Photoevaporative Dispersal of Protoplanetary Disks around Evolving Intermediate-mass Stars

Masanobu Kunitomo1, Shigeru Ida2, Taku Takeuchi3,8, Olja Panić4,9, James M. Miley4,5,6, and Takeru K. Suzuki7

1 Department of Physics, School of Medicine, Kurume University, 67 Asahimachi, Kurume, Fukuoka 830-0011, Japan; kunitomo.masanobu@gmail.com
2 Earth-Life Science Institute, Tokyo Institute of Technology, 2-12-1 Ookayama, Meguro-ku, Tokyo 152-8551, Japan
3 Department of Earth and Planetary Sciences, Tokyo Institute of Technology, 2-12-1 Ookayama, Meguro-ku, Tokyo 152-8550, Japan
4 School of Physics and Astronomy, University of Leeds, Leeds LS2 9JT, UK
5 Joint ALMA Observatory, Alonso de Cordova 3107, Vitacura, Santiago, Chile
6 National Astronomical Observatory of Japan, Alonso de Cordova 3788, 61B Vitacura, Santiago, Chile
7 School of Arts & Sciences, The University of Tokyo, 3-8-1, Komaba, Meguro, Tokyo 153-8902, Japan

Received 2020 April 24; revised 2021 January 9; accepted 2021 January 11; published 2021 March 10

Abstract

We aim to understand the effect of stellar evolution on the evolution of protoplanetary disks. We focus in particular on the disk evolution around intermediate-mass (IM) stars, which evolve more rapidly than low-mass ones. We numerically solve the long-term evolution of disks around 0.5–5 $M_\odot$ stars considering viscous accretion and photoevaporation (PE) driven by stellar far-ultraviolet (FUV), extreme-ultraviolet (EUV), and X-ray emission. We also take stellar evolution into account and consider the time evolution of the PE rate. We find that the FUV, EUV, and X-ray luminosities of IM stars evolve by orders of magnitude within a few million years, along with the time evolution of stellar structure, stellar effective temperature, or accretion rate. Therefore, the PE rate also evolves with time by orders of magnitude, and we conclude that stellar evolution is crucial for the disk evolution around IM stars.

Unified Astronomy Thesaurus concepts: Protoplanetary disks (1300); Pre-main sequence stars (1290); Stellar evolution (1599); Planet formation (1241)

Supporting material: machine-readable table

1. Introduction

So far, the long-term evolution of protoplanetary disks has been mostly investigated by considering viscous accretion and photoevaporation (PE; e.g., Clarke et al. 2001; Alexander et al. 2006a; Gorti et al. 2009; Owen et al. 2010). The PE is a thermally driven disk wind from hot disk atmospheres due to the irradiation of high-energy photons (e.g., Hollenbach et al. 1994), that is, far-ultraviolet (FUV) photons (6–13.6 eV), extreme-ultraviolet (EUV) photons (13.6–100 eV), and X-rays (>100 eV). Most of the previous works, however, do not consider the time evolution of the luminosity of high-energy photons or include all PE mechanisms.

Gorti et al. (2009) investigated the long-term disk evolution considering all PE mechanisms from central stars for the first time. However, they did not consider the temporal evolution of the luminosity of the EUV and X-rays. Moreover, the contribution of each mechanism was not clearly shown. Since the PE rate depends on the UV and X-ray luminosities, it is crucial for disk evolutionary models to adopt realistic models of those luminosities. We also note that Alexander et al. (2004) claimed that FUV and EUV from the stellar photosphere are sensitive to the absorption in the stellar atmosphere.

In this paper, we aim to (i) investigate the long-term disk evolution (i.e., not the dynamical evolution within several Kepler timescales but the disk evolution for millions of years) with realistic FUV, EUV, and X-ray luminosity, considering stellar evolution and the absorption in the stellar atmosphere, and (ii) clarify which mechanism of PE plays a dominant role in dispersing disks.

We focus in particular on the influence of stellar evolution. As we will describe in detail in Section 2, young stars emit UV photons and X-rays through three mechanisms: magnetic activity, accretion shock, and photospheric radiation. Since the magnetic activity originates from the convective motion in the stellar interior, the evolution of the stellar internal structure (i.e., the thickness of the convective envelope, $M_{\text{conv}}$) is important (see Section 2.5). The $M_{\text{conv}}$ value of young stars decreases with time, and a radiative core is developed instead. Moreover, the spectra of photospheric radiation depend on the stellar effective temperature, $T_{\text{eff}}$, and the stellar intrinsic bolometric luminosity, and $c$ is a dimensionless factor that depends on the polytropic index (e.g., $c = 3/7$ for fully convective stars and $3/4$ for radiative stars). Given the weak dependence of the $L_{\ast}$ of pre-main-sequence (pre-MS) stars on $M_{\ast}$ (i.e., roughly $L_{\ast} \propto M_{\ast}^2$ for pre-MS stars), the $\tau_{\text{KH}}$ of higher-mass stars is shorter; thus, the $T_{\text{KH}}$ and $M_{\text{conv}}$ of higher-mass stars evolve more rapidly. Therefore, the PE rate is also expected to change with time, in particular around higher-mass stars.

We note that recent infrared (IR) observations have revealed that the disk evolution around intermediate-mass (IM) stars is different from low-mass stars in the following three respects: the near-IR dust disk lifetime of IM stars is shorter than that of low-mass stars (Hillenbrand et al. 1992; Hernández et al. 2005;
Carpenter et al. 2006; Yasui et al. 2014; Ribas et al. 2015), the \( \text{H}_2 \) gas disk lifetime is also shorter (Kennedy & Kenyon 2009; Yasui et al. 2014), and there is a substantial difference between near- and mid-IR dust disk lifetimes, unlike in low-mass stars (Yasui et al. 2014). Therefore, disk evolution depends on stellar mass. Following the previous studies above, we define IM as stars of mass above 2–5 \( M_\odot \).\(^{10}\) There is also a difference in planetary architectures between low-mass and IM stars (i.e., a lack of close-in planets around IM stars), which may result from the different disk evolution (e.g., Burkert & Ida 2007; Sato et al. 2008; Currie 2009; Kunitomo et al. 2011). To understand these puzzles, as a first step, we investigate the effect of stellar evolution on disk evolution in this paper.

This paper is organized as follows. First, we describe our model of the luminosity of the high-energy photons considering stellar evolution. In Section 3, we describe our physical models of the PE and computation method for simulating the disk evolution. In Section 4, we investigate how the disk evolution is affected by stellar evolution. In Section 5, we describe the caveats of our model. The results are summarized in Section 6.

2. Stellar Evolution

In this section, we first describe the computation methods of the stellar evolution (Section 2.1) and stellar atmosphere (Section 2.2). Using these results and the observational results, we model the evolution of stellar FUV, EUV, and X-ray luminosity (Sections 2.3–2.5).

2.1. Stellar Evolution Calculation

We simulate the stellar evolution using the code MESA (ver. 2258, Paxton et al. 2011; see also Kunitomo et al. 2011, for details). Figure 1 shows the evolutionary tracks of 0.5–5 \( M_\odot \) stars in the Hertzsprung–Russell (H-R) diagram. We assume solar metallicity. We adopt the birthline of Stahler & Palla (2004) in the H-R diagram as an initial condition. This corresponds to the standard scenario of star formation (see Section 5.4). We note that the luminosity of 0.8–2 \( M_\odot \) stars on the birthline is almost the same because of the short thermal timescale, whereas that of >2 \( M_\odot \) stars increases with \( M_\star \) because of deuterium burning (Stahler 1988).

Here we briefly introduce the basic nature of the stellar pre-MS evolution (see, e.g., Kippenhahn & Weigert 1990; Stahler & Palla 2004). From the birthline, young low-mass stars evolve along their Hayashi track, which is almost vertical in the H-R diagram due to the strong temperature dependence of the \( H^\alpha \) opacity (Hayashi 1961). On the Hayashi track, the stars are fully convective and shrink due to the radiative energy loss (i.e., the KH contraction). Since the energy loss results in the increasing internal temperature with time from the virial theorem, and the stellar internal opacity is anticorrelated with temperature (i.e., the Kramers law), the stellar internal opacity decreases with time. Then a radiative core is developed, and a star leaves the Hayashi track. We note that high-mass (>3 \( M_\odot \)) stars are hot enough to have a radiative core from the beginning. Stars evolve on the horizontal Henyey track, and the stellar effective temperature, \( T_\text{eff} \), increases with time. The IM pre-MS stars with a high \( T_\text{eff} \) surrounded by a disk are called Herbig Ae/Be stars (Herbig 1960; van den Ancker et al. 1997). The evolution of stellar structure and \( T_\text{eff} \) is a key ingredient in this work (see Sections 2.2 and 2.5).

For simplicity, we do not consider the \( M_\star \) evolution due to the mass accretion from the disk inner edge to the star or the mass loss via stellar winds (e.g., Suzuki et al. 2013).

2.2. Stellar Spectra and Atmospheric Model

The UV photons are directly emitted from the photosphere of hot IM stars. Those photons, however, are substantially absorbed in the stellar atmosphere; therefore, the stellar spectra deviate from the blackbody (Alexander et al. 2004). Here we quantify the extent of the absorption of UV photons by using a stellar atmospheric model. In this subsection, we describe the method and results.

We used version 13.04 of the Cloudy code, last described by Ferland et al. (2013), to obtain the spectra. We note that the stellar evolution simulations in MESA do not provide stellar spectra. Therefore, we need to independently calculate the stellar absorption using the Cloudy code. We adopt the Atlas grids of Castelli & Kurucz (2003), which are available in \( T_\text{eff} = 3500–50,000 \text{ K} \), in the case of solar metallicity. We assume \( R_\star = 1 R_\odot \) and the Stefan–Boltzmann law gives the bolometric luminosity \( L_\star \). We adopt the stellar surface gravity \( g = 0.33 \, g_\odot \), where \( g_\odot = 10^{4.44} \text{ cm s}^{-2} \) is the surface gravity of the Sun. We note that the results below are not sensitive to the assumed \( g \) value (see Appendix A) or \( R_\star \).

