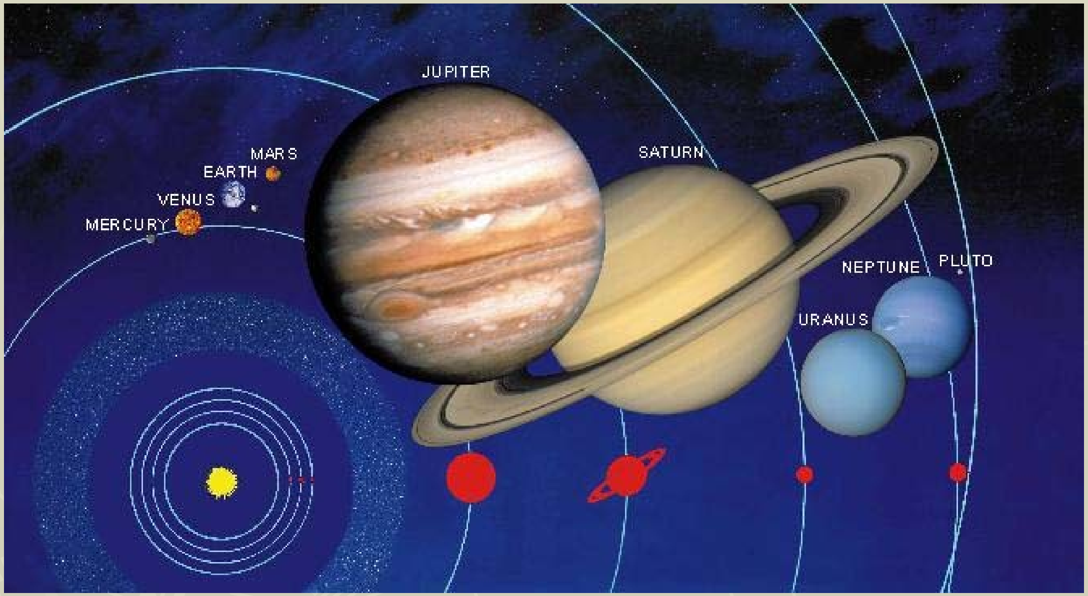
The Solar system:

1) Physical and dynamical properties of planets, in particular giant planets, and their precursors i.e. asteroids and comets.

2) The standard model of planet formation: from circumstellar disks to planets via planetesimals and pebbles.

Bodies in the solar system:

Planets

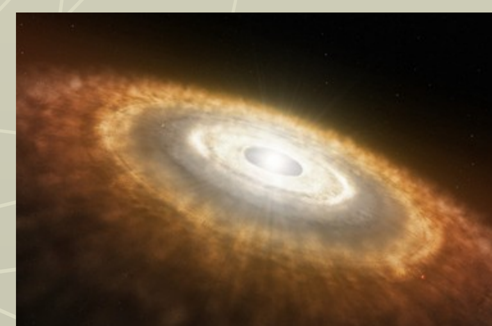


Asteroids

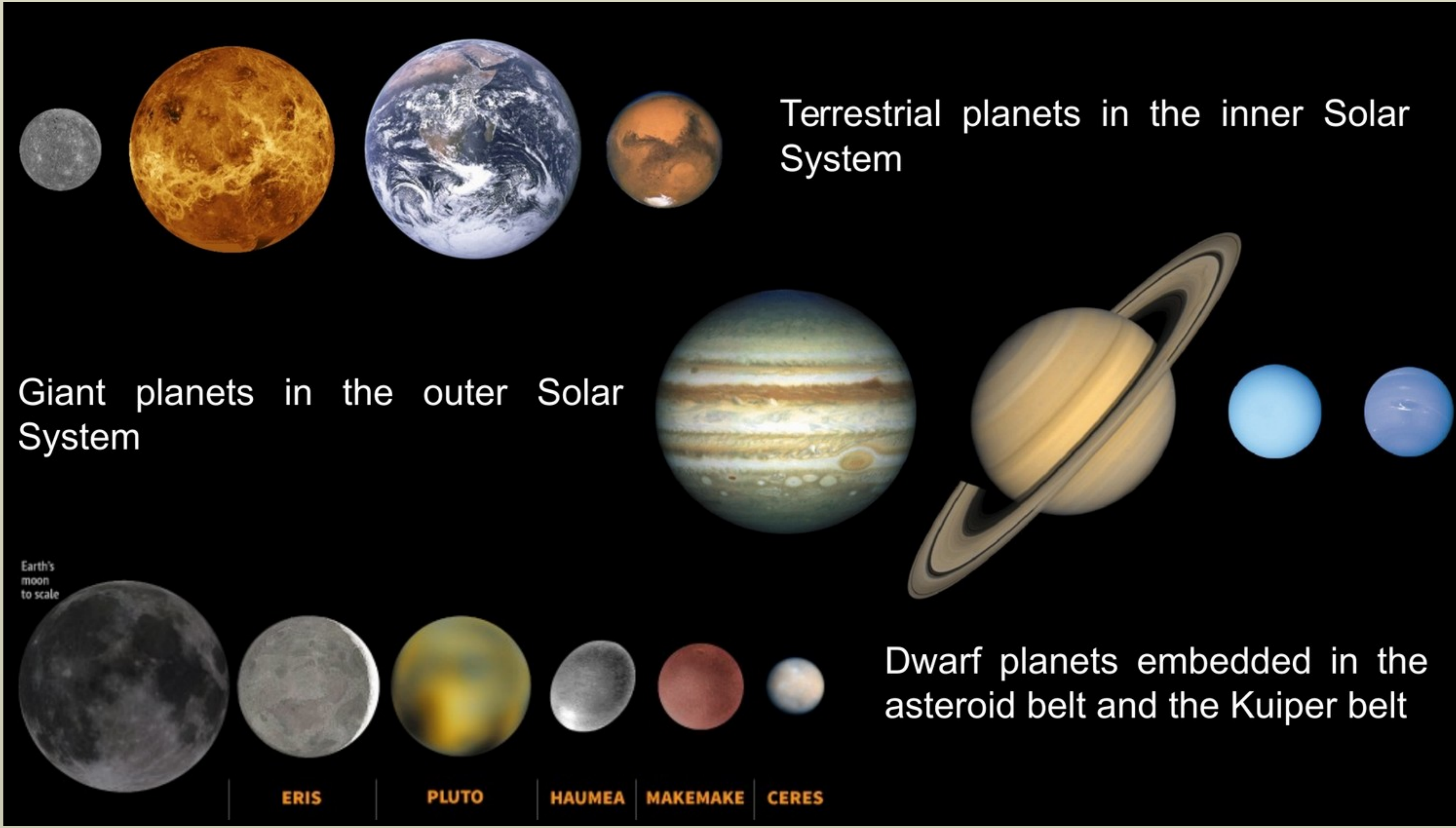
Comets

Moons

Dust particles



Different kinds of planets.



