Star formation

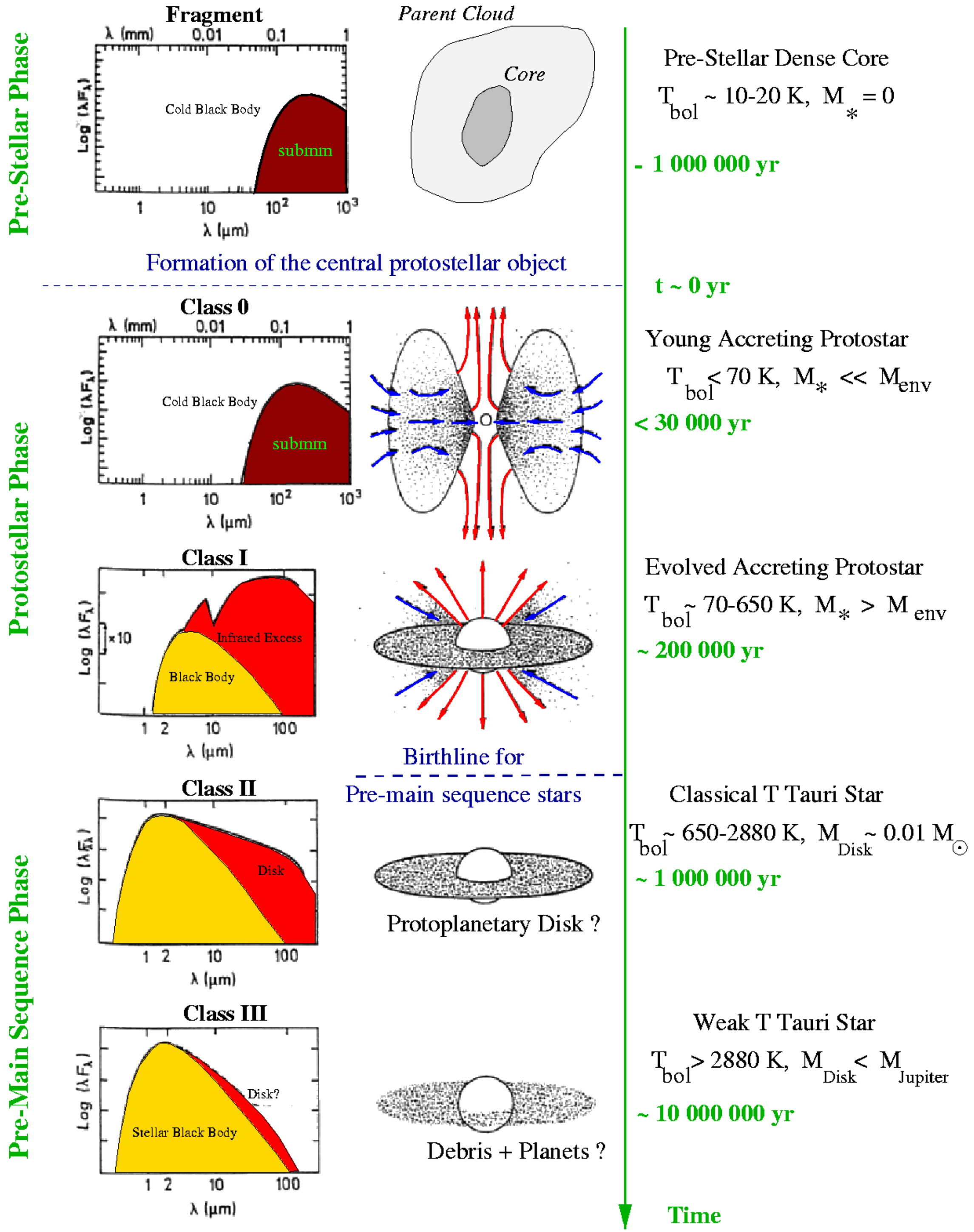
Protostellar accretion disks

Summary of previous lectures and goal for today

Collapse

Protostars
- *main accretion phase*
- *not visible in optical (dust envelope)*

Pre-main-sequence phase
- *(almost) final mass reached*
- *visible in optical (envelope dispersed)*
- *quasi-static contraction*



Formation of disks

So far, no magnetic field and no rotation…

Now: Let's look at rotation! (and magnetic fields)

A consequence of rotation: disks

- disk formation
- disk structure

Environmental impact : jets and outflows

- angular momentum removal

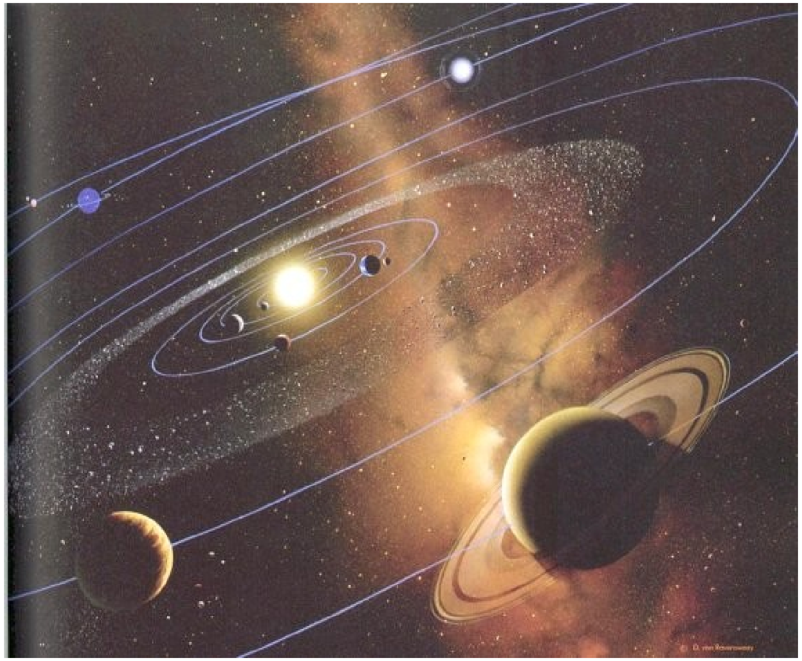
Our solar system

- Planetary orbits are coplanar.

- All planets orbit in the same direction

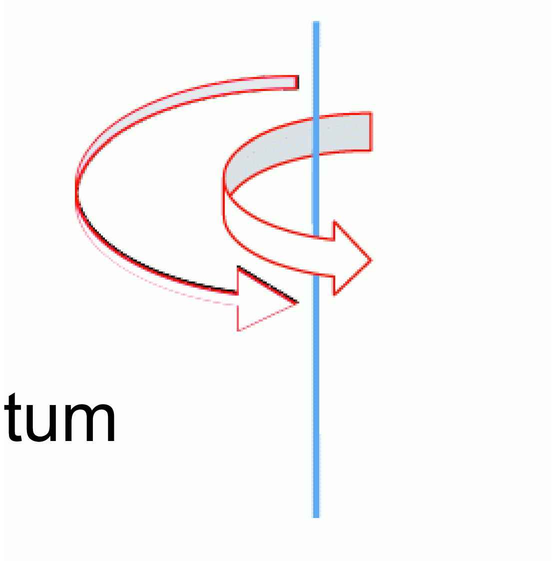
- Sun mass = 98%, sun angular momentum <2%

→ "Nebular hypothesis":
(famous from Kant & Laplace, 18th century, but actually first exposed by Emanuel Swedenborg in 1734)



"Planets are formed from a gaseous protoplanetary disk"

Consequence of rotation



Presence of accretion disks is predicted by theory of infall of rotating clouds

In the simplistic case of rotating infall *without magnetic braking*, the angular momentum of any fluid particle is conserved during infall.

- The angular rotation increases with decreasing distance to the central axis

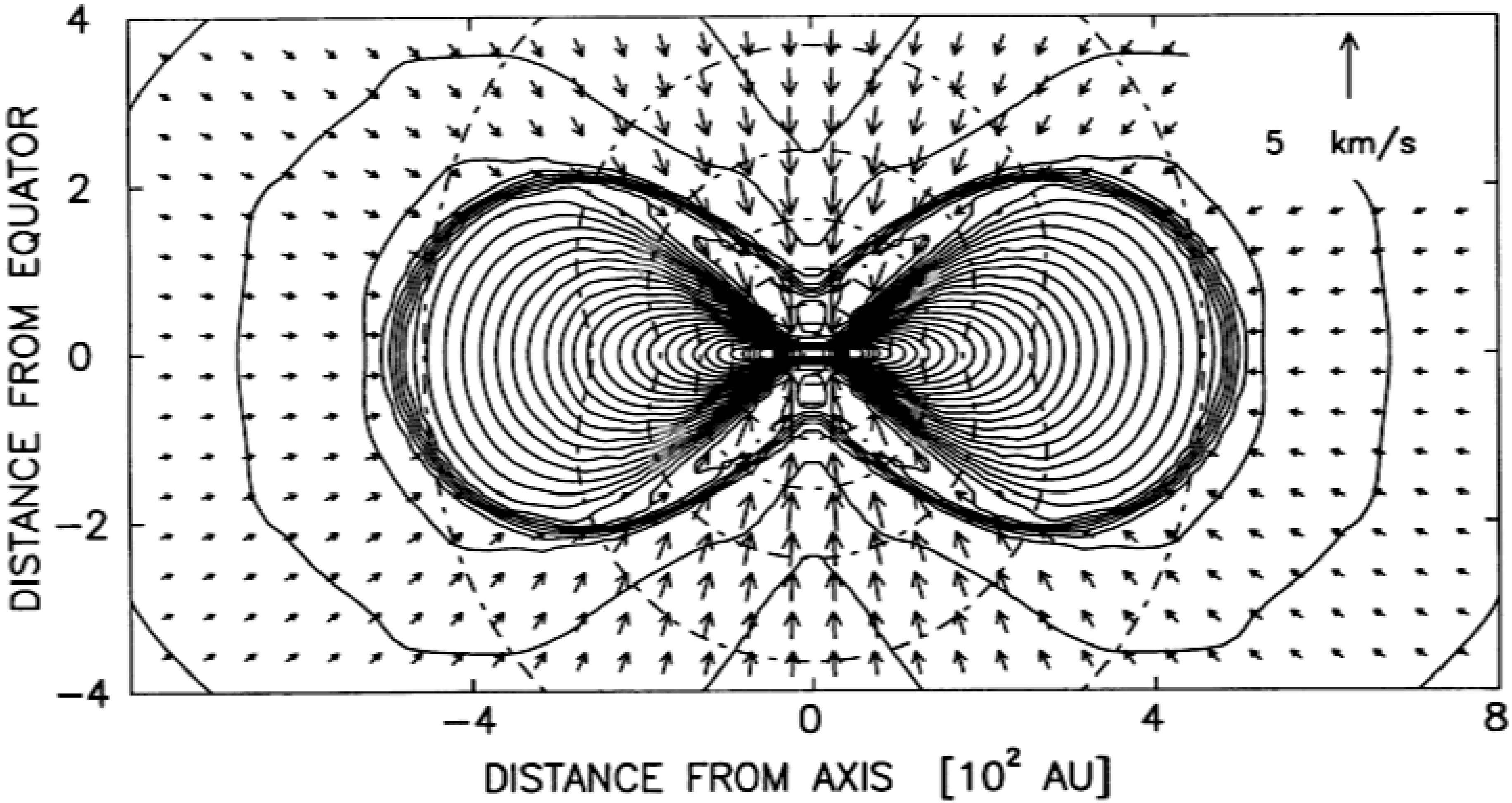
$$\Omega = \Omega_0 \, (r_0/r)^2$$

- Centrifugal force = $m\Omega^2 r \sim r^{-3}$
- Gravitational force $\sim r^{-2}$

→ rotation becomes important close to the central object

The formation of a disk

3-D Radiation-Hydro simulations of disk formation

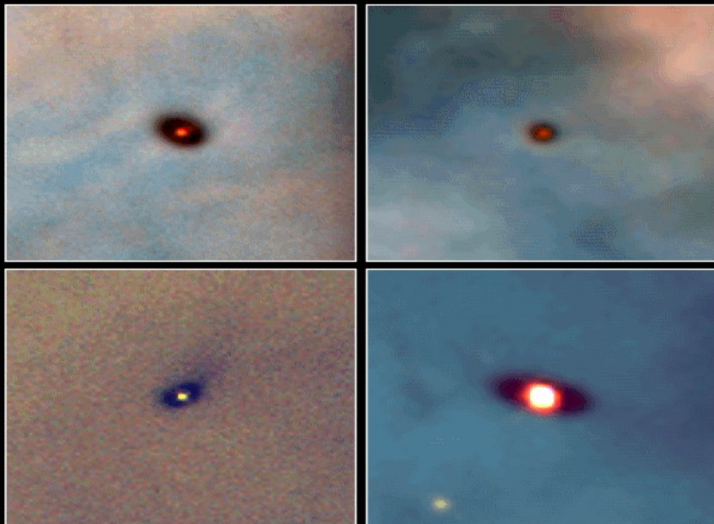


Yorke, Bodenheimer & Laughlin 1993

Observations of disks

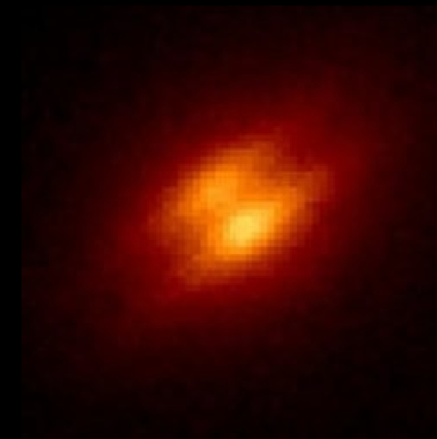
- Direct
  - Optical (silhouette)
    - Not deeply embedded objects
  - Millimeter/Submillimeter
    - Dust
    - Molecular lines
- Indirect
  - SEDs
- Results: Sizes ~ 100 AU

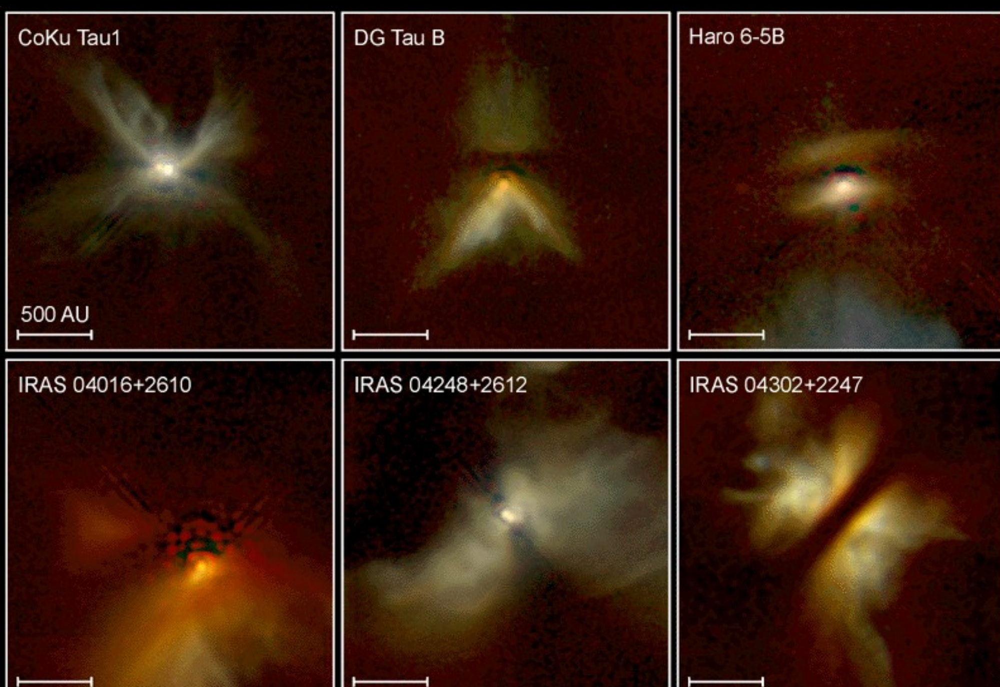
Disks



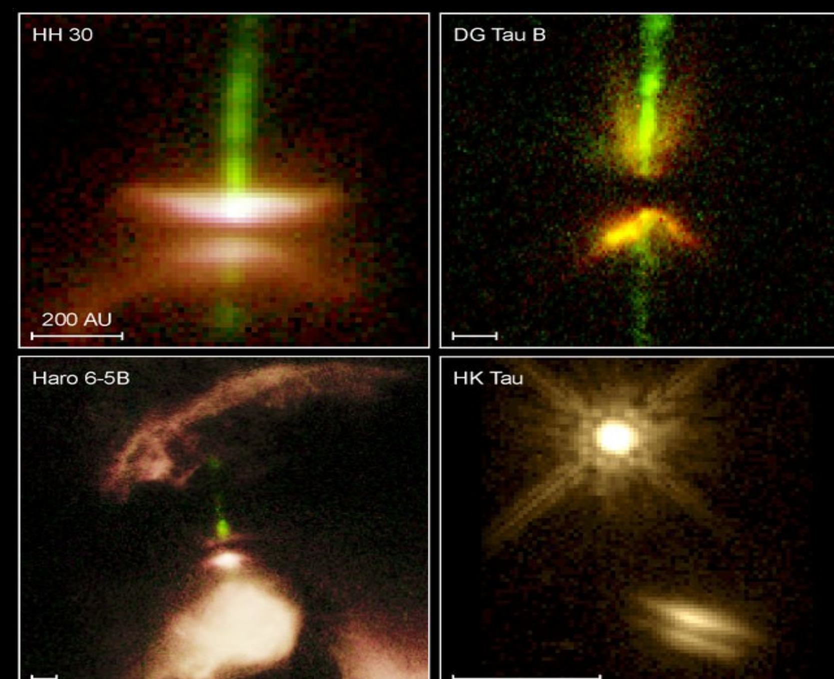
Protoplanetary Disks HST • WFPC2
Orion Nebula
PRC95-45b • ST ScI OPO • November 20, 1995
M. J. McCaughrean (MPIA), C. R. O'Dell (Rice University), NASA

