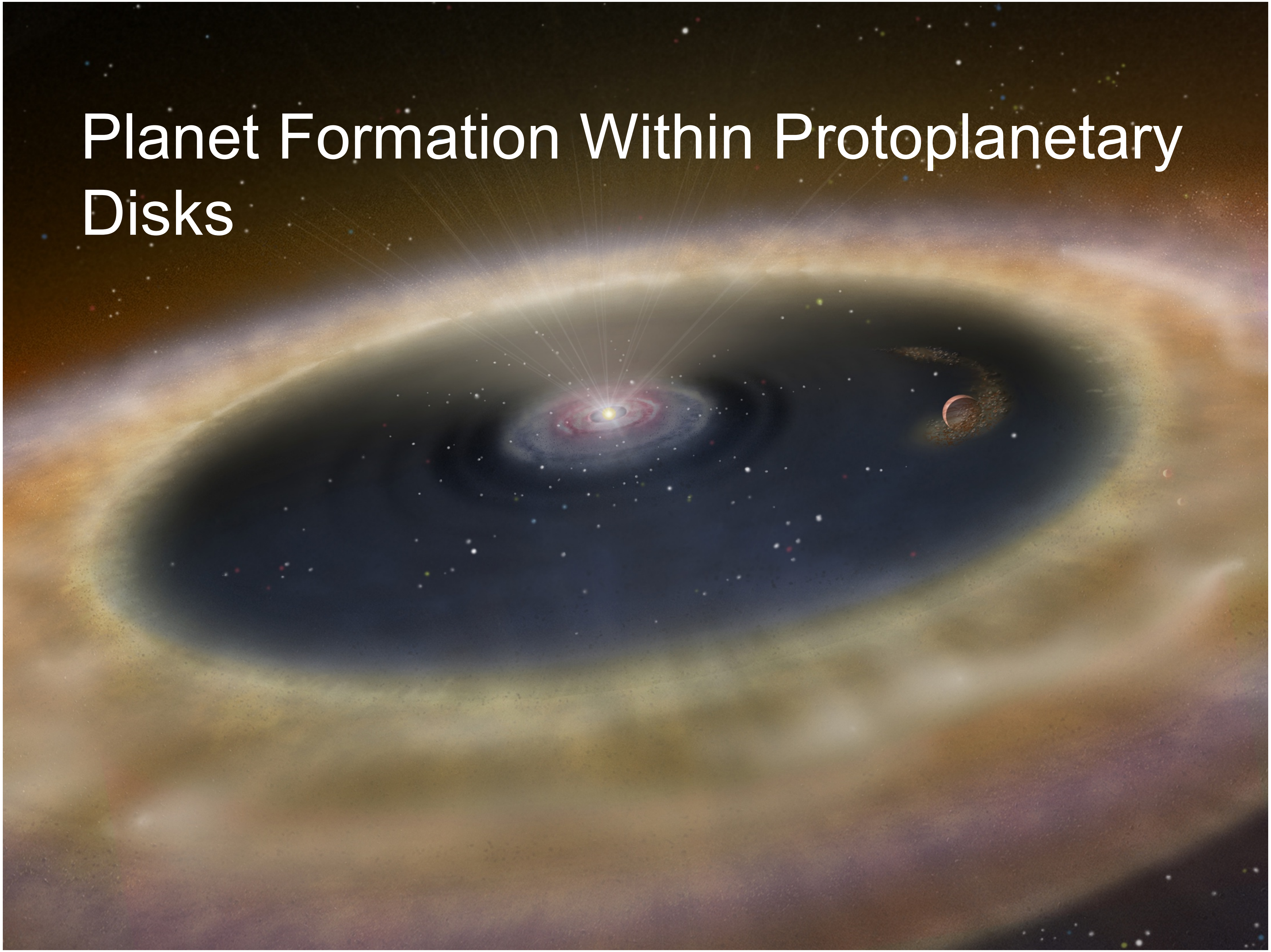
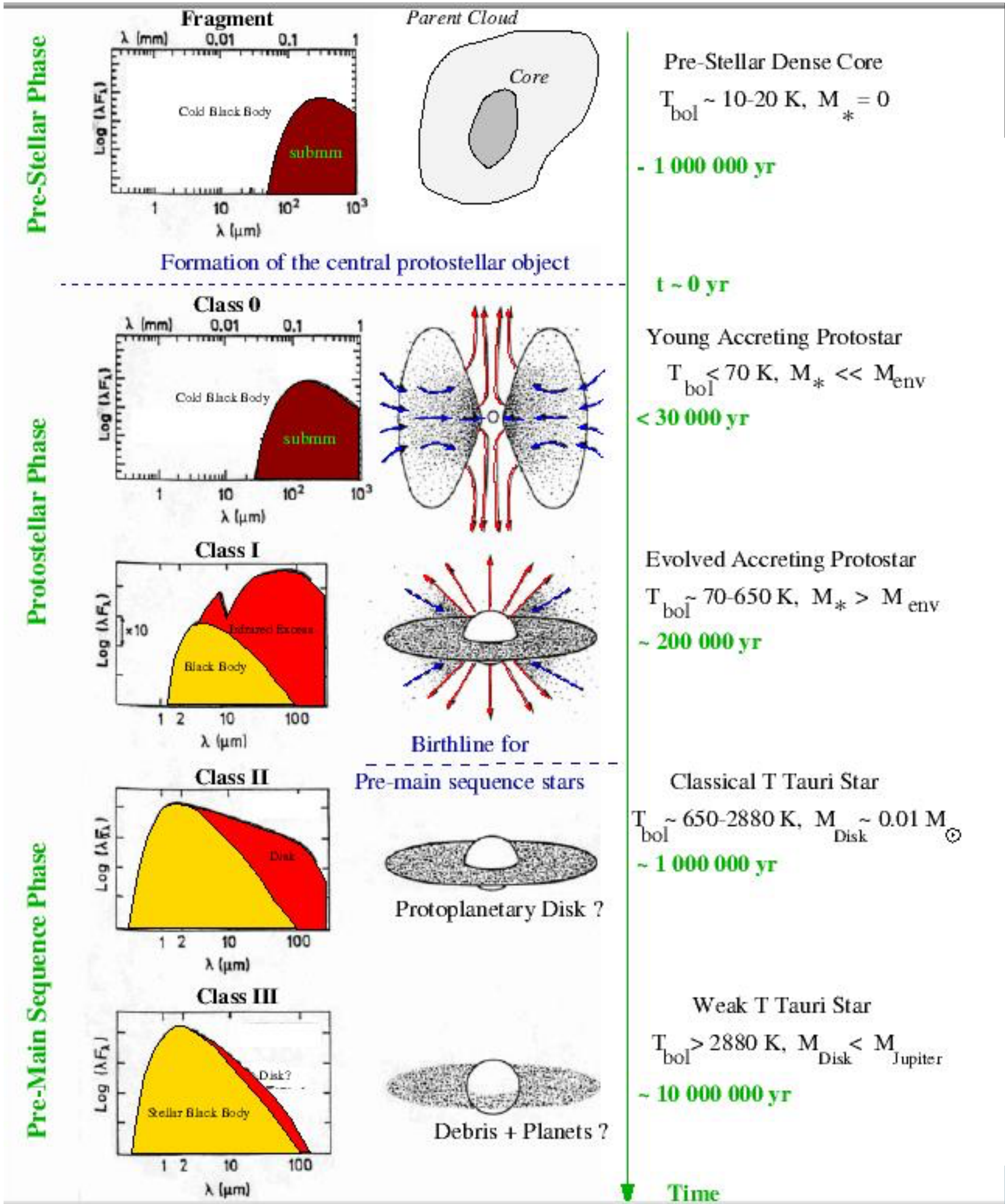
Planet Formation Within Protoplanetary Disks



Full Week Outline

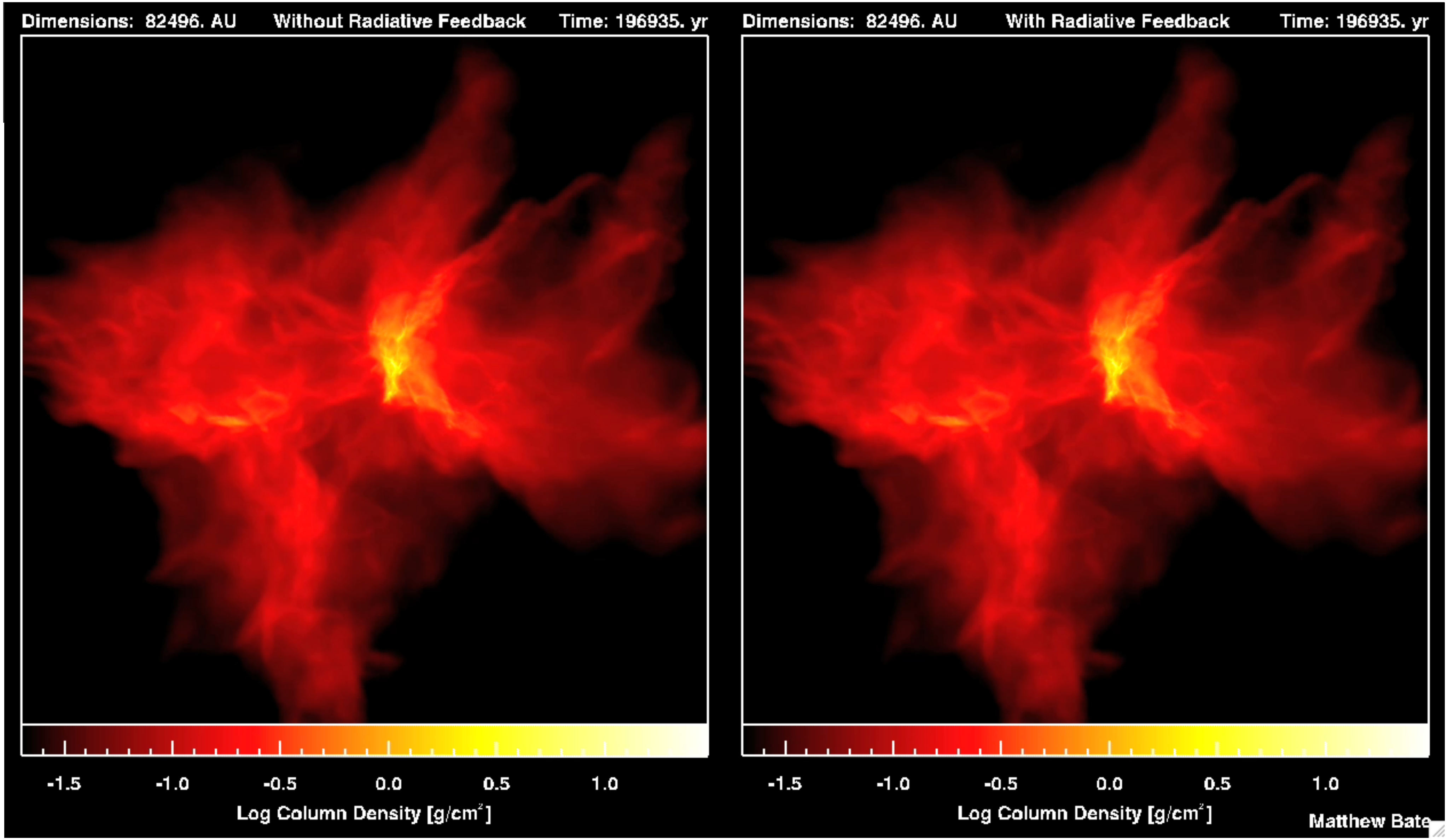
- Big picture intro
- Disk Physics (intro and key references)
  - Hydrostatic equilibrium in a disk plane (Ch. 15 revision)
  - Viscosity (and an intro to α-disk models, Ch. 15 revision)
  - Gravitational instability (Ch. 15 Revision)
  - Minimum mass solar nebula.
  - Radiation transport
  - Grain Growth and gas drag (Mark to cover growth details later)
  - Planetesimal formation (Mark to cover limits to Kelvin-Helmholz instability)
  - Rocky Planet Formation
  - Gas giant formation via core accretion
- Observational Parameter space
- Selected Open Questions
  1. What is a Transition Disk?
  2. What causes accretion in protoplanetary disks?
  3. What is the initial entropy of a giant exoplanet?
  4. What causes the final mass of a giant exoplanet?
- Ask Mark about Dust Traps…



Reminder:

"Textbook" picture of Star and Planet Formation

Reipurth (2005)



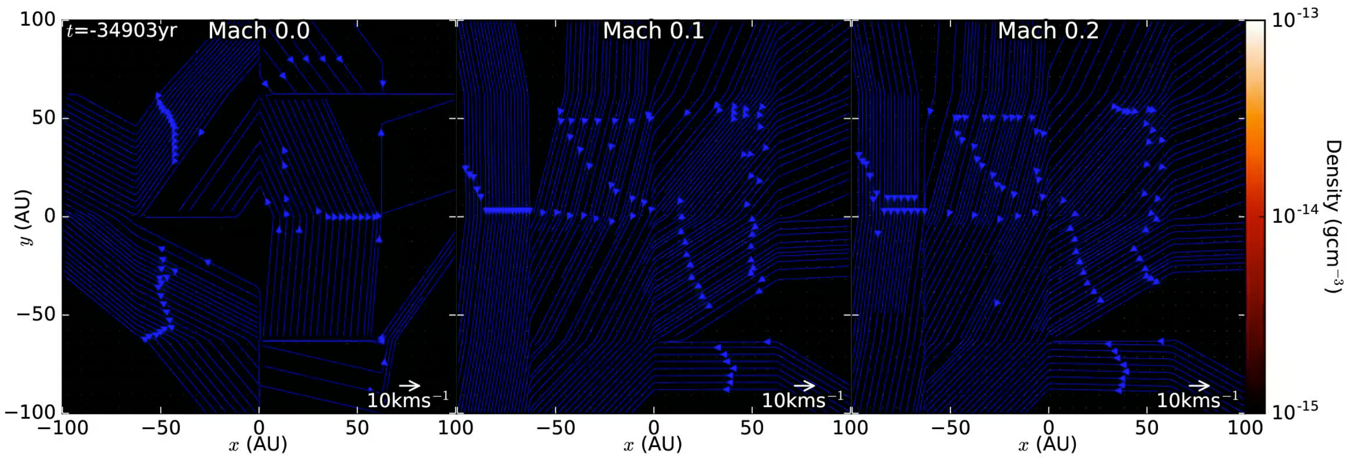
Dynamical picture of star formation

Bate et al (2009)

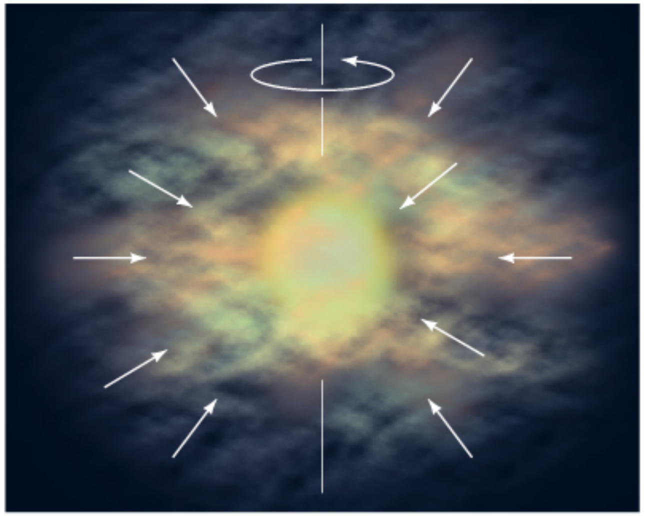
The angular momentum "problem"

