be added to a future version of our model.

5.4.4. Comparing our model to large-scale parametric studies

Our model predictions can be directly integrated into parametric-based modeling of the stellar spin evolution of young stars, including the influence of an accretion disk [e.g., Vasconcelos and Bouvier, 2017]. While Vasconcelos and Bouvier [2017] have conducted a large number of simulations ($\sim 2 \times 10^5$), exceeding the number of our simulations by over two orders of magnitude, their model is parametric and the evolution of the accretion disk and the star-disk interaction is based on simplified assumptions. The disk accretion rate is taken to be independent of stellar mass and the disk lifetime is estimated based on a free parameter. During the disk lifetime, disk-locking is assumed (the stellar rotational period is constant over time) and important aspects like the influence of a stellar magnetic field or an accretion-powered stellar wind are ignored.

Vasconcelos and Bouvier [2017] can present useful guidelines regarding the rotational evolution of low-mass stars. However, none of their models can reproduce all the available observational aspects (e.g., a correlation between stellar mass and rotation period). We can provide mass-dependent disk lifetimes for different disk parameters (e.g., Fig. 5.3) and a relation for the SAM after the disk phase that includes the influence of the stellar magnetic field strength and an accretion-powered stellar wind (Eq. 5.6). Combining our results with the method presented in Vasconcelos and Bouvier [2017], could lead to new constraints for stellar and disk parameters (e.g., the mass-dependent distribution of stellar magnetic strength or mass-dependent APSW efficiency).

5.5. Conclusion and caveats

To investigate the factors that govern the angular momentum distribution of young M dwarfs, we ran more than 500 long-term (up to 10 Myr) simulations of a rotating star and disk and the interaction between them in the innermost disk ($\lesssim 0.05$ AU). This large number of simulations,

spanning a wide parameter space, was possible due to the implicit solver in the code, which circumvents the limitations of using explicit time steps when simulating disk physics on both small (inner edge) and large (outer edge) scales. Our model includes physics in a more self-consistent manner (e.g., pressure gradients and magnetic torques in the inner disk, deviations from the Keplerian velocity or the back-reaction in the disk) compared to previous long-term simulations [e.g., Matt et al., 2010a, Gallet et al., 2019].

The main conclusions of this work are:

- If the initial accretion rate (a proxy for disk mass) is above a critical value $\dot{M}_{\text{crit}} \sim 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$, the accretion disk effectively "erases" the initial spin of the star and the spin and angular momentum (AM) evolves in a convergent manner that is independent of the initial AM. In such cases, the final SAM (SAM) is controlled by stellar and disk parameters, $M_{\star}$, $B_{\star}$, α , and W . Consistent with Armitage and Clarke [1996], we find that sufficiently massive disks can regulate stellar rotation.
- For stellar masses $> 0.3 \text{ M}_{\odot}$, the initial accretion rates $\dot{M}_{\text{init}}$ at $\sim 1 \text{ Myr}$ are comparable to $\dot{M}_{\text{crit}}$. The SAM of these young M dwarfs after the disk phase j_{final} is independent of the initial conditions and scales with the stellar magnetic field strength as $B_{\star}^{-1.1}$. For stellar masses $\lesssim 0.3 \text{ M}_{\odot}$, $\dot{M}_{\text{init}} < \dot{M}_{\text{crit}}$ and the initial stellar rotation periods affect the SAM after the disk phase.
- Stellar dipole field strengths in the range of $\sim 30 \text{ G}$ to $\sim 1 \text{ kG}$ (see Fig. 5.9) are necessary to reproduce the nearly mass-independent SAM distribution reported by Somers et al. [2017] with higher stellar masses requiring stronger field strengths. These results match the range of observed large-scale dipole field strengths reported in the literature. In addition, the positive relation between the stellar mass and the magnetic field strength is needed to explain the mass-independent SAM distribution in the presence of mass-dependent accretion rates and is suggested by theoretical scaling arguments [Browning et al., 2016].
- Our scaling relation is also consistent with (but not robustly tested by) the few available observations of SAM, mass, and magnetic field strength in individual stars.