Figure 1 (a) shows the stellar spectra as a function of wavelength \( \lambda \) in the cases of \( T_\text{eff} = 20,776, 15,097, 10,128, 7951, \) and 3587 K with and without the absorption in the stellar atmosphere. We note that the latter (i.e., the blackbody spectra) is \( \nu \nu B_\nu \), where \( \nu \) is the frequency and \( B_\nu \) is the Planck function. The spectra exhibit strong absorption at the Lyman break and in the EUV range (>13.6 eV); therefore, we should not use the blackbody for \( \Phi_{\text{EUV,ph}} \), as claimed in Alexander et al. (2004). We also find the absorption in the FUV (not only the Ly\( \alpha \) absorption at 1216 Å) in the case with the low \( T_\text{eff} \). We note that we confirmed that the spectrum of a 15,097 K star is almost the same as Figure 1 of Alexander et al. (2004).

\(^{10}\) We note that Hernández et al. (2005) and Ribas et al. (2015) defined >2 \( M_\odot \) stars as Herbig Ae/Be and high-mass stars, respectively, whereas Yasui et al. (2014) defined 1.5–7 \( M_\odot \) stars as IM stars.
We simply assume 50 eV as the average EUV photon energy and $T_{\text{eff}} = 20,776$ K, respectively (i.e., the latter is the blackbody spectrum), in the case of $g = 0.33$ g$_{\odot}$. (Top panel) Stellar spectra in the cases of $T_{\text{eff}} = 20,776, 15,097, 10,128, 7,971$, and 3,587 K (top to bottom). The two vertical lines indicate the wavelengths at 6 and 13.6 eV. In practice, with the stellar spectra, we calculate $\Phi_{\text{EUV}}$, the fraction of the photospheric EUV photon luminosity, of 3–5 $M_{\odot}$ stars (bottom panel).

We set $f_{\text{EUV}} = 0$ where $T_{\text{eff}} < 5 \times 10^3$ K and $f_{\text{EUV}} = 0$ where $< 3.5 \times 10^3$ K. Together with the evolution of $T_{\text{eff}}$ and $L_*$, we obtain the evolution of $\Phi_{\text{EUV}}$ and $L_{\text{EUV,ph}}$.

Figure 2(b) shows the results of $f_{\text{EUV}}$ and $f_{\text{EUV}}$ as a function of $T_{\text{eff}}$. Using the polynomial fitting of Numpy, we obtained the following fitting formulae:

$$\log f_{\text{EUV}} = \sum_{i=0}^{5} a_i (\log T_{\text{eff}})^i$$

(3)

in 5000–50,000 K, where $a_5 = -95.238145$, $a_4 = 1998.2116$, $a_3 = -16728.880$, $a_2 = 69832.410$, $a_1 = -145282.56$, and $a_0 = 120432.67$, and

$$\log f_{\text{FUV}} = \sum_{i=0}^{6} b_i (\log T_{\text{eff}})^i$$

(4)

in the range of $T_{\text{eff}} = 3500–50,000$ K, where $b_6 = 177.14306$, $b_5 = -4452.9922$, $b_4 = 46546.370$, $b_3 = -258939.74$, $b_2 = 808484.59$, $b_1 = -1343172.0$, and $b_0 = 927492.51$.

**Figure 3.** Time evolution of the photospheric FUV luminosity, $L_{\text{FUV,ph}}$, of 1.5–5 $M_{\odot}$ stars (top panel) and the photospheric EUV photon luminosity, $\Phi_{\text{EUV}}$, of 3–5 $M_{\odot}$ stars (bottom panel).

We set $f_{\text{EUV}} = 0$ where $T_{\text{eff}} < 5 \times 10^3$ K and $f_{\text{FUV}} = 0$ where $< 3.5 \times 10^3$ K. Together with the evolution of $T_{\text{eff}}$ and $L_*$, we obtain the evolution of $\Phi_{\text{EUV}}$ and $L_{\text{EUV,ph}}$.

Figure 3 shows the results of the $L_{\text{FUV,ph}}$ evolution of 1.5–5 $M_{\odot}$ stars and the $\Phi_{\text{EUV}}$ evolution of 3–5 $M_{\odot}$. We have combined $T_{\text{eff}}(t)$ and $L_*(t)$ from the stellar evolution simulations (see Section 2.1) and the $f_{\text{EUV}}$ and $f_{\text{FUV}}$ relations (see the solid lines in Figure 2(b)). We find that they abruptly increase by orders of magnitude. Equation (4) shows that $T_{\text{eff}} = 7342$ K is a characteristic temperature; above this temperature, $f_{\text{FUV}}$ exceeds $10^{-2}$. We will show the influence of this rapid increase in the disk evolution in Section 4.

Gorti et al. (2009) investigated disk evolution including $L_{\text{FUV,ph}}$ and $\Phi_{\text{EUV}}$. They adopted the values of MS stars from Parravano et al. (2003): $L_{\text{FUV,ph}} = 3.8 \times 10^{33}$, $2.9 \times 10^{34}$, $1.1 \times 10^{35}$, $4.3 \times 10^{35}$, and $3.3 \times 10^{36}$ erg s$^{-1}$ for 2, 2.5, 3, 4, and 5 $M_{\odot}$ stars, whereas $\Phi_{\text{EUV}} = 2.4 \times 10^{42}$ s$^{-1}$ for a 5 $M_{\odot}$ star (see also Armitage 2000). These values agree well with the values for MS stars in our model (see Figure 3). We also note that Parravano et al. (2003) indirectly verified their models by comparing them with observed interstellar FUV radiation fields.

**2.3. Stellar FUV Luminosity**

Using the results in Sections 2.1 and 2.2 and the observational results, we model the stellar FUV luminosity $L_{\text{FUV}}$. We adopt the same model of $L_{\text{FUV}}$ as Gorti et al. (2009) and assume that $L_{\text{FUV}}$ is the sum of three components,

$$L_{\text{FUV}} = L_{\text{FUV,acc}} + L_{\text{FUV,ph}} + L_{\text{FUV,chr}},$$

(5)

[11] In Parravano et al. (2003), $L_{\text{EUV,ph}}$ of <1.8 $M_{\odot}$ stars and $\Phi_{\text{EUV}}$ of <5 $M_{\odot}$ are not available (see their Table 1).
where \( L_{\text{FUV,acc}} \) originates from the accretion process, and \( L_{\text{FUV,chr}} \) from the stellar chromosphere.

We assume that 4% of the gravitational energy of accreting materials (=\( GM_\ast M_{\text{acc}}/R_\ast \)) is emitted as FUV photons (Calvet & Gullbring 1998); therefore,

\[
L_{\text{FUV,acc}} = 10^{-2} L_\odot \left( \frac{M_\ast}{M_\odot} \right) \left( \frac{R_\ast}{R_\odot} \right)^{-1} \left( \frac{M_{\text{acc}}}{10^{-3} M_\odot \text{yr}^{-1}} \right),
\]

where \( M_{\text{acc}} \) is the mass accretion rate onto the star. Observations also suggest that the \( L_{\text{FUV}} \) of classical T Tauri stars is proportional to \( M_{\text{acc}} \) (e.g., Ingleby et al. 2011). We also adopt the \( L_{\text{FUV,chr}} \) model as

\[
L_{\text{FUV,chr}} = 10^{-3.3} L_\ast
\]

(see Section 3.1 of Alexander et al. 2014, and references therein). We adopt the \( L_{\text{FUV,ph}} \) model in Section 2.2. Because \( L_{\text{FUV,acc}} \) depends on the initial condition and disk evolution, we will show our \( L_{\text{FUV}} \) models in Section 4. We note that all of the components (i.e., \( L_{\text{FUV,acc}}, L_{\text{FUV,ph}}, \) and \( L_{\text{FUV,chr}} \)) are important (see Figure 9).

### 2.4. Stellar EUV Luminosity

The origin of EUV photons from pre-MS stars remains unclear because interstellar hydrogen atoms easily absorb EUV and it is difficult to observationally measure their \( \Phi_{\text{EUV}} \). In this paper, we consider EUV from the stellar corona and photosphere and assume that \( \Phi_{\text{EUV}} \) is the sum of the \( \Phi_{\text{EUV,cor}} \) and \( \Phi_{\text{EUV,ph}} \), respectively as

\[
\Phi_{\text{EUV}} = \Phi_{\text{EUV,cor}} + \Phi_{\text{EUV,ph}}.
\]

We simply adopt \( \Phi_{\text{EUV,cor}} = 10^{31} \text{s}^{-1} \) in this paper (see Section 5.4). We adopt the \( \Phi_{\text{EUV,ph}} \) model in Section 2.2.

### 2.5. Stellar X-Ray Luminosity

Stellar X-rays are emitted from the hot corona by magnetic activity. Although the accretion onto the star may also contribute to the X-ray luminosity \( L_X \) (see, e.g., Kastner et al. 2002, 2004), in this paper, we neglect this possibility for simplicity (see Section 5.4). We model the evolution of \( L_X \) based on the following two observed features.

First, observations have suggested that \( L_X \) depends on the stellar Rossby number. The \( L_X \) of rapid rotators is known to be a function of \( L_x \); that is, the fractional X-ray luminosity \( (R_X = L_X/L_* \) is constant at around \( 10^{-3} \) (e.g., Vilhu & Rucinski 1983). Most T Tauri stars rotate rapidly and have this relation (so-called “saturation”; Flaccomio et al. 2003; Preibisch et al. 2005; Telleschi et al. 2007). On the other hand, the \( R_X \) of IM stars or slow rotators is more complex. Since the dynamo efficiency depends on both the rotation period and the depth of the convective zone, Mangeney & Praderie (1984) and Nayes et al. (1984) introduced the Rossby number, which is the ratio of the rotational period to the convective turnover timescale \( (\tau = P_{\text{rot}}/\tau_{\text{conv}}) \), as an indicator of the X-ray activity. Wright et al. (2011) combined the observed data of both saturated and unsaturated stars and derived the following empirical formula: \( R_X = \min[10^{-3.13}, 5.3 \times 10^{-6} \tau_6^{-2.7}] \). The threshold value of the saturation is \( \tau_{\text{sat}} = 0.16 \).