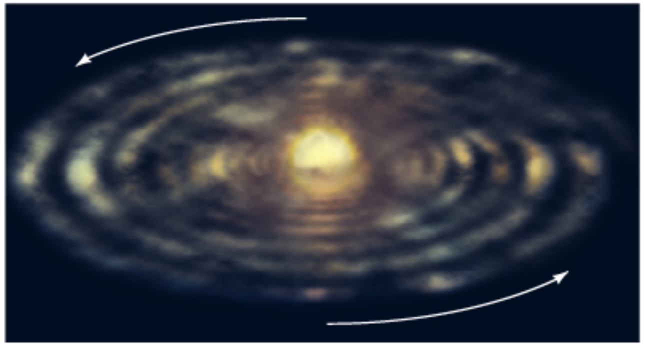
- The Galaxy has turbulence on all scales, so gas clouds always start off rotating slowly.
- During a collapse over a factor of 1 million in distance, the cloud spins faster and faster, and eventually has an aspect ratio < 0.1 (thin disk)

$$v \propto r^{-1} \qquad v_{\mathrm{Kep}} \propto r^{-1/2}$$



Kuruwita and Federrath (2019)





The Challenge…

- We know planet formation occurs in disks, which have a variety of disk initial conditions, compositions and stellar types.
- We know that there are a wide variety of exoplanet architectures
- How do we link these, and answer questions like: *Is our planetary system unusual?* and *What pre-determines if a planetary system is likely to be habitable?*

Physics of Disks

Scaling Parameters

- It is traditional to avoid the use of GM where possible, and replace T with the square of the *isothermal* sound speed c_{is}^2

$$c_{\mathrm{is}} = \sqrt{\frac{k_B T}{m}}$$

$$\Omega = \sqrt{\frac{GM_*}{\varpi^3}} \qquad v_\phi = \sqrt{\frac{GM_*}{\varpi}} \qquad = \sqrt{\frac{k_B T}{\mu u}}$$

- Disks modelling is in cylindrical co-ordinates, with surface density Σ.

$$\mathbf{x} \sim (\varpi, z, \phi)$$

Dynamics

- Equation of motion:

$$\frac{d\mathbf{v}}{dt} = \underbrace{-\frac{1}{\rho}\nabla P}_{\text{gas pressure}} + \underbrace{\mathbf{F}_{\text{drag}}}_{\text{drag force}} + \underbrace{\mathbf{F}_{\text{grav}}}_{\text{gravity}} + \underbrace{\mathbf{F}_{\text{visc}}}_{\text{viscosity}}$$

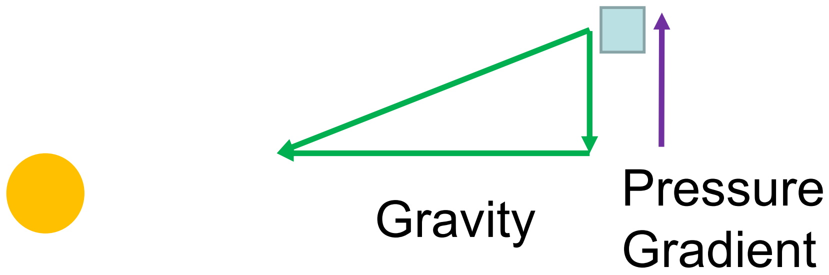
- **Viscosity**: drive gas evolution in disk [viscous accretion disk]
- **Gravity**: – keeps things rotating
  – "holds" disks together [hydrostatic eq: $F_{grav}=F_{pres}$]
  – massive self-grav → spirals → fragment → giants?
- **Drag**: – couples dust and gas (depends on grain size)
  – [huge!] effect on dust dynamics
- **Ordered magnetic field**. Not above. Will discuss briefly at end.

Disk Vertical Structure

- *Thin* disks can be approximated as having a separable radial *surface* density profile and a vertical density profile.
- If we approximate temperature as being uniform with *z*, the vertical structure is Gaussian. Derivation is ~10 mins (try it), balancing the vertical pressure gradient with the vertical component of the Gravitational force.

$$\rho = \rho_0 \exp\left(\frac{-z^2}{2H^2}\right) \qquad \frac{H}{\varpi} = \frac{c_{\mathrm{is}}}{v_\phi}$$

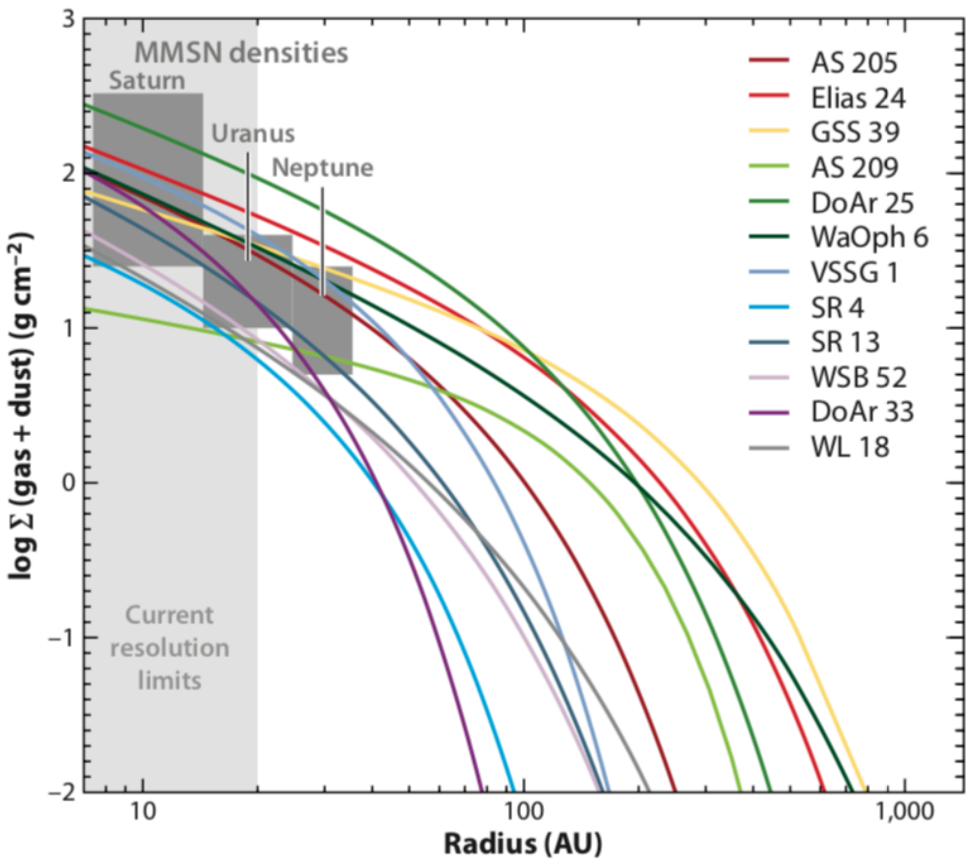
Shakura and Sunyaev (1973)
Armitage (2017)



Disk Radial Structure

- Radial disk structure depends on viscous disk evolution.
- Kinematic gas viscosity is far too low (evolutionary timescales > Gyr)

$$\tau_\nu \approx \frac{\varpi^2}{\nu} = \frac{\varpi^2 \rho}{\mu}$$



Shakura and Sunyaev (1973)
Armitage (2017)
Williams and Cieza (2011)

Disk Radial Structure

- Common to consider an "α-disk" prescription, where we put:

$$\nu = \alpha c_{\rm is} H = \alpha \frac{c_{\rm is}^2}{\Omega}$$

- Some physical motivation for α, (e.g. Shakura and Sunyaev 1973) but not for it being constant. Caused by e.g. turbulence or magneto-rotational instability.
- For a steady-state disk with power law Σ and T, this predicts:

$$p + q = -\frac{3}{2} \text{ for } \Sigma(r) \propto r^p \text{ and } T(r) \propto r^q.$$

- E.g. $\Sigma \sim r^{-1}$ for $T \sim R^{-1/2}$

$\varpi \sim r$

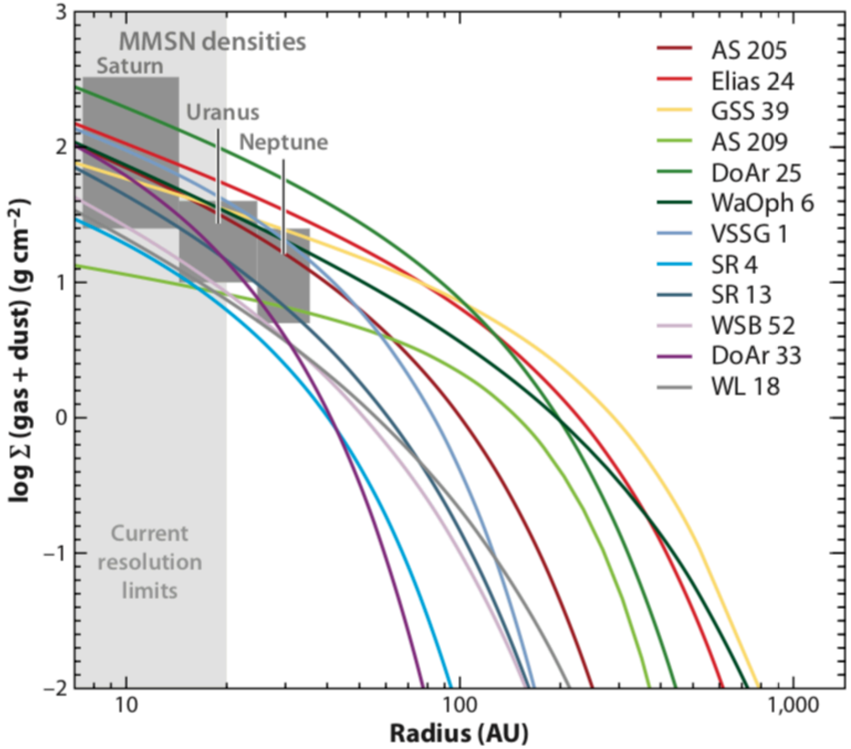
e.g. Dullemond lecture notes

Minimum Mass Solar Nebula (MMSN)

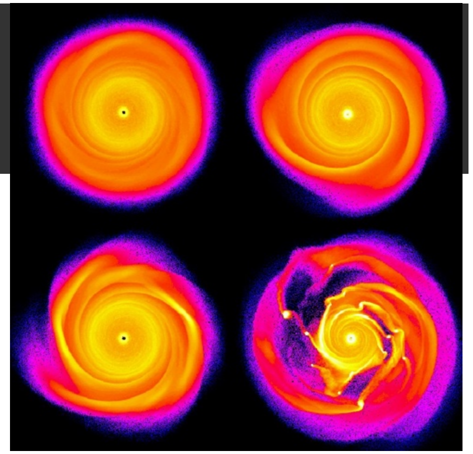
- Based on the amount of solid material in the solar system planets, and a solar dust to gas ratio, there was ~0.01M$_{sun}$ when the planets formed.

$$\Sigma = \Sigma_0 \varpi_0^{-3/2}$$
$$\Sigma_0 \approx 1700 \text{ g cm}^{-2}$$
$$\varpi_0 = \varpi / \text{AU}$$
$$H = 0.03 \varpi_0^{5/4} \text{ AU}$$

- Any planet formation theory has to ask: *Does it work for the MMSN?*



Gravitational Instability (revision)



- The Jean's mass could be derived by considering the *dispersion relation* of sinusoidal density fluctuations in 3D:

$$\Delta\rho \propto e^{i(kx-\omega t)} \qquad \omega^2 = c_{\rm is}^2 k^2 - 4\pi G \rho_0$$

- In 2D, with a natural epicyclic frequency $\kappa \sim \Omega$, we have:

$$\omega^2 = c_{\rm is}^2 k^2 - 2\pi G \Sigma k + \kappa^2$$

- This is unstable ($\omega^2 < 0$) if the Toomre Q parameter:

$$Q = \frac{c_{\rm is}\kappa}{\pi G \Sigma} \lesssim 1$$

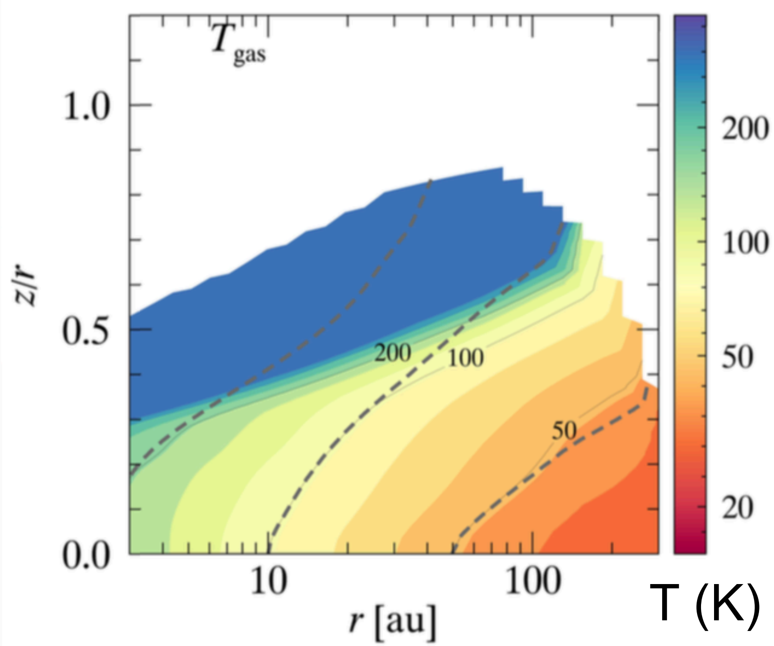
Kratter and Lodato (2016)

Australian National University

Radiation Transport

- Unlike typical assumptions by hydro modellers, real disk temperatures are determined by *viscous heating* and *stellar irradiation.*
- For optically-thin grey dust, we have: $T_{\rm thin,grey} = \left(\frac{L_*}{16\pi\sigma_{\rm SB}}\right)^{1/4} r^{-1/2}$
- Shadowed material has a *cooler* and *steeper* temperature profile.
- Small grains have a *warmer* and *shallower* temperature profile.
- Gas heated by UV or X-rays in low density surface regions can be very warm.
- Gas and different dust species are decoupled through most the disk.