MBM12 3C
Jayawardhana et al. 2002





Young Stellar Disks in Infrared HST • NICMOS
PRC99-05a • STScI OPO
D. Padgett (IPAC/Caltech), W. Brandner (IPAC), K. Stapelfeldt (JPL) and NASA



Disks around Young Stars HST • WFPC2
PRC99-05b • STScI OPO
C. Burrows and J. Krist (STScI), K. Stapelfeldt (JPL) and NASA

β Pic

GG Tau – the ring world

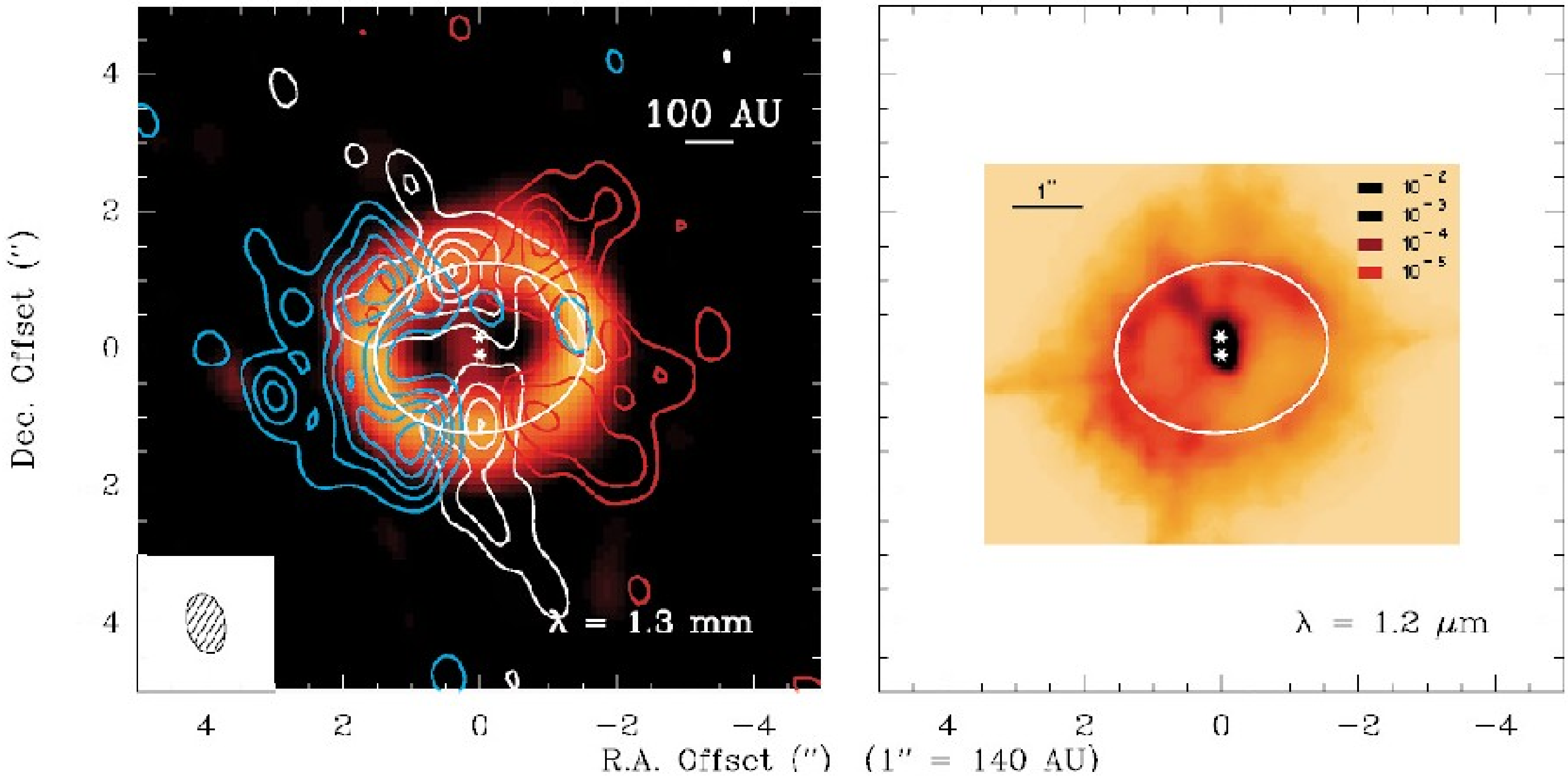
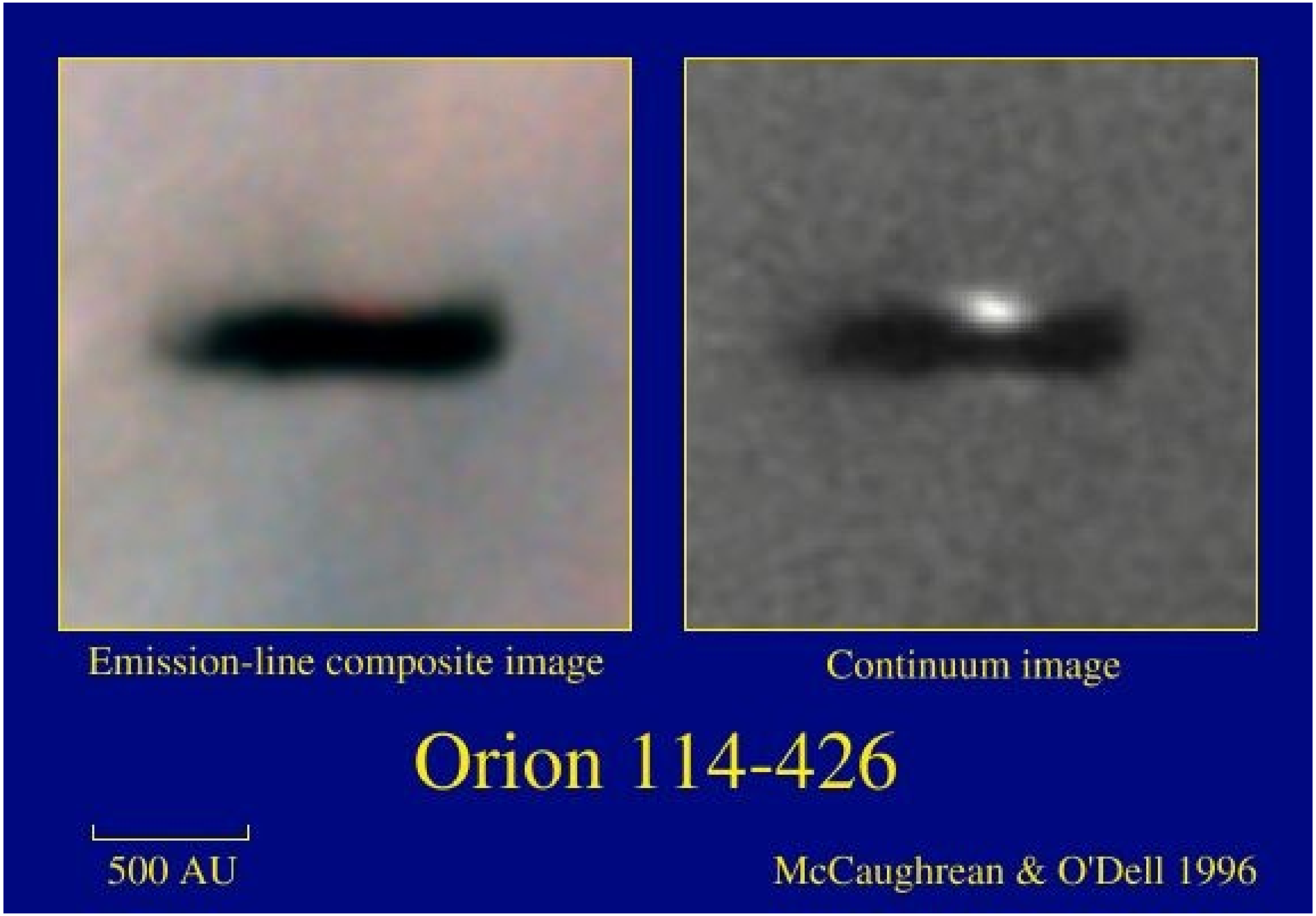
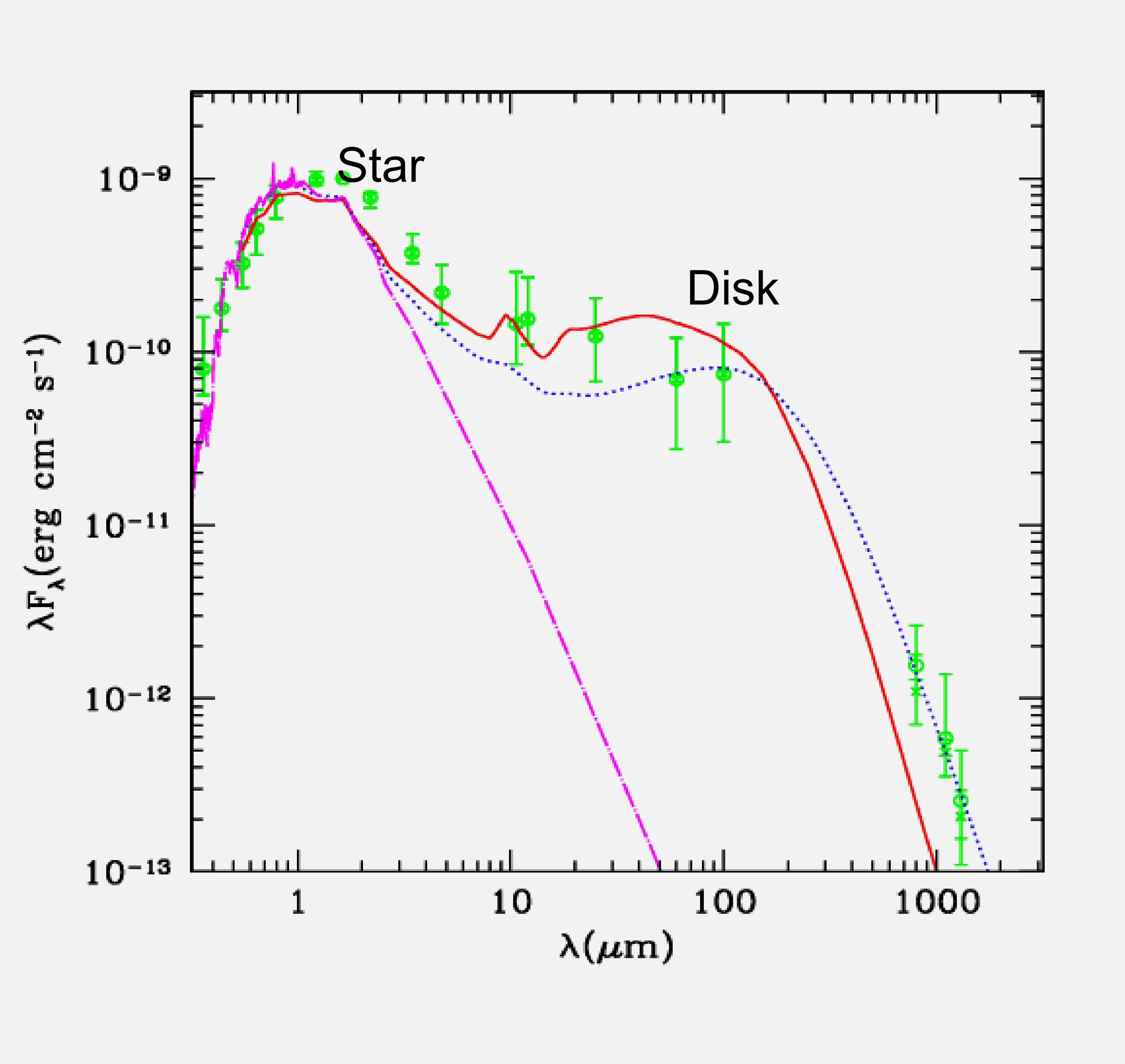


Fig. 6a and b. Comparison of the **a** ^{13}CO J=2–1 line emission at V = 5.55 (blue contours), 6.30 (white contours) and 7.05 km s^{-1} (red contours), overlaid on a false colour image of the 1.4 mm continuum emission and **b** J-band emission resolved by RRNGJ. The white ellipse indicates the ring average radius.

Disk evaporation: proplyds



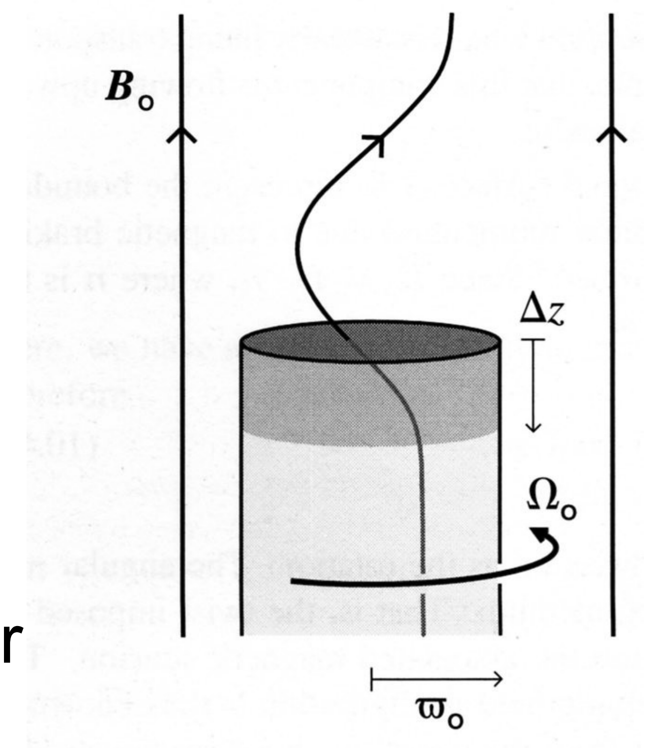
SEDs



Star
Disk
10^{-9}
10^{-10}
10^{-11}
10^{-12}
10^{-13}
$\lambda F_\lambda (\mathrm{erg\ cm^{-2}\ s^{-1}})$
1
10
100
1000
$\lambda(\mu m)$

Magnetic braking

- Existence of an azimutal component of **B.**
- Field line twisting due to rotation (**B** is anchored on large scales).
- This creates a torque which acts against the rotation.
  - Transport of angular momentum by torsional Alfvén waves
- Close to the center, **B** and matter decouple because density increases and thus ionization decreases.
- **B** is not anymore “frozen”.
- Magnetic braking only efficient in outer disk.



Consequence of rotation

Matter that falls towards the center “misses” the center because of the centrifugal forces.

If a particle with specific angular momentum l (=angular momentum per unit mass) falls in and ends up on a circular orbit while maintaining its angular momentum, then the radius r of the orbit is:

$$r = l^2 / GM$$

In the inside-out infall model, all material arriving at the center at a certain time started from the same radius r_0.