Terrestrial planets: close to the star, rocky and metallic, high density.

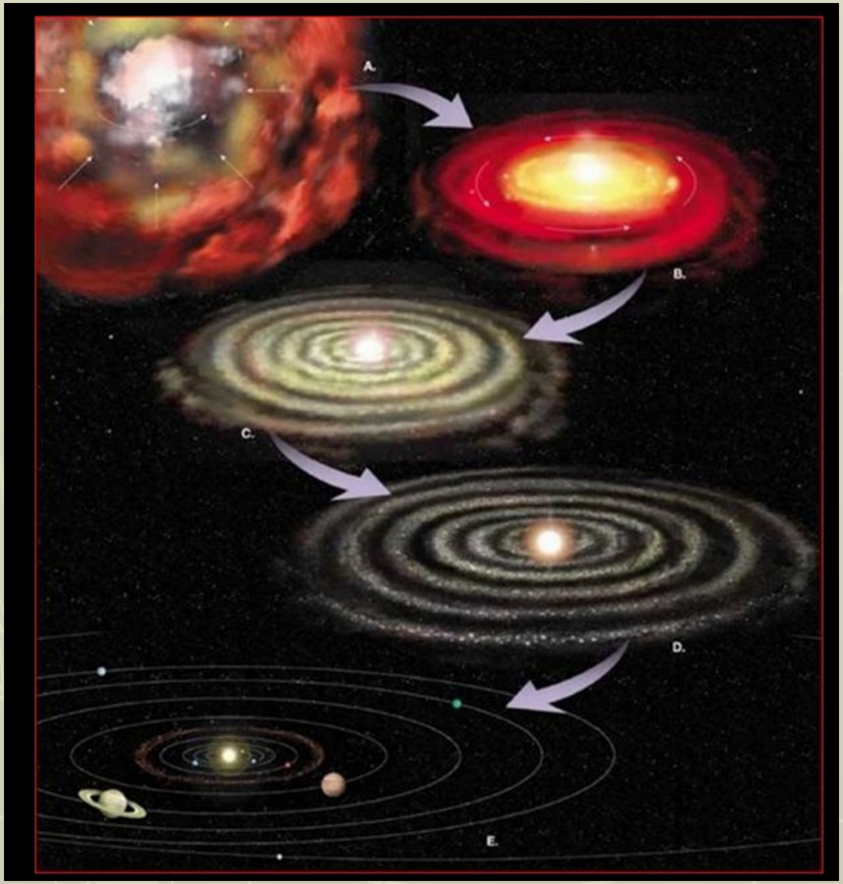
Pianeta	Terra	Mercurio	Venere	Marte
Diametro equatore(Km)	12756	4878	12104	6794
Densità (kg/m')	5517	5500	5250	3933
Mass a(Terra=1)	1	0.055	0.815	0.107
Gravità (Terra=1)	1	0.38	0.903	0.38
Velocità di fuga (km/s)	11.2	4.3	10.36	5.03
Distanza dal sole (UA)	1.00000	0.38710	0.72333	1.52369
Distanza dal sole (10^6 Km)	149.6	57.9	108.2	227.9
Periodo orbitale (anni)	1	0.241	0.615	1.88
Velocità orbitale (Km/sec)	29.79	47.89	35.03	24.13
Temperature superficie (C)	10	127	457	63

Giant and icy planets: gas, ice and rock. Far from the star.

Pianeta	Giove	Saturno	Urano	Neptuno	
Diametro equatore (Km)	142796	120000	50800	48600	
Densità (kg/cm(3))	1330	706	1270	1700	
Massa (Terra=1)	318.832	95.162	14.536	17.139	
Gravità (Terra=1)	2.643	1.159	1.11	1.21	
Velocità di fuga(km/s)	60.22	32.26	22.5	23.9	
Distanza dal sole (UA)	5.20248	9.53884	19.18184	30.05798	
Distanza dal sole (10^6 Km)	778.3	1427	2869	4497.1	
Periodo orbitale (anni)	11.867	29.461	84.013	164.793	
Velocità orbitale (Km/sec)	13.06	9.64	6.81	5.43	
Temperature superficie (C)	-148	-178	-213	-213	