Fedele (2016)



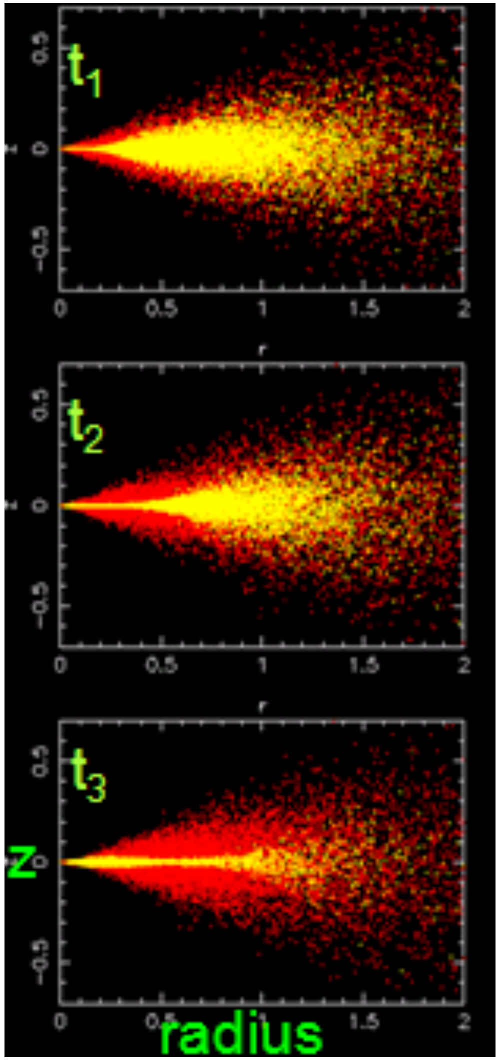
Grain Growth, Settling and Gas Drag
(details next week)

- When grains grow, they settle to the midplane in the absence of strong turbulence.
- Governed by the *stopping time* of dust, which is characterised in the Epstein regime as:

$$\tau_s = \frac{\rho_m s}{\rho v_{\rm th}}$$

s: grain radius, ρ_m: material density, v_{th}: mean thermal velocity (near sound speed)

- Wherever $\Omega\tau_s >> 1$, dust is strongly coupled to the gas. If $\Omega\tau_s$ <<1, dust is weakly coupled to the gas.
- Boundary at 0.1mm to 1m depending on gas density and orbital radius.



Grain Growth, Settling and Gas Drag
(details next week)

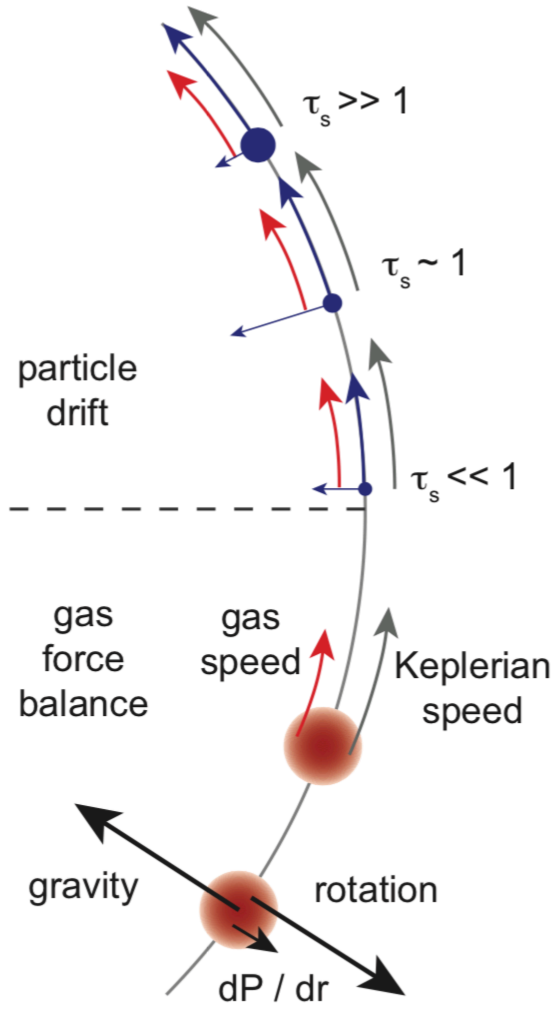
- In addition to Keplerian terms, gas velocity **u** has an equation of motion term:

$$\frac{\partial \mathbf{u}}{\partial t} = \ldots - \frac{c_{\rm is}^2}{\rho}\nabla\rho$$

- Dust velocity **w** doesn't have this term, but does have:

$$\frac{\partial \mathbf{w}}{\partial t} = \ldots - \frac{1}{\tau_s}(\mathbf{w} - \mathbf{u})$$

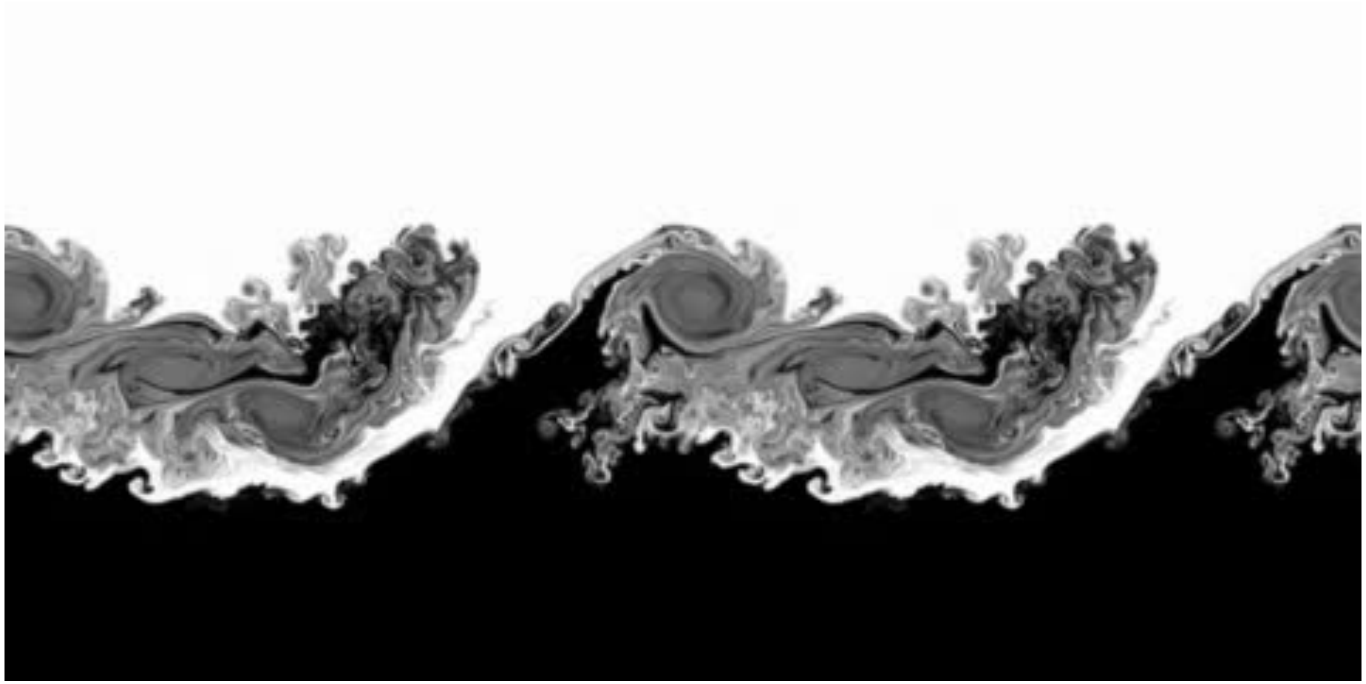
- Gas has the opposite term, modified by the gas to dust ratio.
- This means that gas flows at sub-Keplerian speeds (by a fraction ~$(H/r)^2$), and intermediate size particles experience lossy drag, moving towards high pressure regions (e.g. inner disk)



Youdin (2007), Armitage (2017)

Kelvin-Helmholz Instability
(details next week)

- When dust slows down gas in the midplane, you can get a Kelvin-Helmholz instability



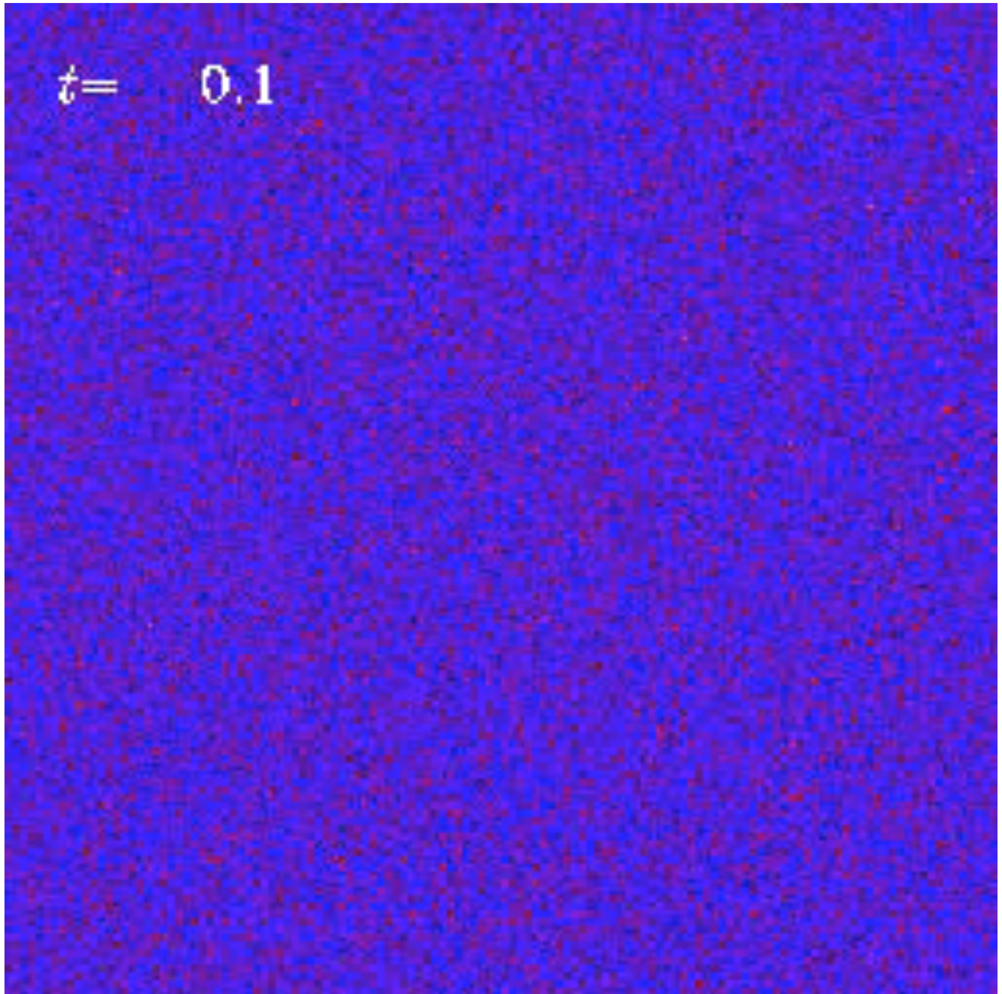
Wikipedia movie

Streaming Instability

- More generally, when dust and gas have equal-ish densities, this produces a streaming instability.
- With a slightly enhanced m-size dust density at radius ϖ, gas will slow down, then dust migrating inwards will “bunch up”, ehnancing the instability.
- (see Mike’s holiday movie of a waterfall)
- https://www.youtube.com/watch?v=1ubjiIXHAYE (29 minute mark)

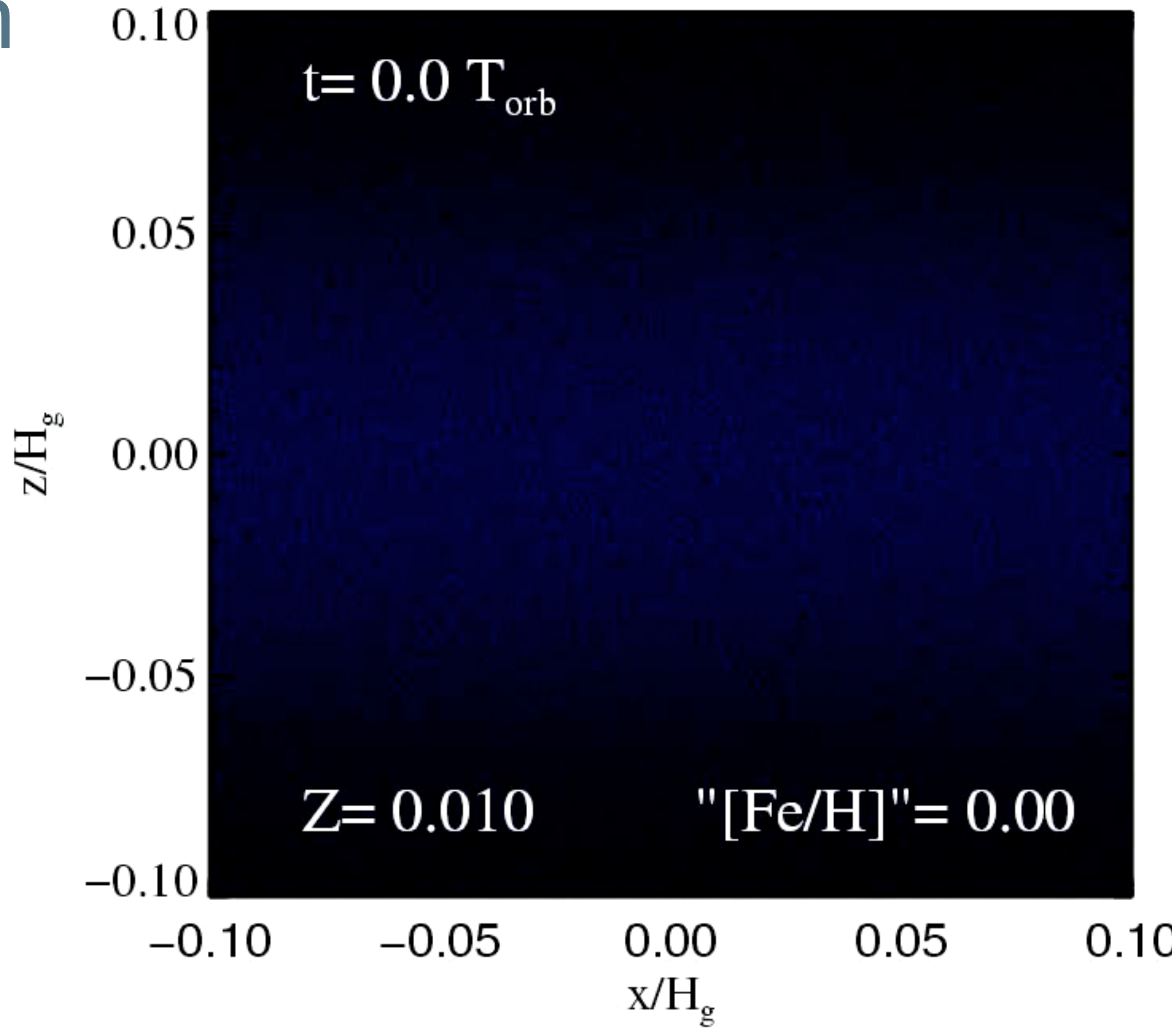
Streaming Instability: Pure Aerodynamic Concentration