Second, the \( L_X \) of pre-MS IM stars depends strongly on their age. Hamaguchi et al. (2005) and Huenemoerder et al. (2009) reported that young IM stars on or leaving their Hayashi track have a high \( R_X \) (~\( 10^{-3} - 10^{-4} \)). On the other hand, the older counterparts, Herbig Ae/Be stars, have smaller values of \( R_X \) ranging from \( 10^{-5} \) to \( 10^{-7} \) according to observations (Zinnecker & Preibisch 1994; Hamaguchi et al. 2005; Hamidouche et al. 2008; Stelzer et al. 2009). The strong dependence of the \( L_X \) of IM stars on age (or \( T_{\text{eff}} \)) is shown in Flaccomio et al. (2003), Hamaguchi et al. (2005), and Gregory et al. (2016). Flaccomio et al. (2003, see their Figure 9) showed that the median value of \( L_X \) of 2–3 \( M_\odot \) stars decreases at around a few million years by orders of magnitude. We note that this is consistent with recent observations by Villebrun et al. (2019), who have suggested that young IM stars possess magnetic fields, whereas most (90%–95%) Herbig Ae/Be stars do not. Therefore, we assume that the evolution of the \( L_X \) of IM stars can also be modeled with the Rossby number; the increase of the Ro number with time results in the \( R_X \) decrease. Although the physical origin of the weak X-ray emission of Herbig Ae/Be stars is still under debate, we impose a lower limit to \( R_X = 10^{-7} \) even if \( R_X > 4.35 \equiv R_{\text{floor}} \).

Considering the above observational constraints, we model the \( L_X \) evolution as follows:

\[
L_X = \text{max} \left[ \min(10^{-3.13}, 5.3 \times 10^{-6} \tau_6^{-2.7}), 10^{-7} \right] L_\ast.
\]

We note that the choice of the lower limit of \( R_X = 10^{-7} \) has little impact on disk evolution.

To compute \( L_X \) with Equation (9), we need the evolution of \( \tau_{\text{conv}} \) and \( P_{\text{rot}} \). From the mixing-length theory (Cox & Giuli 1968), \( \tau_{\text{conv}} \) in the stellar interior can be estimated as

\[
\tau_{\text{conv}} = \left[ \frac{M_{\text{conv}}(R_* - R_{\text{conv}})^2}{3L_*} \right]^{1/3},
\]

where \( M_{\text{conv}} \) is the mass in the convective envelope and \( R_{\text{conv}} \) is the radius at the base of the envelope (Zahn 1977; Rasio et al. 1996; Villaver & Livio 2009).

The \( P_{\text{rot}} \) value of stars younger than several million years (corresponding to the disk lifetime) ranges from 1 to 10 days and is almost constant with time, probably due to the star–disk locking (Rebull et al. 2004; Bouvier 2008; Gallet & Bouvier 2013). Therefore, we set the fiducial value of \( P_{\text{rot}} \) to be 3 days and investigate the influence of its variation in Section 5.2.

Figures 4(a) and (b) show the time evolution of \( L_\ast \) and \( R_X \), respectively. Figure 5 shows our model of the \( L_X \) evolution combining Equation (9) and Figure 4. One might be skeptical about our prescription of \( L_X \). We compare our model of \( L_X \) over time with observational data in Gregory et al. (2016, see their Figure C2). Figure 6 shows that our model of the \( L_X \) evolution is in good agreement with the data in Gregory et al. (2016). The observed \( L_X \) data show that 0.5–1 \( M_\odot \) stars have a gradual decrease for \( \sim 10 \) Myr, whereas \( \geq 1.5 \) \( M_\odot \) stars have a decrease by orders of magnitude. Our model captures such features, and the \( L_X \) values and the timing of decrease are also reproduced. We have also confirmed that our model is consistent with Flaccomio et al. (2003), Güdel (2004), and the Sun.13 We admit, however, that Figure 6 shows that the \( L_X \) values of 1–1.5 \( M_\odot \) in our model are several times larger

---

12 We note that we should be careful with the contribution by an unresolved binary star, but Hamidouche et al. (2008) ruled out this possibility with an 80% confidence level.

13 The Sun has \( M_{\text{conv}} = 0.025 M_\odot \) and \( R_{\text{conv}} = 0.713 R_\odot \) (Bahcall et al. 2005) and therefore \( \tau_{\text{conv,0}} = 13.9 \) days. Combining this with \( P_{\text{rot}} \approx 26.9 \) days, \( R_{\text{sat}} \approx 194.9 \). Equation (9) with \( R_{\text{sat}} \approx 194.9 \) gives \( R_X = 5.9 \times 10^{-7} \), which is consistent with the observed solar value, \( R_X = 10^{-7} - 10^{-6} \) (Judge et al. 2003).
than the median value of the observed \( L_X \). Moreover, the observed \( K_X \) has a large scatter (\sim 1 \text{ dex}; e.g., Preibisch et al. 2005). We will investigate the impact of the larger/smaller value of \( L_X \) in Section 5.2.

Our stellar evolutionary models described in Section 2 are provided in Table 1.

### 3. Physical Model and Computation Method of Disk Evolution

We simulate the time evolution of protoplanetary disks including the effects of viscous accretion and time-dependent PE (Section 3.1). We adopt the PE models from the literature (Section 3.2), considering stellar evolution on the pre-MS (see Section 2). The criterion for the disk dispersal is described in Section 3.3. The numerical method and computational settings are summarized in Section 3.4.

#### 3.1. Basic Equations

We solve the one-dimensional diffusion equation for the surface density profile (e.g., Lynden-Bell & Pringle 1974; Clarke et al. 2001),

\[
\frac{\partial \Sigma}{\partial t} = \frac{3}{R} \frac{\partial}{\partial R} \left[ \sqrt{R} \frac{\partial}{\partial R} \left( \nu_{\text{vis}} \Sigma \sqrt{R} \right) \right] - \Sigma_{\text{PE}}(R, t),
\]

where \( \Sigma \) is the surface density, \( t \) is the time, \( R \) is the distance from the central star, \( \nu_{\text{vis}} \) is the viscosity, and \( \Sigma_{\text{PE}} \) is the PE rate, under the cylindrical coordinates \((R, \phi, z)\).

We adopt the viscosity model of Shakura & Sunyaev (1973), \( \nu_{\text{vis}} = (2/3) c_s^2 / \Omega \), where \( c_s \) is the sound speed at the disk midplane and \( \Omega \) is the angular velocity. We neglect the disk self-gravity and pressure gradient force and adopt \( \Omega = \sqrt{G M / R^3} \), where \( G \) is the gravitational constant.

For the midplane temperature \( T_{\text{mid}} \), we consider both the viscous heating and stellar irradiation, following Kunitomo et al. (2020, see their Section 2.2), which is based on Suzuki et al. (2016). Since in this paper, we consider the \( L_\star \) evolution (see Section 2), the disk temperature in the entire region evolves with time because both viscous heating and stellar irradiation change with time. We define \( c_s^2 = k_B T_{\text{mid}} / (\mu m_H) \), where \( k_B \) is the Boltzmann constant, \( \mu = 2.34 \) is the mean molecular weight, and \( m_H \) is the atomic mass unit.

#### 3.2. PE Models

In this paper, we consider the PE driven by the irradiation from a central star (so-called “internal PE”) and do not consider the external irradiation by a nearby massive star (e.g., Adams et al. 2004; Haworth & Clarke 2019).

So far, a number of studies have been performed on the internal PE (e.g., Hollenbach et al. 1994; Font et al. 2004; Ercolano et al. 2008; Gorti & Hollenbach 2009; Tanaka et al. 2013; Komaki et al. 2020). We also refer the reader to recent reviews such as Alexander et al. (2014), Gorti et al. (2016), and Ercolano & Pascucci (2017). We consider the PE driven by FUV, EUV, and X-rays, and we adopt their mass-loss rates from the literature. We assume that the dominant heating source among the three at the wind launching region determines the mass-loss rate \( \dot{\Sigma}_{\text{PE}} \), and therefore

\[
\dot{\Sigma}_{\text{PE}}(R, t) = \max [\dot{\Sigma}_{\text{FUV}}(R, t), \dot{\Sigma}_{\text{EUV}}(R, t), \dot{\Sigma}_X(R, t)],
\]

where \( \dot{\Sigma}_{\text{FUV}}, \dot{\Sigma}_{\text{EUV}}, \) and \( \dot{\Sigma}_X \) are the PE rate driven by FUV, EUV, and X-rays, respectively. We note that one might think that \( \dot{\Sigma}_{\text{PE}} \) can be proportional to the total energy deposited in the disk atmosphere, and therefore \( \dot{\Sigma}_{\text{PE}} = \dot{\Sigma}_{\text{FUV}} + \dot{\Sigma}_{\text{EUV}} + \dot{\Sigma}_X \).

We have confirmed that the two expressions of \( \dot{\Sigma}_{\text{PE}} \) make little difference in the results (\(<8\% \) in disk lifetime) because one process among the three almost always dominates.

The PE rate has two regimes: one is for primordial disks, and the other is for disks with an inner hole. In the latter, the outer disk is directly irradiated, and therefore the PE profile is changed (so-called “direct PE”). We consider both regimes.