If the cloud core is initially in uniform rotation Ω, then the specific angular momentum at r_0 varies with angle θ from the rotation axis:

$$l = \Omega\, r_0^2 \sin\theta$$

Consequence of rotation

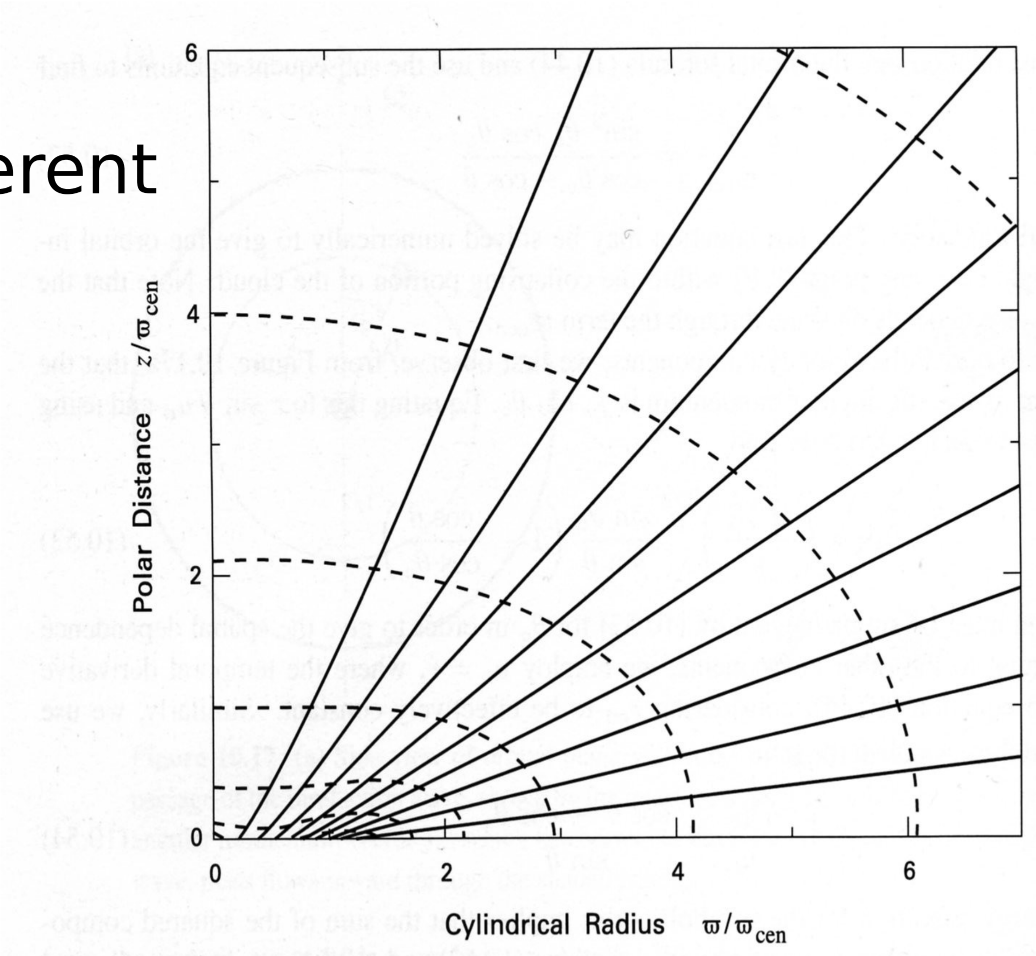
Infalling material from different directions will arrive at the midplane at different radii.

Maximum radius
  "centrifugal radius"

(for $\sin\theta = 1$)

$$r_{cen} = r_0^4\,\Omega^2 / GM$$

→ scale of the protostellar disk



Fluid stream lines (*solid curves*) and isodensity contours (*dashed curves*) within a collapsing rotating cloud.

Orbits

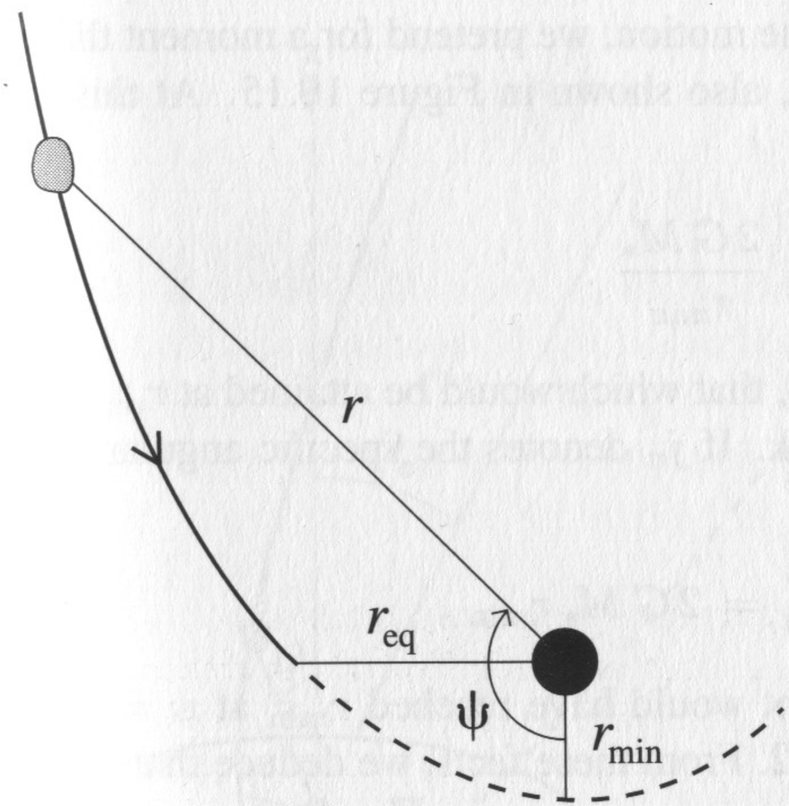
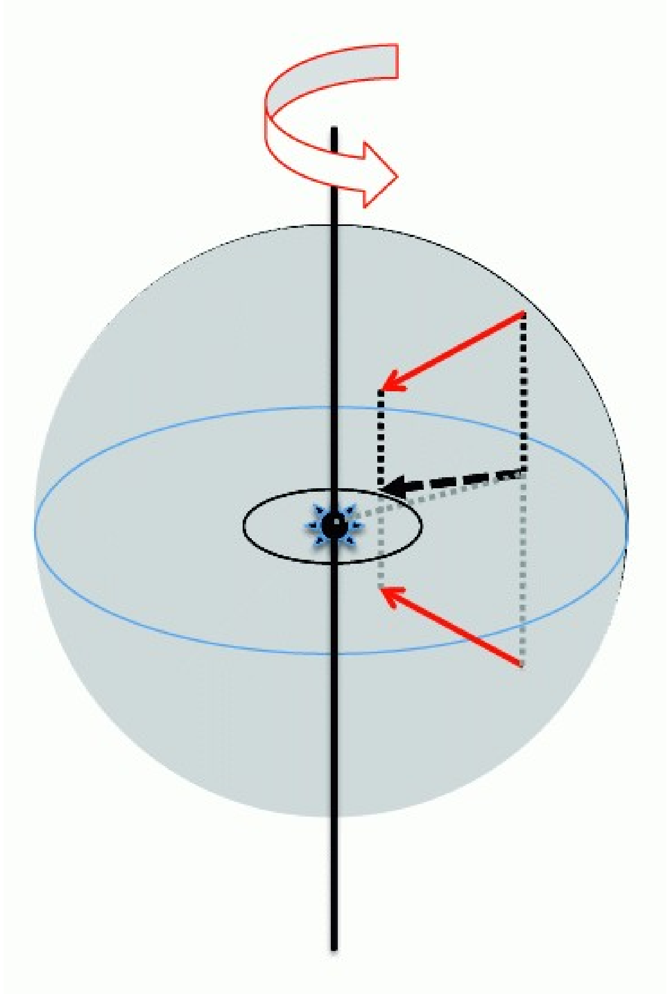


Figure 10.15 Parabolic orbit in rotating infall. A fluid element, with instantaneous polar coordinates (r, ψ) in the orbital plane, falls into the radial distance $r_{\rm eq}$, where it impacts the disk. If the disk were absent, the element would have reached the smaller distance $r_{\rm min}$ before swinging back out.

- Mass does not reach Keplerian perihel, but closes radius where it hits the disk

Rotating infall

Axisymmetry + equatorial symmetry



• material falling from above and below the equatorial plan meet on the equator through a shock which dissipates the vertical kinetic energy.

• If cooling efficient: material accumulates and forms a disk

Most of the matter will continue towards the center, whereas some angular momentum is transported outwards by viscosity.

Typical sizes: >100 AU

Streamlines in disks

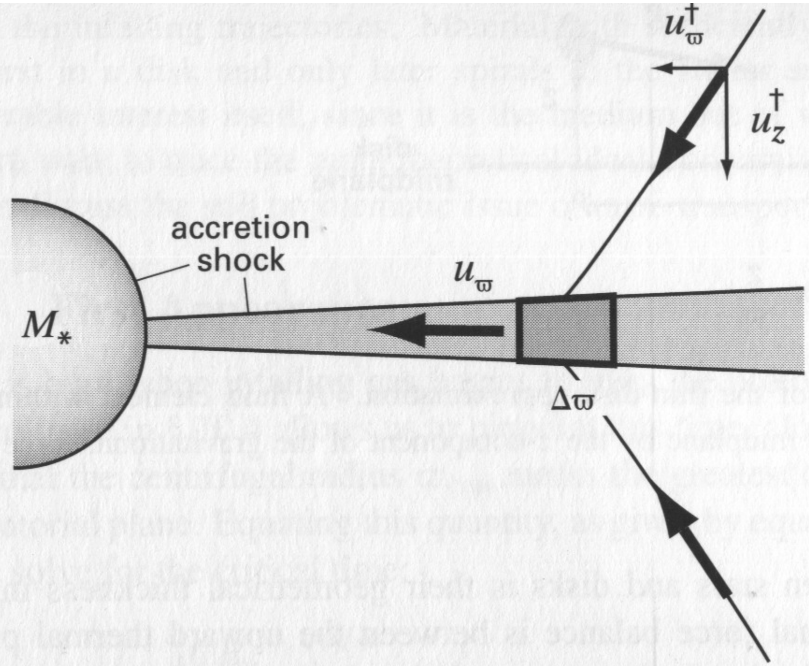


Figure 11.11 Collision of streamlines during rotating, protostellar collapse. The velocity of any infalling element consists of vertical and horizontal components. An element within the disk acquires a horizontal velocity u_ϖ from impact of the external streams. The accretion shock now covers both the protostellar and disk surfaces.

- The accretion shock occurs on the surface of the disk as well as on the protostar
- Radial infall in the disk

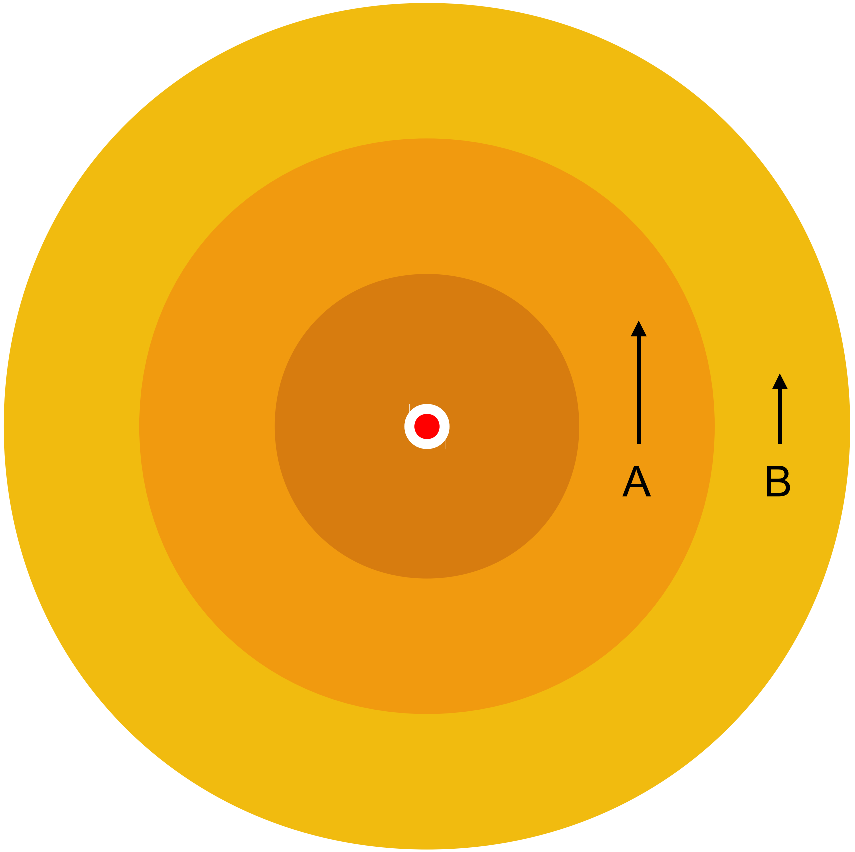
The angular momentum problem

- Angular momentum of 1 $M_{\odot}$ in 10 AU disk: 3×10^{49} m^2/s

Angular momentum of 1 $M_{\odot}$ in 1 $R_{\odot}$ star: << 6×10^{47} m^2/s
(=breakup-rotation-speed)

- Original angular momentum of disk = 50x higher than maximum allowed for a star
- Angular momentum is strictly conserved!
- Two possible solutions:
  - Torque against external medium (via magnetic fields?)
  - Very outer disk absorbs all angular momentum by moving outward, while rest moves inward. Need *friction through viscosity*!

Angular momentum transport



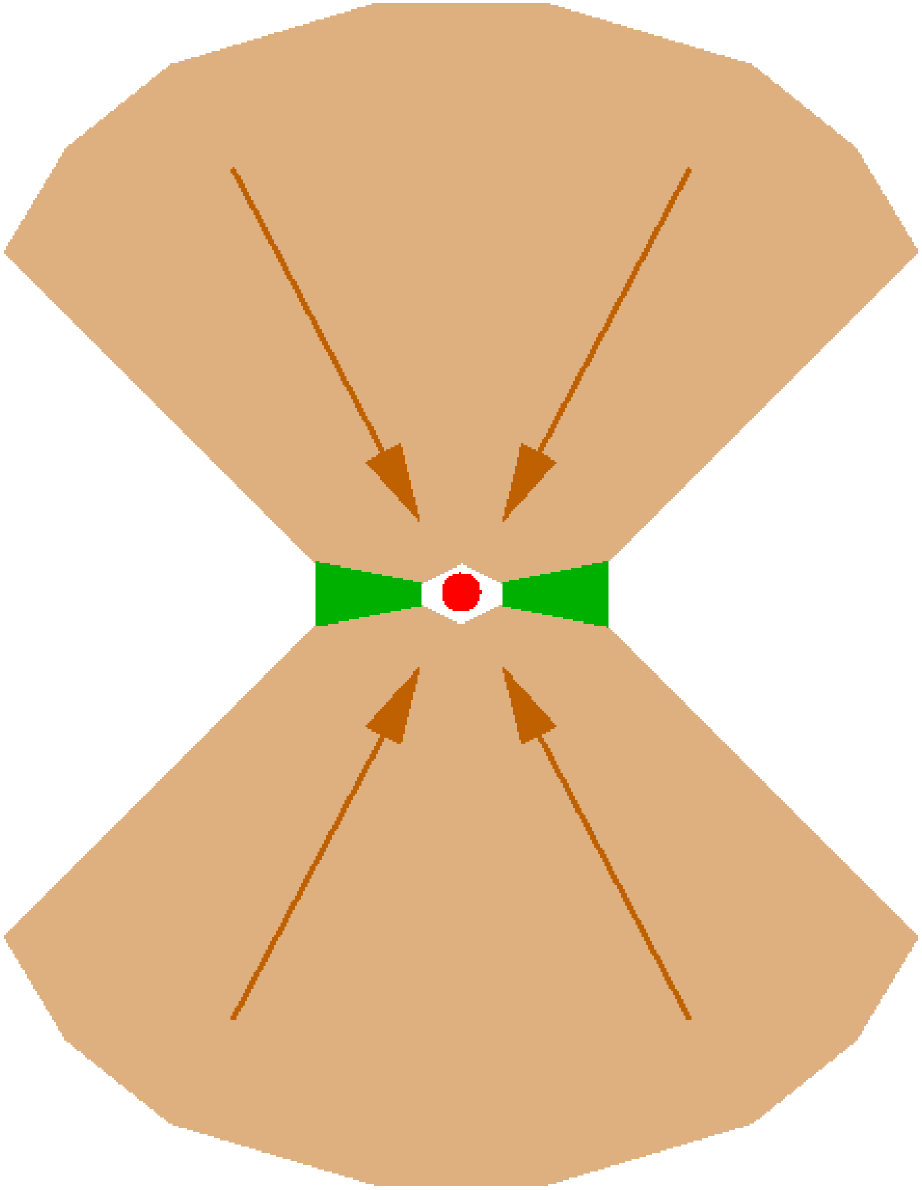
- Ring A moves faster than ring B.
- Friction between the two will try to slow down A and speed up B.
- Angular momentum is transferred from A to B.