Youdin and Goodman (2005) then Johansen (2007…)



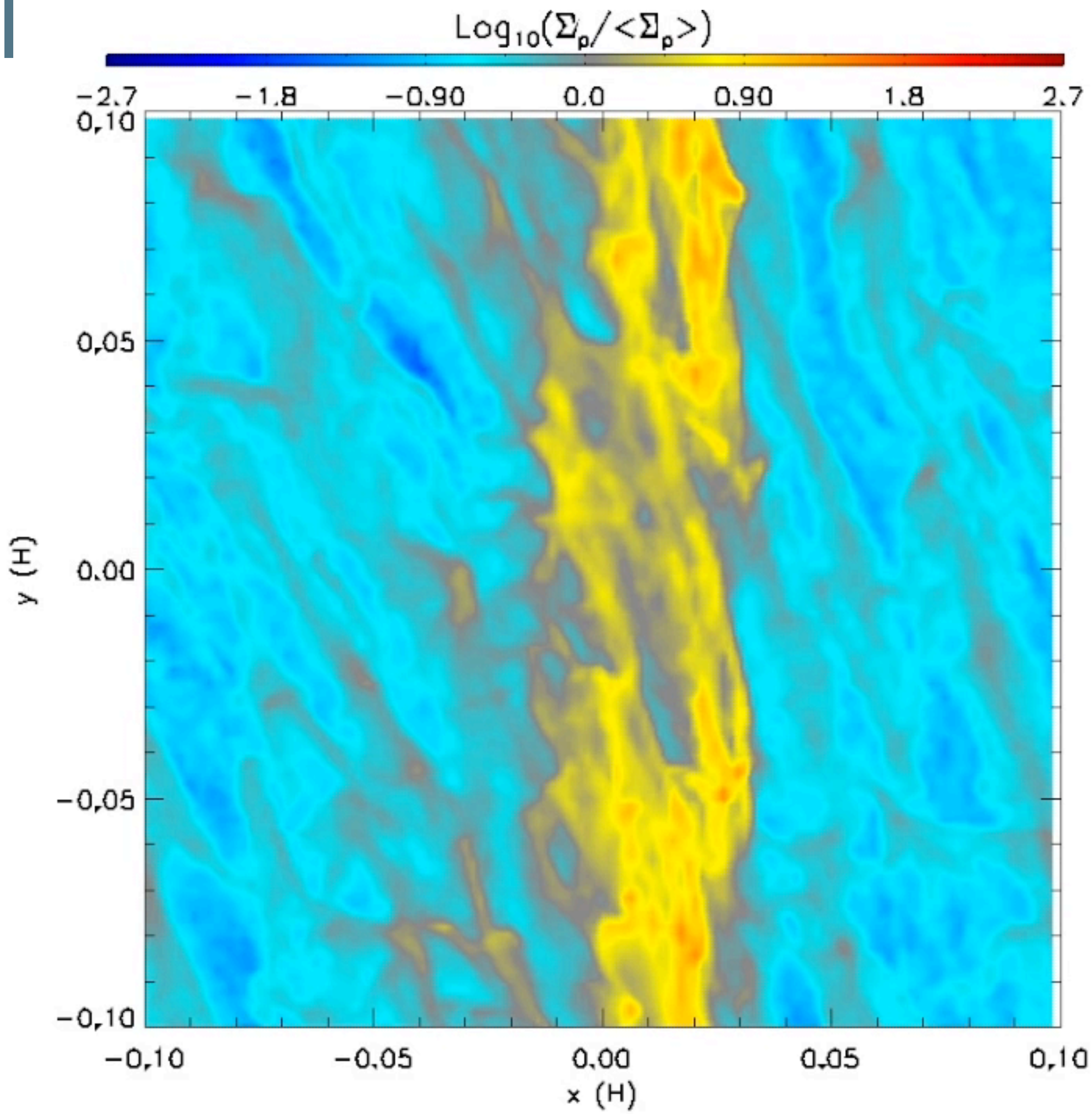
Streaming Instability: With sedementation

Johansen, Youdin and Mac Low (2007)



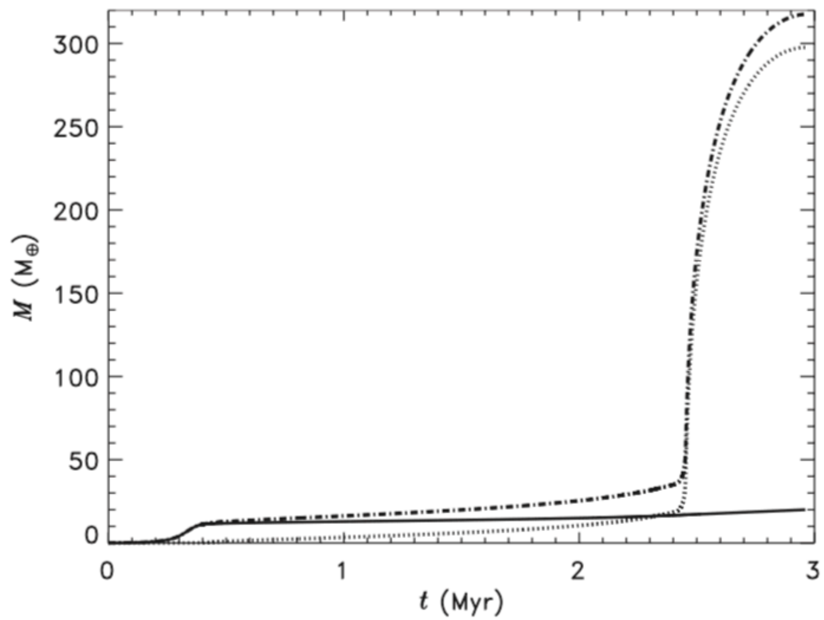
Streaming Instability: Direct Planetesimal Formation?

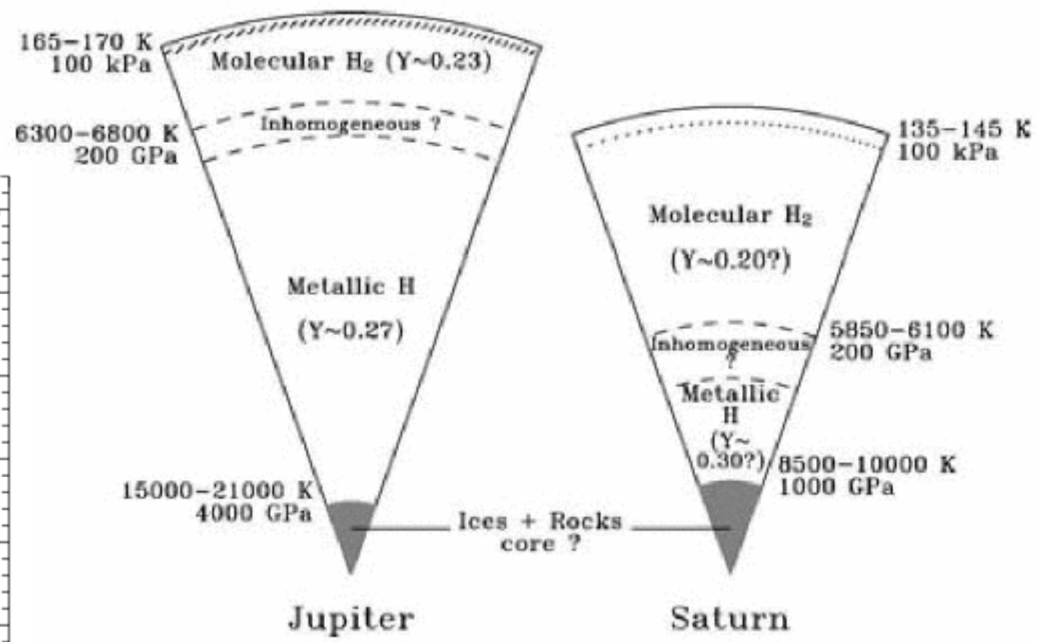
Simon et al (2016)



Planet Formation Theory 1: Core-accretion.

- "Core Accretion": Dust grains grow larger and larger, caught in local pressure maxima, forming rocky planetesimals and then large planets.
- Only works close to a star, and best at the *snow line* at 5-10 AU.





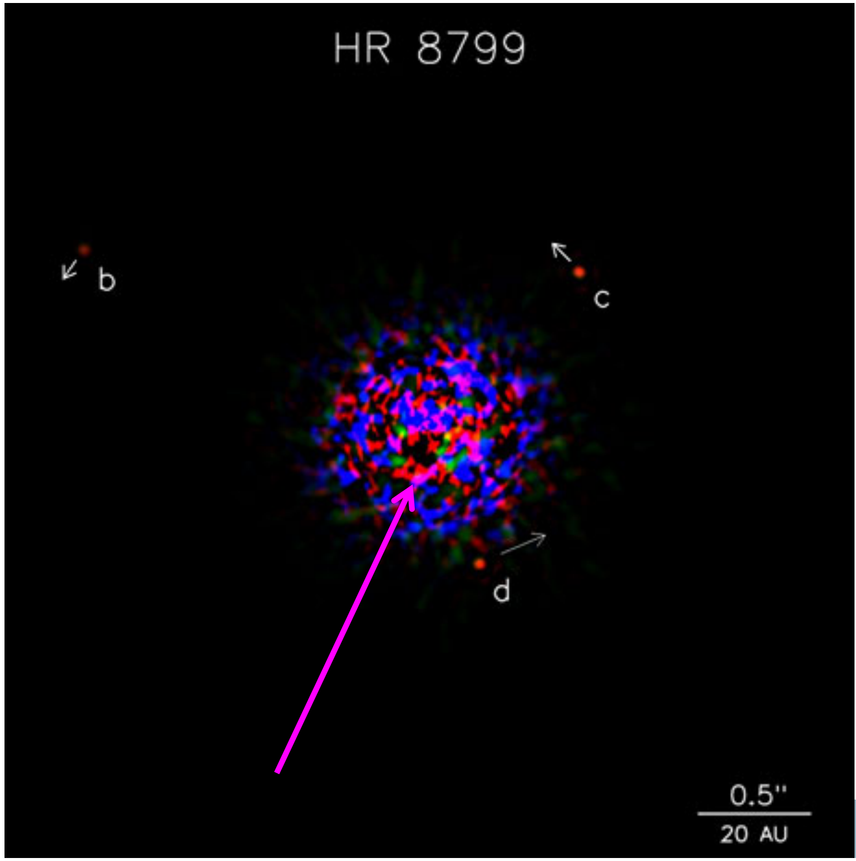
Guillot (2005)

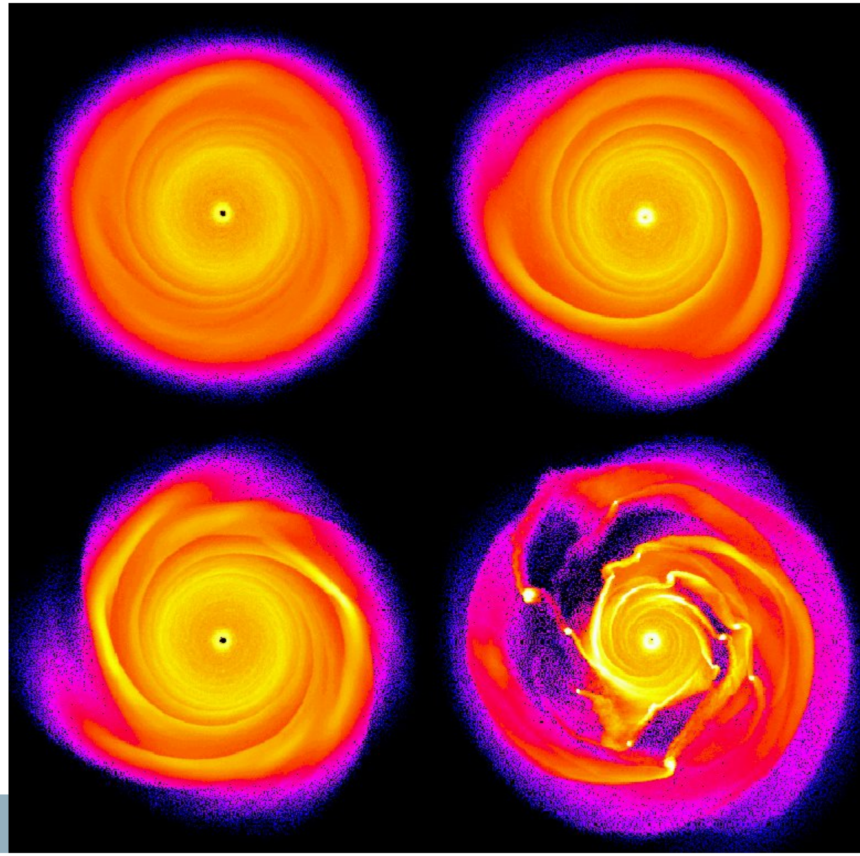
Lissauer (2009), in D'Angelo (2010)

Planet Formation Theory 2: Gravitational Instability.

- If a disk is massive and cold enough (low Toomre Q), spiral density waves form and the disk fragments.
- Works best far from a star (e.g. >10 AU), so *may* explain systems like HR 8799

e.g. Boss (1997)





“Solid” Bodies

- Anything in astrophysics (i.e. an object governed by self-gravity) has characteristic timescales, pressure and velocities defined by its mass and radius.
- For solid bodies, thinking of this characterisation by density and radius is more useful.

$$\Omega_{\rm body} = \sqrt{\frac{GM}{r^3}}$$
$$= 1.9\,{\rm hr}^{-1}\left(\frac{\rho}{1\,g\,cm^{-3}}\right)^{1/2}$$

$$v_{\rm body} = R \times \Omega_{\rm body}$$
$$= 0.53\,m\,s^{-1}\left(\frac{\rho}{1\,g\,cm^{-3}}\right)^{1/2}\left(\frac{R}{1\,km}\right)$$

$$p_{\rm body} = \frac{2}{3\pi}G\rho^2R^2$$
$$= 14\,{\rm Pa}\left(\frac{\rho}{1\,g\,cm^{-3}}\right)^2\left(\frac{R}{1\,km}\right)^2$$

Planetesimal

- IAU: *A planetesimal is a solid object arising during the accumulation of planets whose internal strength is dominated by self-gravity and whose orbital dynamics is not significantly affected by gas drag. This corresponds to objects larger than approximately 1 km in the solar nebula.*
- Is this really true? What is the internal pressure of an object of comet-like density (0.6 g/cm^3) and 1km radius? How does this compare to the pressure needed to make a snowball with your hand?
- What about liquid subsurface water?

"Tutorial" Question…

Consider a planetesimal with an ^{26}Al fraction X =10^{-7}(applicable to canonical numbers with an ice/rock mixture) an Aluminium half life of 1.06 Myr and an ^{26}Al heat production of 1.5 x 10^{13} J/kg. 1) For a heat capacity of 1000 J/kg/K, 3 x 10^{5} J/kg to melt the ice and a thermal conductivity of 0.2 W/m/K, find:

a) The temperature that the planetesimal would reach with no heat loss.

b) Assuming equilibrium between heat output and loss, find the surface temperature at 10 and 100km.

c) For 10 and 100km radii, determine if equilibrium is a reasonable assumption.