We adopt the same \( \dot{\Sigma}_{\text{EUV}} \) model as in Kunitomo et al. (2020), \( \dot{\Sigma}_{\text{EUV}} \) for primordial disks in Alexander & Armitage (2007) and that for the direct PE in Alexander et al. (2006b).
The total mass-loss rates for the EUV PE in both regimes are

\[
M_{\text{EUV,p}} = 1.6 \times 10^{-10} M_\odot \text{ yr}^{-1} \left( \frac{\Phi_{\text{EUV}}}{10^{41} \text{s}^{-1}} \right)^{1/2} \left( \frac{M_*}{1 M_\odot} \right)^{1/2} \tag{13}
\]

and

\[
M_{\text{EUV,d}} = 1.3 \times 10^{-9} M_\odot \text{ yr}^{-1} \left( \frac{\Phi_{\text{EUV}}}{10^{41} \text{s}^{-1}} \right)^{1/2} \left( \frac{R_{\text{hole,EUV}}}{3 \text{ au}} \right)^{1/2}, \tag{14}
\]

where \(\Phi_{\text{EUV}}\) is the EUV photon luminosity and \(R_{\text{hole,EUV}}\) is the inner hole size for the EUV. We assume the aspect ratio \(h/R = 0.05\) in Equation (14) (see Alexander et al. 2006b), where \(h = \sqrt{2} G M_* / \Omega\) is the gas scale height.\(^ {14}\) The \(\Sigma_{\text{EUV}}\) profile of primordial disks has a peak at \(\approx 1\) au \((M_\odot/M_*\odot)\). We refer the reader to Alexander & Armitage (2007) for the full formula of \(\Sigma_{\text{EUV}}\) (see also Figure 7).

As for the X-ray PE rate, the prescription in Owen et al. (2012) has been widely used. In the case of a 1 \(M_\odot\) star with \(L_X = 10^{30} \text{ erg s}^{-1}\), \(\Sigma_X\) has a peak value \((\approx \Sigma_{X,0} = 5.1 \times 10^{-12} \text{ g cm}^{-2})\) at 2.5 au \((M_\odot/M_*\odot)\)\(^ {15}\), decreases with radius as \(R^{-2}\), and has a sharp cutoff at several tens of astronomical units. The cutoff is, however, not seen in the recent study by Picogna et al. (2019, see their Figure 5). For the primordial disks, we adopt

\[
\Sigma_X = \Sigma_{X,0} \left( \frac{L_X}{10^{30} \text{ erg s}^{-1}} \right) \left( \frac{R}{2.5 \text{ au}} \right)^{-2}, \tag{15}
\]

outside 2.5 au \((M_\odot/M_*\odot)\). In the inner region, the disk gas is gravitationally bound to the disk and does not flow out (i.e., \(\Sigma_X = 0\)). We note that we neglect the weak dependence on stellar mass \((\propto M_*^{-0.068})\) in the original \(\Sigma_X\) profile in Owen et al. (2012). We note that the X-ray PE rate in Owen et al. (2012) has recently been called into question; the radiation-hydrodynamic (RHD) simulations with self-consistent thermochemistry by Wang & Goodman (2017) and Nakatani et al. (2018a) disagree with the results in Owen et al. (2012), and therefore Equation (15) may overestimate the X-ray PE rate (see also a pioneering study by Gorti & Hollenbach 2009). We will discuss this issue in Section 5.4.

For the \(\Sigma_X\) of the direct PE, we adopt the model in Owen et al. (2012, see their Appendix B), which peaks at the inner edge of the outer disk. The total mass-loss rate is

\[
M_{X,d} = 4.8 \times 10^{-9} M_\odot \text{ yr}^{-1} \left( \frac{L_X}{10^{30} \text{ erg s}^{-1}} \right)^{1.14} \left( \frac{M_*}{M_\odot} \right)^{-0.148}, \tag{16}
\]

where the subscript “d” stands for the direct PE.

We need to define the inner hole sizes for the direct PE for the EUV and X-rays. We also modify \(\Sigma_{\text{EUV}}\) and \(\Sigma_X\) for the direct PE to avoid numerical problems using “smoothing functions.” We refer readers to Kunitomo et al. (2020, see their Section 2.4) for the full details of these prescriptions.

As for the \(\Sigma_{\text{FUV}}\) model, no formula is available to date in the literature. We construct the \(\Sigma_{\text{FUV}}\) model as a function of \(R\) and the stellar FUV luminosity \(L_{\text{FUV}}\), based on the results in Gorti & Hollenbach (2009) and Wang & Goodman (2017). The latter performed RHD simulations, whereas the former conducted calculations using a hydrostatic model.

Gorti & Hollenbach (2009) investigated the dependence of \(\Sigma_{\text{FUV}}\) on \(L_{\text{FUV}}\) (see models F10, S, and F0.1 in their Figure 4) around a 1 \(M_\odot\) star. The \(\Sigma_{\text{FUV}}\) value beyond 4 au changes by about 1 order of magnitude by varying \(L_{\text{FUV}}\) by an order of magnitude. Therefore, we assume that \(\Sigma_{\text{FUV}} \propto L_{\text{FUV}}\) and the FUV PE mass loss occurs beyond 4 au \((M_\odot/M_*\odot)\). We note that 4 au corresponds to the critical radius for \(\approx 2000\) K gas around a 1 \(M_\odot\) star (Liffman 2003). The gas heated by FUV is much

\[^{14}\text{We note that the factor of } \sqrt{2} \text{ is sometimes not included. We include it following Kunitomo et al. (2020).}\]

\[^{15}\text{The peak of the X-ray PE (at 2.5 au for a 1 \(M_\odot\) star) is farther than that of the EUV PE (at 1 au) because the X-ray PE is launched from the atomic layer} (\approx 3000–5000 \text{ K}), \text{ whereas the EUV PE is from the } 10^7 \text{ K layer (Alexander et al. 2014).}\]
### Table 1
Stellar Evolutionary Models

| \(M_\ast\) \(\text{[}M_\odot\text{]}\) | \(\log t\) \(\text{[}\text{yr}\text{]}\) | \(R_\ast\) \(\text{[}R_\odot\text{]}\) | \(L_\ast\) \(\text{[}L_\odot\text{]}\) | \(T_{\text{eff}}\) \(\text{[K]}\) | \(M_{\text{conv}}\) \(\text{[}M_\odot\text{]}\) | \(R_{\text{conv}}\) \(\text{[}R_\odot\text{]}\) | \(\tau_{\text{conv}}\) \(\text{[days]}\) | \(L_X\) \(\text{[erg s}^{-1}\text{]}\) | \(\Phi_{\text{EUV,ph}}\) \(\text{[1 s}^{-1}\text{]}\) | \(L_{\text{FUV,ph}}\) \(\text{[erg s}^{-1}\text{]}\) |
|---|---|---|---|---|---|---|---|---|---|---|
| 0.5 | 0.00 | 4.537E+00 | 4.883E+00 | 4.032E+03 | 5.000E+00 | 5.113E-02 | 1.387E+00 | 1.391E+31 | 0.000E+00 | 1.162E+26 |
| 0.5 | 1.00 | 4.536E+00 | 4.686E+00 | 3.991E+03 | 4.998E+00 | 5.113E-02 | 1.387E+02 | 1.335E+31 | 0.000E+00 | 8.965E+25 |
| 0.5 | 2.00 | 4.536E+00 | 4.686E+00 | 3.991E+03 | 4.998E+00 | 5.113E-02 | 1.387E+02 | 1.334E+31 | 0.000E+00 | 8.964E+25 |

**Note.** Evolutionary models of young \(0.5-5\ M_\odot\) stars. A portion is shown here for guidance regarding its form and content. (This table is available in its entirety in machine-readable form.)
cooler than that by EUV, which is \( \approx 10^4 \) K (see also Nakatani et al. 2018b).

The \( \Sigma_{\text{FUV}} \) profile in Gorti & Hollenbach (2009) is a complex function of \( R \) (see their Figure 4), whereas Wang & Goodman (2017) claimed that \( 2\pi R^2 \Sigma_{\text{FUV}} \) is almost constant (see their Figure 8). Wang & Goodman (2017, see their Section 5.2) confirmed that the difference arises from how to estimate the mass-loss rate; the sonic point is different between the hydrodynamic simulations in Wang & Goodman (2017) and the hydrostatic models in Gorti & Hollenbach (2009). The \( \Sigma_{\text{FUV}} \propto R^{-2} \) profile seems energetically reasonable. Therefore, from the results in Wang & Goodman (2017), we assume that \( \Sigma_{\text{FUV}} \propto R^{-2} \) and \( \Sigma_{\text{FUV}} = 10^{-12} \) g cm\(^{-2}\) (\( \equiv \Sigma_{\text{FUV},0} \)) at 4 au around a 1 \( M_\odot \) star. As a result, we adopt the following \( \Sigma_{\text{FUV}} \) profile: in the outer disk beyond 4 au\((M_*/M_\odot) \),

\[
\Sigma_{\text{FUV}} = \Sigma_{\text{FUV},0} \left( \frac{L_{\text{FUV}}}{10^{31.7} \text{ erg s}^{-1}} \right) \left( \frac{R}{4 \text{ au}} \right)^{-2},
\]

and in the inner disk \((R < 4\text{ au}) (M_*/M_\odot)\), \( \Sigma_{\text{FUV}} = 0 \).

Figure 7 shows examples of the \( \Sigma_{\text{PE}} \) profiles of two cases; one is around a 3 \( M_\odot \) star with \( L_{\text{FUV}} = 10^{34} \) erg s\(^{-1}\), \( \Phi_{\text{EUV}} = 10^{31} \) erg s\(^{-1}\), and \( L_\text{X} = 3 \times 10^{38} \text{ erg s}^{-1} \), and the other is around a 1 \( M_\odot \) star with \( L_{\text{FUV}} = 10^{32} \) erg s\(^{-1}\), \( \Phi_{\text{EUV}} = 10^{31} \) erg s\(^{-1}\), and \( L_\text{X} = 3 \times 10^{31} \text{ erg s}^{-1} \). As described in Section 2, these luminosities evolve with time, and therefore the PE rate varies with time.

3.3. Disk Dispersal Criterion

In this study, we define the time when the disk mass, \( M_{\text{disk}} \), decreases down to \( 10^{-8} M_{\text{d,ini}} \) as the disk lifetime, \( t_{\text{disk}} \), where \( M_{\text{d,ini}} \) is the initial disk mass. Here we take a numerical factor \( 10^{-8} \), but \( t_{\text{disk}} \) is insensitive to it, if it is \( \lesssim 10^{-4} \).