Specific angular momentum for a Keplerian disk:

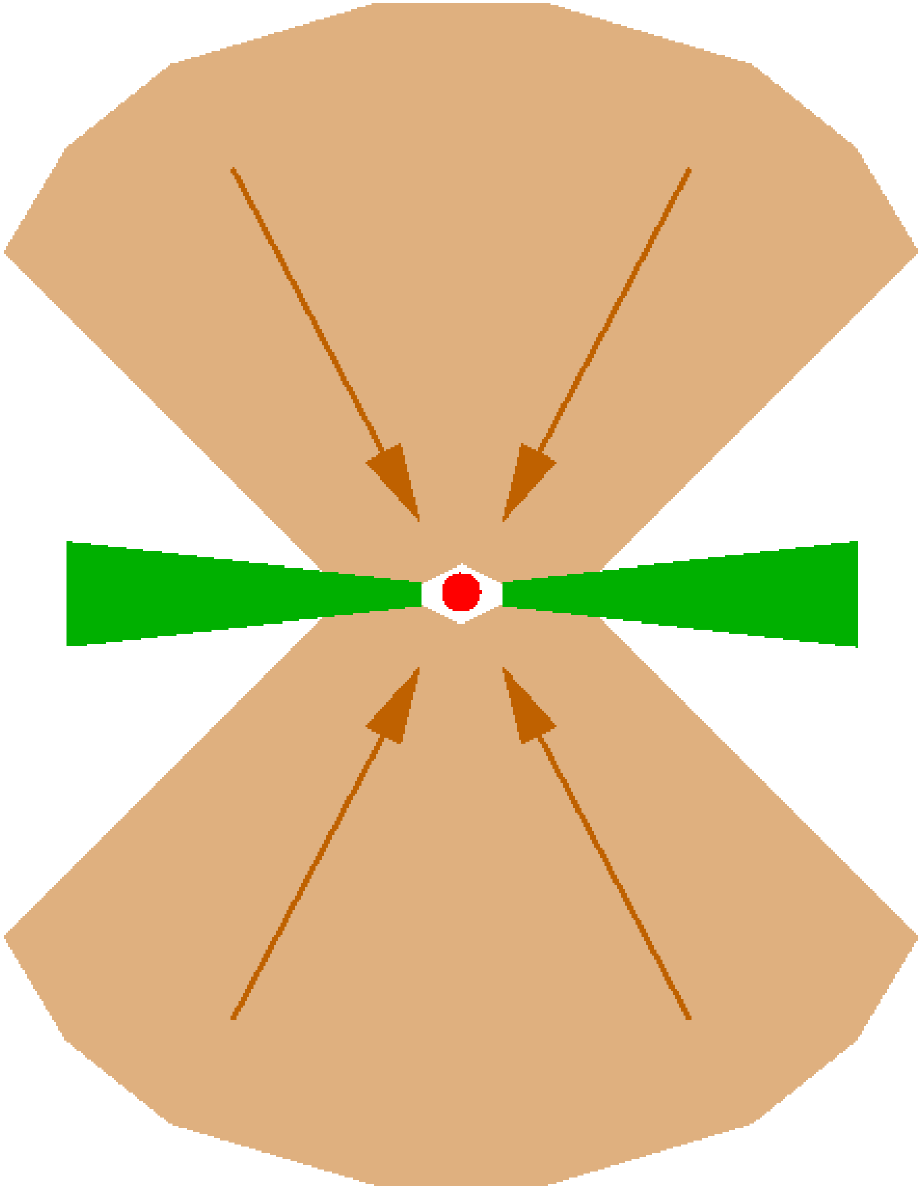
$$l = r v_f = r^2 \Omega_K = \sqrt{GM\, r}$$

So if ring A looses angular momentum, but is forced to remain on a Kepler orbit, it must move inward! Ring B moves outward, unless it, too, has friction (with a ring C, which has friction with D, etc.).

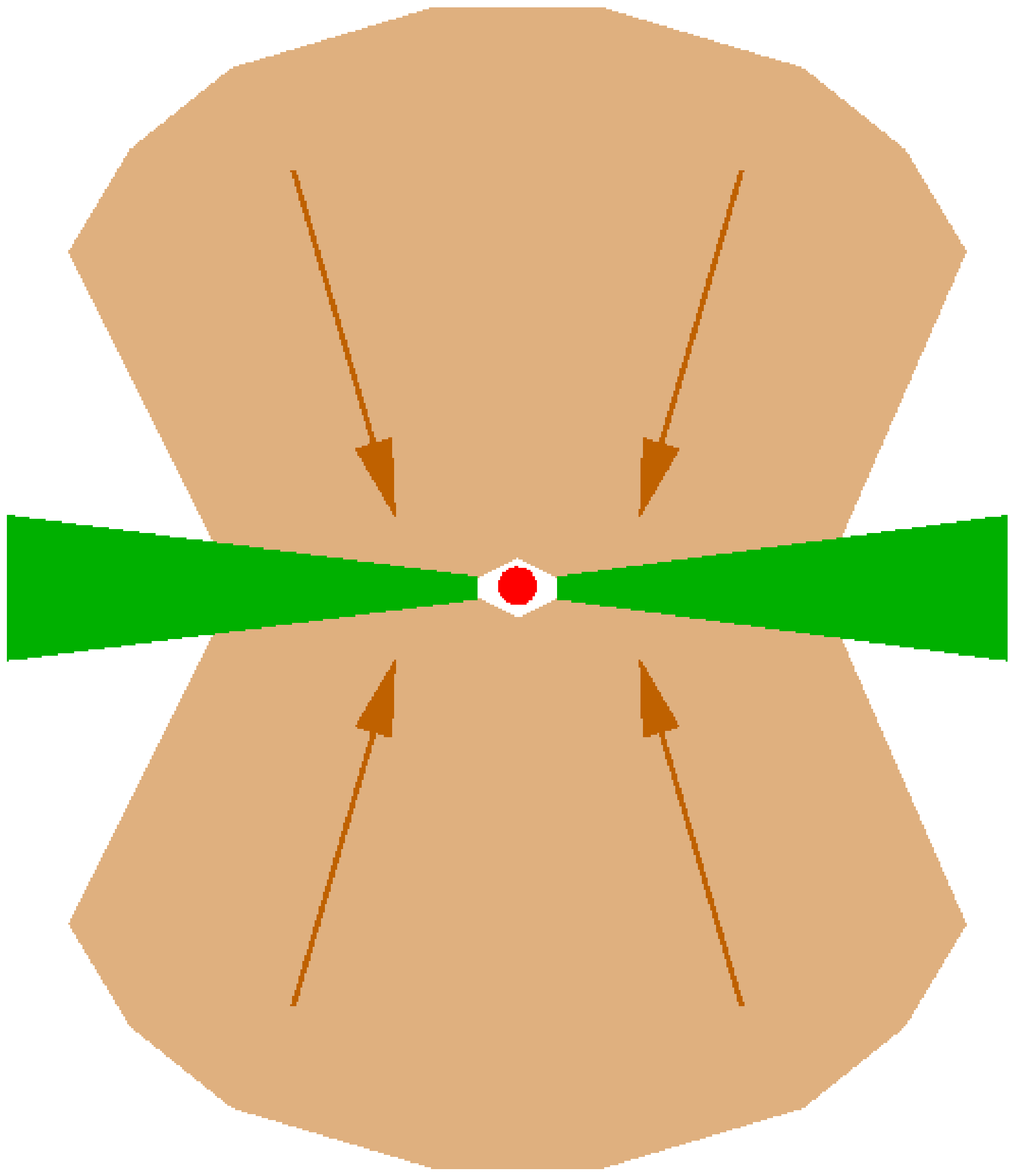
Formation & spreading of disk



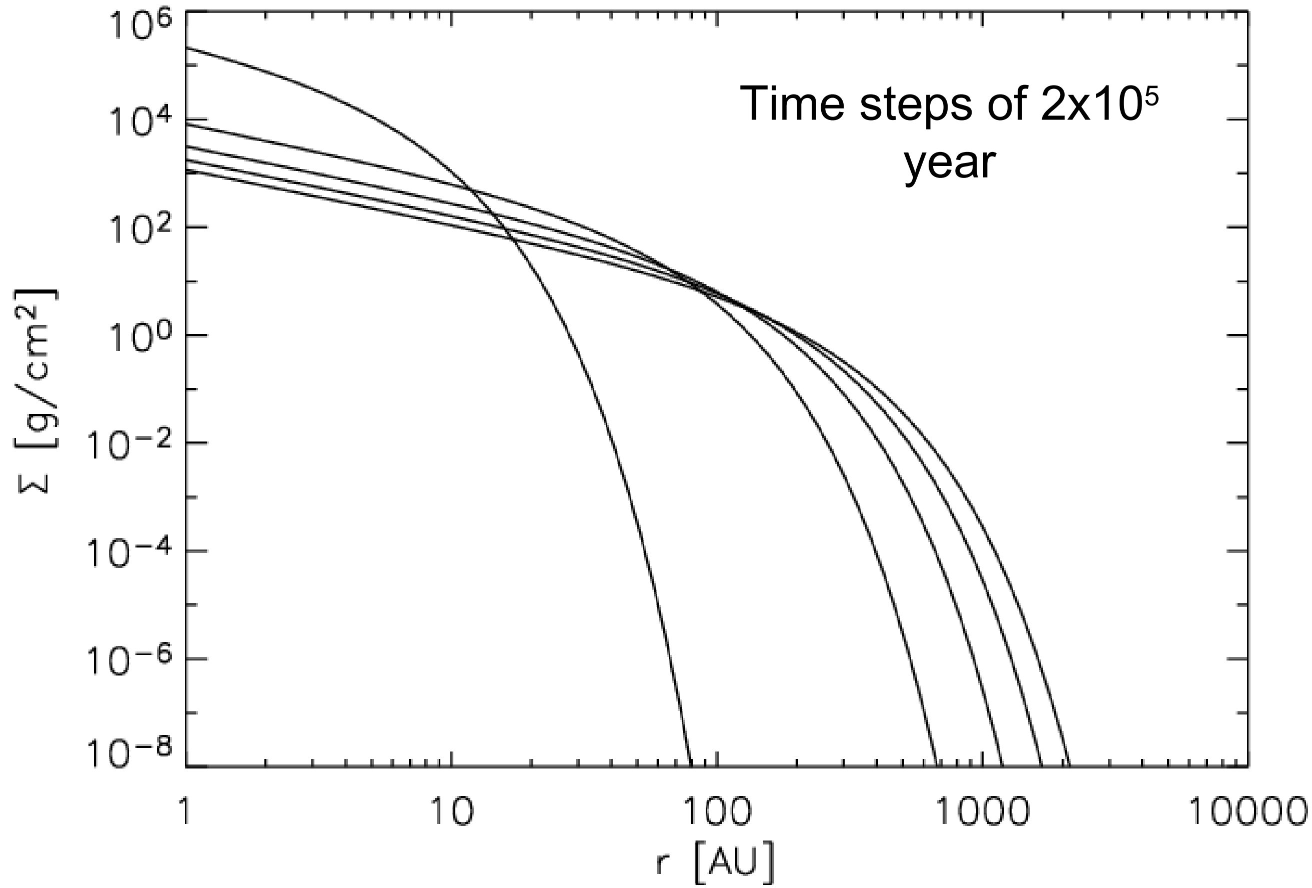
Formation & spreading of disk



Formation & spreading of disk



Non-stationary (spreading) disks



Lynden-Bell & Pringle (1974), Hartmann et al. (1998)

$$r_{cen} \approx \frac{c_s \Omega_0^2 t^3}{16} = 0.3\ AU \left(\frac{T}{10\,K}\right)^{1/2} \left(\frac{\Omega_0}{10^{-14}\,s^{-1}}\right)^2 \left(\frac{t}{10^5\,yrs}\right)^3$$

Birth of the disk

Critical time determined by $r_{cen} = R_*$

$$t_0 = \left(\frac{16 R_*}{\Omega_0^2 v_s}\right)^{1/3} = 3\times 10^4 [yr] \left(\frac{R_*}{3_\odot}\right)^{1/3} \left(\frac{\Omega_0}{10^{-14} s^{-1}}\right)^{-2/3} \left(\frac{v_s}{0.3\, km\, s^{-1}}\right)^{-1/3}$$

the mass of the protostar is $M_{¿} = \dot{M} t_0$

leading to initial stellar mass:

$$M_0 = \left(\frac{16 R_* v_s^8}{G^3 \Omega_0^2}\right)^{1/3} = 0.2\, M_\odot \left(\frac{R_*}{3_\odot}\right)^{1/3} \left(\frac{\Omega_0}{10^{-14} s^{-1}}\right)^{-2/3} \left(\frac{v_s}{0.3\, km\, s^{-1}}\right)^{8/3}$$

Disk evolution

- Viscuous angular mo-mentum transport in a fully evolved accretion disk enables efficient accretion
- Accretion fades with outer supply

→ birth of PMS

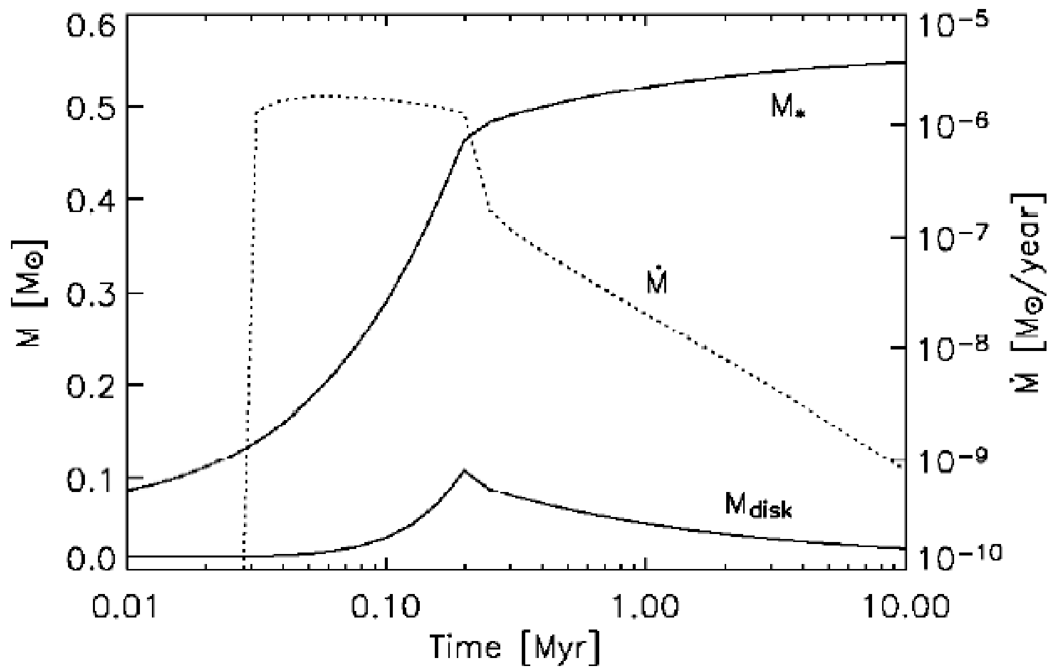


Fig. 1.— Evolution of various disk and star quantities as a function of time after the onset of collapse of the cloud core (after *Hueso and Guillot*, 2005). Solid lines: stellar mass (upper) and disk mass (lower). Dotted line: accretion rate *in the disk*. In this model the disk is formed at $t \simeq 0.03$ Myr, causing the jump in the dotted line at this point. The collapse phase is finished by 2×10^5 years.

Hueso & Guillot (2005)

Disk evolution

- Spiral patterns expands with t^3
- At time $t_1 = 1.43\ t_0$, the streamlines start missing the protostar (at a radius of 0.34 r_{cen})
- boundary between (tenuous) outer disk and (dense) inner disk with almost circular orbits

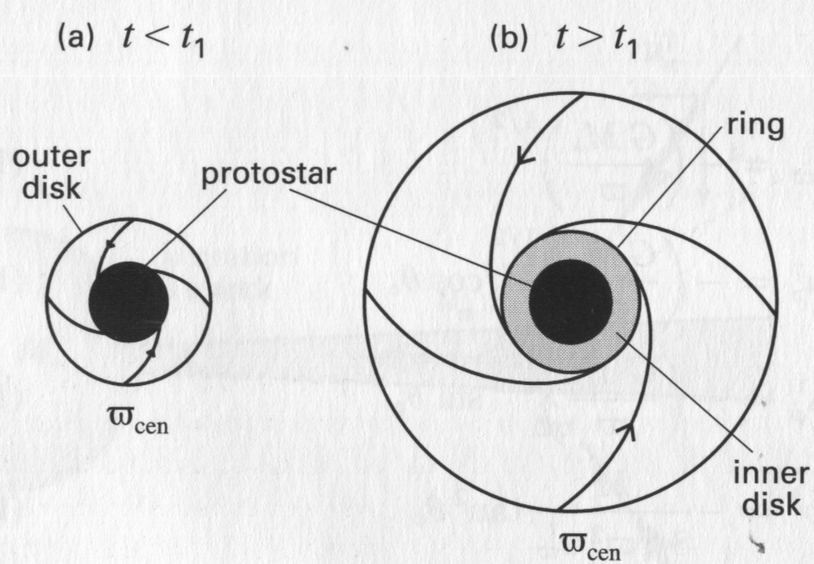


Figure 11.12 Early expansion of a protostellar disk. (a) Before time t_1, curved streamlines from the outer disk impact the protostar directly. (b) After t_1, these streamlines converge to form a dense ring, which transfers mass to an inner disk surrounding the star. At all times, the outer disk boundary is the growing centrifugal radius ϖ_{cen}.