Our study points to future directions in advancement in both modeling and observations but should be considered preliminary because of important caveats and limitations of our model.

Model dimensionality: Our one-dimensional model does not resolve the vertical structure of the disk or include vertical gradients as shear. Thus, certain mechanisms such as for example vertical shear instabilities [e.g., Urpin and Brandenburg, 1998, Nelson et al., 2013, Latter and Kunz, 2022] and vertically resolved “dead” zones [e.g., Jankovic et al., 2021, Delage et al., 2022] which could be important for the transport of mass and angular momentum cannot be included. An expansion to two dimensions, including the vertical dimension, would allow such phenomena to be included.

Early disk evolution: Phenomena relevant to the earlier ($\lesssim 1$ Myr) phase (Class 0 and Class 1 phases) of stellar evolution of the star-disk system are not included in the TAPIR code. The effects of a more massive disk, accretion outbursts, and in-falling material from the parent cloud core could be included in future studies. Vorobyov et al. [2017b] show that unsteady accretion during this early phase can affect later stellar evolution.

Premain sequence contraction: In the current TAPIR version the contraction of the central star is described by the simplified Kelvin-Helmholtz contraction. Due to unsteady accretion, which can occur even in later disk stages [e.g., HBC 722, Kóspál et al., 2016], the contraction can deviate significantly from the Kelvin-Helmholtz relation. A future improvement could be combining the TAPIR code with the predictions of a full stellar evolution code [e.g., MESA Paxton et al., 2019a]. The combination of a stellar evolution code with our current model would further allow a more self-consistent treatment of the evolution of the stellar magnetic field strength, assuming the magnetic energy density scales with the kinetic energy of the convective flow [e.g., Browning et al., 2016].

Stellar multiplicity: In the current study, we only consider single stars. Binary or multiple systems tend to influence the evolution of the accretion disk [e.g., Akeson et al., 2019, Zagaria et al., 2021]. The influence of a binary companion on the angular momentum of a star will be studied in future work.

Transport of mass and angular momentum: Our model relies on a simple (constant α -parameter) prescription for viscous transport of mass and angular momentum in the disk. α is almost certainly not constant with time and location in the disk, and other mechanisms such as disk winds and photo-evaporation, which cannot be described with this parameterization, are important for disk evolution [e.g., Königl et al., 2011, Roquette et al., 2021]. These are currently

not included but could, in principle, be added in some future version of our model.

5.6. Appendix

5.6.1. Influence of MEs for different configurations

The torque due to magnetospheric ejections (MEs) is shown for different configurations in Fig. 5.11 compared to a reference model. The reference model (solid black line) has a magnetic field strength of $B_\star = 0.5$ kG, an initial rotation period of $P_{\text{init}} = 10$ days, an initial accretion rate of $\dot{M}_{\text{init}} = 1 \times 10^{-8} M_\odot \text{ yr}^{-1}$, a viscosity parameter $\alpha = 0.01$, and a stellar mass of $M_\star = 0.3 M_\odot$. We increase the strength of the magnetic field to $B_\star = 1.0$ kG (model " $B_\star$ high," solid red line), decrease the initial rotation period to 1 day (model " P_{init} fast," solid blue line), and reduce the initial accretion rate to $\dot{M}_{\text{init}} = 3 \times 10^{-9} M_\odot \text{ yr}^{-1}$ (model " $\dot{M}_{\text{init}}$ low," solid yellow line). The transfer between the spin-up and spin-down regime is marked with a horizontal black dashed line. For comparison, the spin-down torque due to the APSW, Γ_W , for model $B_\star$ high is shown as a dashed red line.