PLANET FORMATION

The standard model developed for the Solar System



1) Gas+dust cloud collapses into a protostar

2) A circumstellar disk of gas and dust forms around the star

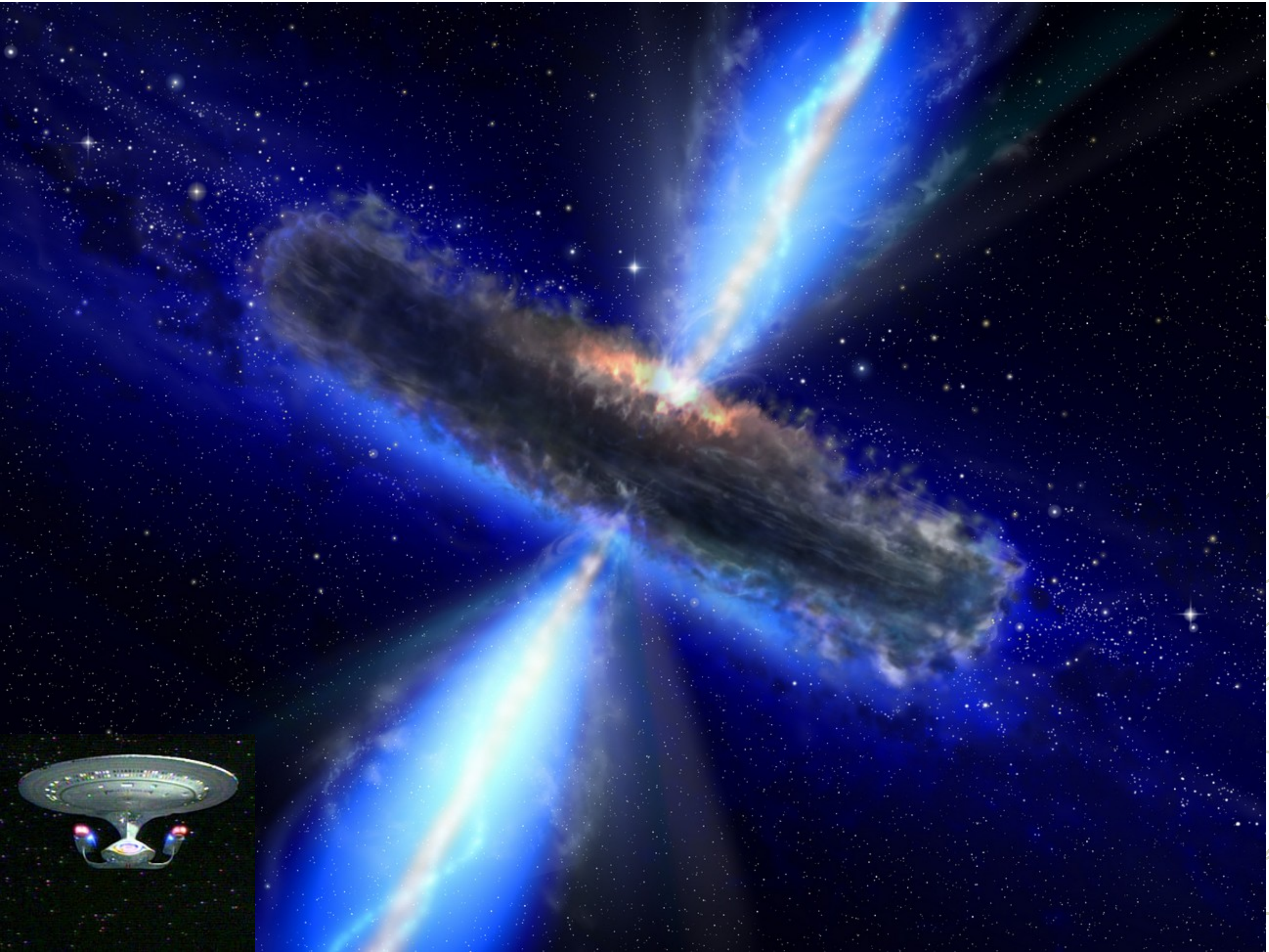
3) From the dust, pebbles and planetesimals (D>1-100 km) grow

4) Collisions and gravitational accumulation into planet embryos

5) Terrestrial planet formation and core of giants.

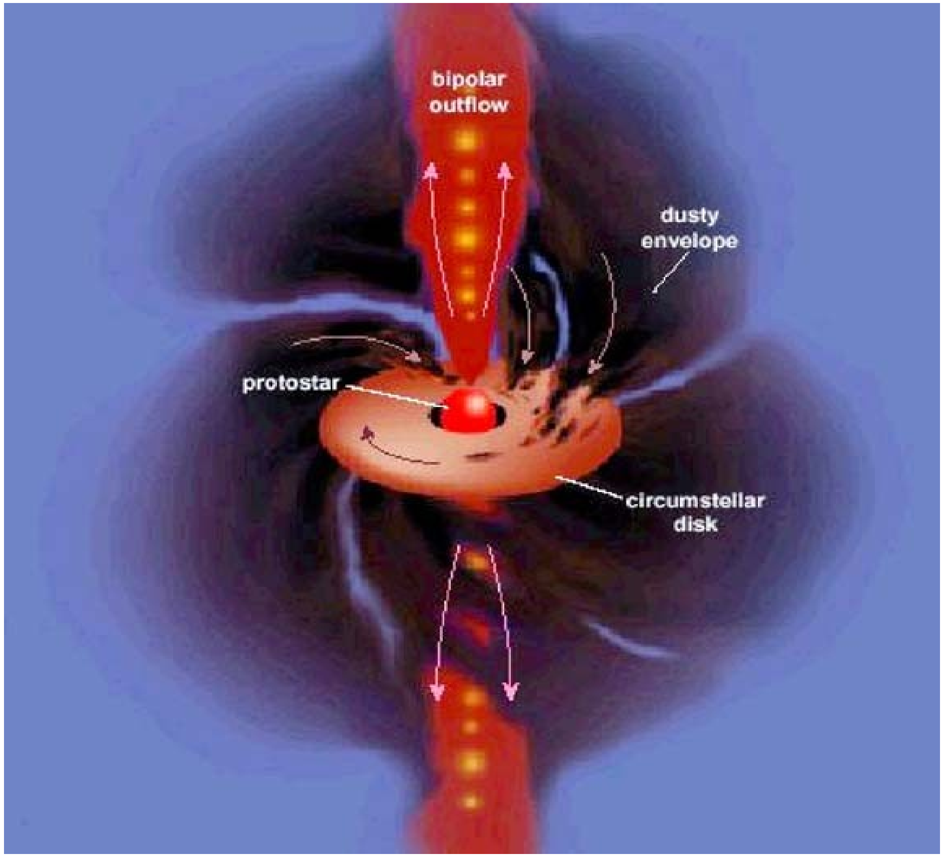
6) Final infall of gas on the core: giant planet.

1. Circumstellar disks

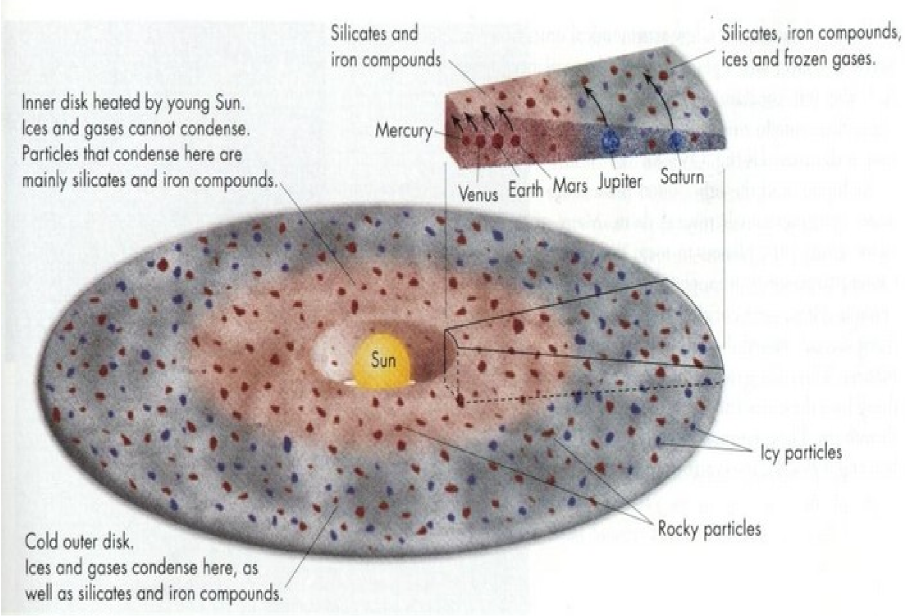


Circumstellar disks: gas + dust (~1/100 ratio)

The initial evolution of the disk is quite turbulent (gas and dust infall from the envelope, jets...) but it is followed by....