We note that Kimura et al. (2016) and Kunitomo et al. (2020) measured the inner disk lifetime when the optical depth of the inner disk (i.e., the IR-emitting region) becomes unity. Considering the fact that the IR is emitted by dust grains that are not modeled in this study (see Section 5.4), here we measure \( t_{\text{disk}} \) using \( M_{\text{disk}} \). However, we note that the inner disk lifetime using the optical depth is almost the same as \( t_{\text{disk}} \) in this study, because an entire disk disperses quickly once a gap opens (see Figure 8(a)).

3.4. Numerical Method

We numerically solve Equation (11) using the time-explicit method based on Kunitomo et al. (2020). The calculation domain ranges from 0.01 to 10\(^3\) au. The grid size is in proportion to \( \sqrt{R} \), and the number of mesh points is 2000. The zero-torque boundary condition is imposed at both the inner and outer boundaries. We measure \( M_{\text{acc}} \) at the innermost cell. We stop calculations when the disk is completely dispersed.

We adopt the self-similar solution (Lynden-Bell & Pringle 1974) as an initial surface density profile given by

\[
\Sigma(R, t=0) = \frac{M_{\text{d,ini}} \exp(-R/R_1)}{2\pi R_1^2} R/R_1.
\]

The characteristic radius \( R_1 \) represents the location outside which the \( e^{-1} \) of the disk mass resides.

We choose input parameters to reproduce observational constraints as summarized in Table 2. First, from the observed relation that disk masses are proportional to \( M_* \) (e.g., Williams & Cieza 2011; Andrews et al. 2013; Mohanty et al. 2013; Pascucci et al. 2016), we adopt

\[
M_{\text{d,ini}} \propto M_*,
\]

The proportionality factor ranges from 0.001 to 0.1. Given that this value decreases with time, we start calculations with a massive disk, \( M_{\text{d,ini}} = 0.1 M_* \), (i.e., from the early phase). We note that the quantity of \( M_{\text{d,ini}}/M_* \) does not change the qualitative results on the disk lifetimes.

Second, following Gorti et al. (2009), we adopt

\[
\alpha \propto M_*,
\]

in order to reproduce the observed relation \( M_{\text{acc}} \propto M_*^2 \) (e.g., Calvet et al. 2004; Muzerolle et al. 2005). We assume that magnetorotational instability (MRI; Velikhov 1959; Chandrasekhar 1961; Balbus & Hawley 1991) is the source of the turbulent viscosity, and we adopt \( \alpha = 10^{-2} (M_*/M_\odot) \).

Equation (20) is derived with the following assumptions: steady-state accretion \( (\dot{M}_{\text{acc}} = 3\pi \Sigma \nu_{\text{visc}}) \), constant \( R_1 \) with \( M_* \) and Equation (19) (therefore \( \Sigma \propto M_* \)), optically thin disk temperature (see, e.g., Equation (6) of Kunitomo et al. 2020), \( L_* \propto M_*^2 \) (see \( L_* \) at 1 Myr in Figure 4(a) or Siess et al. 2000, and Keplerian \( \Omega \).

Finally, we adopt the initial disk radius \( R_1 = 50 \) au. Andrews et al. (2010) measured dust disk radii from millimeter-wavelength observations and found that they range from 14 to 200 au and peak at \( \approx 30 \) au (see their Figure 3). Considering that recent studies have suggested that gas disks are likely to be larger than dust disks (e.g., Ansdell et al. 2018), we adopt \( R_1 = 50 \) au in this paper. Andrews et al. (2010) did not find a clear correlation between the disk radius and \( M_* \) (see also Ansdell et al. 2018; Long et al. 2019). Although Andrews et al. (2018) recently suggested a weak correlation with \( M_* \) in this paper, we adopt the constant \( R_1 \) with \( M_* \) for simplicity.
4. Results

4.1. Overview of Disk Evolution

In this subsection, we show the disk evolution around a 3 $M_{\odot}$ star with the fiducial settings listed in Table 2. In our results, $t=0$ corresponds to the time when stars appear on their birthline. We consider the three PE mechanisms: FUV, EUV, and X-rays. In the four panels of Figure 8, we show the evolution of (a) the $\Sigma$ profile, (b) the $T_{\text{mid}}$ profile, (c) $M_{\text{acc}}$ and the mass-loss rates, and (d) the time-integrated accreted or lost mass. We define $M_{\text{FUV}} \equiv \int 2\pi R \Sigma_{\text{FUV}} \, dR$ and $M_{\text{EUV}} \equiv \int 2\pi R \Sigma_{\text{EUV}} \, dR$ (see Equations (15) and (17)). Both are integrated over the entire computation domain. The time-integrated accreted mass is $M_{\text{acc}} \equiv \int M_{\text{acc}} \, dt$, and the total mass lost by the PE is $M_{\text{PE}} \equiv \int \Sigma_{\text{PE}} \, dR \, dt$ (see Equation (12)). We note that we have checked the mass conservation in our simulations: $M_{\text{lim}} = M_{\text{disk}}(t) + M_{\text{acc}}(t) + M_{\text{PE}}(t)$ with a precision of $<10^{-10}$.

The qualitative behavior of the evolution in Figure 8 is the same as the results in previous works (e.g., Clarke et al. 2001; Alexander et al. 2006a; Gorti et al. 2009; Owen et al. 2010): (i) the $M_{\text{disk}}$ decreases with time due to viscous accretion, (ii) a gap is created when and where the accretion rate decreases down to the PE rate, (iii) the inner disk depletes in the viscous timescale at the gap, and (iv) after the dispersal of the inner disk, the outer disk is directly irradiated and also quickly dispersed. The gap opens at $\sim 15$ au, slightly outside the peak of $\Sigma_{\text{FUV}}$ (see Section 3.2). We note that the period of the phase (iii) is consistent with the viscous timescale, $\tau_{\text{vis}}$, at the gap given by

$$\tau_{\text{vis}} \equiv \frac{R^2}{\nu_{\text{vis}}} = 0.07 \text{Myr} \left( \frac{R}{15\text{ au}} \right)^{1/2} \left( \frac{T_{\text{mid}}}{100\text{ K}} \right)^{-1} \left( \frac{M_{\odot}}{3} \right)^{1/2} \left( \frac{\alpha}{0.03} \right)^{-1}. \quad (21)$$

We note that the nonsmooth $T_{\text{mid}}$ profile in Figure 8(b) results from the nonlinear function of the opacity (see Kunitomo et al. 2020).

Figure 8(c) shows that the mass-loss rates evolve with time, unlike the previous studies. Although the X-ray PE rate, $M_{\text{X}}$, is high (a few $10^{-7} \ M_{\odot} \text{ yr}^{-1}$) in the early phase, $M_{\text{X}}$ decreases by more than 3 orders of magnitude between 0.4 and 0.8 Myr. This is induced by stellar evolution; at this phase, a 3 $M_{\odot}$ star develops a large radiative core. Ro increases, and therefore $L_{\text{X}}$ and $M_{\text{X}}$ decrease (see Figures 4(b) and 5). Instead,
$M_{\text{FUV}}$ rapidly increases by more than 1 order of magnitude at $\approx 1$ Myr. This is because, after $\approx 1$ Myr, $T_{\text{eff}} > 7300$ K and $L_{\text{FUV,ph}} > 10^{-2}$; that is, the stellar surface becomes hot enough to emit FUV from the photosphere. We stress that although Gorti et al. (2009) already found that the rapid disk dispersal around IM stars is induced by the PE driven by photospheric UV, they did not consider stellar evolution (see Section 4.2). Since the rapid increase of $M_{\text{FUV}}$ has a strong impact on the disk evolution, we claim that stellar evolution is crucial for the disk dispersal around IM stars.

Figure 9 shows the evolution of $L_{\text{FUV}}$ in the cases of $M_*=1$ and $3 M_\odot$. The disk evolution around a $1 M_\odot$ star is shown in Appendix B. In the $1 M_\odot$ case, $L_{\text{FUV,acc}}$ dominates in almost the entire phase, which is consistent with observations (see, e.g., Ingleby et al. 2011) and previous theoretical study (Gorti et al. 2009). Along with the decrease in $M_{\text{acc}}$, $L_{\text{FUV}}$ decreases with time, and in the late phase, $L_{\text{FUV,ph}}$ dominates. In the case of $3 M_\odot$ stars, however, although $L_{\text{FUV,acc}}$ dominates in the early phase, $L_{\text{FUV,ph}}$ rapidly increases by orders of magnitude as $T_{\text{eff}}$ increases at $\approx 1$–1.5 Myr. We note that, in the $4 M_\odot$ case, the switch occurs at $4 \times 10^4$ yr, and in the $5 M_\odot$ case, $L_{\text{FUV,ph}}$ always dominates.

We note that the initial value of the $L_{\text{FUV,acc}}$ of the $3 M_\odot$ star is $\approx 1$ order of magnitude larger than that of $1 M_\odot$. This is because we adopt the initial condition to reproduce the observed relation $M_{\text{acc}} \propto M_*^2$ (see Section 3.4).

### 4.2. Importance of Stellar Evolution

In Section 4.1, we showed that the photospheric FUV radiation has a dominant role in the disk dispersal around a $3 M_\odot$ star. We again note that it had already been found by Gorti et al. (2009), and as an update from their study, we considered the stellar evolution. To illustrate its importance, we performed the same simulation of Figure 8 but without the time evolution of $L_{\text{FUV,ph}}$, $\Phi_{\text{FUV}}$, and $L_\chi$ as in Gorti et al. (2009). We adopt $L_{\text{FUV,ph}} = 1.1 \times 10^{35}$, $\Phi_{\text{FUV}} = 1.0 \times 10^{39}$, and $L_\chi = 5.0 \times 10^{28}$ erg s$^{-1}$ following Gorti et al. (2009); see also Section 2.1). Figure 10 shows that $M_{\text{FUV}}$ is kept high from the beginning; therefore, the disk disperses much earlier than the case in Figure 8. As shown in Figure 9, the $L_{\text{FUV}}$ of a pre-MS $3 M_\odot$ star should be much lower than that of an MS star, but, in approaching the MS, should suddenly increase by orders of magnitude. This time evolution has a strong impact on the disk lifetime. We note that Equation (5) is adopted, but $L_{\text{FUV,ph}}$ always dominates; therefore, $L_{\text{FUV}}$ is almost constant with time.