Disk evolution

Mass transport rate:

- Standard 2-D disk models provide a constant accretion rate across the disk

- Spirals lead to spatial and temporal discontinuities

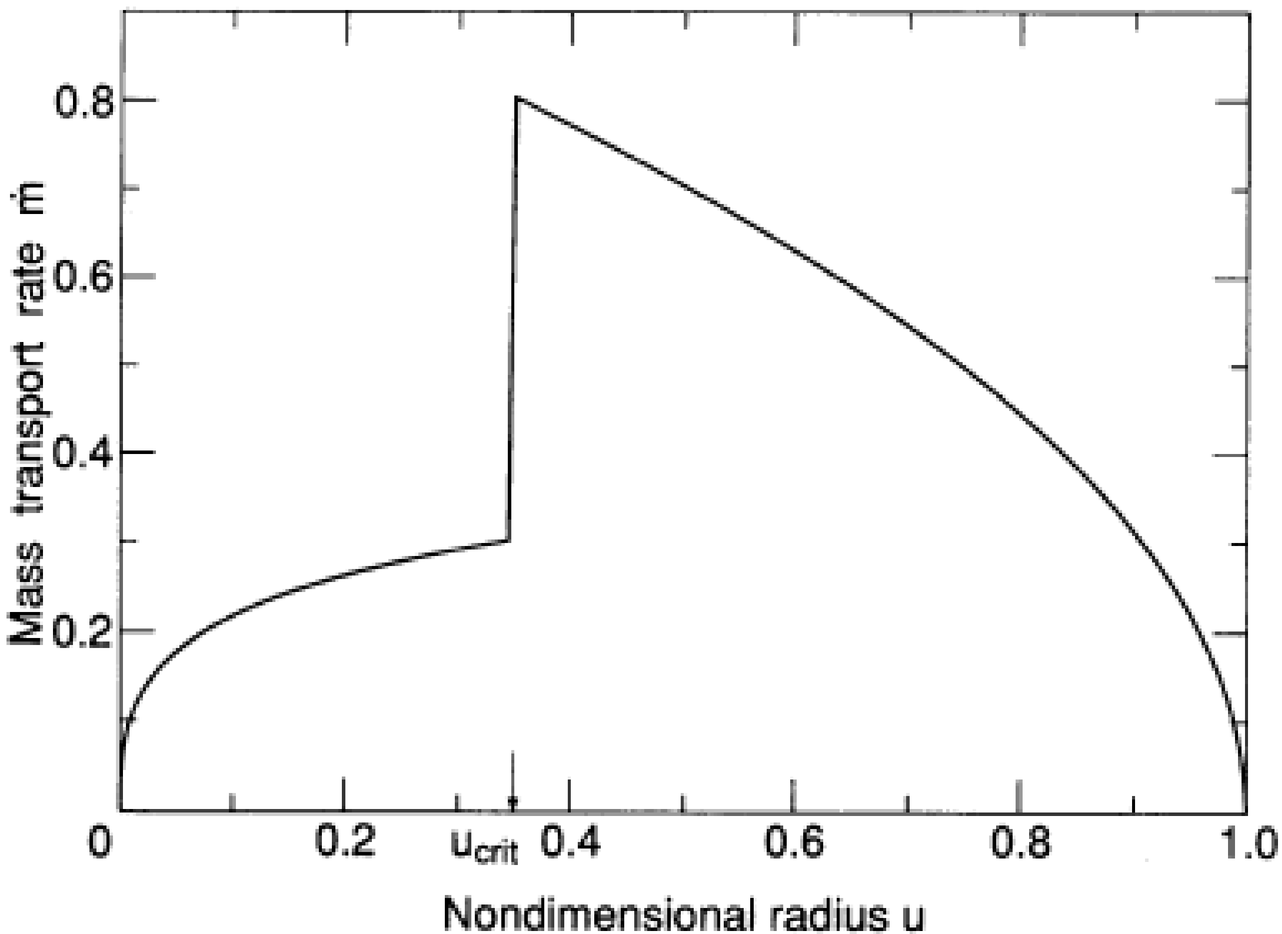


FIG. 11.—Mass transport rate in the disk. Displayed are both $\dot{m}_{\rm inner}$ and $\dot{m}_{\rm outer}$, as derived in the text. Note the discontinuity in the transport rate at $u = u_{\rm crit}$, where the ring is being built up. As in Fig. 10, the true inner boundary is at $u = u_*$.

Accretion in an inhomogeneous disk

Temporal evolution of mass accretion rate:

FU Orionis outbursts:

• from spiral inflow channel

•Clumpy disk

→ still not fully understood

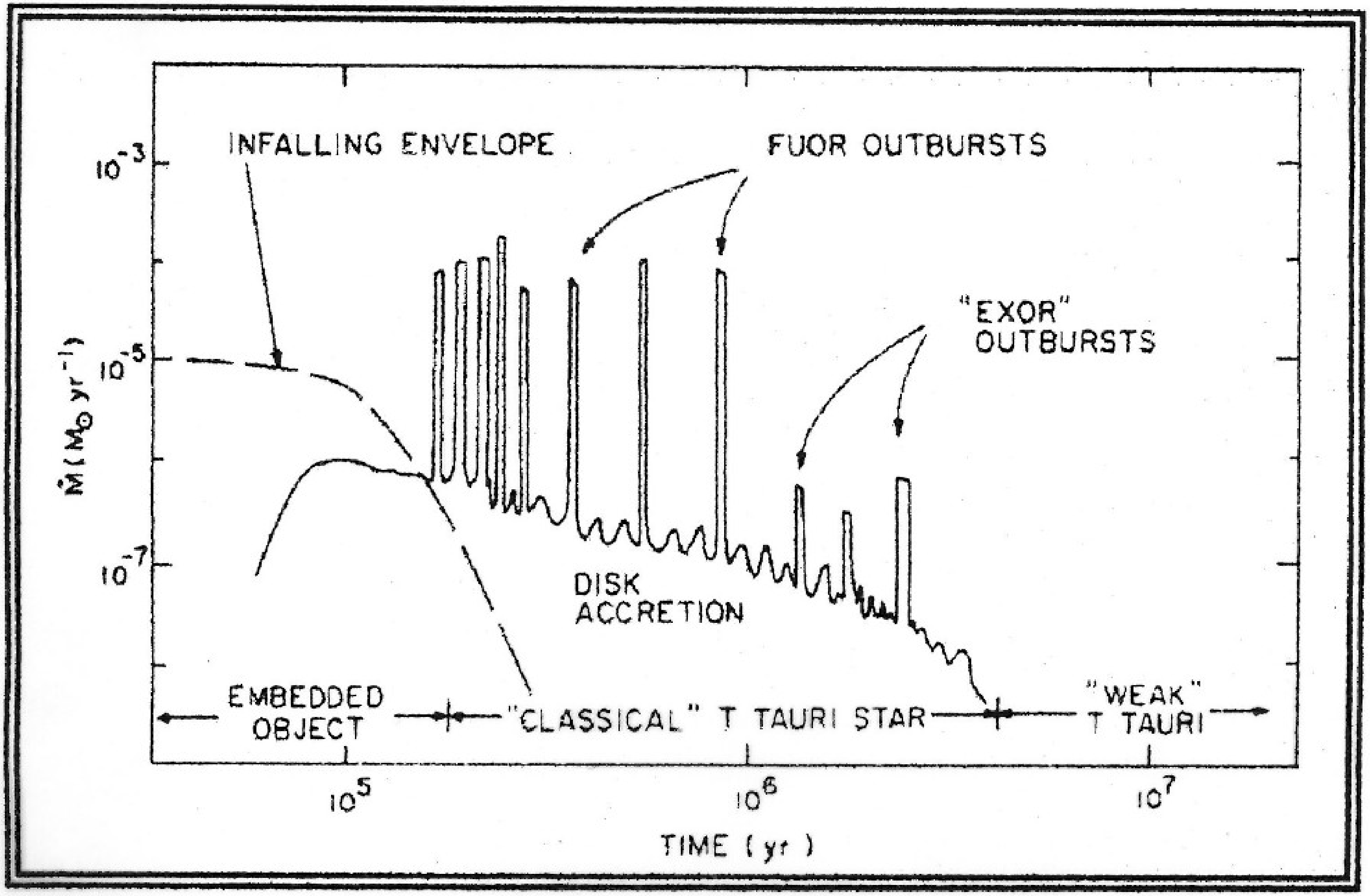


Abb.12 Der zeitliche Verlauf der Akkretionsrate

FU Orionis outburst

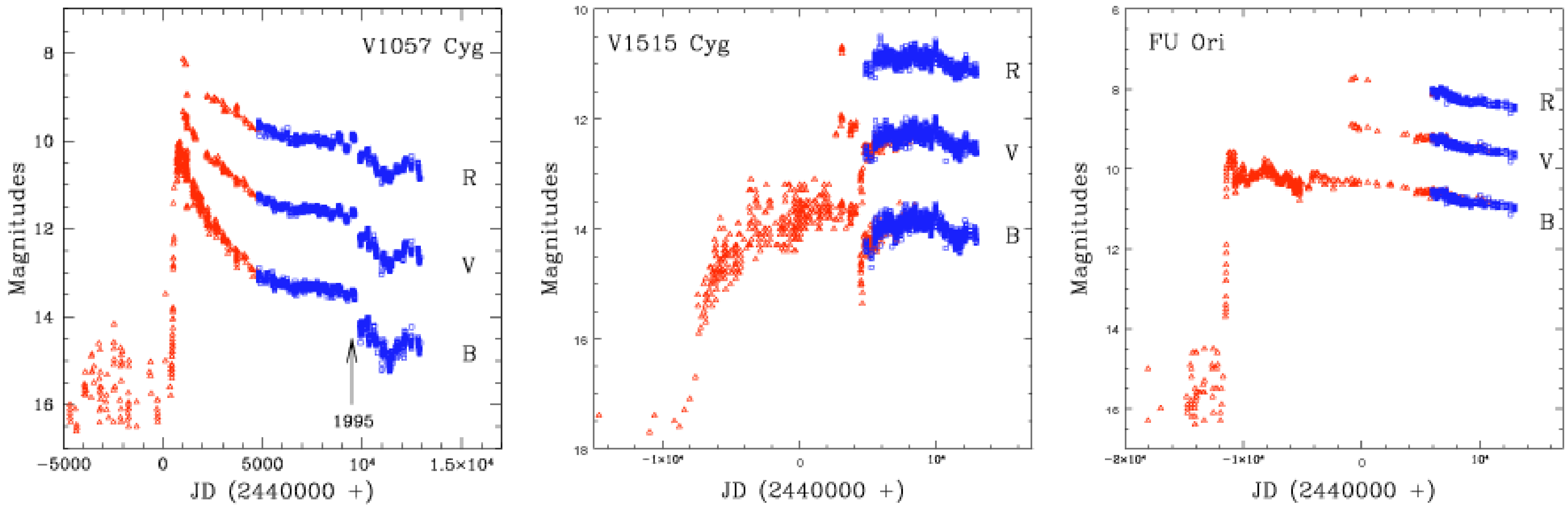


Figure 1. Light curves of V1057 Cyg (left), V1515 Cyg (middle) and FU Ori (right). Squares indicate our new data, while triangles indicate historical data taken from Mendoza (1971), Rieke, Lee & Coyne (1972), Schwartz & Snow (1972), Bossen (1972), Welin (1975, 1976), Landolt (1975, 1977), Kolotilov (1977), Kopatskaya (1984) and Ibragimov & Shevchenko (1988).

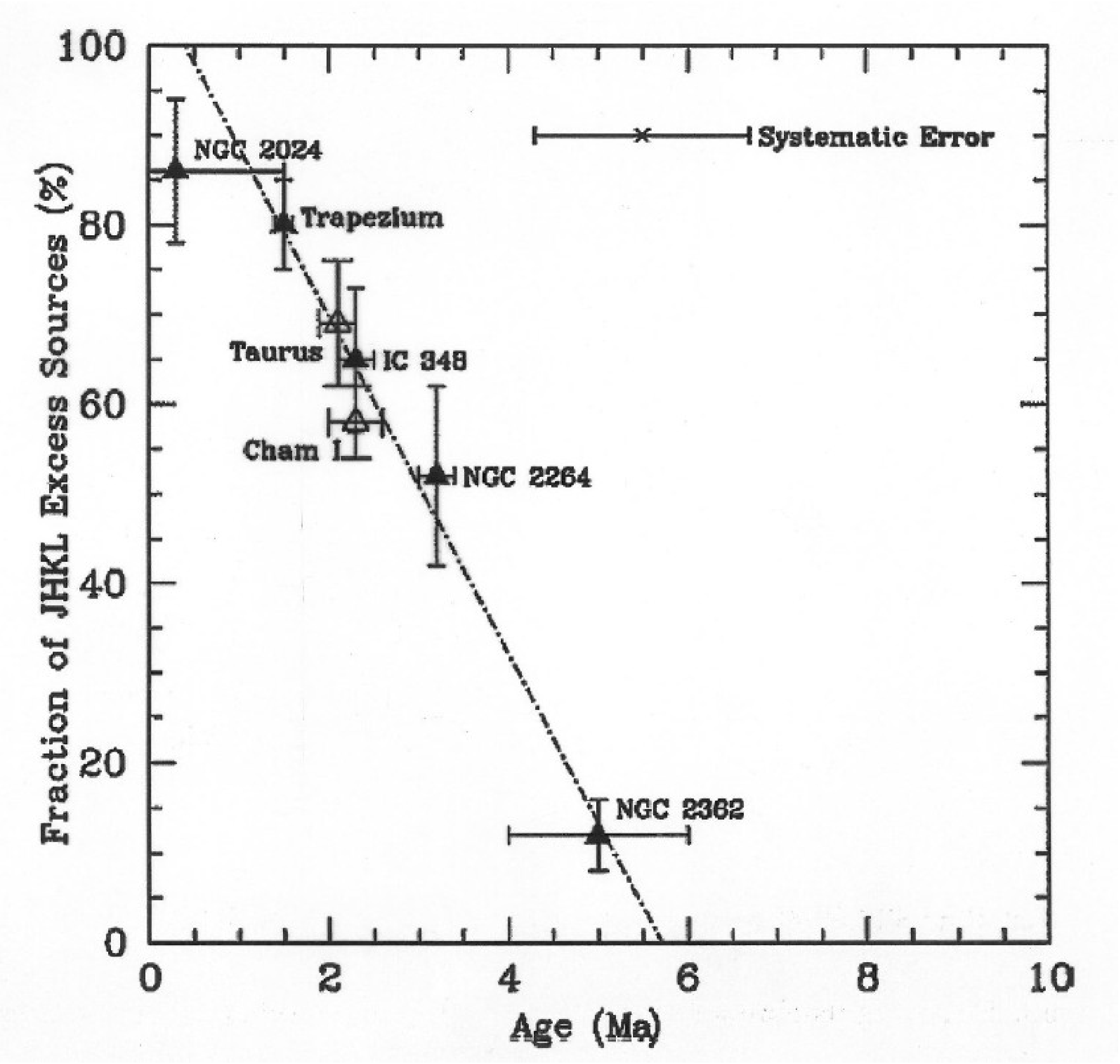
- Mass accretion rates varies from 10^{-7} $M_\odot$/yr to 10^{-4} $M_\odot$/yr
- Origin: mass pileup in disk – relation to spirals unclear
- Thermal or flow instabilites probably trigger outbursts

Disk evolution

Disk dispersal:

Observations of IR excess show that the disk life time is only a few million years.

Sets the time scale for
• Angular momentum transfer
• Planet formation



Haisch et al. 2000

Angular momentum transfer

Angular momentum:

Accretion from the Keplerian disk adds angular momentum to the pre-main sequence star.

It takes 10^8a to remove the angular momentum by stellar radiation.

→ Young stars rotate much faster than main sequence stars.

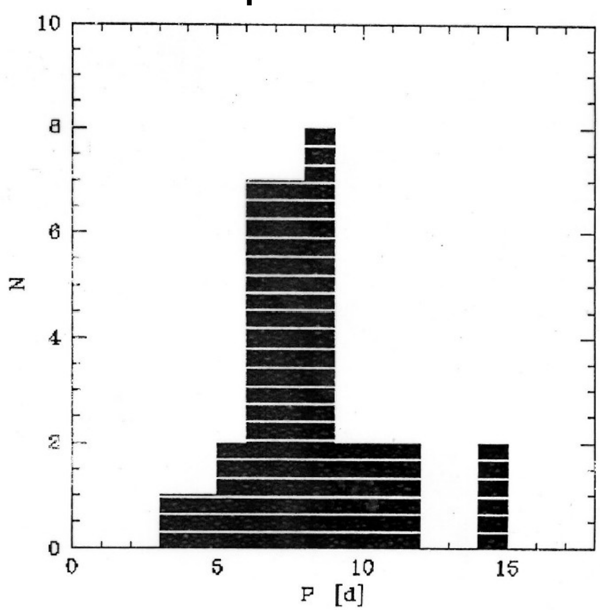


Fig. 36. The distribution of rotation periods for 34 classical T Tauri stars. (From Bouvier et al. [13])