….. a more quiet state



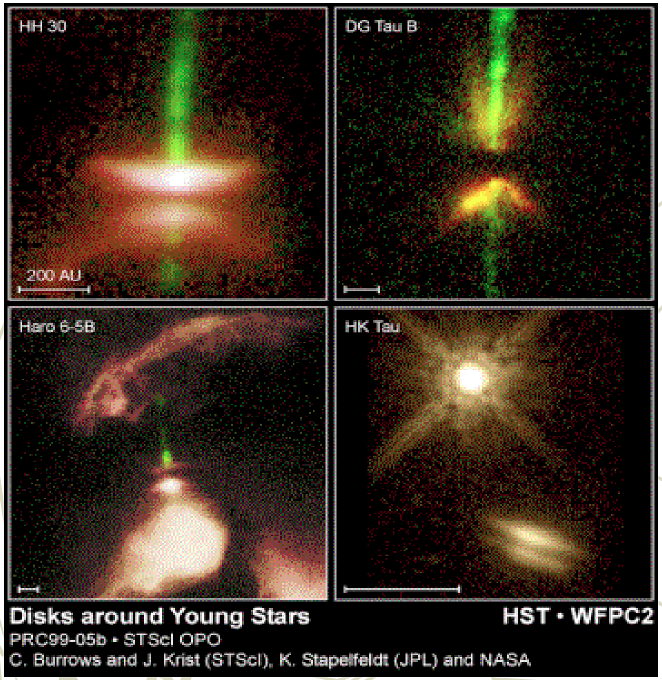
$M_d \sim 0.003 – 0.3\ M_S$

Detection methods:

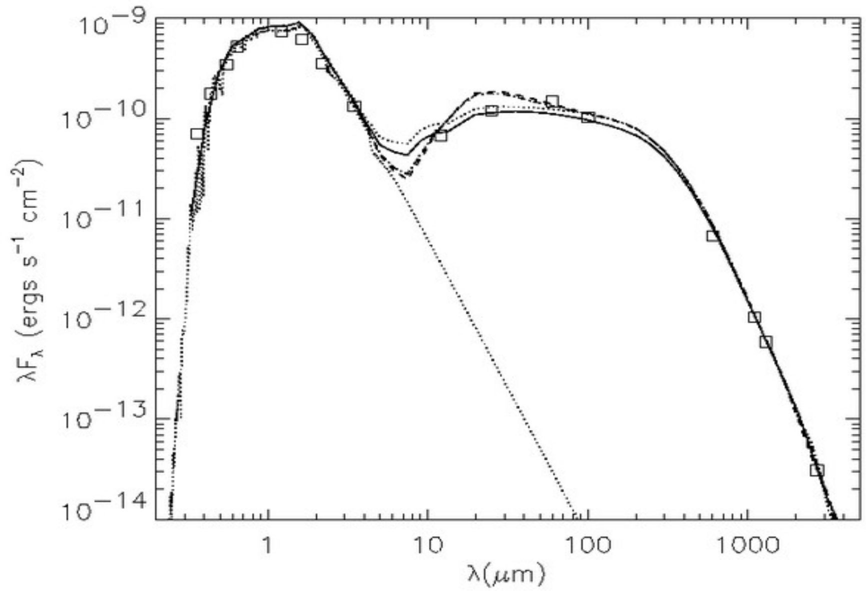
1) Infrared excess in the star spectrum

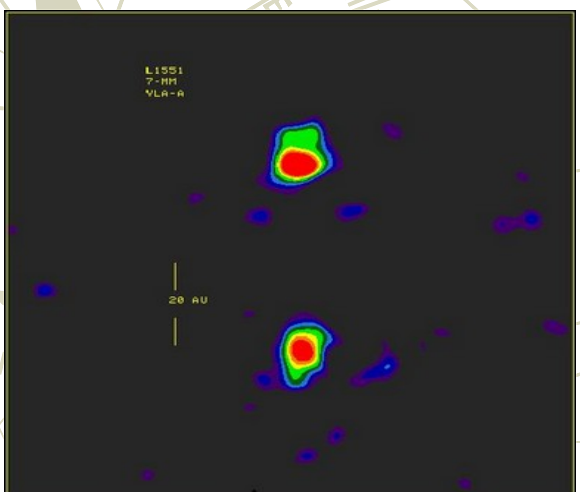
2) Direct imaging (HST, ALMA, SPHERE)

HST images of circumstellar disks around T-Tauri stars

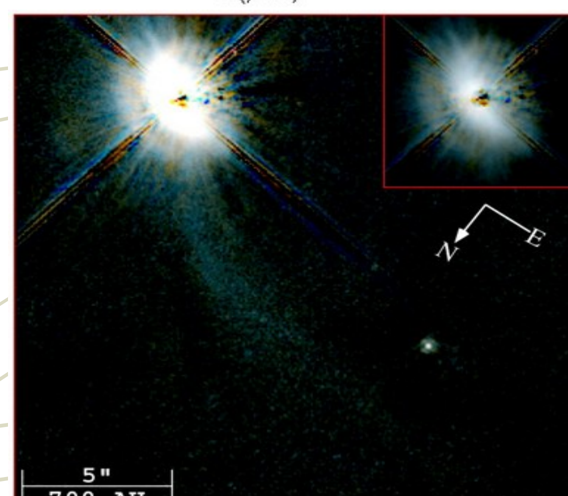


GM Aur disk. Top: infrared excess. Bottom: coronograpic image from HST. R~ 300 AU $M=0.01\ M_{S.}$ The disk is truncated inside, the SED has a dip.