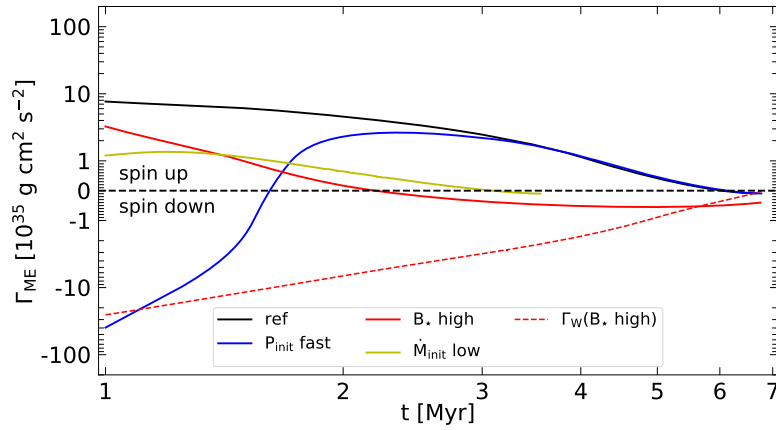


Figure 5.11.: Evolution of the torque due to magnetospheric ejections (MEs) for different configurations (see text). The transfer between the spin-up and spin-down regime is marked with a horizontal black dashed line. For comparison, the spin-down torque due to the APSW, Γ_W , for model $B_\star$ high is shown as a dashed red line.

5.6.2. Flat spectrum sources

Observed accretion rates of flat spectrum sources used in Fig. 5.3.

Table 5.3.: Stellar masses and accretion rates of flat-spectrum sources with $M_{\star} < 1 M_{\odot}$ used in Fig. 5.3. In addition, the method for determining the accretion rate ($\text{Method}_{M_{\star}}$) is given. If only one reference is given it contains stellar mass and the accretion rate.

Name	$M_{\star} [M_{\odot}]$	$\dot{M}_{\star} [10^{-8} M_{\odot} \text{ yr}^{-1}]$	$\text{Method}_{M_{\star}}$	Ref.
V846Ori	0.55	12.0	$\text{CaII}/\text{Pa}\beta/\text{Bry lines}$	Caratti o Garatti et al. [2012]
J05390536-0711052	0.55	13.0	$\text{CaII}/\text{Pa}\beta/\text{Bry lines}$	Caratti o Garatti et al. [2012]
IRAS 05379-0815	0.73	7.5	$\text{CaII}/\text{Pa}\beta/\text{Bry lines}$	Caratti o Garatti et al. [2012]
J05413033-0840177	0.56	10.0	$\text{CaII}/\text{Pa}\beta/\text{Bry lines}$	Caratti o Garatti et al. [2012]
J03290895+3122562	0.15	1.1	$\text{Pa}\beta/\text{Bry lines}$	Fiorellino et al. [2021]
J03292044+3118342	0.28	2.1	$\text{Pa}\beta/\text{Bry lines}$	Fiorellino et al. [2021]
IRAS 04385+2550	0.59	12.0 ^a	SED fit ^a	Güdel et al. [2007], Ribas et al. [2020]
DG Tau	0.50	25.0	U-band excess	Güdel et al. [2018], Alonso-Martínez et al. [2017]

^a SED fitting with an artificial neural network using the D'Alessio Irradiated Accretion Disk models [DIAD, see Ribas et al., 2020].

5.6.3. Observational data for stars after the disk phase

Stellar observational data used in Fig. 5.9 and Fig. 5.10 is shown in Table 5.4.

Table 5.4.: Stellar observational data used in Fig. 5.9 and Fig. 5.10. The respective uncertainties are given when available. Missing data is represented by "—".