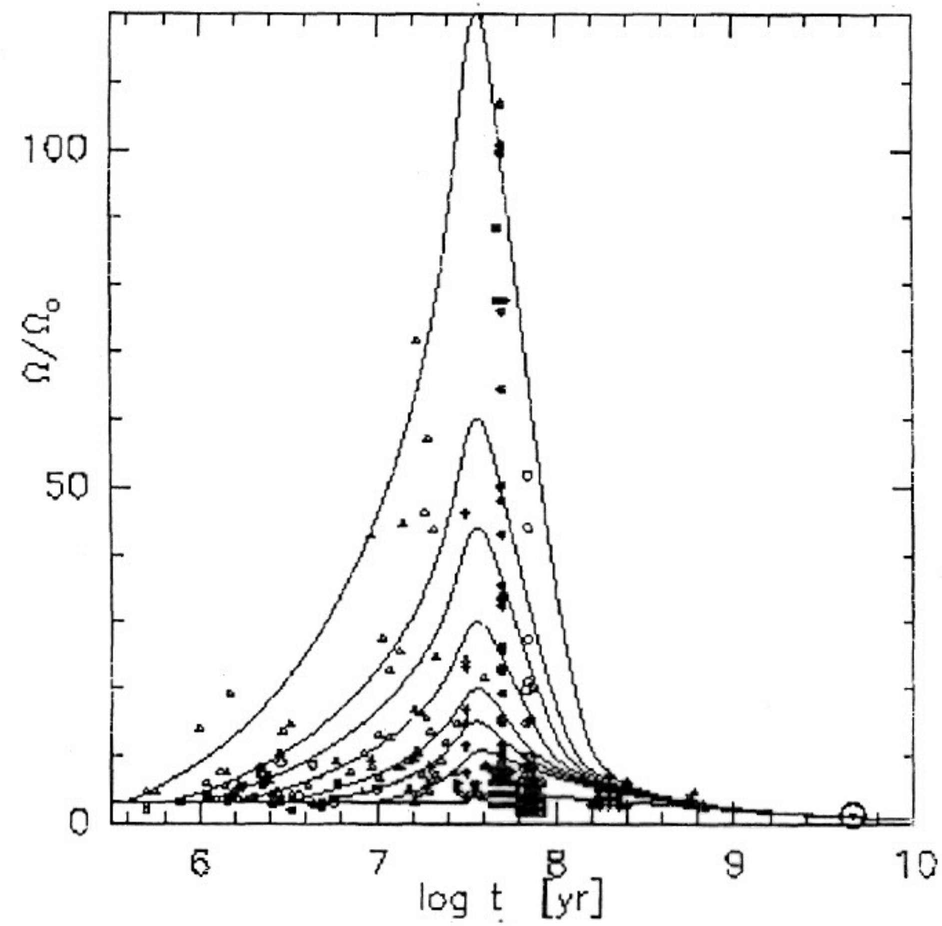
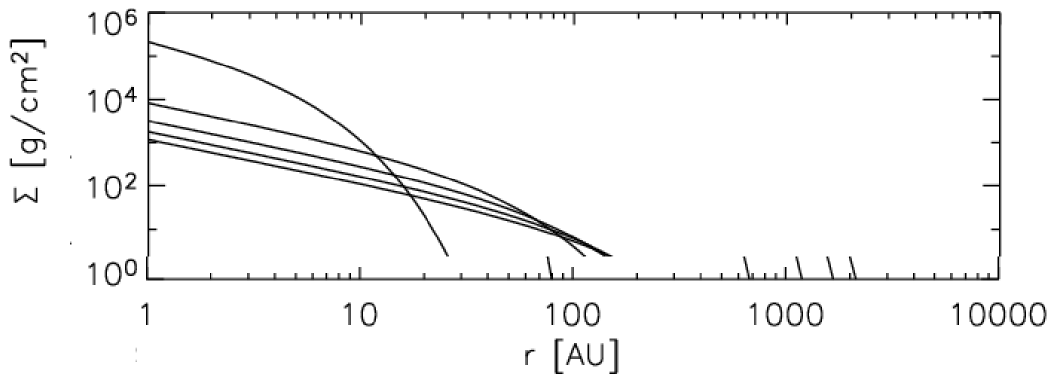


Fig. 37. The evolution of the angular velocity Ω for a 1 $M_{\odot}$ star as a function of time for various disk life times. The uppermost curve is for stars with short lived disks (4×10^5 yr), while the lowest one is for long-lived disks (30×10^6 yr). The observed angular velocities in a variety of T Tauri stars and open clusters are shown by the various symbols. (From Bouvier et al. [10])

Typical TTau rotation at 1/10 of disruption velocity

The density structure

Consider density scaling in inner main part:



- Self-similar collapse in 2-D would provide $\sigma \sim r^{-1.5}$
- Viscous disk:
  - Density structure determined by flow of material: $$\sigma = \frac{\dot{M}}{3\pi\nu}$$
  - Assumes constant flow of material throw the disk (continuity)
  - Viscosity ν determines mass and angular momentum transfer

The disk viscosity

Molecular viscosity: $\nu = \langle v_T \rangle \lambda_{\mathrm{free}} \approx c_s \lambda_{\mathrm{free}}$

•λ_{free} = mean free path of molecule

Typical disk (at 1 AU): $N = 1\times10^{14}$ cm^{-3}, T=500 K, L=0.01AU

$\sigma_{H_2} \approx \pi(1\text{Å})^2 \rightarrow \lambda_{\mathrm{free}} = 1/(N\sigma) \approx 32\,\mathrm{cm}$

$\rightarrow \quad \nu = 730\,\mathrm{m}^2/\mathrm{s}$

- Extremely low
- Would allow a maximum accretion rate of 3 10^{-12} $M_\odot$/a for a 1$M_\odot$, 100AU disk
- Actual viscosity must be > 10^6 times higher
  → provided by turbulent, not thermal motion

The disk viscosity

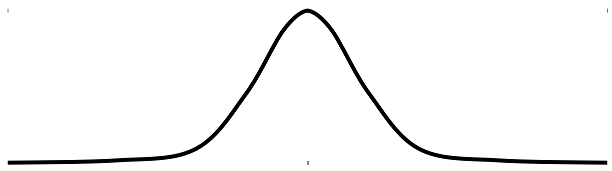
Turbulent viscosity: Shakura-Sunyaev model

- Dissipation provided at scale of minimum eddy size
- Characteristics: scale height h, sound speed c_s
- Additional "efficiency" factor α

$$\nu = \alpha c_s h \qquad \alpha = 0.001 \ldots 0.1$$

- Scale height of disk:
  - Solution of hydrostatic balance in z-direction
  - $\rho(z) = \rho_0 \exp\left(-\frac{z^2}{2h^2}\right)$ with $h = \sqrt{\frac{kTr^3}{\mu m_p G M_x}} = \frac{c_s^2}{\Omega_{Kepler}}$

A Gaussian!

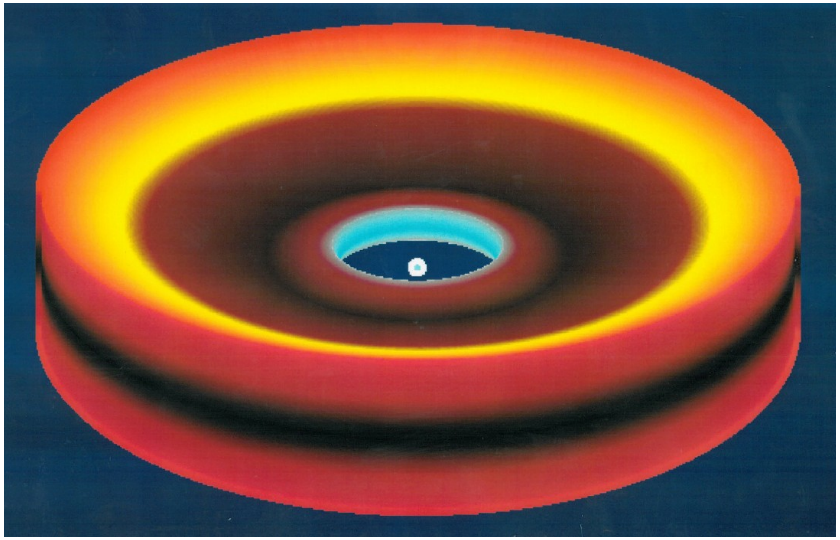


The density structure

Disk height: $\Omega_{\mathrm{Kepler}} \propto r^{-3/2}$

$\rightarrow$ $h \propto r^{3/2}$ for constant temperature

$\rightarrow$ "Flared disk"



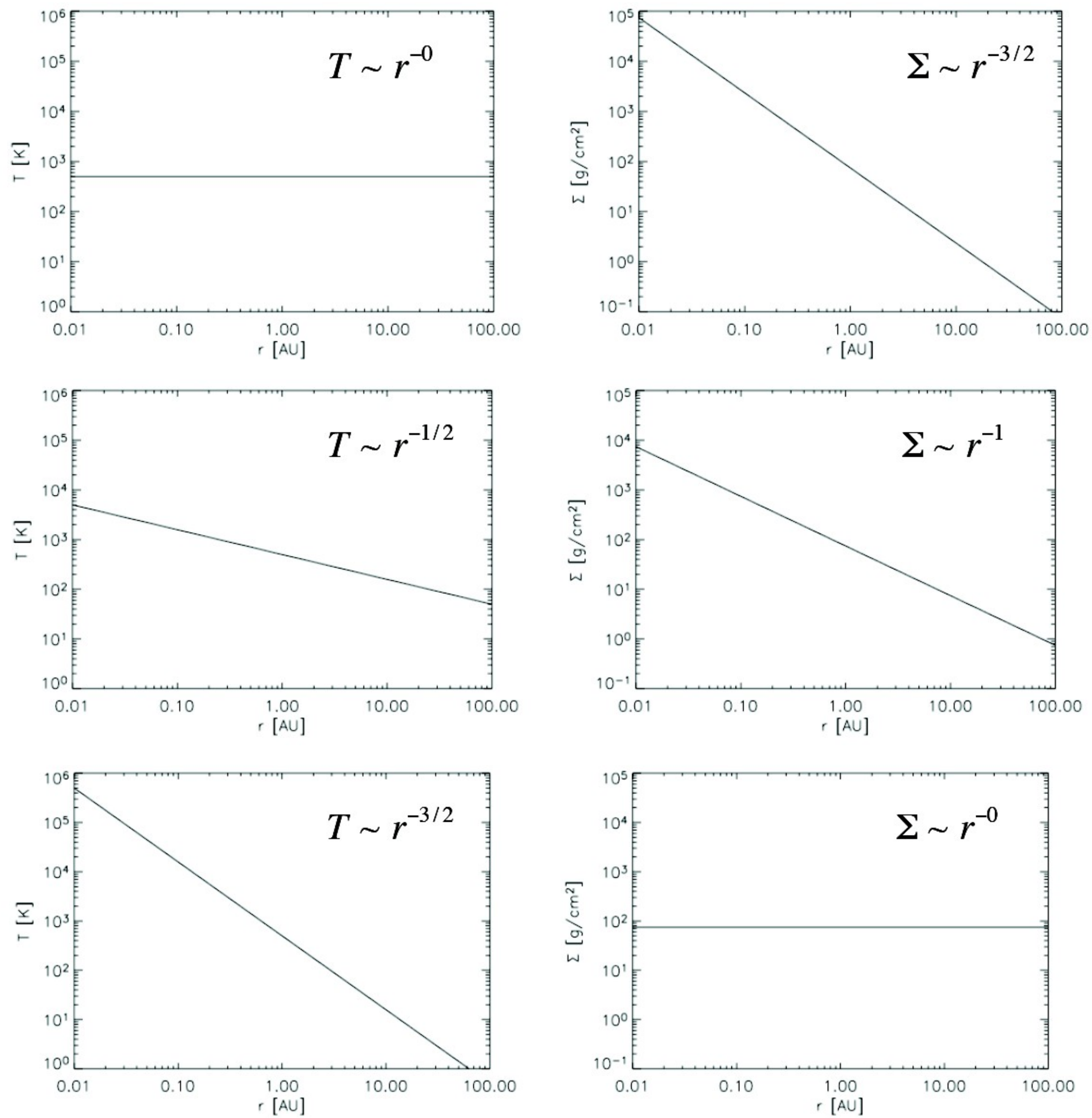
Flared disk with inner rim from heating

- Resulting surface density $\sigma = \frac{\dot{M}}{3\pi\nu} \propto r^{-3/2}$

  - However, this ignores all heating processes by assuming constant c_s.

The density structure

Temperature structure
→ disk height
→ density structure:



The temperature structure

2 Heating processes:

- Viscous heating from the friction within the disk
  - Only relevant for massive, optically thick disks
- Irradiation of the surface from the protostar/PMS

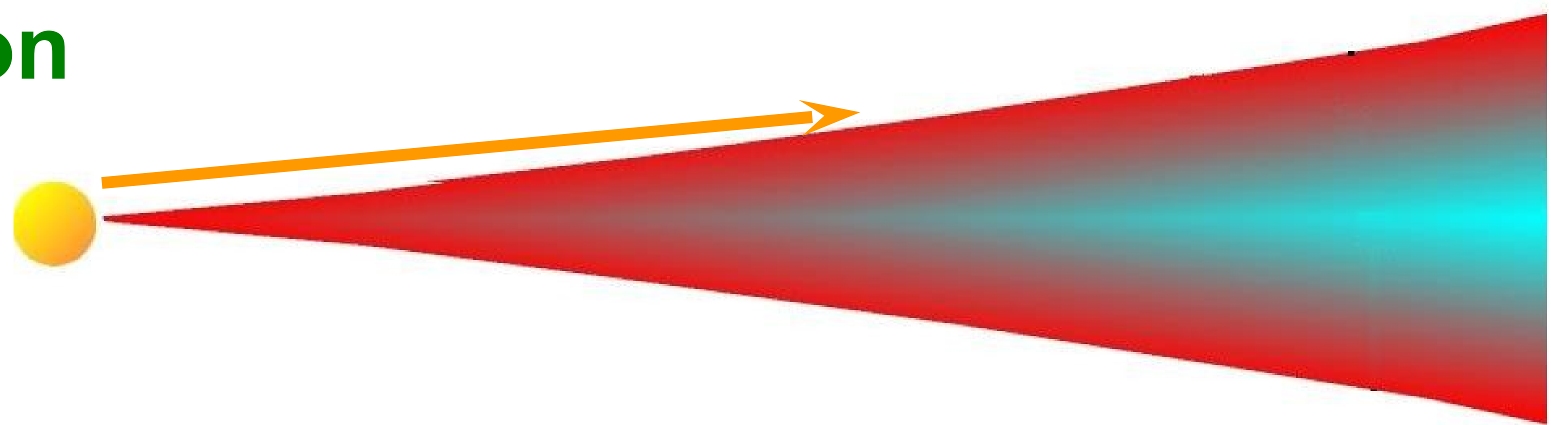
Viscous heating:

- Dissipation in shear flow: $$\dot{E} = \nu\sigma \left| r \frac{\partial \Omega}{\partial r} \right|^2$$
- Keplerian orbits: $$\frac{\partial \Omega}{\partial r} = -\frac{3}{2}\sqrt{\frac{GM_\star}{r^{5/2}}}$$
- Radiative cooling in optically thick disk: $$\dot{E} = -2\sigma_S T^4$$

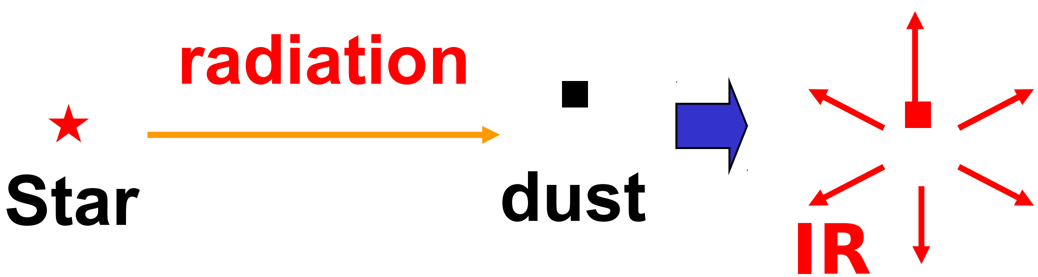
$$\rightarrow \quad T = \left(\frac{9GM_\star\nu\sigma}{8\sigma_S r^3}\right)^{1/4} = \left(\frac{3GM\star\dot{M}}{8\pi\sigma_S r^3}\right)^{1/4} \propto r^{-3/4}$$

Radiative heating

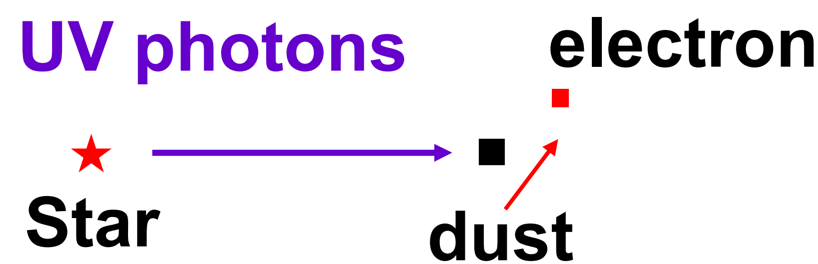
Irradiation



Dust heating



Gas heating



Irradiation from the central star heats gas & dust in disks

- $T \sim r^{-1/2}$ for optically thick disk
- $T \sim r^{-2/(4+\beta)}$ for optically thin disk with $\dot{E} = -2\sigma_S T^{4+\beta} \frac{Q_0 \sigma}{2}$

Temperature Structure

Verification by observations:

- Observable spectrum from unresolved disk
  - Determine spectral index $\nu S_\nu(\nu) \propto \nu^\alpha$
  - Most T Tau disks: $\alpha \approx 0$
- Theoretical spectrum of optically thick disk:

$$S_\nu(\nu) = \int_{R_{in}}^{R_{out}} \frac{2h\nu^3}{c^2 \left[\exp\left(h\nu/kT\right) - 1\right]} \frac{2\pi R}{D} \frac{dR}{D}$$

  - With $T \sim r^{-p}$

$$\rightarrow \nu S_\nu(\nu) \propto \nu^{4-2/p}$$

    - Most disks must have $p = ½$
    - Contribution of viscous heating negligible
    - Support of Flared disk model with large surface