Disk in a binary system L1551-IRS5 by Rodriguez et al. (1993) imaged in the radio to see the dust.



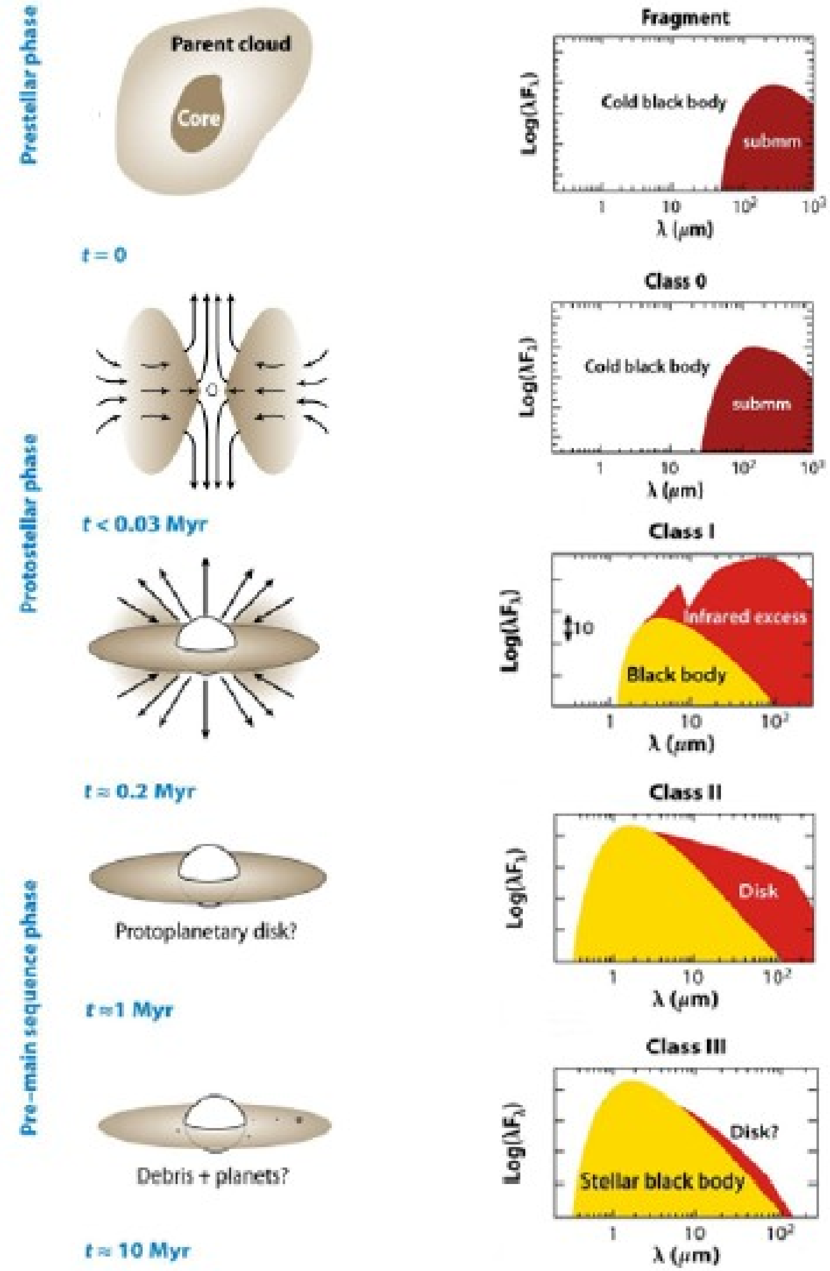
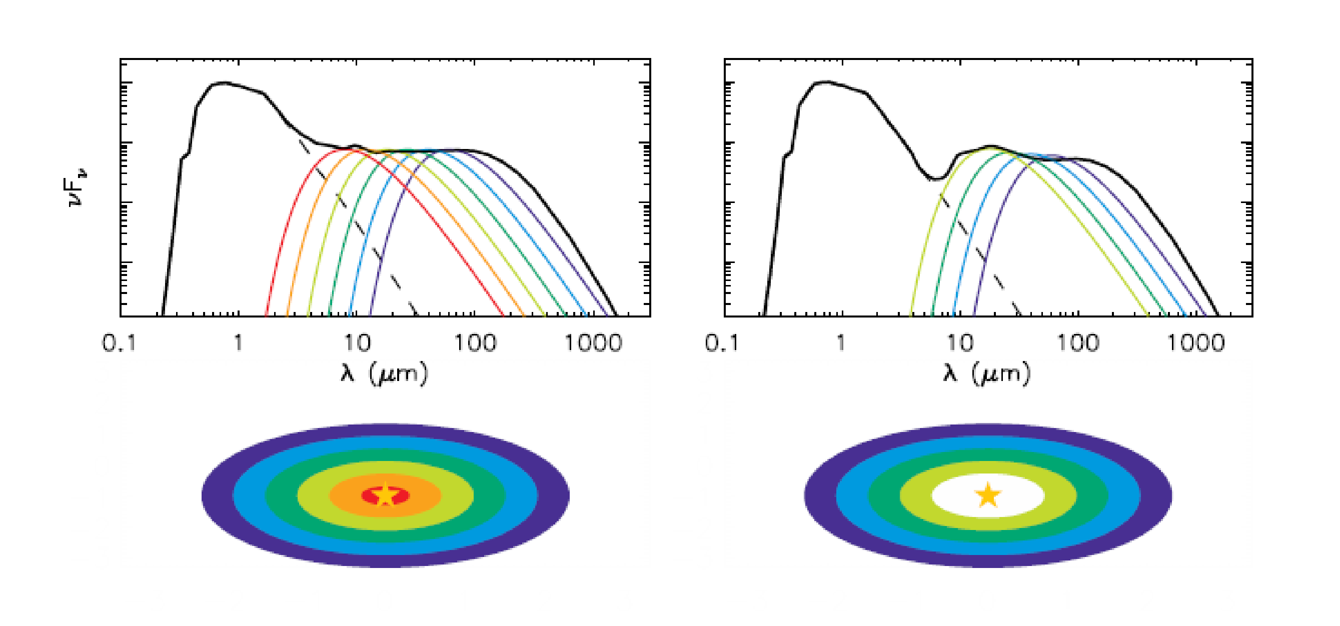


Figure 1.1: Cartoon image illustrating the star and formation process (left) with the expected SEDs (right). A dense molecular cloud collapses under its own gravity and, due to its initial rotation, forms a disc which carries most of the angular momentum of the primordial nebula and slowly accretes onto the central star. Image adapted from André (2002).