Better Planetesimal Discussion

- When the escape velocity (~v_{body}) is more than the gas velocity difference to Keplarian, growth is inevitable. E.g. "pebble accretion".
- Accretion continues only until all material within an orbital radius of the *Hill Sphere* is removed. This is a protoplanet.

$$R_H = \varpi \left(\frac{M_p}{3M_*} \right)^{1/3}$$

- Until a size of >100km, accretion energy in this gas-damped environment is not very important, but radioactive decay is (cryovolcanism).
- When the escape velocity (~v_{body}) exceeds the local sound speed c_g, a (hydrogen-rich) gas atmosphere is stable.

“Tutorial” Question

In core accretion, one of the key requirements for a planetesimal being stable and being able to grow is for the escape velocity to be more than the maximum of the local turbulent velocity or the difference between gas and large solid body velocities. Assume that this value is 30m/s, and that the density of a planetesimal is 1 g/cm^3. What is the minimum radius according to this criteria that means the planetesimal can grow?

Now for gas accretion, the relevant velocity is ~1km/s. How large does a protoplanet have to be to retain a Hydrogen atmosphere?

Isolation Mass

D'Angelo (2010)

- A protoplanet can accrete all "nearby" solids up to a radius difference: $\Delta\varpi = bR_H \approx 4R_H$
- If this applies to both sides of the planetary orbit, then there is a maximum mass of accreted solids:

$$M_{\rm iso} \sim 4\pi\varpi b R_H \Sigma_s$$

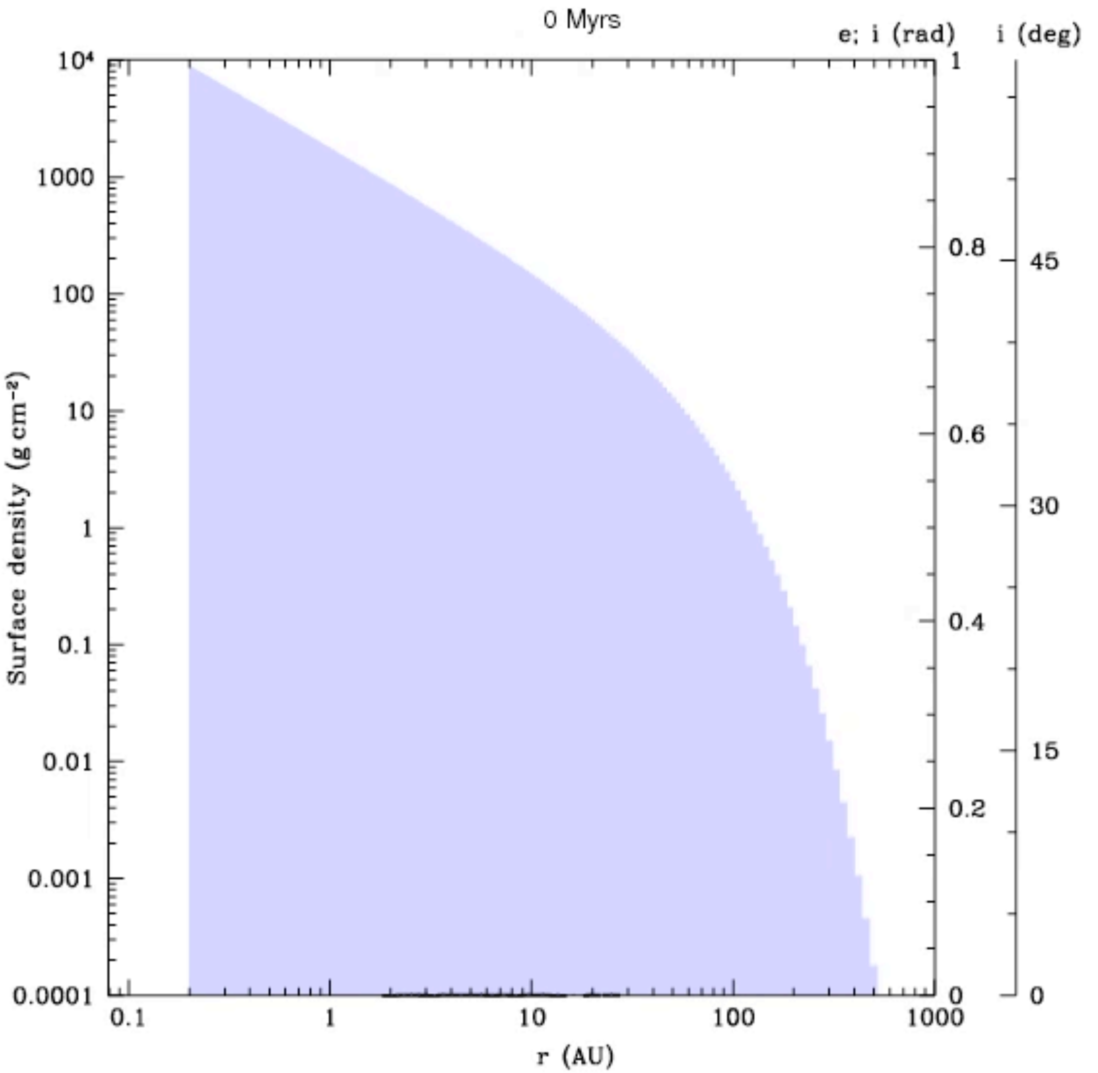
- Substituting in the value of the Hill radius gives:

$$M_{\rm iso} \sim \sqrt{\frac{(4\pi\varpi^2 b\Sigma_s)^3}{3M_*}}$$

- For gas to dust of 200:1, we have ~10g/cm^2 at 1au and an isolation mass of only ~0.1 $M_{\rm earth}$...

Side Comment: Ogliarchic Growth

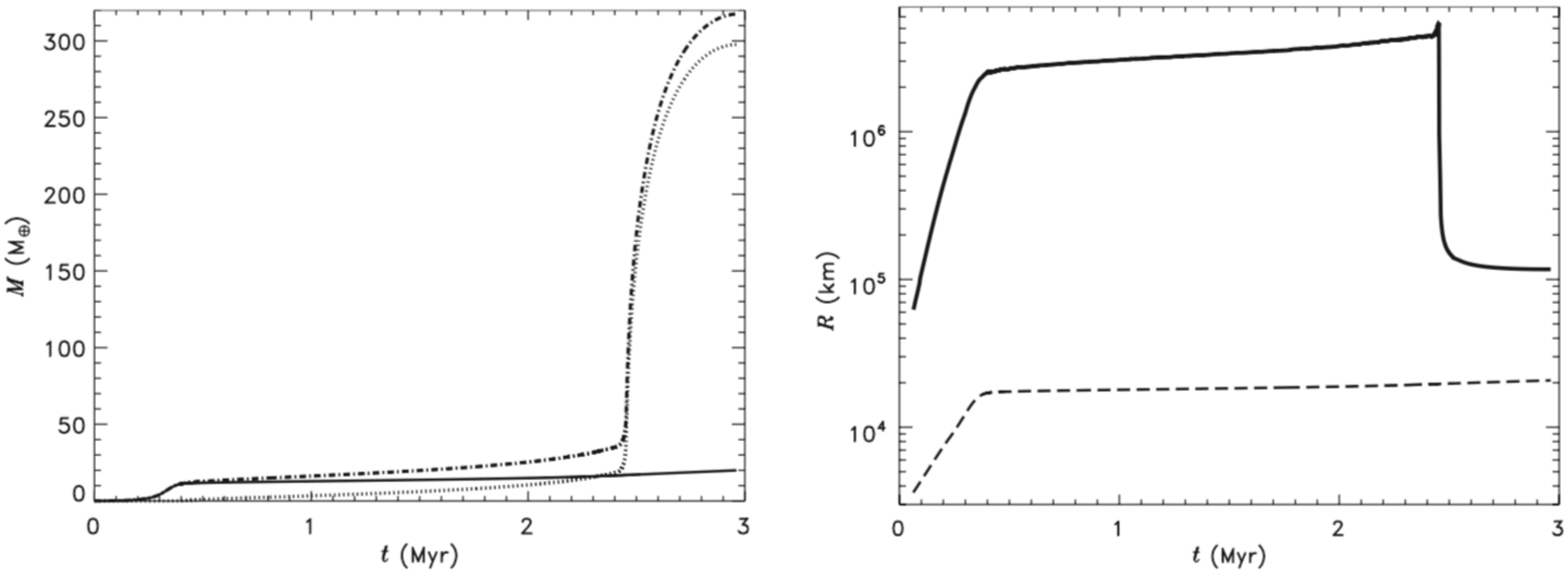
- There are 2 key ways to get beyond the isolation mass:
  - “Ocliarchic” growth, where larger protoplanets collide with smaller protoplanets, accreting them. Easier with no gas.
  - Moving the “pebbles” though the gas – pebble accretion.
- [The movie on the right, Thommes et al. (Science, 2008) was about a model of ogliarchic growth for giant planets, but hopefully you get the idea]



Gas Accretion

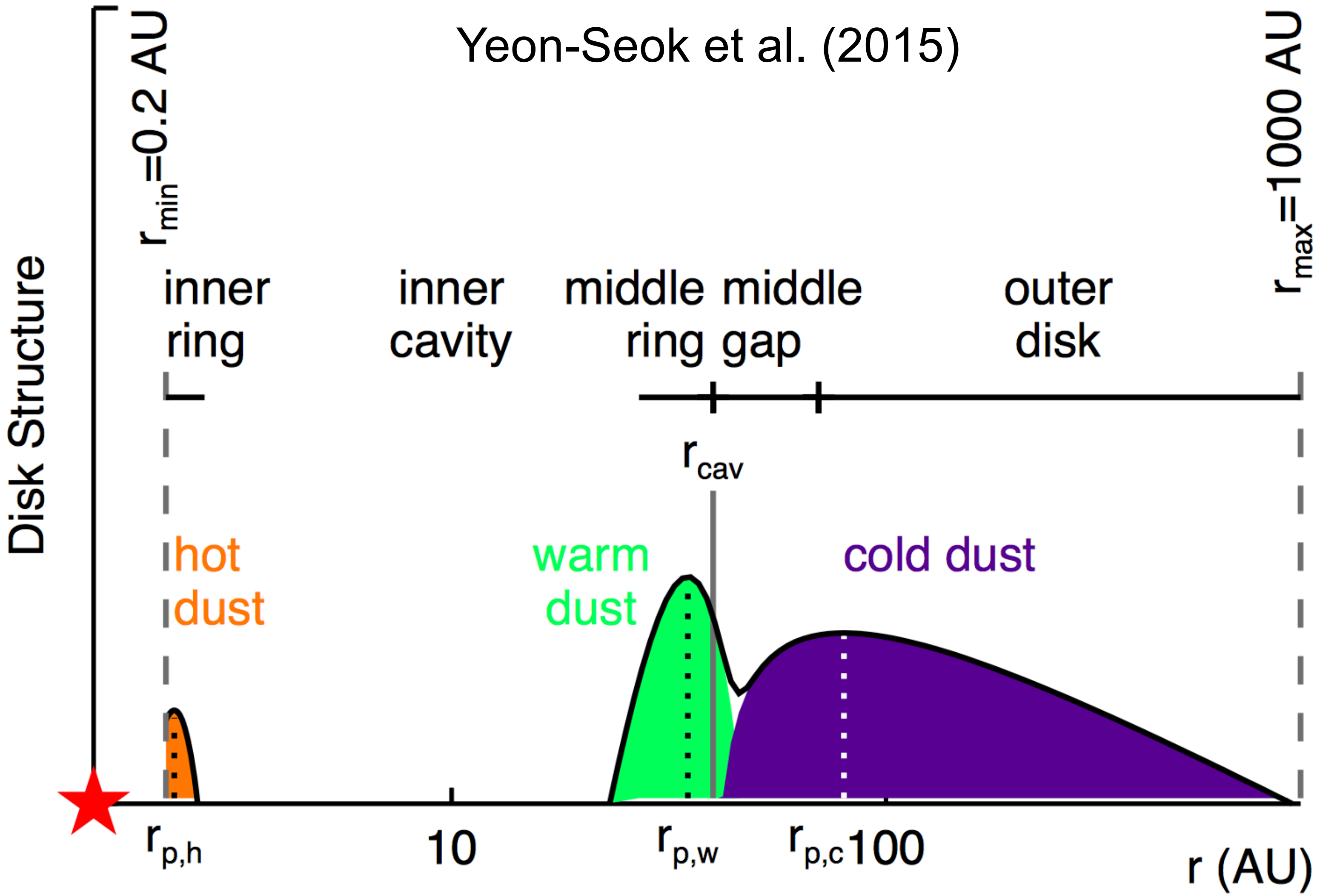
D'Angelo (2010), Lissauer (2009)

- For a long time, gas can slowly accrete with the planet in *equilibrium* with the surrounding disk medium.
- Gas can only accrete as the atmosphere cools and contracts, which is hampered by solid accretion and e.g. Al-26 decay.
- Eventually, such an atmosphere becomes unstable and collapses. After this point, any gas near the planet accelerates on to the planet and is accreted – runaway growth of a giant planet, clearing the disk!