The fact that a disk is dispersed mainly by the PE driven by $L_{\text{FUV,ph}}$ is the same in both cases in Figures 8 and 10. However, for a realistic disk evolution model around IM stars, we claim that stellar evolution is one important ingredient.

We note the difference between the results in Figure 10 and Gorti et al. (2009); even though the $L_{\text{FUV,ph}}$, $\Phi_{\text{FUV}}$, and $L_\chi$ values are the same, the disk lifetimes differ by about 1 order of magnitude (0.2 and 4 Myr, respectively). We speculate that the difference probably originates from the absorption of high-energy photons in disk winds from an inner disk. Gorti et al. (2009), see their Section 2.4.1) considered this effect, whereas we do not. This effect can suppress the PE rate in the early phase. We will discuss this issue in Section 5.4. Nevertheless, our claim that the time-dependent $L_{\text{FUV,ph}}$ is important for disk evolution is still valid.

### 4.3. Disk Lifetime

We perform a suite of disk evolution simulations around 0.5–5 $M_\odot$ stars as in Section 4.1. We find that $\tau_{\text{disk}}$ decreases with increasing $M_*$ (Figure 11(a)).

To understand which mechanism plays the dominant role, we also perform three sets of simulations: (i) with only the FUV PE, (ii) with only the X-ray PE, and (iii) without the EUV PE. The other settings are the same as the fiducial runs (see Table 2). In the models where we do not include the FUV and EUV PE (the “only X” model in Figure 11), the disk lifetime around $\gtrsim 3 M_\odot$ stars increases significantly, while any combination of mechanisms that includes FUV causes similarly short lifetimes for $\gtrsim 3 M_\odot$ stars. These results clearly illustrate that disks around $\gtrsim 3 M_\odot$ stars are dispersed mainly by the FUV PE.

Figure 11(b) shows the evolution of $M_{\text{disk}}$ around a $3 M_\odot$ star. After the X-ray PE becomes less effective at 0.4 Myr, it takes time for the FUV PE to become strong at 1.0 Myr, and then the disks quickly disperse if the FUV PE is considered.

Figure 11(a) also shows the time when stars reach $R_0 = R_{\text{crit}}$ (i.e., $L_\chi = 10^{-3.13} L_*$; the maximum value of $L_\chi$), $L_{\text{eff}} = 7342$ K (i.e., $L_{\text{FUV,ph}} = 10^{-2} L_*$; as an indicative timescale for $M_{\text{FUV}}$ to increase). These timescales decrease with $M_*$, This is because higher-mass stars have a shorter KH timescale $\tau_{KH}$ (see Equation (1)) and therefore develop a radiative core and have a hotter photosphere more rapidly. We note that the $T_{\text{eff}}$ of stars with
Figure 11. Top panel: disk lifetime as a function of $M_*$ in cases with the FUV, EUV, and X-ray PE (black solid line), only the FUV PE (red dashed line), only the X-ray PE (blue dotted line), and the FUV and X-ray PE (green dotted-dashed line). The two thin black dotted lines show the time when $R_0 = R_{\text{out}}$ (left; i.e., $L_{\text{X}}/L_\star = 3 \times 10^{-13}$) and $T_{\text{eff}} = 7342$ K (right; i.e., $L_{\text{FUV,ph}}/L_\star = 10^{-7}$). The cyan shaded region illustrates the phase in which the X-ray PE dominates, whereas in the magenta shaded region, the FUV PE dominates. Bottom panel: time evolution of disk mass, $M_{\text{disk}}$, around a 3 $M_\odot$ star.

less than 1.6 $M_\odot$ never reaches 7342 K in the pre-MS and MS phases; therefore, $L_{\text{FUV,ph}}$ is always below $10^{-2} L_\star$.

In the cases with FUV, the disks around $\gtrsim 2 M_\odot$ stars disappear after $L_{\text{FUV,ph}}$ reaches $10^{-2} L_\star$. In the case with only the FUV PE, the disk lifetime around $\sim 1.5$–3 $M_\odot$ stars is almost the same as the timescale to reach $L_{\text{FUV,ph}} = 10^{-2} L_\star$. Therefore, if the X-ray PE is less effective, the disk lifetime around IM stars is determined by the stellar evolution. We note that even though the $L_{\text{FUV}}$ of 4 and 5 $M_\odot$ stars becomes luminous in the early phase, it takes time for $M_{\text{acc}}$ to decrease and the disks to disperse. On the other hand, the disk lifetime around $\gtrsim 1 M_\odot$ stars in the case with only FUV exceeds 30 Myr. This is because the $L_{\text{FUV}}$ of low-mass stars is dominated by $L_{\text{FUV,acc}}$, which is self-regulated; $L_{\text{FUV,acc}}$ decreases along with decreasing $M_{\text{acc}}$ over time. Therefore, the PE mainly by $L_{\text{FUV,acc}}$ does not open a gap.

If we compare the cases with and without the X-ray PE, one finds that the disks around $\lesssim 2.5 M_\odot$ stars disperse mainly by the X-ray PE. The influence of the EUV PE on $t_{\text{disk}}$ is negligible in the entire mass range. Therefore, under the current settings, $\gtrsim 3 M_\odot$ stars disperse mainly by the FUV PE, whereas $\lesssim 2.5 M_\odot$ stars disperse by the X-ray PE. However, we note that, although we adopt the X-ray and EUV PE rates from the literature in this study, they are still under debate (see Section 5.4). If our X-ray PE rate is overestimated, then the realistic $t_{\text{disk}}$ should be in between the $t_{\text{disk}}$ of the fiducial case and that of the “only FUV” case. Nevertheless, the importance of the $L_{\text{FUV,ph}}$ evolution around IM stars is not affected by the uncertainty of the X-ray PE model.

On the high-mass side ($\gtrsim 3 M_\odot$), $t_{\text{disk}}$ decreases with $M_*$ because of the shorter $\tau_{\text{rot}}$, as described above. Here we explain why we obtain the same trend on the low-mass side. The $t_{\text{disk}}$ value is almost the same as the timescale of the gap opening, which occurs when $M_{\text{acc}}$ decreases down to $M_{\text{PE}}$ (see Section 4.1). Both have a similar dependence on $M_*$, because the input parameter $\alpha$ to reproduce the observed relation $M_{\text{acc}} \propto M_*^\alpha$ (Section 3.4). Around low-mass stars, the X-ray PE dominates, and therefore $M_{\text{PE}} \approx M_X$. We adopt the X-ray PE model based on Owen et al. (2012), which is in proportion to $L_X$. Figure 5 shows that $L_X$ is roughly proportional to $M_*^{1.6}$ in the case of 1 Myr old low-mass stars. Since both $M_{\text{acc}}$ and $M_{\text{PE}}$ have a similar correlation with $M_*$, the gap-opening timescale is determined by the timescale for $M_{\text{acc}}$ to decrease, that is, the viscous timescale $\tau_{\text{vis}}$ (Clarke et al. 2001). Given that $\tau_{\text{vis}} \propto M_*$ (see Section 3.4) and that we assume $\tau_1$ does not correlate with $M_*$, $\tau_{\text{vis}} \propto M_*^{-1}$. Therefore, $M_{\text{acc}}$ decreases faster around high-mass stars and $t_{\text{disk}}$ decreases with $M_*$. We note that for this correlation, Equation (20) is essentially important because this gives the relation $\tau_{\text{vis}} \propto M_*$ (see discussions in Section 5.3).

5. Discussion

5.1. Comparison with Observations

In this subsection, we compare our results with observations. Here we focus only on the gas disk lifetime (see Section 1 for dust disk lifetimes); recent Hα observations have revealed that the gas disk lifetime around IM stars is shorter than that of low-mass stars (Kennedy & Kenyon 2009; Yasiu et al. 2014). This is consistent with our results of $t_{\text{disk}}$ in Figure 11(a). We again stress that the realistic $L_{\text{FUV,ph}}$ model with stellar evolution is crucial for this trend on the high-mass side, whereas the $L_X$ and $\alpha$ models are important on the low-mass side. Since we have not explored the dependence on the input parameters and the PE models are still under debate, we limit ourselves to focusing only on the qualitative results in this study. We leave the quantitatively detailed discussions for future studies.

5.2. Dependence on the Variety in X-Ray Luminosity

We have found that disks around $\lesssim 2.5 M_\odot$ stars are dispersed mainly by the X-ray PE, and therefore $t_{\text{disk}}$ depends on $L_X$. Observations have revealed that the stellar $L_X$ has a large variety. Although in this paper, we have adopted the empirical relation of Wright et al. (2011, see Equation (9)), the observed data of $R_X$ (see, e.g., Preibisch et al. 2005) exhibit a variety by a factor of 4.5 ($\sim 0.65$ dex). Moreover, although we have assumed $P_{\text{rot}} = 3$ days, the observed rotational period of pre-MS stars has a variety from $\sim 1$ to 10 days (see Section 2.5). In this section, we explore the influence of these varieties on the results of $t_{\text{disk}}$.