The infrared excess changes according to the disk evolution. As the disk evolves, the density and temperature profiles are different.

The infrared excess gives valuable information on the density and temperature profiles of disks.



Transition disk

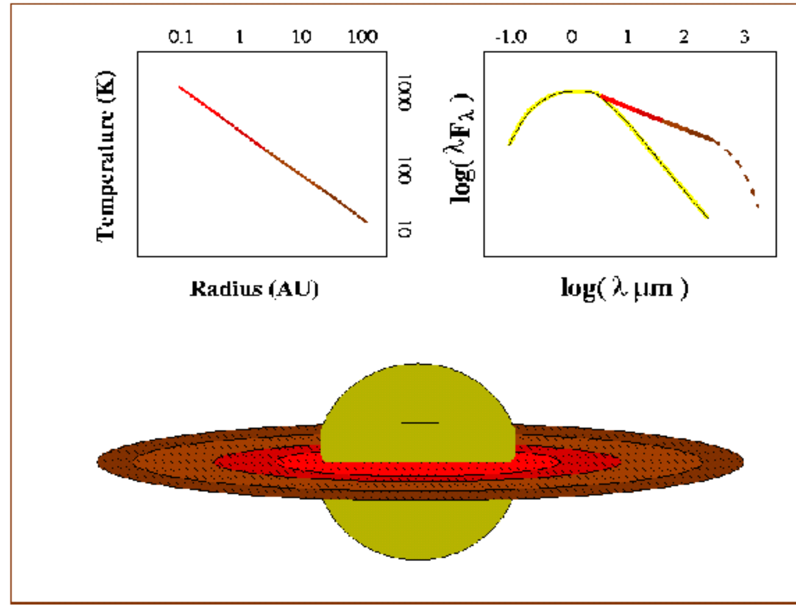
Standard power law disk model **Beckwith et al. (1990)**

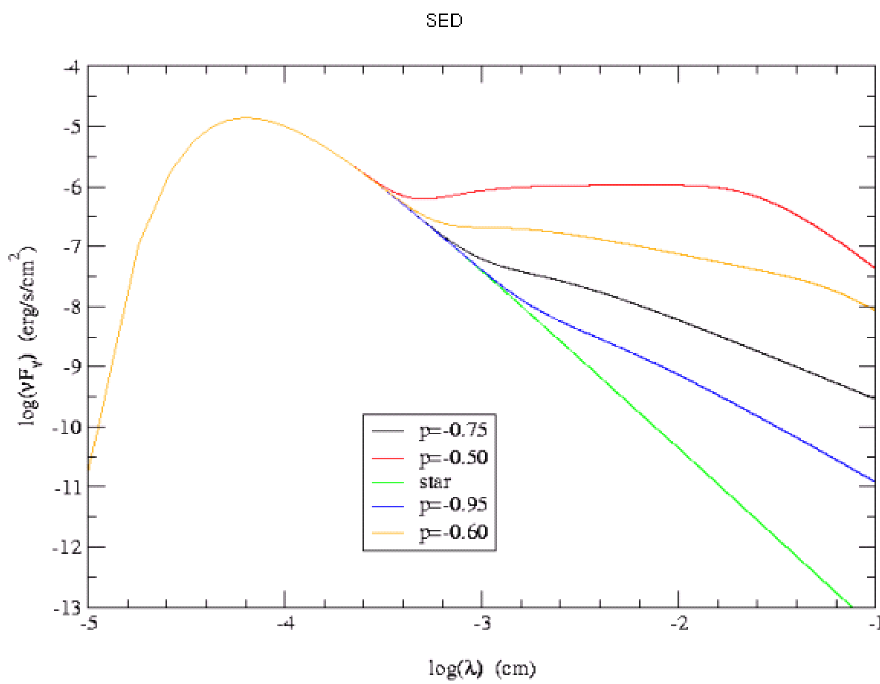
$$\Sigma = \Sigma_0 \left(\frac{R}{R_0} \right)^q$$

$$\Sigma_0 \approx 10^2 - 10^4 \ g/cm^2$$

$$T(R) = T_0 \left(\frac{R}{R_0} \right)^p$$

$$T_0 \approx 1000 - 2000 \ K$$





Solar type stars, R_0 = 1 AU

q ~ -1/2, -1, -3/2 …..
p ~ -1/2, -3/4, -3/7…..