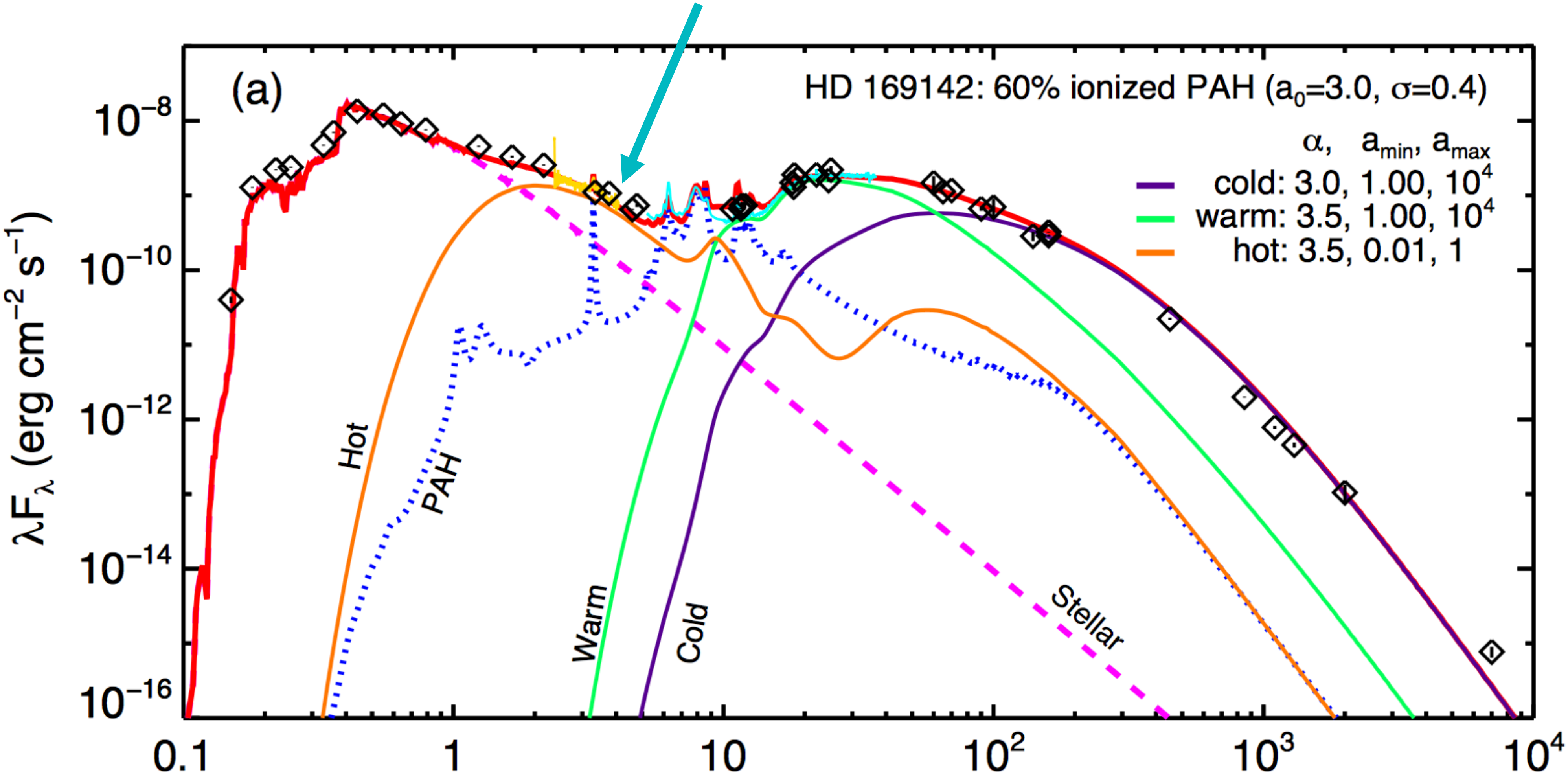
Observational Parameter Space

Example: HD 169142: Literature model



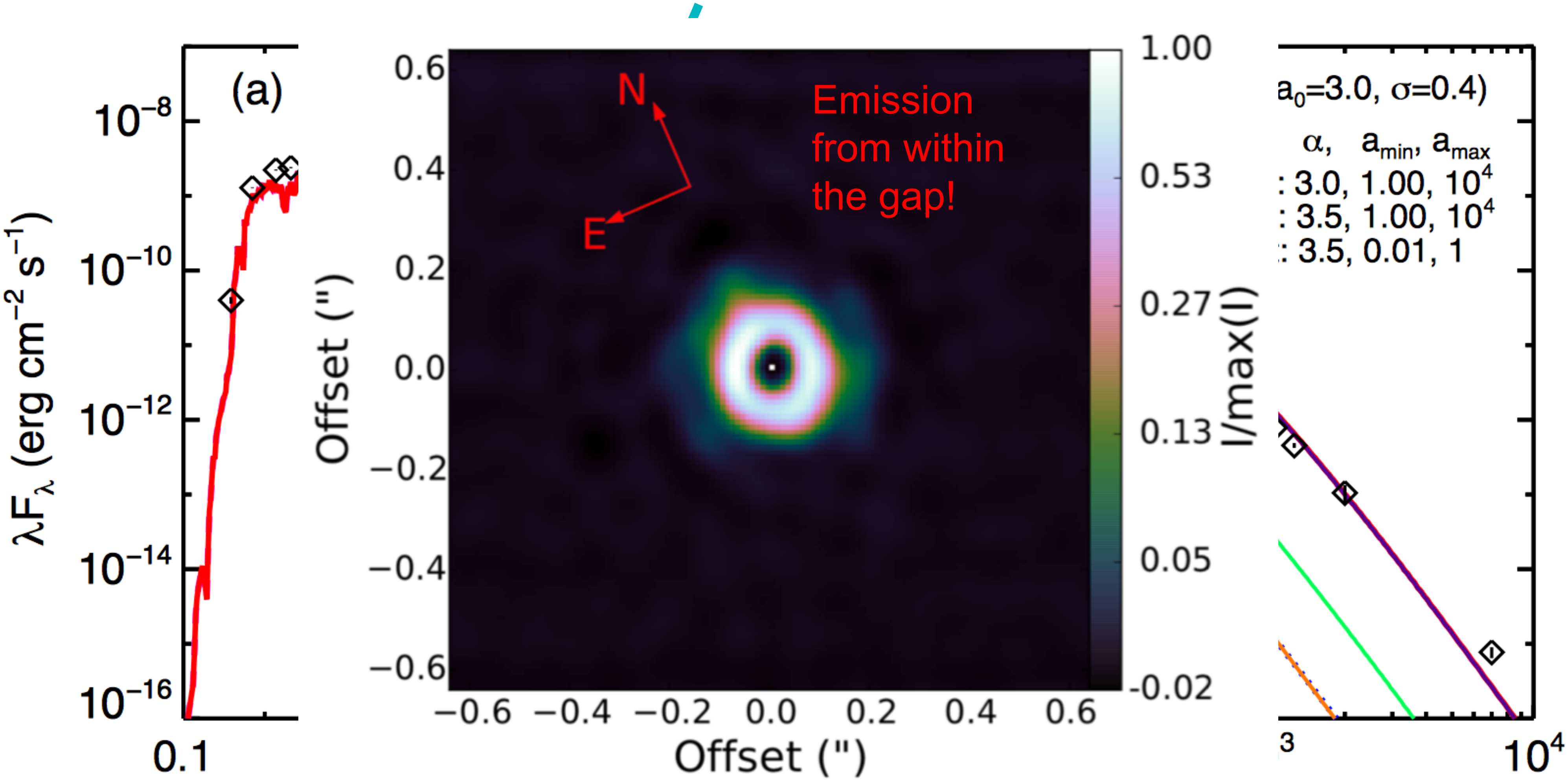
HD 169142: SED models this is based on

- 3.7 microns dominated by inner disk + star



HD 169142: SED models this is based on

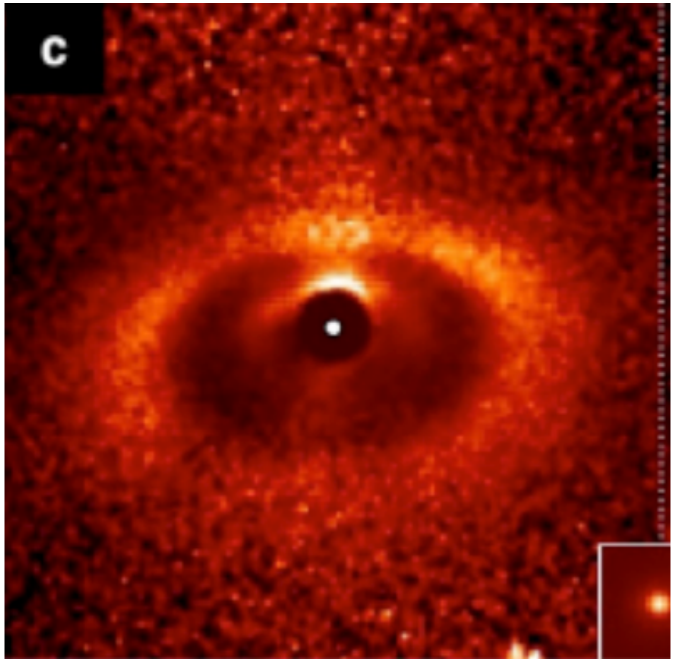
- 3.7 microns dominated by ~~inner disk + star~~



Observational Limitations

- ALMA gets down to about 3.5 AU resolution at best – similar to adaptive optics observations on large telescopes.
- Contrast is a major limitation at optical and near-infrared wavelengths (usually polarized intensity only)
- Spatial resolution and sensitivity

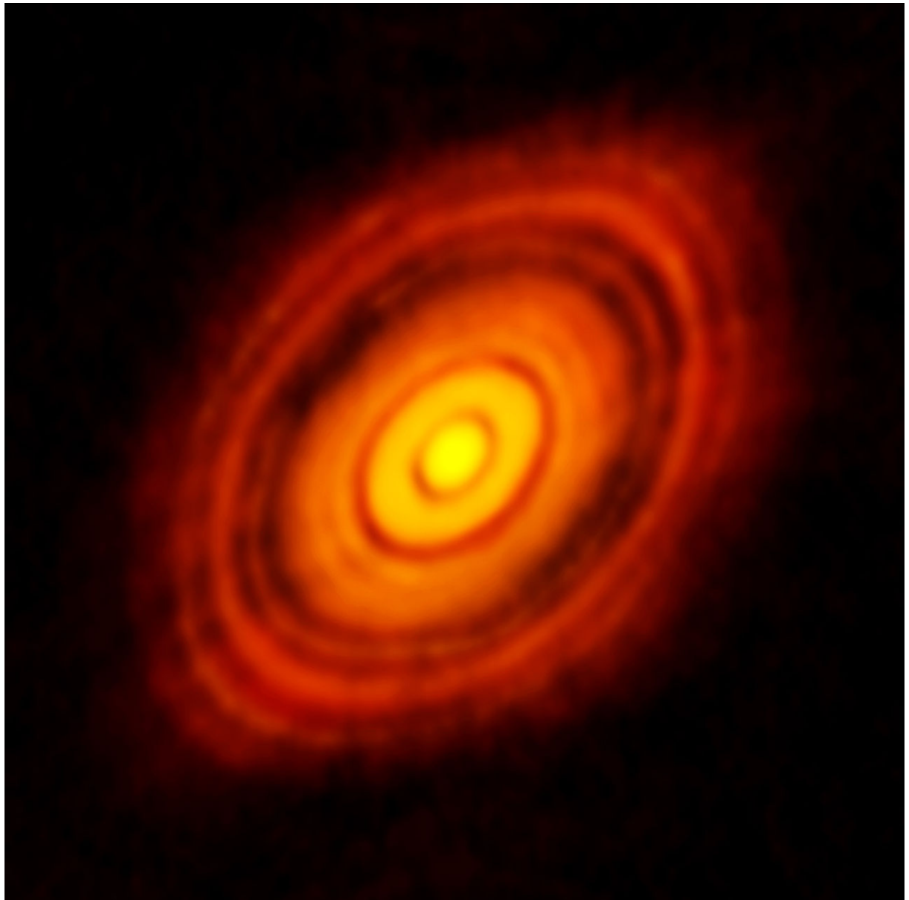
limitations at longer wavelengths



15 AU

LkCa 15

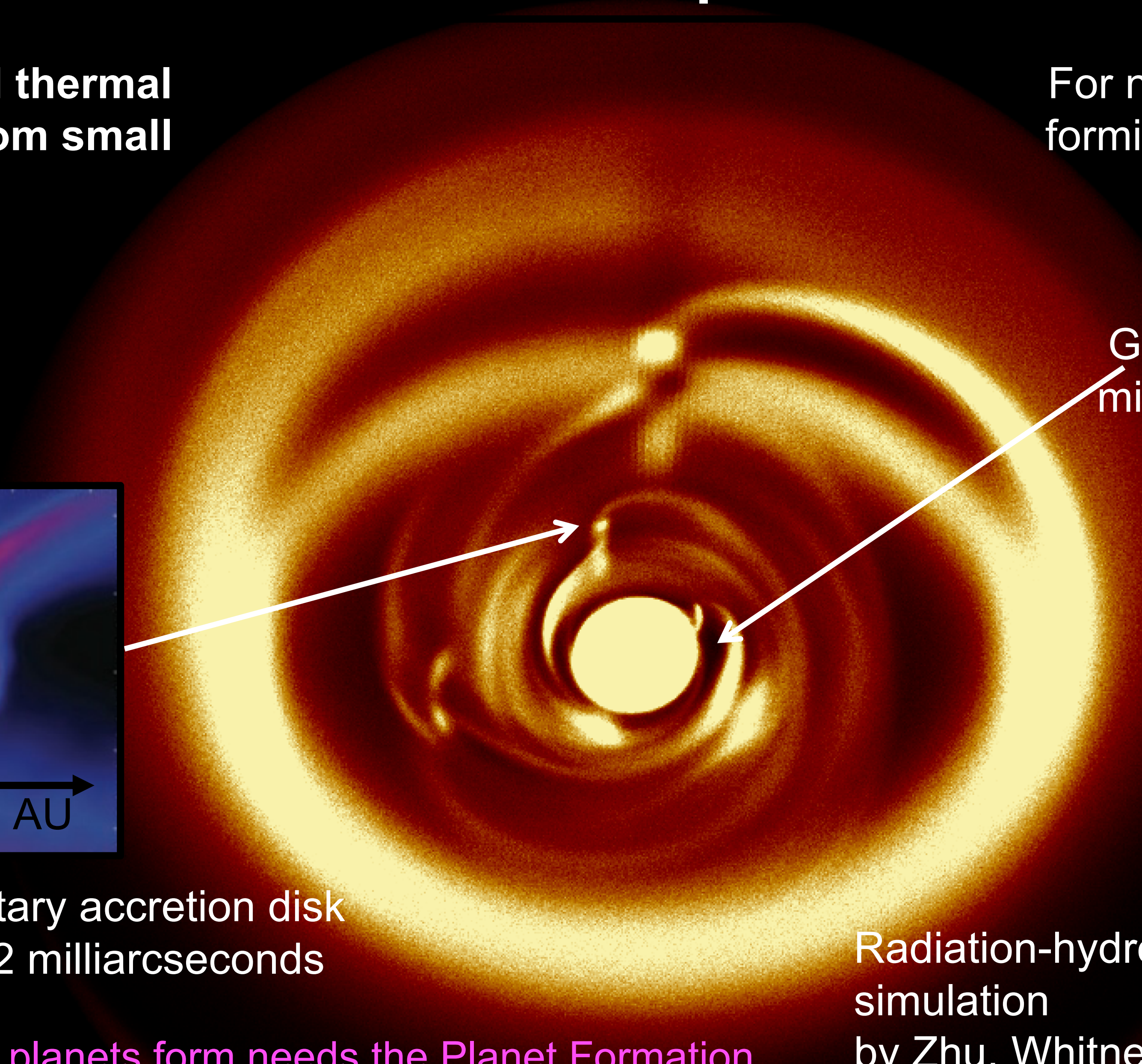
HL Tau



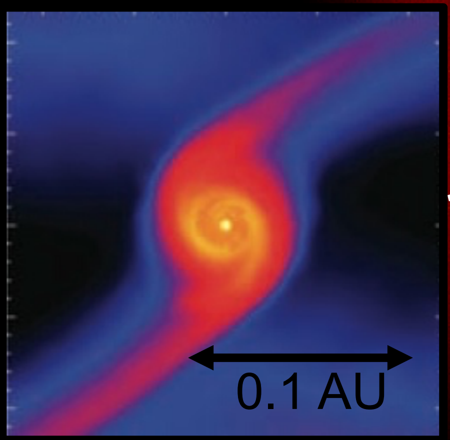
What are the relevant spatial scales?