Disk SED

Fit of the spectral energy distribution in the Rayleigh-Jeans domain

- Dominated by central layer with high density
- Optically thick emission

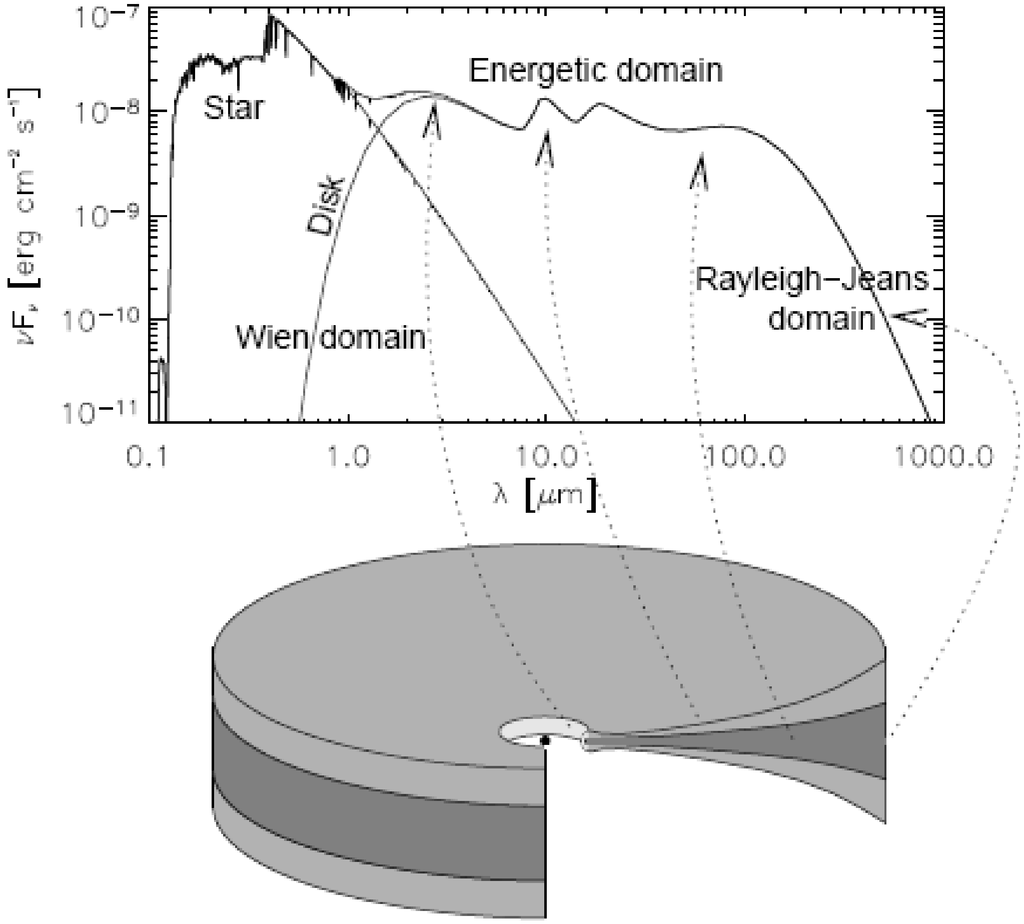


Fig. 2.— Build-up of the SED of a flaring protoplanetary disk and the origin of various components: the near-infrared bump comes from the inner rim, the infrared dust features from the warm surface layer, and the underlying continuum from the deeper (cooler) disk regions. Typically the near- and mid-infrared emission comes from small radii, while the far-infrared comes from the outer disk regions. The (sub-)millimeter emission mostly comes from the midplane of the outer disk. Scattering is not included here.

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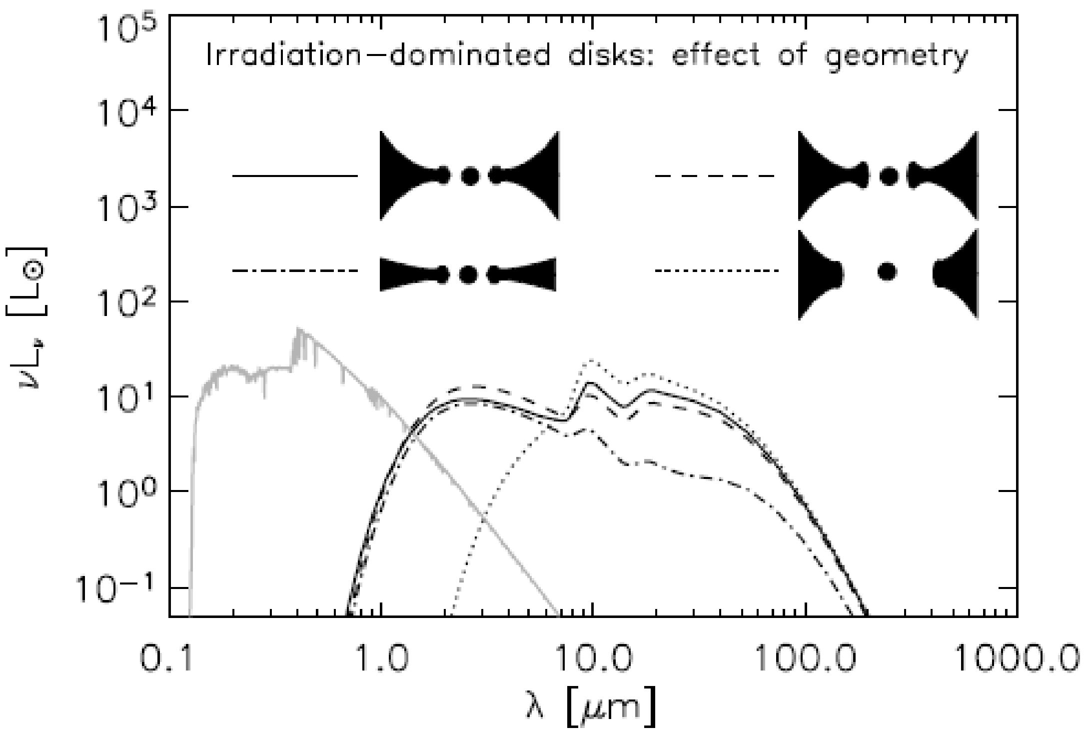
Disk SED

Flared disk with inner rim:

- Temperature increase at inner boundary produces additional pressure
- Inner rim is "puffed-up"
- Shadowed region behind rim



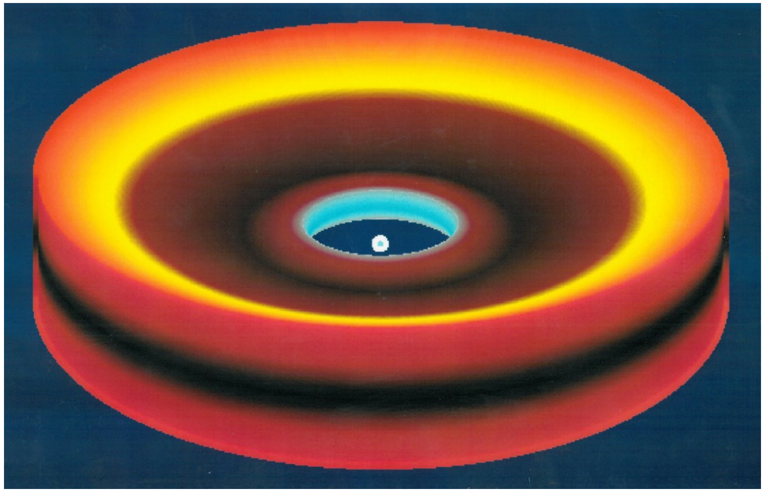
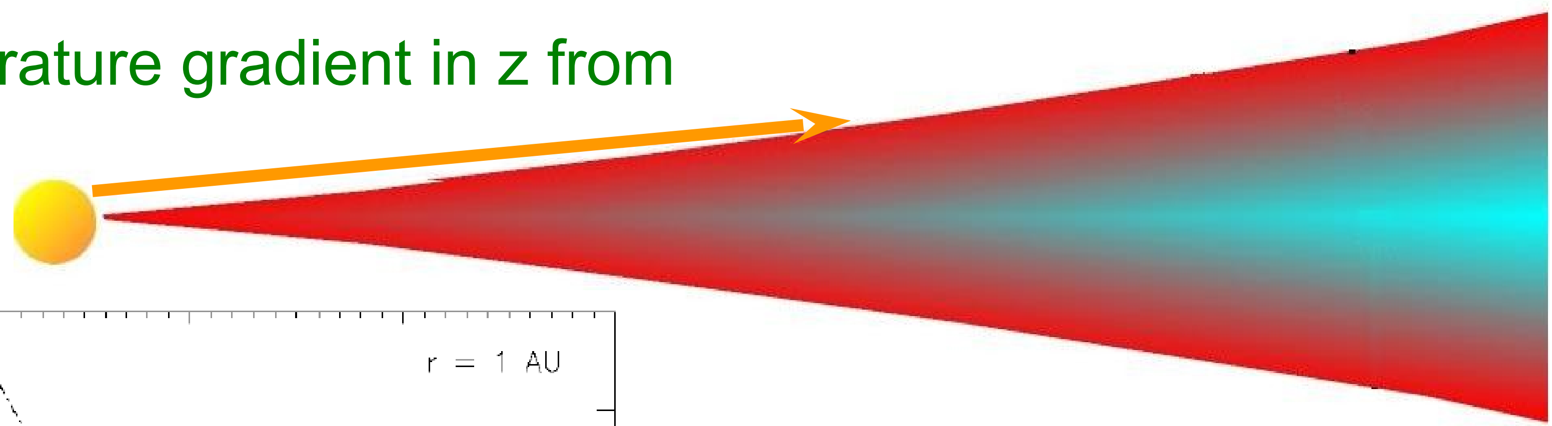


Fig. 5.— Overall SED shape for *non-accreting* disks with stellar irradiation, computed using the 2-D radiative transfer tools from *Dullemond & Dominik* (2004a). The stellar spectrum is added in grey-scale. Scattered light is not included in these SEDs. Solid line is normal flaring disk with inner dust rim; dashed line is when the rim is made higher; dot-dashed line is when the flaring is reduced (or when the disk becomes 'self-shadowed'); dotted line is when the inner rim is at 10× larger radius.

Temperature Structure

Strong temperature gradient in z from irradiation:



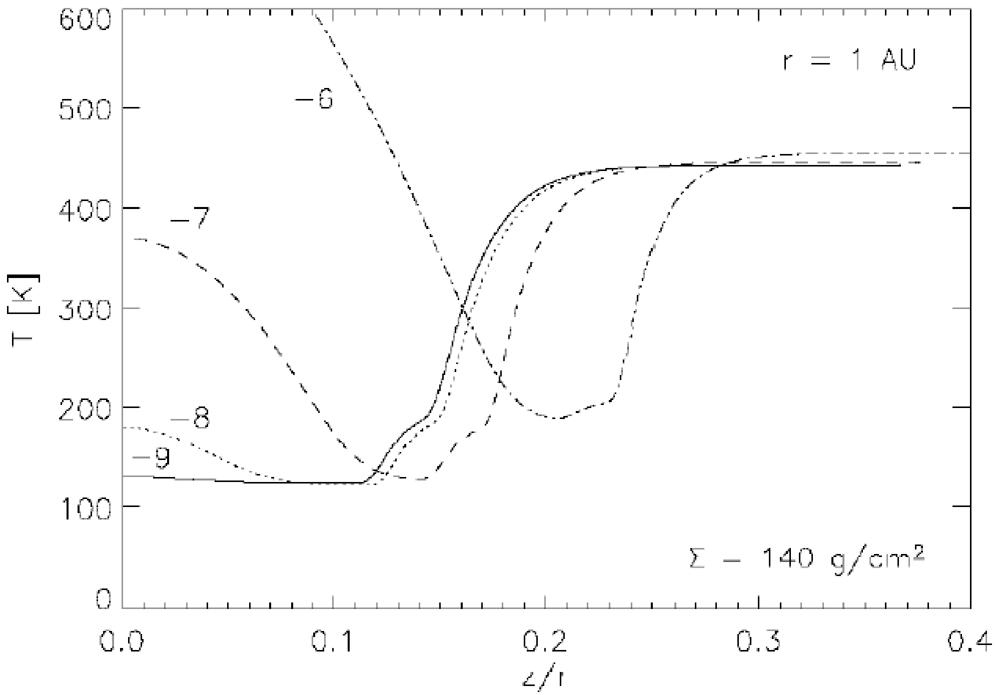
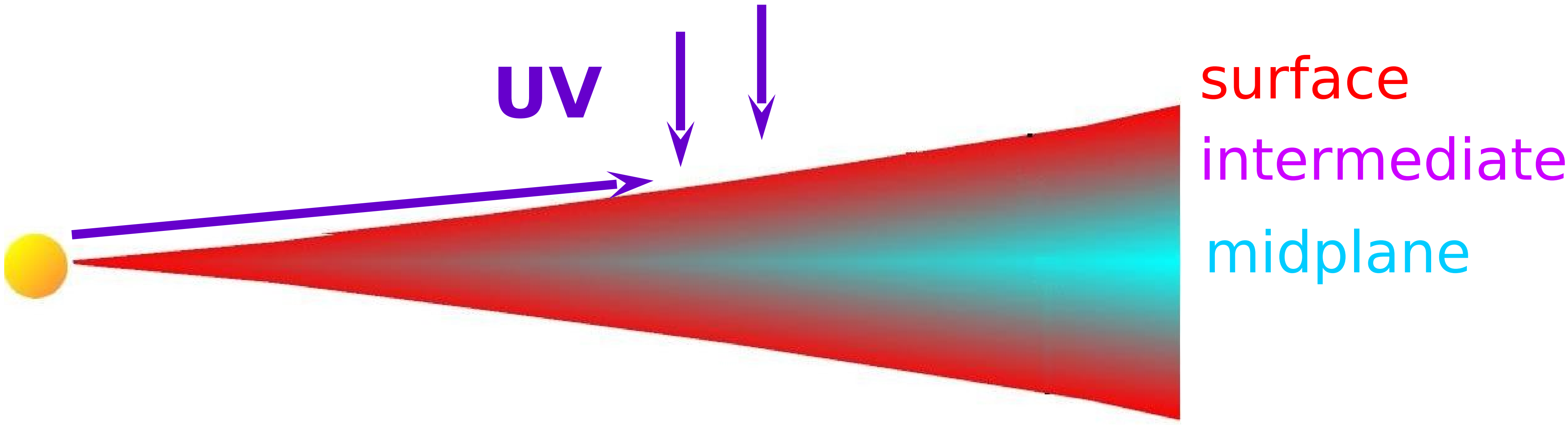


Fig. 4.— Vertical temperature distribution of an irradiated α-disk at 1 AU, for a fixed Σ (chosen to be that of a disk model with $\dot{M} = 10^{-8}\, M_\odot/\mathrm{yr}$ for $\alpha = 0.01$), but varying $\dot{M}$, computed using the models of *D'Alessio et al.* (1998). The labels of the curves denote the 10-log of the accretion rate in $M_\odot$/yr.

Temperature of central layer dominated by overall gas density and accretion flow

Surface temperature determined by local gas cooling

Chemical Structure of PPDs



Surface layer : $n \sim 10^{4-5} cm^{-3}$, $T > 50K$

Photochemistry

Intermediate : $n \sim 10^{6-7} cm^{-3}$, $T > 40K$

Dense cloud chemistry

Midplane : $n > 10^{7} cm^{-3}$, $T < 20K$

Freeze-out

Disk chemical structure

Y. Aikawa et al.: Warm molecular layers in disks 625

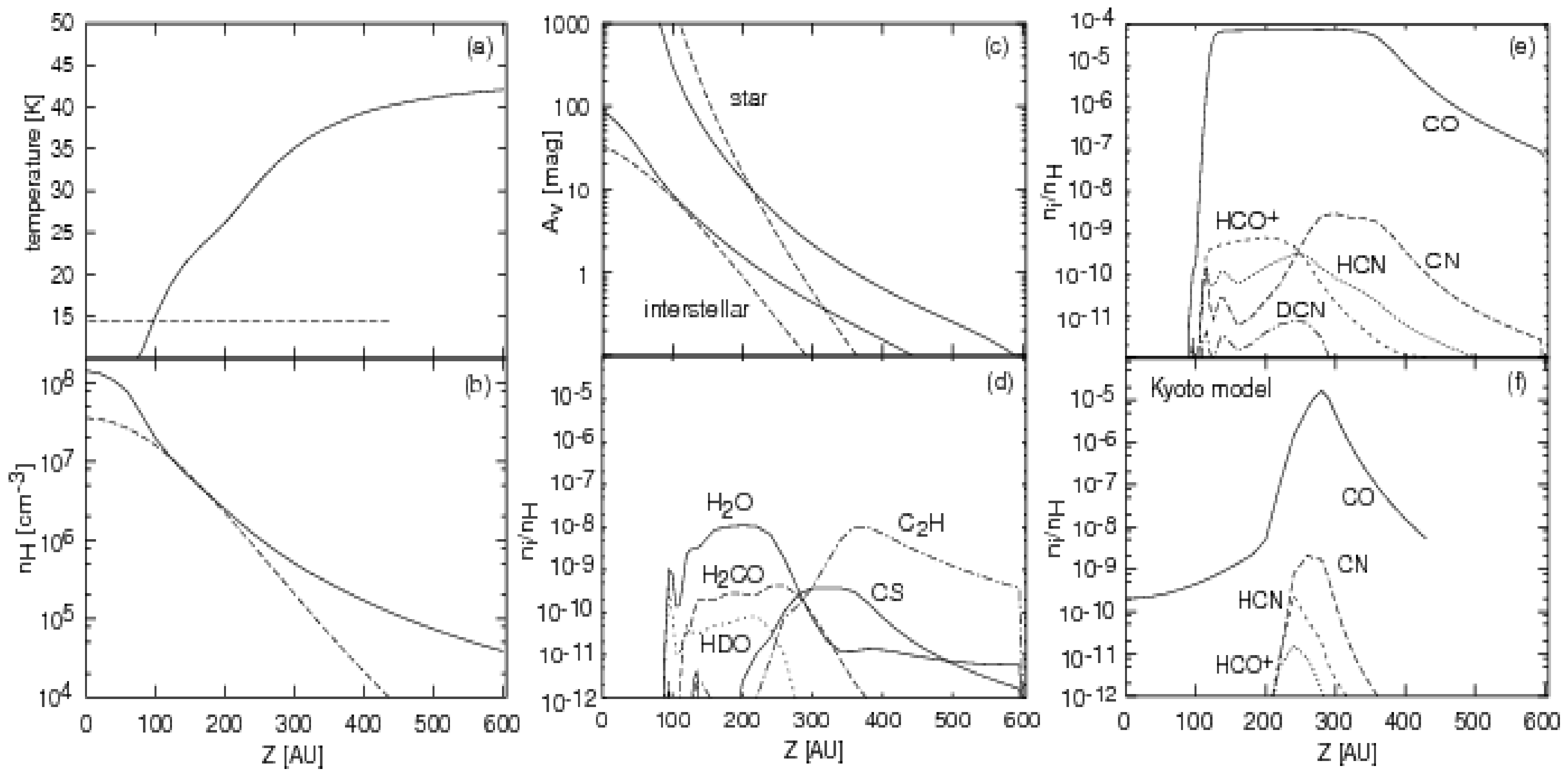


Fig. 2. Vertical distributions at R = 373 AU of **a)** temperature, **b)** density ($n_{\rm H} \equiv 2n({\rm H_2}) + n({\rm H})$), **c)** attenuation of the interstellar radiation ($A_{\rm V}^{\rm IS}$) and stellar radiation ($A_{\rm V}^{\rm star}$), **d–e)** molecular abundances in the D'Alessio et al. model, and **f)** molecular abundances in the Kyoto model with $S = 0.03$. In panels **a–c)**, the physical parameters of the D'Alessio et al. and Kyoto models are shown via *solid* and *dashed* lines, respectively. In panels **d–f)**, the disk age is assumed to be $t = 1.0 \times 10^6$ yr.