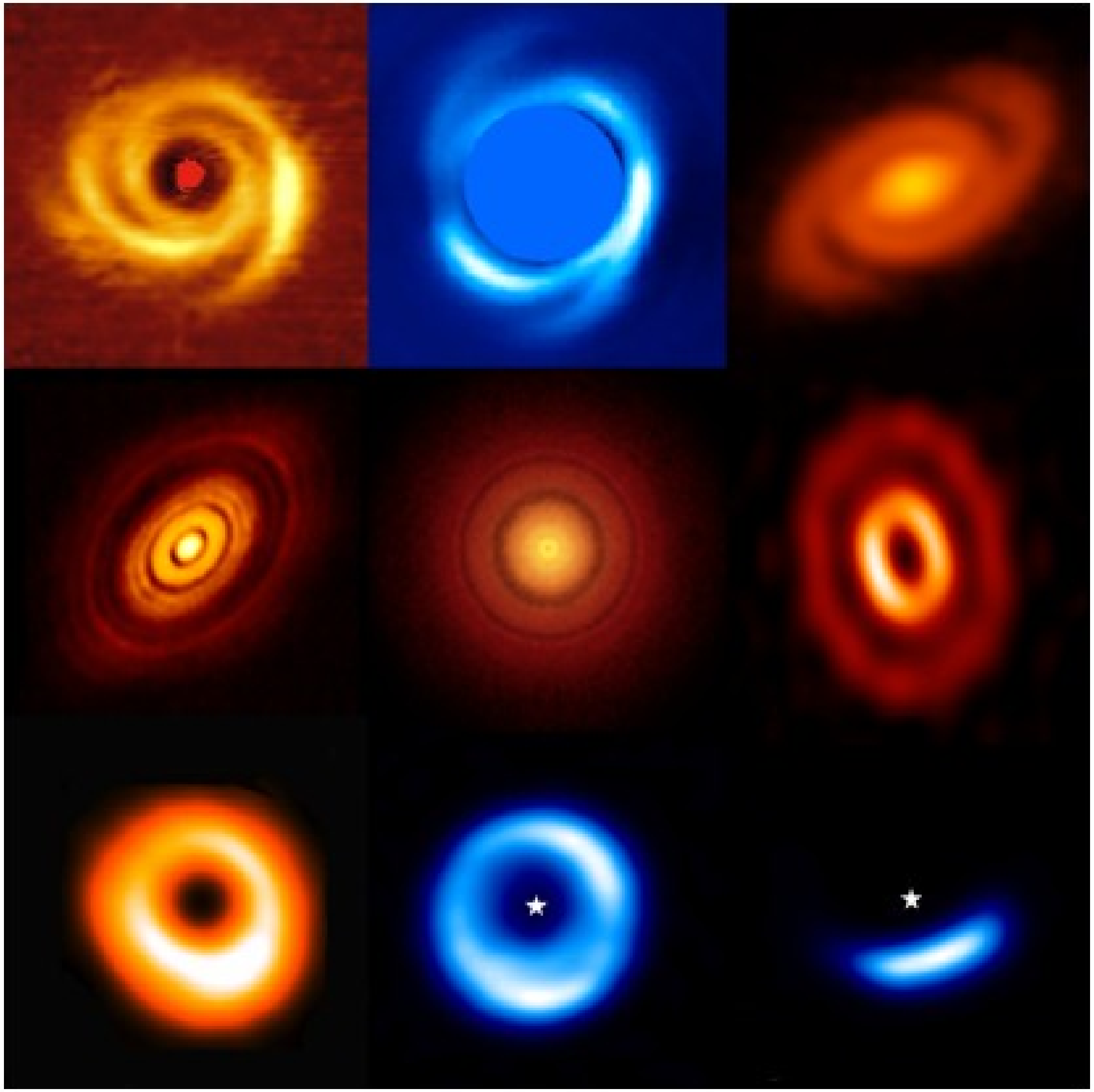
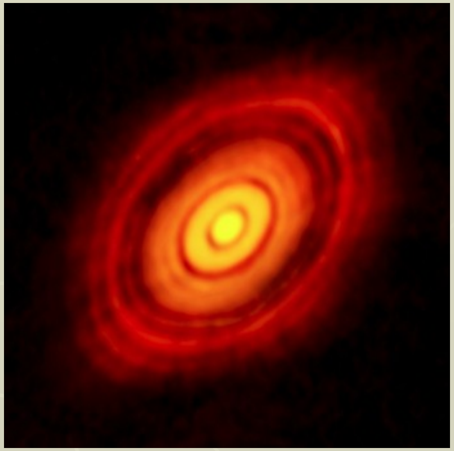


Figure 3: Gallery of disc substructures. Spiral structures are observed in polarized scattered light of (top left) HD 135344B (Garufi et al. 2013), (top center) HD 100453 (Wagner et al. 2015) and (top right) in ALMA band 6 of Elias 2-27 (Pérez et al. 2016). Gap-like features are detected in thermal dust emission at (sub-)millimetre wavelengths in (middle left) HL Tau (ALMA Partnership et al. 2015), (middle center) TW Hydrae (Andrews et al. 2016) and (middle right) HD 97048 (van der Plas et al. 2016). Horseshoe are observed in thermal dust emission at (sub-)millimetre wavelengths in (bottom left) HD 135344B (Pérez et al. 2014), (bottom center) SR 21 (Pinilla et al. 2015) and (bottom right) Oph IRS 48 (van der Marel et al. 2013). All the images are on different scale.

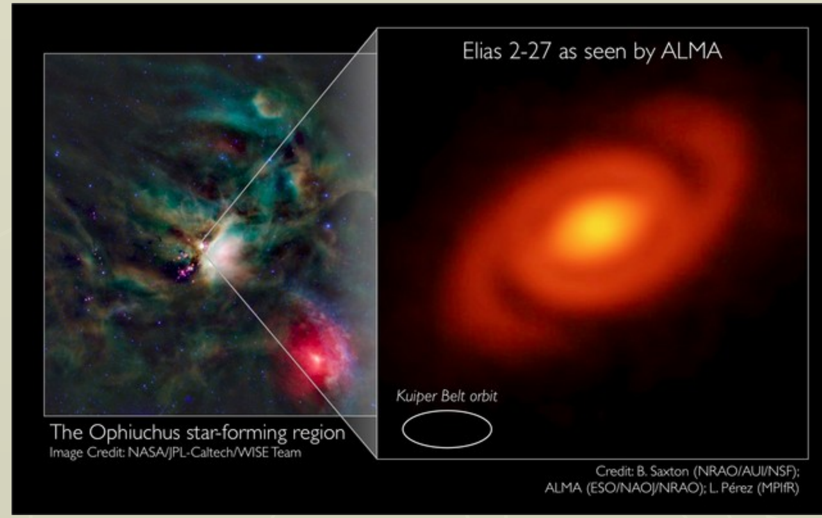
And then there was ALMA ….

The circumstellar disk zoo…..

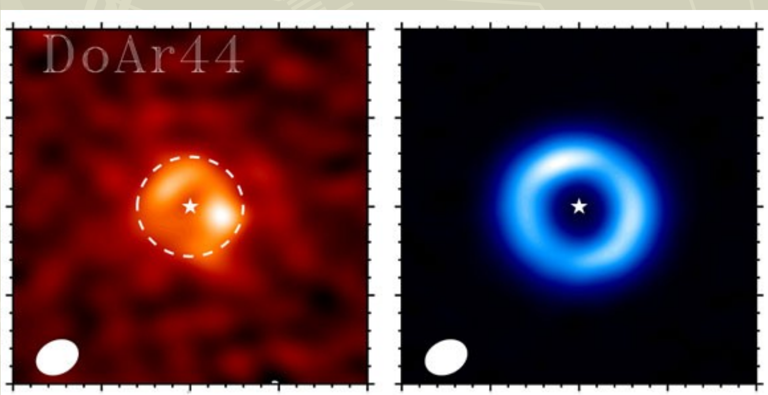
..gaps (HL Tau) ….