Figures 12(a) and (b) show the $L_X$ evolution of 3 and 1 $M_\odot$ stars, respectively. We consider the cases with $P_{\text{rot}} = 1$ and 10 days and $L_X$ multiplied or divided by a factor of 4.5. We find that 1 $M_\odot$ stars develop a radiative core at $\approx 10$ Myr, and until then, pre-MS stars are in the saturated regime irrespective of $P_{\text{rot}}$, whereas it happens for 3 $M_\odot$ stars in the early ($\approx 0.4$ Myr) phase. We note that Tu et al. (2015) claimed that the $L_X$ of 1 $M_\odot$ MS stars has a large variety depending on the $P_{\text{rot}}$. This is because the $\tau_{\text{conv}}$ of 1 $M_\odot$ MS stars is short enough for their
$L_X$ to depend on $P_{\text{rot}}$ (see also Equation (9)). However, our results show that the $L_X$ of pre-MS 1 $M_\odot$ stars does not depend on $P_{\text{rot}}$ until $\approx$10 Myr.

Figure 12(c) shows $t_{\text{disk}}$ with different $L_X$ models. Here we adopt fiducial settings other than $L_X$. We find that the variation in $P_{\text{rot}}$ has little impact on $t_{\text{disk}}$. On the other hand, if we change $L_X$ by a factor of 4.5, $t_{\text{disk}}$ changes by up to 1 dex. The variation of $L_X$ has a larger impact on $t_{\text{disk}}$ around lower-mass stars. Therefore, for the detailed comparison with observed disk fractions with time, we need to consider the $L_X$ variation as claimed by Kimura et al. (2016).

The trend of $t_{\text{disk}}$ with $M_*$ depends on different $L_X$ models: $t_{\text{disk}}$ decreases with increasing $M_*$ in the low-$L_X$ case, whereas the $t_{\text{disk}}$ of <3 $M_\odot$ stars is almost constant in the high-$L_X$ case. For the former, the reason is the same as the fiducial case (i.e., the shorter $\tau_{\text{vis}}$; see Section 4.3). For the latter, $M_X \gg M_{\text{acc}}$ from the beginning, and therefore the gap-opening timescale ($\approx t_{\text{disk}}$) is determined by $\tau_{\text{PE}}(R_{\text{gap}}) = \Sigma/\Sigma_{\text{PE}}$, where $R_{\text{gap}}$ is the radius where the PE opens a gap. Below, we briefly show that $\tau_{\text{PE}}(R_{\text{gap}})$ is insensitive to $M_*$, First, $R_{\text{gap}} \propto M_*$ because the location of the peak of $\Sigma_X$ is proportional to $M_*$ (see Section 3.2). Since we assume $\Sigma \propto R^{-1}M_*$ as an initial condition, the initial $\Sigma$ at $R_{\text{gap}}$ does not depend on $M_*$. Second, $\Sigma_{\text{PE}}(R_{\text{gap}}) \propto \Sigma_X \propto L_X R_{\text{gap}}^{-2}$ (see Equation (15)), where $L_X \propto M_\odot^{1.6}$ but $R_{\text{gap}}^{-2} \propto M_\odot^{-2}$. These two opposite effects make the peak $\Sigma_X$ value almost constant with $M_*$.

Therefore, the $\tau_{\text{PE}}$ (and thus $t_{\text{disk}}$) of <3 $M_\odot$ stars is insensitive to $M_*$ in the high-$L_X$ case.

5.3. Dependence on the Variety in Viscosity

We have adopted $\alpha \propto M_*$ to reproduce the observed relation ($M_{\text{acc}} \propto M_*^{2.7}$; see Section 3.4), but the physical origin of this relation is still unclear. In addition, the absolute value of $\alpha$ is also under debate. As a fiducial value, we adopt a relatively large $\alpha$ value ($=10^{-2}(M_*/M_\odot)$) assuming that the disks are turbulent. However, recent observations (e.g., Pinte et al. 2016; Flaherty et al. 2017) and theoretical studies (see, e.g., Turner et al. 2014, and references therein) have suggested a low $\alpha$ (e.g., $\lesssim 10^{-3}$ from the observations).

To explore the dependence of $t_{\text{disk}}$ on the $\alpha$ model, we simulate disk evolutions with $\alpha = 10^{-2}$ (i.e., constant with $M_*$) and $\alpha = 10^{-3}(M_*/M_\odot)$ (i.e., 10 times lower than the fiducial model). Figure 13 shows that the decreasing $t_{\text{disk}}$ with $\alpha$ on the high-mass side ($\geq 3 M_\odot$) remains even if we adopt a different $\alpha$ model because the rapid increase of $L_{\text{FUV,ph}}$ has a dominant role.

We note that the variety in $\alpha$ affects the $t_{\text{disk}}$ values; a lower $\alpha$ by a factor of 10 results in a larger $t_{\text{disk}}$ by a factor of $\approx 3$ (as shown in Figure 11 of Gorti et al. 2009). We also note that if $\alpha$ is constant with $M_*$, the $t_{\text{disk}}$ value is also constant with $M_*$ ($\approx 2$ Myr) in the range $M_* \leq 3 M_\odot$. Therefore, to compare theoretical $t_{\text{disk}}$ values with observations, it is crucial to understand the origin of the relation $M_{\text{acc}} \propto M_*^2$ and constrain the absolute value of $\alpha$ in protoplanetary disks.

5.4. Model Caveats

In this subsection, we describe the caveats on the PE models, evolution of dust disks, magnetohydrodynamic (MHD) winds, and variations of input parameters.

We point out two issues on the PE models. First, although we adopt the X-ray PE model by Owen et al. (2012), their $M_X$ is higher than that of recent RHD simulations with a self-consistent thermochemistry by Wang & Goodman (2017) and Nakatani et al. (2018a). Therefore, although our results suggest that the disks around $\lesssim 2.5 M_\odot$ stars disperse mainly by the X-ray PE (Section 4.1), the $t_{\text{disk}}$ of $\lesssim 2.5 M_\odot$ stars may be underestimated. Future works should investigate the long-term disk evolution with the updated X-ray PE rate. Second, the PE may be suppressed in particular in the early phase in the outer region due to the absorption of high-energy photons. These photons can be shielded by dense gas, such as accretion flows onto the star (Alexander et al. 2004), inner disk winds (Bai 2017; Takasao et al. 2018), stellar winds (Hollenbach et al. 2000), and dust grains in the disk atmosphere (Nakatani et al. 2018b). If the high-energy photons are shielded, the PE rate can decrease by orders of magnitude, and the PE profiles can also be changed (see also Section 4.2).

There are two issues in the luminosity and spectra of stellar high-energy photons. First, in this paper, we have used a simple
model of $\Phi_{EUV}$, but this is quite uncertain (Section 2.4). Bouret & Catala (1998) suggested that Herbig Ae/Be stars have $\Phi_{EUV} \sim 10^{33} - 10^{45}$ using an indirect estimation. Although the EUV PE has a marginal effect on the disk evolution in our results, we expect that future works constrain the $\Phi_{EUV}$ of IM stars more precisely. Second, the hardness of the X-ray spectra of young stars remains a matter of debate. Some observations have suggested that the X-ray spectra of accreting stars may be softer (e.g., Kastner et al. 2002, 2004). Gorti et al. (2009) showed that a softer X-ray spectrum results in a larger PE rate even with the same $L_X$ (see their Figure 9). Future studies should investigate the influence of the evolution of the X-ray hardness on the disk evolution.

The uncertainties and varieties in the PE models above would be important for some observational results. Although most IM stars have a shorter inner disk lifetime (see Section 5.1), some have a long disk lifetime (e.g., Pančić et al. 2008; Fedele et al. 2017; Booth et al. 2019; Miley et al. 2019; Muro-Arena et al. 2020). These long-lived disks may have a lower PE rate. Since most of these long-lived disks around Herbig stars are well studied due to the relative ease of detecting their large bright disks, there are a lot of existing high-quality data for theoretical models to be compared with. Theoretical models should be compared in detail with and explain these observations in future.

We stress the importance of the dust disk evolution, which is not considered in this paper. Previous studies have found that gas and dust disk lifetimes can differ (see, e.g., Takeuchi et al. 2005; Alexander & Armitage 2007; Gorti et al. 2015; Owen & Kollmeier 2019). Since IR observations trace the small dust grains, we need to simulate the long-term evolution of gas and dust to compare theoretical models with IR observations. The number of dust grains in the disk atmosphere may also affect the FUV PE rate (Gorti et al. 2015; Nakatani et al. 2020). However, the motion and evolution of dust grains are quite complicated; we need to consider a number of effects, such as radial drift (Adachi et al. 1976), gas pressure gradient (Taki et al. 2016, 2020), coagulation, fragmentation and collisional cascade (Kobayashi & Tanaka 2010), and entrainment in the PE or MHD disk winds (Gorti et al. 2015; Miyake et al. 2016; Franz et al. 2020). Future studies with the dust and stellar evolution around IM stars are needed to investigate the realistic lifetimes of dust disks.

In this paper, we have not included MHD disk winds, but recently much attention has been paid to them (e.g., Suzuki & Inutsuka 2009; Fromang et al. 2013; Lesur et al. 2013; Bai & Stone 2013a; Bai 2017; Wang et al. 2019). The MHD winds carry away not only mass but also angular momentum (so-called wind-driven accretion; Bai & Stone 2013b; Bai 2016; Suzuki et al. 2016). Kunitomo et al. (2020) claimed that the MHD and PE winds have different roles (see also recent radiation-MHD simulations by Wang et al. 2019; Rodenkirch et al. 2020; Gressel et al. 2020), and both winds and the wind-driven accretion should be considered for a realistic disk evolution, in particular for disks with weak turbulence. We will investigate the long-term disk evolution around IM stars including both winds in our next paper.

We have not varied the input parameters in this paper. The variety of the initial disk condition, $M_{d,\text{ini}}$ and $R_1$, should be related to the properties of parental clouds using a disk formation model (Takahashi et al. 2013; Kimura et al. 2016). For a detailed comparison with the observations of disk fractions over time, we need Monte Carlo simulations covering the variety of input parameters ($M_{d,\text{ini}}$, $R_1$, and $\alpha$; Alexander & Armitage 2009; Kimura et al. 2016).