Mid-infrared thermal emission from small dust grains

For nearby star-forming regions, d~140pc



Gaps 5AU ~35 milliarcseconds



Circumplanetary accretion disk
0.03 AU = 0.2 milliarcseconds

Really seeing planets form needs the Planet Formation Imager (ground or space-based)
http://planetformationimager.org/

Radiation-hydrodynamics simulation
by Zhu, Whitney & Dong
(Kraus et al. 2014, Ayliffe & Bate 2009)

1) What is a Transition Disk?

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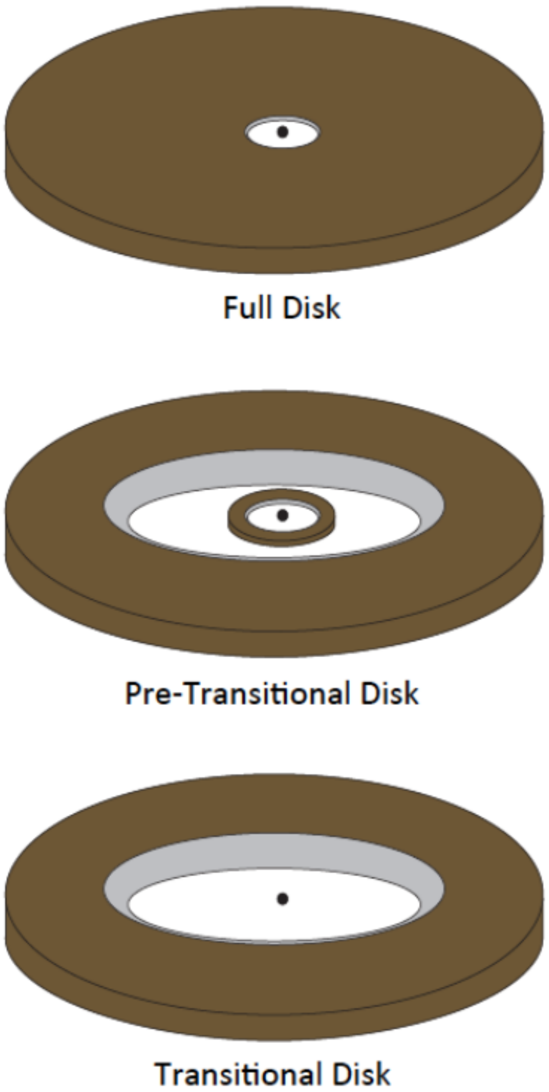
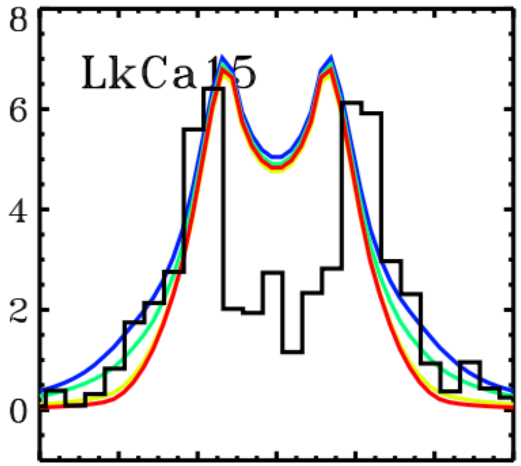
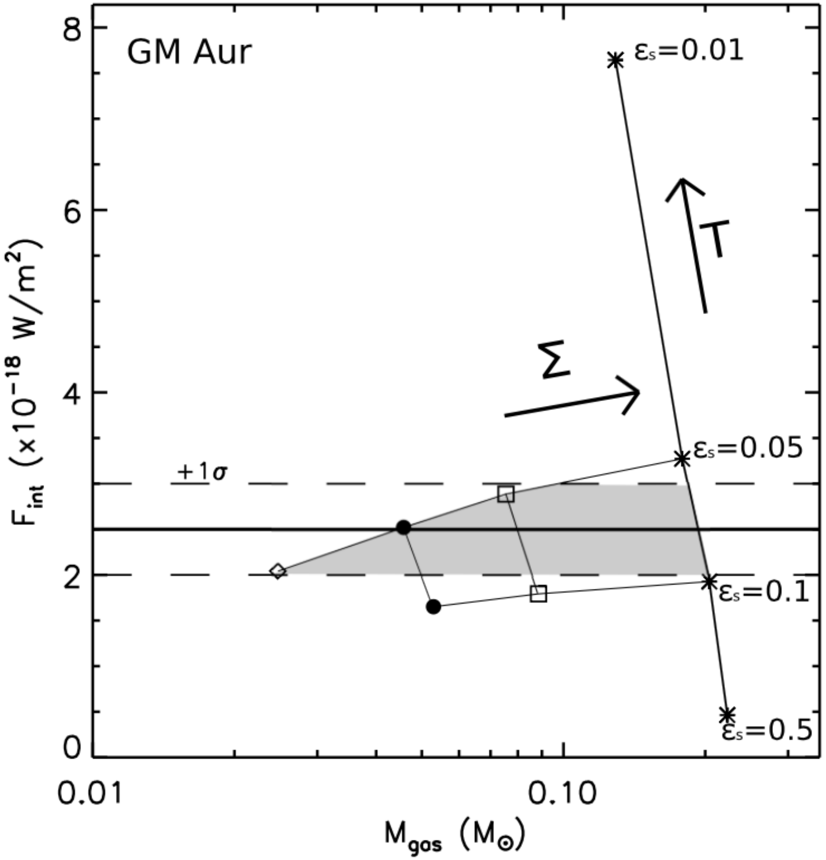
Transitional Disks…



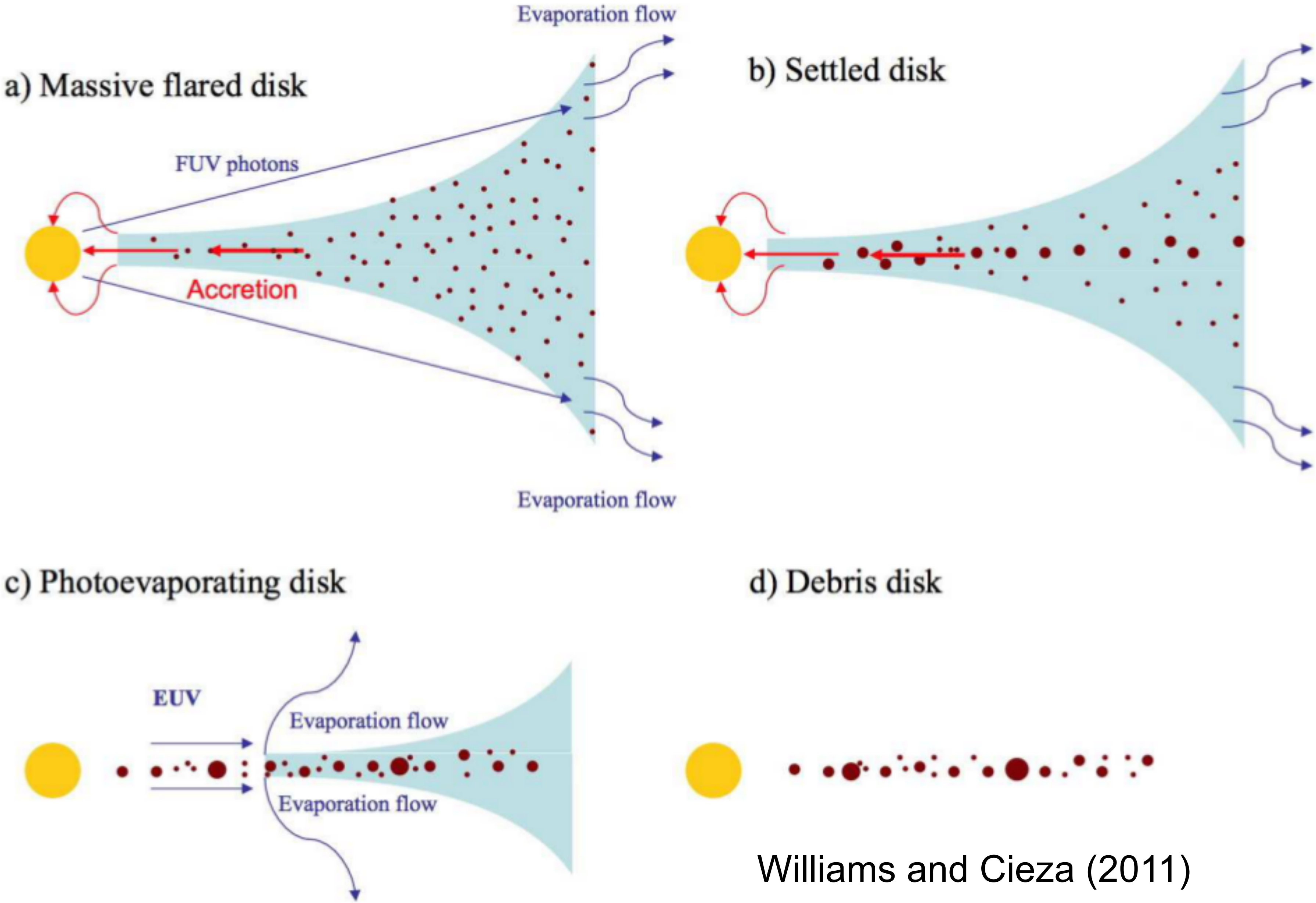
A sign of planet formation, or not?

Espaillat (2014, left), van der Marel (2016, below) and McClure (2017, right)





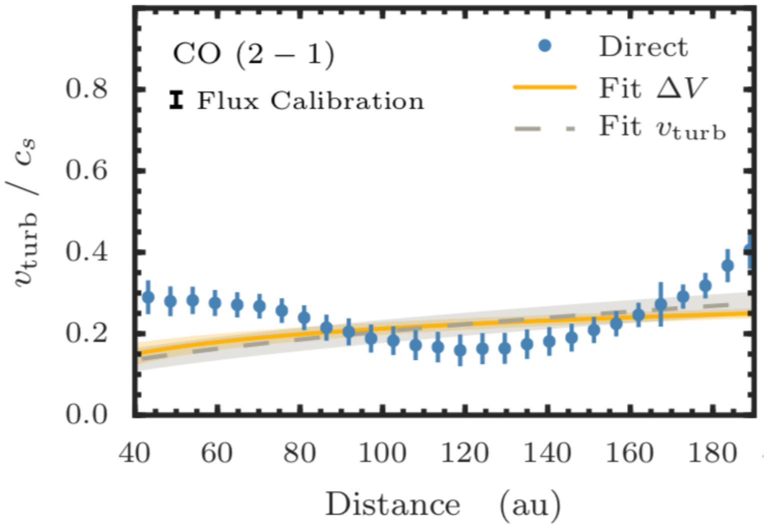
Grain Settling and Photoevaporation



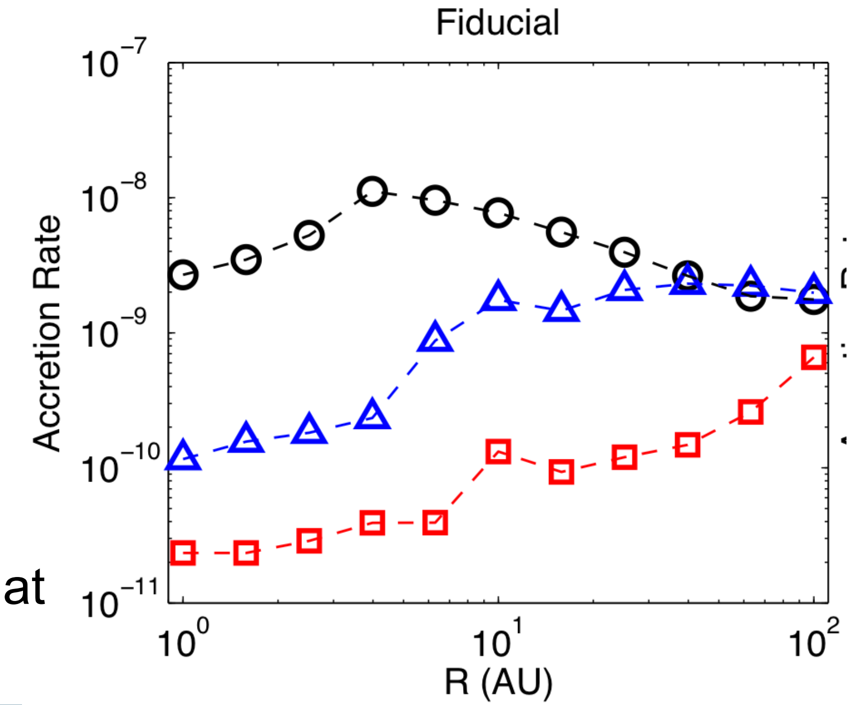
Williams and Cieza (2011)

2) What Causes Accretion in Disks?

- Youdin says that streaming instability doesn't work to produce angular momentum transport (with caveats). But it certainly mixes the disk.
- Turbulent velocities are observed to be small (~50m/s max) in outer disks. But ~0.2 c_s is almost enough….



Teague (2016) TW Hya



Right: Bai (2011) has ~10 Myr accretion timescale at MMSN

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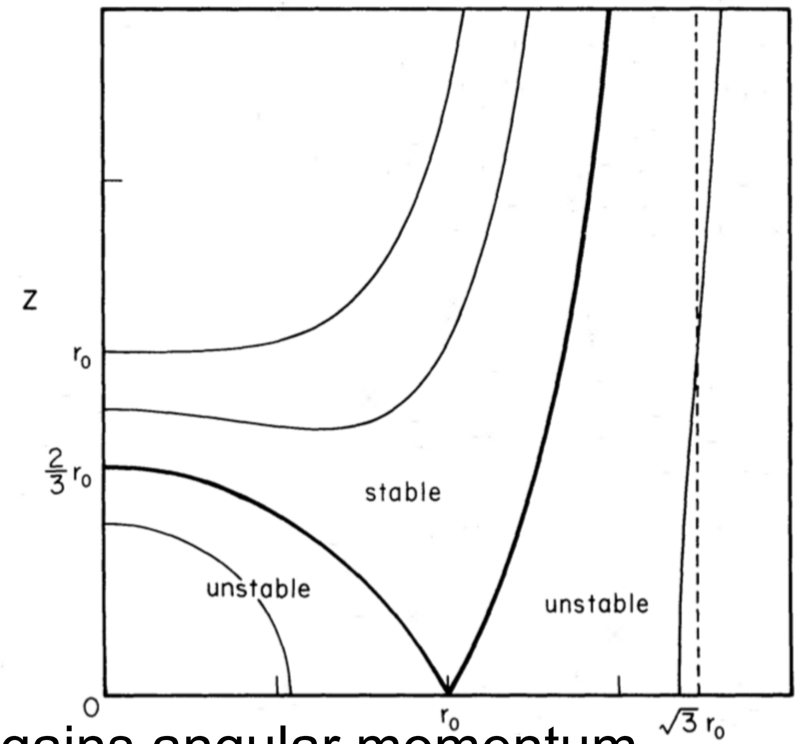
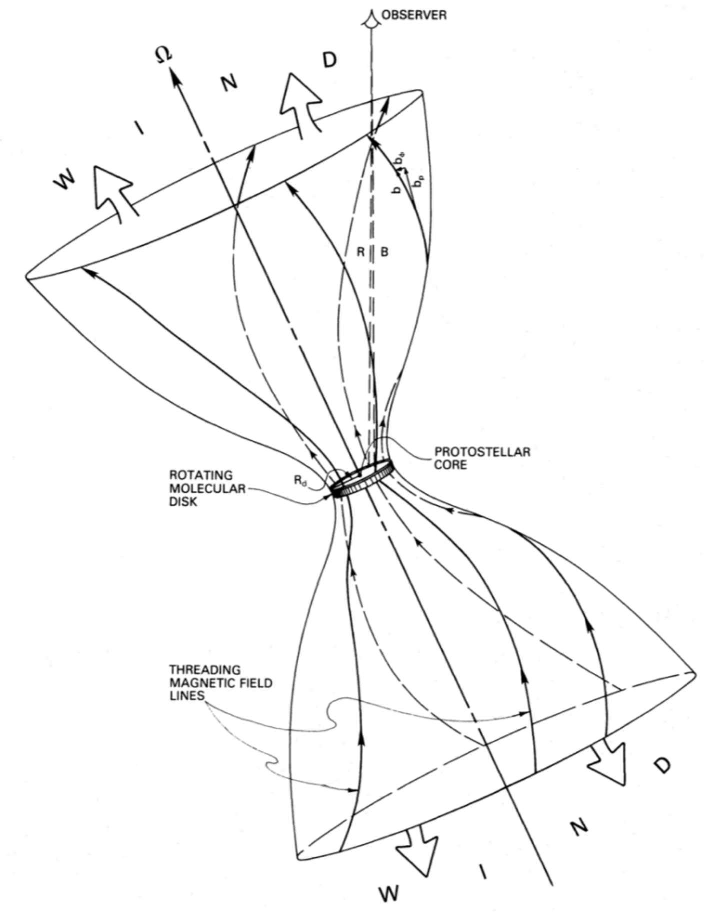
Outflows and Angular Momentum Loss

- In the ideal MHD approximation, mass can only move along magnetic field lines, like a bead on a wire.
- Where field lines make an angle of less than 60 degrees to the disk plane, matter accelerates outwards.