Name	Age [Myr]	$M_{\star}$ [$M_{\odot}$]	$R_{\star}$ [$R_{\odot}$]	$P_{\star}$ [d]	B_I [G]	B_V [G]	B_{QU} [G]
AU Mic	23 ± 1 (1)	0.50 ± 0.03 (2)	0.75 ± 0.03 (2)	4.86 ± 0.01 (2)	2310 ± 200 (1)	88 (1)	2000 ± 200 (1)
WW Psa	27 ± 7 (3)	0.19 ± 0.04 (3)	0.70 ± 0.13 (3)	2.37 ± 0.01 (3)	2750 ± 250 (3)	—	—
TX Psa	27 ± 7 (3)	0.11 ± 0.03 (3)	0.52 ± 0.10 (3)	1.09 ± 0.01 (3)	2750 ± 250 (3)	—	—
V1124 Ori ^a	1.2 (4)	0.26 (4)	2.15 (4)	8.18 (4)	2090 ± 340 (4)	—	—
V1123 Ori ^a	0.8 (4)	0.35 (4)	2.59 (4)	7.64 (4)	2510 ± 410 (4)	—	—
V1568 Ori ^a	1.3 (4)	0.40 (4)	2.37 (4)	5.54 (4)	1420 ± 230 (4)	—	—
Upp Sco 55 ^b	5 (5)	0.10 (5)	0.52 (5)	2.90 ± 0.30 (5)	2300 ± 1300 (5)	—	—
TWA 9B	10 (6)	0.30 (6)	0.91 (6)	3.98 (6)	3100 ± 200 (6)	—	—
TWA 8A	11 ± 5 (7)	0.45 ± 0.10 (7)	0.80 ± 0.20 (7)	4.58 ± 0.01 (7)	6000 ± 500 (7)	720 (7)	—
GJ 490B	35-350 (8)	0.33 (9)	0.34 (9)	0.54 (9)	3200 ± 400 (9)	680 (9)	—
V2247 Oph ^a	1.0 (10)	0.35 (10)	2.00 (10)	3.40 (10)	2700 ± 100 (11)	90 (10)	—
TWA 7	17 ± 10 (12)	0.62 ± 0.03 (12)	0.80 ± 0.10 (12)	5.00 ± 0.01 (12)	2700 ± 200 (11)	50 (12)	—

^a Despite the young age of those stars, they are identified as NTTS or weak accretors and thus, included in the list. ^b $P_{\star}$ calculated from *vsini* data. References: (1) Kochukhov and Reiners [2020]; (2) Plavchan et al. [2020]; (3) Messina et al. [2017]; (4) Yang and Johns-Krull [2011]; (5) Reiners et al. [2009]; (6) Yang et al. [2008]; (7) Hill et al. [2019]; (8) Kluttsch et al. [2014]; (9) Phan-Bao et al. [2009]; (10) Donati et al. [2010]; (11) Laval et al. [2019], (12) Nicholson et al. [2021].

Part III.

Conclusions and outlook

6. Conclusions and outlook

The formation and evolution of the star-disk system involve a variety of complex physical processes that range over several spatial and temporal scales. To understand these processes, the star-disk system is approached by various theoretical models, and the results are compared to available observations. In the course of this thesis, it was demonstrated that a long-term evolution model, which couples the stellar and disk components, can provide new insights into a young star-disk system with respect to the evolution of the disk, the stellar structure, and the stellar spin:

- The stellar spin evolution affects the evolution of the protoplanetary disk. Disks surrounding fast-rotating stars are located closer to the star, are considerably hotter, and are subject to more and longer episodic outburst events compared to disks around slow-rotating stars.
- The evolution of protoplanetary disks and the rotation of the star are influenced by stellar metallicity. Disks in low-metallicity environments are heated more efficiently and have shorter lifetimes compared to solar metallicities. Furthermore, low-metallicity stars rotate faster compared to their solar-metallicity counterparts due to smaller stellar radii, different contraction rates, and shorter disk lifetimes.
- The amount of angular momentum removed from the star during the first one million years significantly depends on the interaction between the star and the disk. During this early evolutionary stage, the accretion history, the amount of energy added to the stellar interior, and the presence of winds have the strongest influence on the stellar evolution. In combination with different magnetic field strengths, the observed distribution of rotation periods ranging from $\lesssim 1$ day to ~ 10 days can be matched.
- For later evolutionary stages (Class II and later), the stellar spin evolution depends signific-

antly on the magnetic field strength for stellar masses $\gtrsim 0.3 M_{\odot}$. The amount of energy added to the star during the accretion process and the spin distribution at an age of ~ 1 Myr have no notable influence on the stellar spin evolution. The spin evolution of smaller stellar masses during the Class II and Class III phases, on the other hand, does depend on the initial spin distribution.