Structure of disks

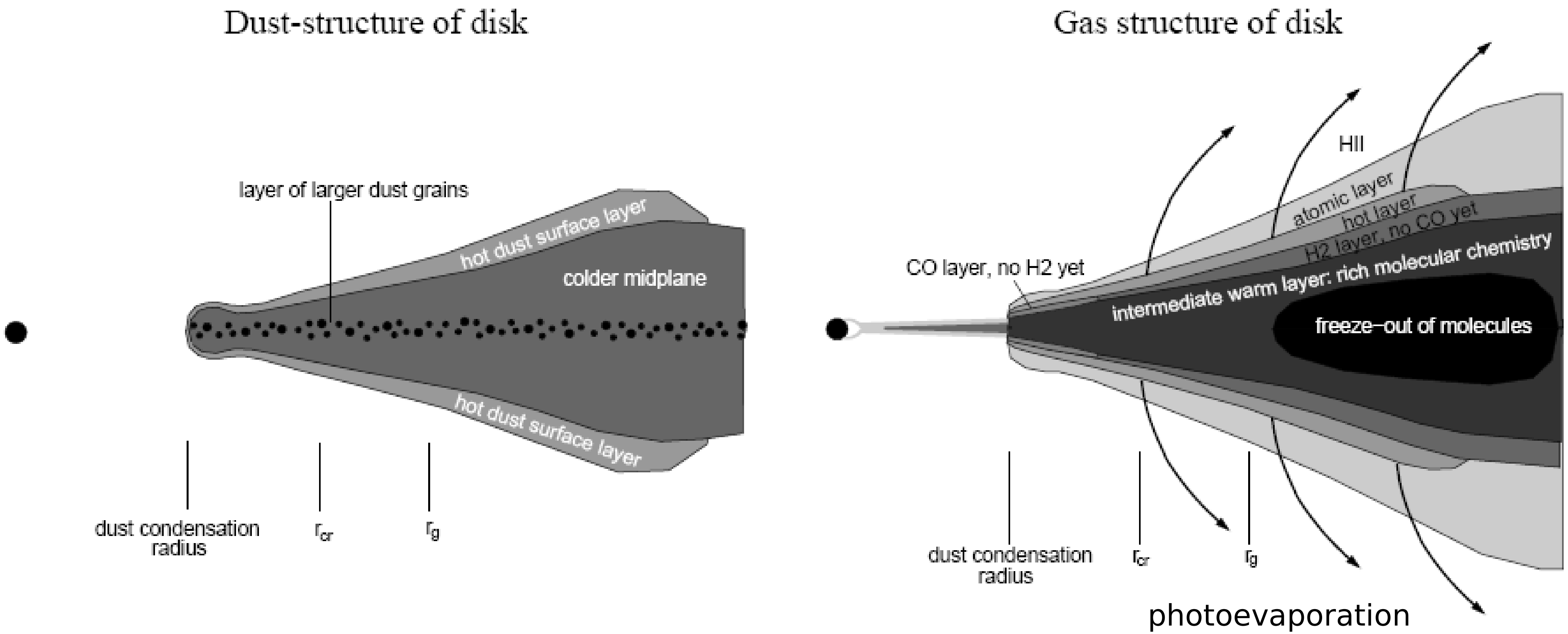


Fig. 13.— Pictograms of the structure of a flaring protoplanetary disk, in dust (left) and gas (right).

Photoevaporation

- Dust evaporates close to the central star
  - Material still flows throw the gap onto the central object

- Evaporation from the surface of the disk into a disk wind/outflow if

$$c_s > v_{esc} \approx \left(\frac{GM_\star}{r}\right)^{1/2}$$

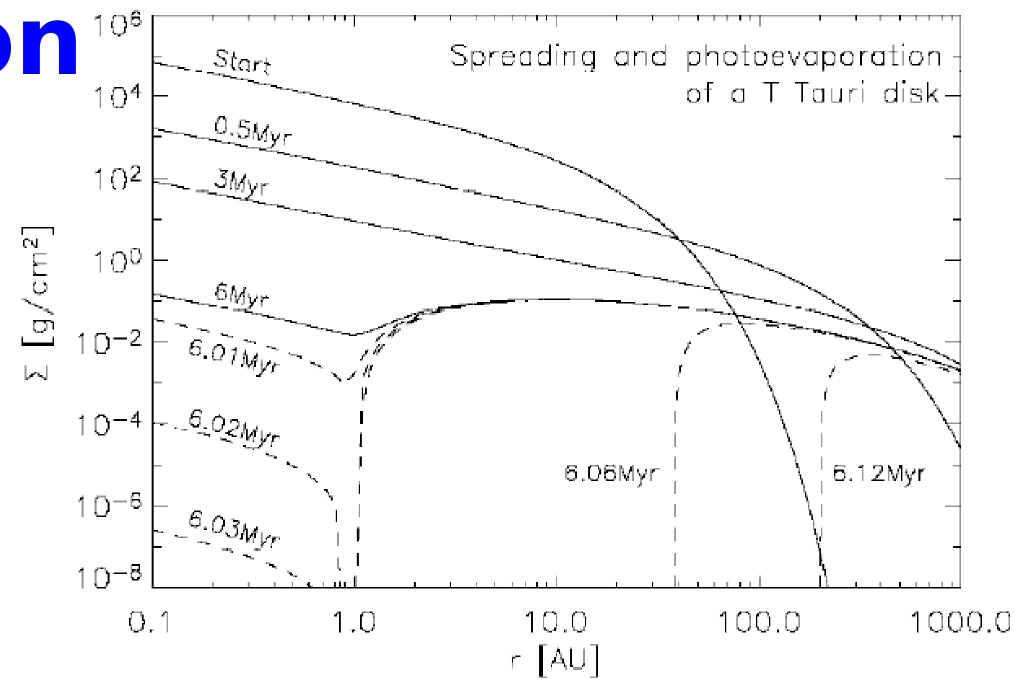


Fig. 11.— Evolution of the surface density of a EUV-photoevaporating disk (Figure adapted from *Alexander et al.*, 2006b). This simulation starts from a given disk structure of about 0.05 $M_\odot$ (marked with 'Start' in the figure). Initially the disk accretes and viscously spreads (solid lines). At $t = 6 \times 10^6$ yr the photoevaporation starts affecting the disk. Once the EUV-photoevaporation has drilled a gap in the disk at ~ 1 AU, the destruction of the disk goes very rapidly (dashed lines).The inner disk quickly accretes onto the star, followed by a rapid erosion of the outer disk from inside out. In this model the disk viscosity spreads to $>$ 1000 AU; however, FUV-photoevaporation (not included) will likely truncate the outer disk.

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Disk wind

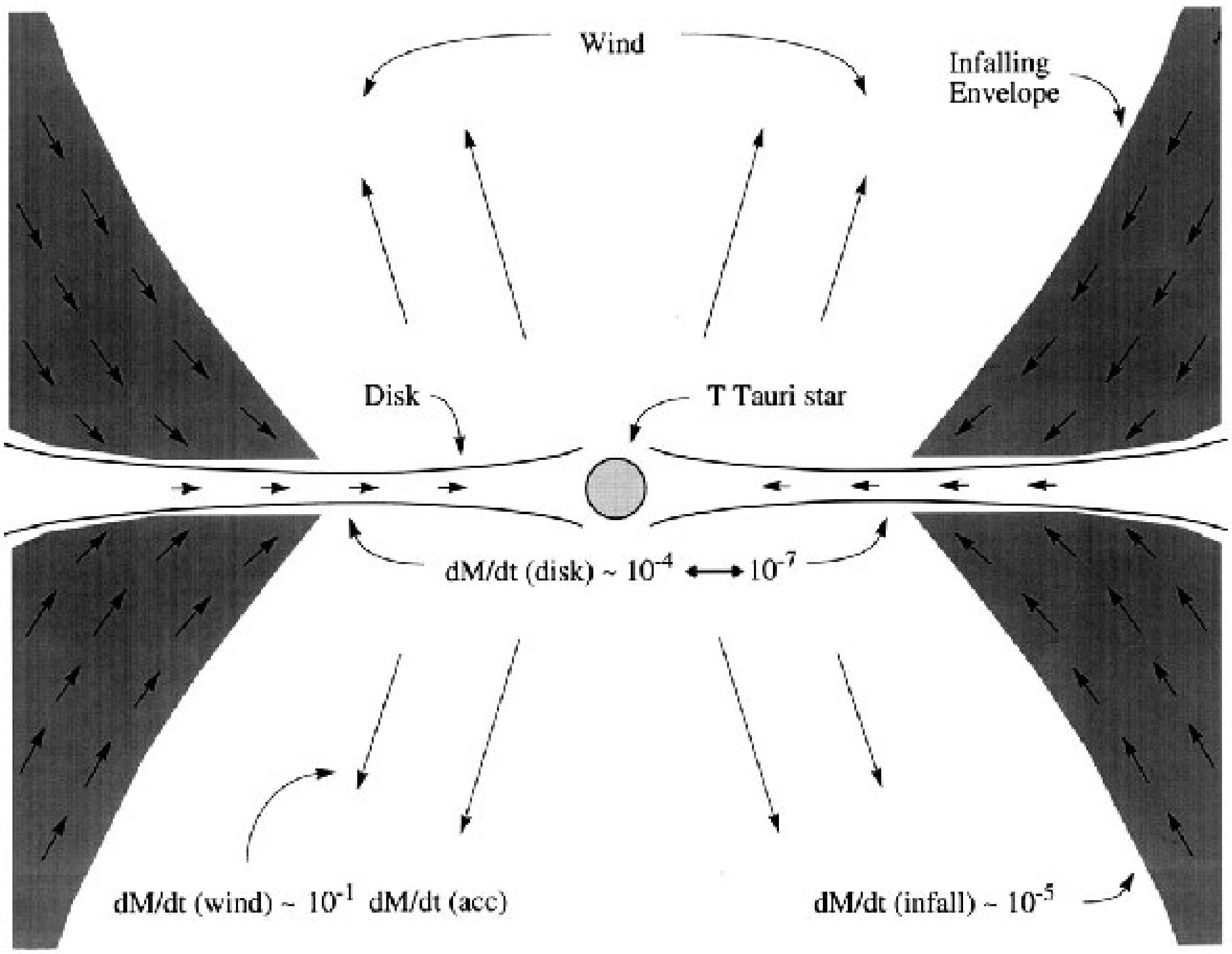
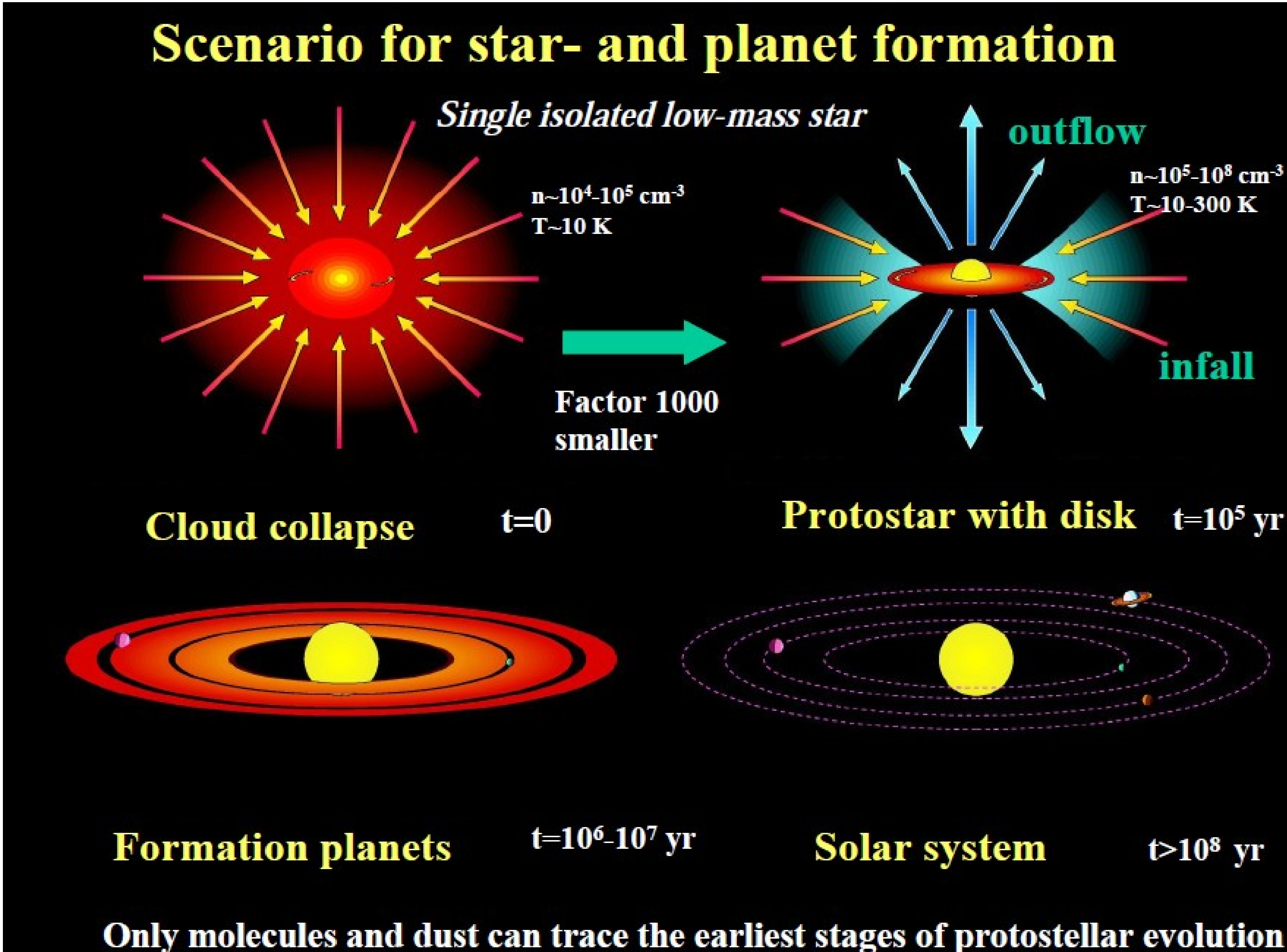


Figure 1 Schematic picture of FU Ori objects. FU Ori outbursts are caused by disk accretion increasing from $\sim 10^{-7}\ M_\odot\ \mathrm{yr}^{-1}$ to $\sim 10^{-4}\ M_\odot\ \mathrm{yr}^{-1}$, adding $\sim 10^{-2}\ M_\odot$ to the central T Tauri star during the event. Mass is fed into the disk by the remanant collapsing protostellar envelope with an infall rate $\lesssim 10^{-5}\ M_\odot\ \mathrm{yr}^{-1}$; the disk ejects roughly 10% of the accreted material in a high-velocity wind.

→ Leads us to outflows


Scenario for star- and planet formation
Single isolated low-mass star
n~10^4-10^5 cm^-3
T~10 K
outflow
n~10^5-10^8 cm^-3
T~10-300 K
infall
Factor 1000
smaller
Cloud collapse
t=0
Protostar with disk
t=10^5 yr
Formation planets
t=10^6-10^7 yr
Solar system
t>10^8 yr
Only molecules and dust can trace the earliest stages of protostellar evolution