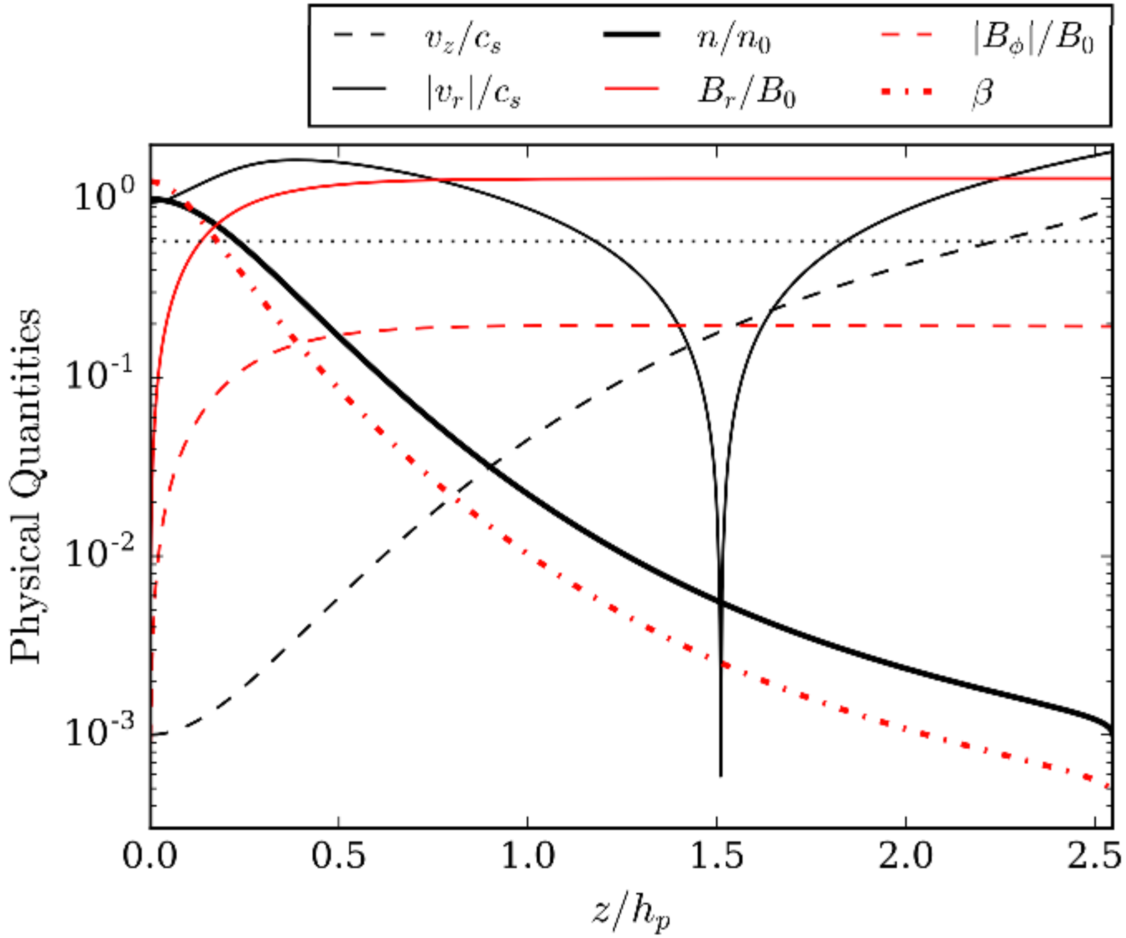
Blandford and Payne (1982)



Pudritz and Norman (1986)

- The outflow gains angular momentum, while the star or disk loses angular momentum.
- Controversy about where disks are launched from (disk, X-point), with few observational constraints due to angular resolution requirements.

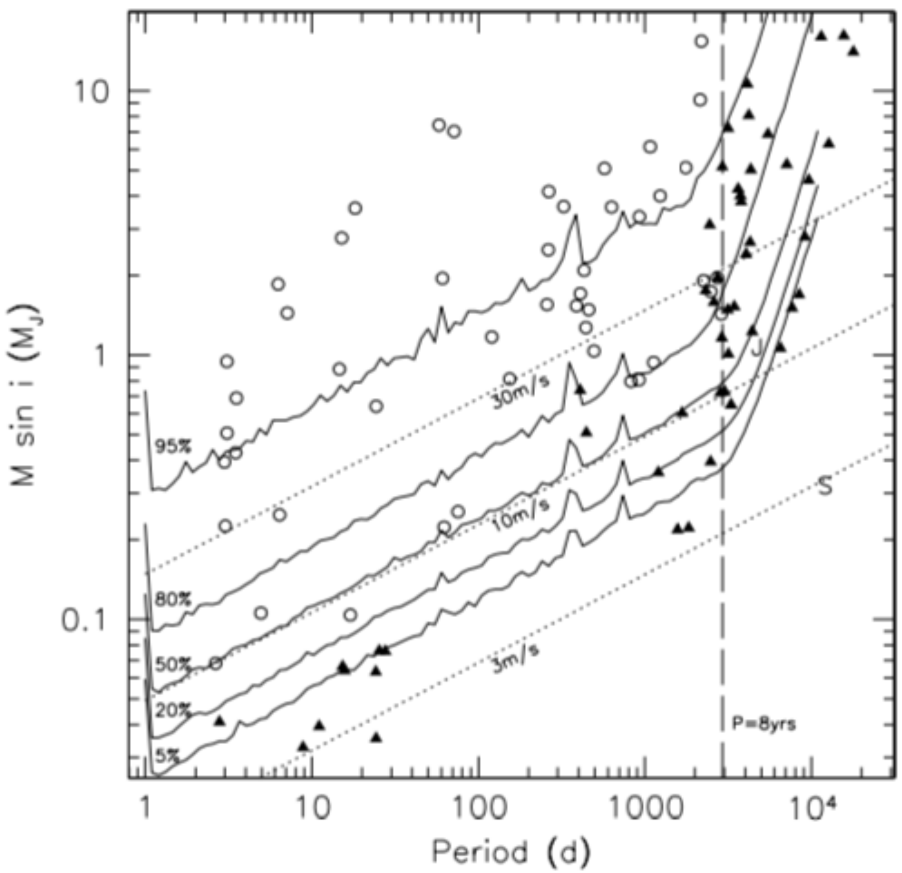
- But gas isn't sufficiently ionized fpr ideal MHD!
- Ordered magnetic fields mean that winds can still drive accretion and viscosity isn't needed. Probably.



Wang and Goodman (2017): Konigl and Wardle winds drive accretion *at the sound speed*.

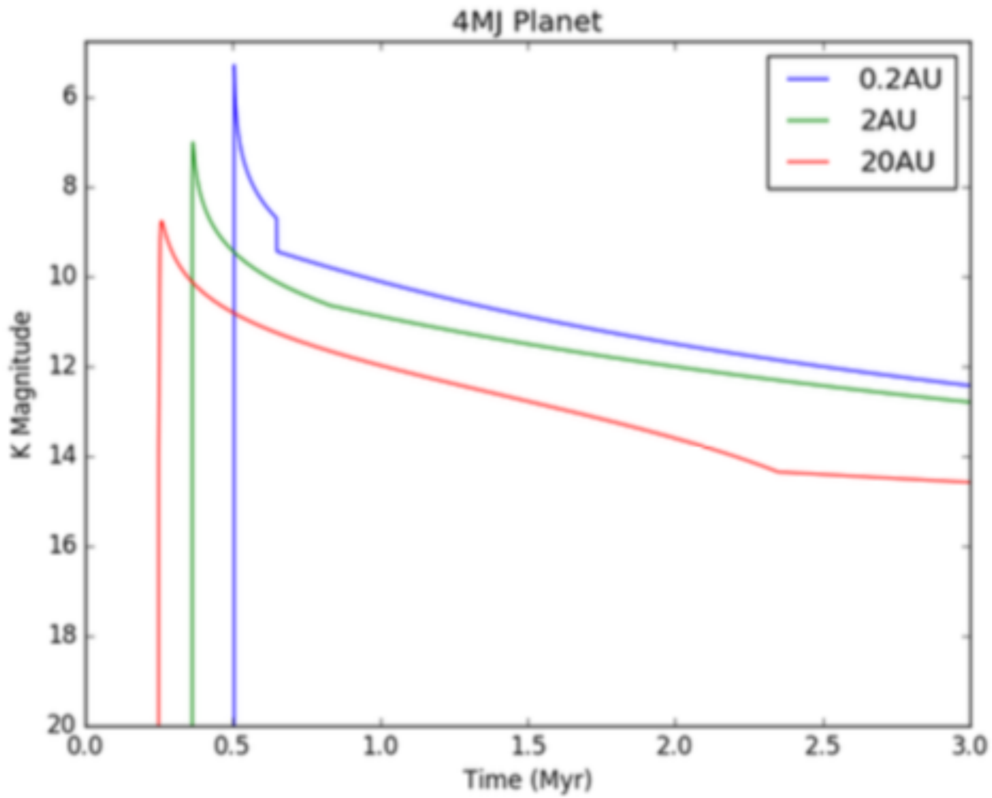
3) Do planetesimals retain their ice, and if not, is the ice line still important?

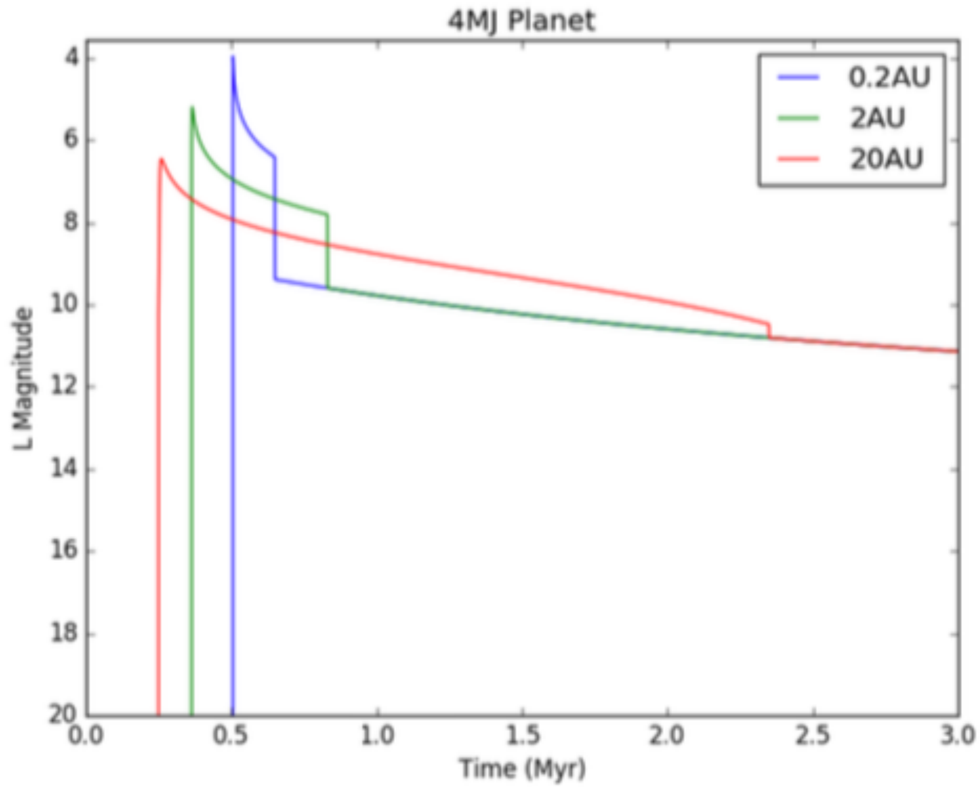
- The "Ice Line" was based on an unusually warm back-of-the-envelope idea of direct stellar illumination.
- Even Cumming (2008) showed that there were not many ~2AU Jupiters suddenly turning up.
- The lack of a follow-up paper by the obvious groups (and verbal conversations) shows that the next paper is a non-detection paper…



4) What is the initial entropy of Giant Exoplanets?

- Hot-start versus cold-start.
- Alex working on this but a long way to go…





4) What determines the final mass of a giant planet? (post hydrodynamic collapse)

- Initial modelling papers are pretty unsatisfactory… they give just a few example calculations designed to match e.g. our Jupiter (Lissauer…)

Australian National University

Full Week Outline

- Big picture intro
- Disk Physics (intro and key references)
  - Hydrostatic equilibrium in a disk plane (Ch. 15 revision)
  - Viscosity (and an intro to α-disk models, Ch. 15 revision)
  - Gravitational instability (Ch. 15 Revision)
  - Minimum mass solar nebula.
  - Radiation transport
  - Grain Growth and gas drag (Mark to cover growth details later)
  - Planetesimal formation (Mark to cover limits to Kelvin-Helmholz instability)
  - Rocky Planet Formation
  - Gas giant formation via core accretion
- Observational Parameter space
- Selected Open Questions
  1. What is a Transition Disk?
  2. What causes accretion in protoplanetary disks?
  3. What is the initial entropy of a giant exoplanet?
  4. What causes the final mass of a giant exoplanet?
- Ask Mark about Dust Traps…