Finally, we discuss the variety of stellar evolution. Although in this paper, we adopted the birthline based on the standard star formation scenario, recent studies have shown that the luminosity of the birthline depends on star formation processes (such as the variety in the entropy of accreting materials or deuterium abundance; see Baraffe et al. 2009; Hosokawa et al. 2011; Tognelli et al. 2015; Kunitomo et al. 2017; Kuffmeier et al. 2018). Stellar $T_{\text{eff}}$ depends on the metallicity and mixing-length parameter $\alpha_{\text{MLT}}$, a lower metallicity or larger $\alpha_{\text{MLT}}$ results in a higher $T_{\text{eff}}$ (Kippenhahn & Weigert 1990). Although in this paper, we have adopted the solar metallicity and $\alpha_{\text{MLT}} = 2.0$, the varieties of these parameters affect the $T_{\text{eff}}$ evolution and therefore the $L_{\text{FUV,ph}}$ and $\Phi_{EUV,ph}$ evolution.

5.5. Implications for Planet Formation

The disk evolution models have important implications for planet formation. Since planets form and evolve in a protoplanetary disk, their characteristics may reflect the disk properties. For example, the orbital configuration of planets around IM stars is different from low-mass stars; there is a paucity of close-in planets around $\gtrsim 2 M_\odot$ stars (e.g., Sato et al. 2008). One possible origin is the different disk evolution; the rapid disk dispersal may hinder planets from migrating inward (e.g., Burkert & Ida 2007; Currie 2009; Kunitomo et al. 2011). Radial velocity surveys have revealed that the occurrence rate of detected giant planets depends upon $M_*$ (e.g., Johnson et al. 2010; Reffert et al. 2015). The mass fraction and/or composition of planet atmospheres can give an indication as to when or where the planet was formed in a disk (Guillot & Hueso 2006; Ogihara et al. 2020; Miley et al. 2021). We expect that our disk evolution models also lead to the understanding of planet formation processes around IM stars.

6. Summary and Conclusions

We investigated the long-term disk evolution around 0.5–5 $M_\odot$ stars by considering the viscous accretion; the PE

\footnote{Although Kunitomo et al. (2011) described that $\alpha_{\text{MLT}} = 1.5$, this was a typo. In standard solar models (see, e.g., Serenelli et al. 2009), $\alpha_{\text{MLT}} = 2.0$ is suggested.}
mass loss by stellar FUV, EUV, and X-rays; and stellar evolution. We started calculations from the early phase and initial conditions with a compact \((R_1 = 50 \text{ au})\) and massive \((M_{d,\text{ini}} = 0.1 M_\odot)\) disk.

We found that the nature of the emission of stellar high-energy photons changes with time; low-mass stars strongly emit X-rays until the typical disk lifetime (i.e., several million years), whereas the X-ray luminosity of higher-mass stars decreases and instead, their FUV luminosity rapidly increases due to stellar evolution (e.g., at around 1 Myr in the case of 3 \(M_\odot\) stars). The critical mass is \(\sim 2.5 M_\odot\) because the KH timescale becomes comparable to the disk dispersal timescale. Therefore, the effect of stellar evolution is not negligible, as assumed in previous works, and should be considered for realistic disk evolution models around IM stars.

We are grateful for the simulation results of star formation provided by Dr. Steven W. Stahler. We are also grateful to Drs. Kei E. I. Tanaka, Chikako Yasui, Masahiro Ikoma, Taishi Nakamoto, Hideko Nomura, Philip J. Armitage, Richard D. Alexander, Jaehan Bae, Kenji Hamaguchi, Shinsuke Takasao, Shu-ichiro Inutsuka, and Hiroshi Kobayashi for fruitful discussions and comments. We appreciate the constructive comments of the anonymous referee, which helped us to improve this paper. M.K. and S.I. thank the University of Leeds for the financial support through the International Mobility Fund and hospitality during their stay in Leeds. This work was supported by JSPS KAKENHI grant Nos. 12J09296, 23244027, 15H02065, 16H02160, 17H01105, 17H01153, and 20K14542. The work of O.P. is funded by the Royal Society Dorothy Hodgkin Fellowship. J.M.M. is supported through the University of Leeds Doctoral Scholarship.

**Appendix A**

**Dependence of Photospheric UV Luminosity on the Stellar Surface Gravity**

In Section 2.2, we derived the empirical formulae of photospheric FUV and EUV luminosities in the case of \(g = 0.33 g_\odot\). We note that there is a variety in \(g\log)\) of 0.5–5 \(M_\odot\) pre-MS stars; from 0.1 to 10 Myr, it ranges from 2.7 to 4.3. Figure 14 shows the weak dependence of \(f_{\text{EUV}}\) and \(f_{\text{FUV}}\) on \(g\). We find that the difference of \(f_{\text{FUV}}\) from the fiducial case with 0.33 \(g_\odot\) is at most 13%, but that of \(f_{\text{EUV}}\) is up to a factor of 3. In this paper, we neglect this weak dependence for simplicity.

![Figure 14](image-url)

*Figure 14.* The \(f_{\text{EUV}}\) (left panel) and \(f_{\text{FUV}}\) (right panel) with varying \(g = 3\) (red double dotted–dashed lines), 1 (green dotted–dashed lines), 0.33 (fiducial; blue solid lines), 0.1 (purple dashed lines) and 0.01 (cyan dotted lines) \(g_\odot\).*
Appendix B
Disk Evolution around Low-mass Stars

In this Appendix, we show the disk evolution around low-mass stars in our model. Since the X-ray PE is a matter of debate (see Section 5.4), it should be noted that the results may be updated in future work.

Figure 15 shows the disk evolution around a 1 $M_\odot$ star. The qualitative behavior of the surface density evolution is the same as the 3 $M_\odot$ star case (Section 4.1). However, unlike the case of IM stars (Figure 8), $M_X$ is always larger than $M_{FUV}$ and $M_{UV}$ (see, however, the caveats in Section 5.4). This is because the $L_X$ of $\approx 1$ $M_\odot$ young stars is in the saturated regime and therefore as large as $\sim 10^{29}$–$10^{31}$ erg s$^{-1}$. Therefore, most materials are lost by either accretion or the X-ray PE.

We note that, as described in Kunitomo et al. (2020, see their Section 4.4), we see the gradual decrease of $M_X$ over 3 Myr, but the qualitative behavior described above is the same as the cases with constant $L_X$ in the previous works (e.g., Owen et al. 2010). This is expected from the long KH timescale of low-mass stars (see Section 1).

ORCID iDs
Masanobu Kunitomo https://orcid.org/0000-0002-1932-3358
Shigeru Ida https://orcid.org/0000-0002-9676-3891
Olja Panić https://orcid.org/0000-0002-6648-2968
James M. Miley https://orcid.org/0000-0002-1575-680X
Takeru K. Suzuki https://orcid.org/0000-0001-9734-9601

References
Adachi, I., Hayashi, C., & Nakazawa, K. 1976, PThPh, 56, 1756
Adams, F. C., Hollenbach, D., Laughlin, G., & Gorti, U. 2004, ApJ, 611, 360
Alexander, R., Pascucci, I., Andrews, S., Armitage, P., & Cieza, L. 2014, in In Protostars and Planets VI, ed. H. Beuther et al. (Tucson, AZ: Univ. of Arizona Press), 475
Alexander, R. D., & Armitage, P. J. 2007, MNRAS, 375, 500
Alexander, R. D., & Armitage, P. J. 2009, ApJ, 704, 989
Alexander, R. D., Clarke, C. J., & Pringle, J. E. 2004, MNRAS, 348, 879
Alexander, R. D., Clarke, C. J., & Pringle, J. E. 2006a, MNRAS, 369, 229
Alexander, R. D., Clarke, C. J., & Pringle, J. E. 2006b, MNRAS, 369, 216
Andrews, S. M., Rosenfield, K. A., Kraus, A. L., & Wilner, D. J. 2013, ApJ, 771, 129
Andrews, S. M., Terrell, M., Tripathi, A., et al. 2018, ApJ, 865, 157
Andrews, S. M., Wilner, D. J., Hughes, A. M., Qi, C., & Dullemond, C. P. 2010, ApJ, 723, 1241
Ansdell, M., Williams, J. P., Trapman, L., et al. 2018, ApJ, 859, 21
Armitage, P. J. 2000, A&A, 362, 968
Bahcall, J. N., Basu, S., Pinsonneault, M., & Serenelli, A. M. 2005, ApJ, 618, 1049
Bai, X.-N. 2016, ApJ, 821, 80
Bai, X.-N. 2017, ApJ, 845, 75
Bai, X.-N., & Stone, J. M. 2013a, ApJ, 767, 30
Bai, X.-N., & Stone, J. M. 2013b, ApJ, 769, 76
Balbus, S. A., & Hawley, J. F. 1991, ApJ, 376, 214
Baraffe, I., Chabrier, G., & Gallardo, J. 2009, ApJ, 702, L27
Booth, A. S., Walsh, C., Ilee, J. D., et al. 2019, ApJL, 882, L31
Bouret, J.-C., & Catala, C. 1998, A&A, 340, 163
Bouvier, J. 2008, A&A, 489, L53
Burkert, A., & Ida, S. 2007, ApJ, 660, 845
Calvet, N., & Gullbring, E. 1998, ApJ, 509, 802
Calvet, N., Muzerolle, J., Briceño, C., et al. 2004, AJ, 128, 1294
Carpenter, J. M., Mamajek, E. E., Hillenbrand, L. A., & Meyer, M. R. 2006, ApJL, 651, L49
Castelli, F., & Kurucz, R. L. 2003, in IAU Symp. 210, Modelling of Stellar Atmospheres, ed. N. Piskunov, W. W. Weiss, & D. F. Gray (Cambridge: Cambridge Univ. Press), A20
Chandrasekhar, S. 1961, Hydrodynamic and Hydromagnetic Stability (Oxford: Clarendon)
Clarke, C. J., Gendrin, A., & Sotomayor, M. 2001, MNRAS, 328, 485
Cox, J., & Giuli, R. 1968, Principles of Stellar Structure (New York: Gordon and Breach), 401

Figure 15. Same as Figure 8 but around a 1 $M_\odot$ star.