Due to the complexity of the star-disk system, certain stages and regions cannot be coupled self-consistently with the current numerical models presented in this work. For example, the model for the inner disk region is suited for simulations starting from the Class II stage. At earlier stages, large disk masses violate certain approximations, like the thin disk approximation [e.g., Ragossnig et al., 2020]. To solve this problem, the current model, which is restricted to the radial dimension, has to be extended in the vertical direction. The detailed modeling of the vertical direction would allow precise simulations of the accretion process along magnetic field lines, the evolution of a large-scale disk magnetic field, photo-evaporation, or disk winds to be included. Such an improved inner disk model could be combined with the stellar evolution code [e.g., MESA, Paxton et al., 2011, 2019a] at the inner boundary and with a multi-dimensional model of the outer disk [e.g., Vorobyov et al., 2019]. Thus, the results presented in this thesis motivate the advancement of the current numerical models to obtain a closed, self-consistent picture of the formation and evolution of a star-disk system even during the early stages. Additionally, the coupled model presented in this thesis could serve as a realistic boundary condition for different applications, such as chemical evolution models of protoplanetary disks [e.g., Rab et al., 2022].

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A. Kurzfassung

Um die Entstehung von Planeten und des Lebens selbst zu verstehen, muss die Evolution von protoplanetaren Scheiben, der Wiege junger Planeten, verstanden werden. Eine Vielzahl von Beobachtungen von protoplanetaren Scheiben und Planetensystemen wirft wichtige Fragen auf, die mit aktuellen theoretischen Modellen nicht beantwortet werden können. Um weitere Einblicke in die Evolution von jungen Stern-Scheiben-Systemen zu gewinnen, werden verschiedene numerische Methoden auf eine selbstkonsistente Weise gekoppelt, die eine kombinierte Behandlung der stellaren und Scheibenkomponente ermöglicht. Für jede Komponente des Stern-Scheiben-Systems wird ein bestimmtes Modell verwendet. Die stellare Komponente wird mit einem ausgefeilten, langfristigen Stellar-Evolutionsmodell modelliert. Um die inneren Scheibenregionen und die Ergebnisse der Stern-Scheiben-Interaktion zu simulieren, wird ein implizites, eindimensionales Scheibenmodell verwendet, das wichtige physikalische Mechanismen wie den Einfluss des stellaren Magnetfeldes auf die Scheibe einschließt. Eigenschaften der äußeren Scheibenregionen basieren auf einem detaillierten, mehrdimensionalen Modell. Das gekoppelte Stern-Scheiben-Modell zeigt, dass unterschiedliche stellare Rotationsperioden oder Metallizitäten die Temperaturstruktur, die dynamische Evolution und die Lebensdauer der Scheibe erheblich beeinflussen können. Gleichzeitig wird die stellare Drehungsevolution und die Verteilung von Rotationsperioden durch die Akkretionsrate und die Entwicklung der Scheibe beeinflusst. Die innere Scheibe spielt in dem Modell eine Schlüsselrolle, da diese Region die Stern-Scheiben-Interaktion definiert. Diese Ergebnisse zeigen, dass eine gekoppelte Behandlung der stellaren und Scheibenentwicklung wichtig ist, um diese komplexen Stern-Scheiben-Systeme zu verstehen. Eine weitere Verbesserung des kombinierten Stern-Scheiben-Modells, insbesondere der innersten Scheibenregion, wird zusätzliche Einblicke in bisher unerforschte Prozesse liefern.

B. Appendix

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