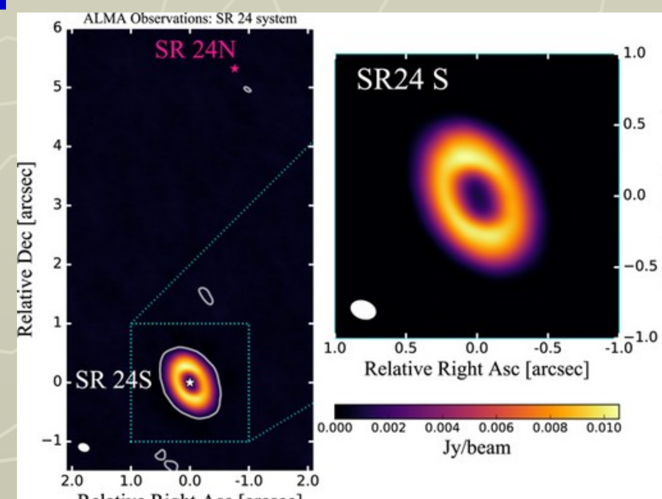
… spirals...



..different gas and dust distribution...

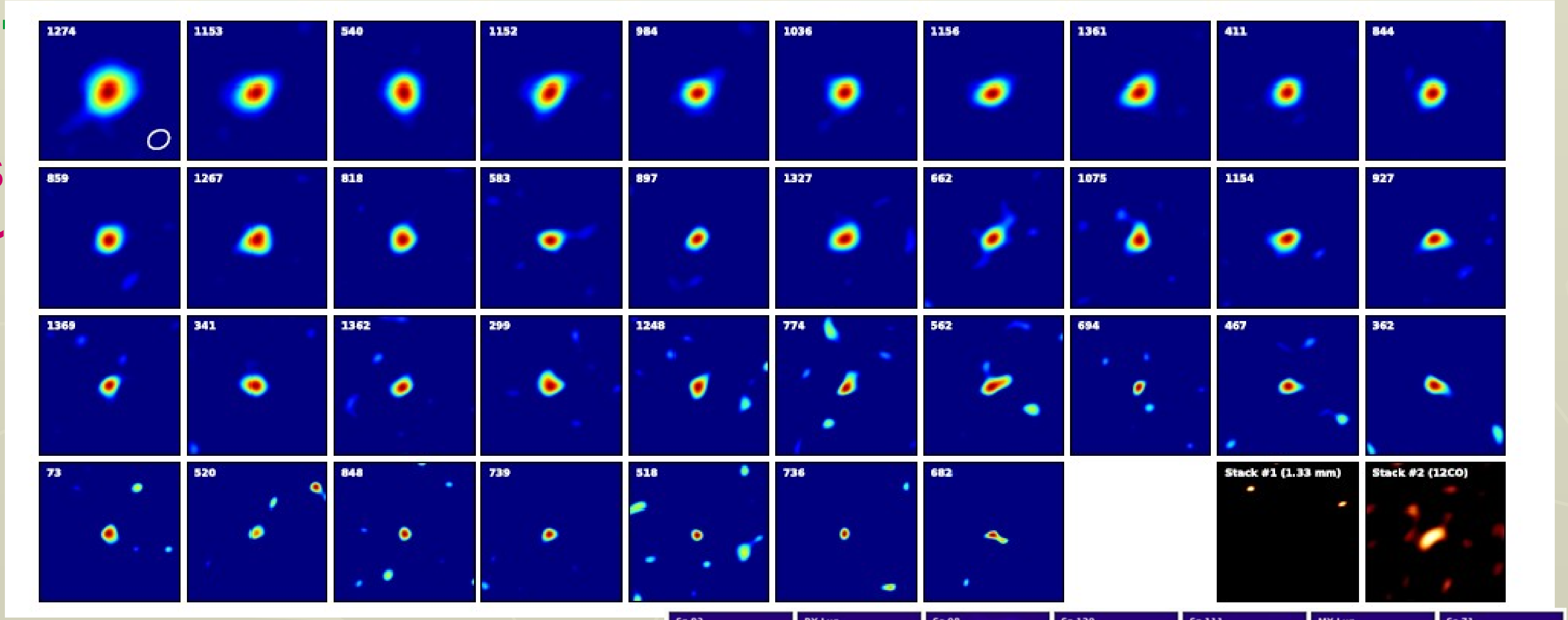


...transition disks with inner gas+dust depletion…..

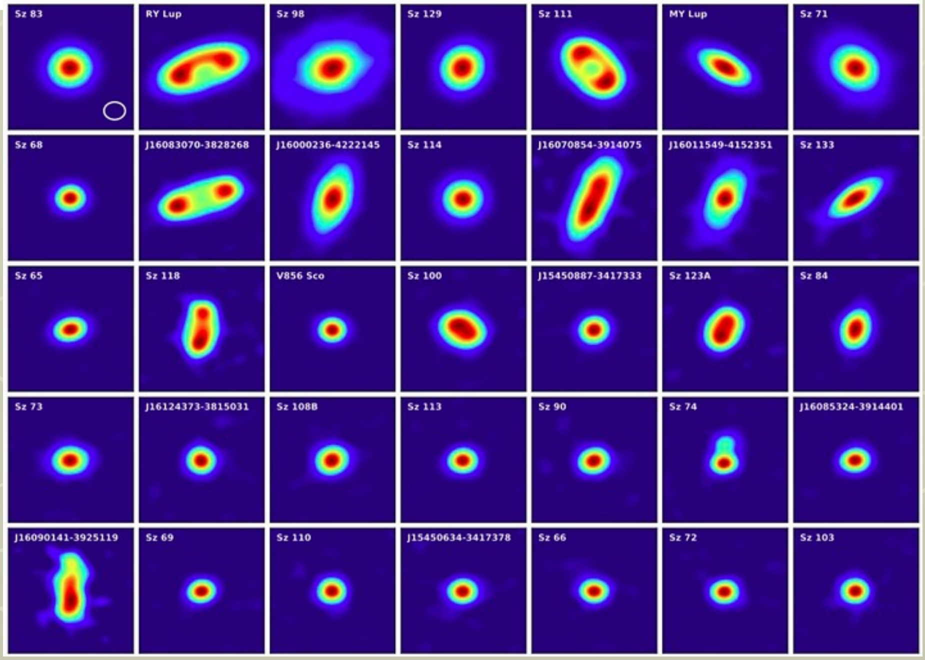


Disk demography with ALMA..

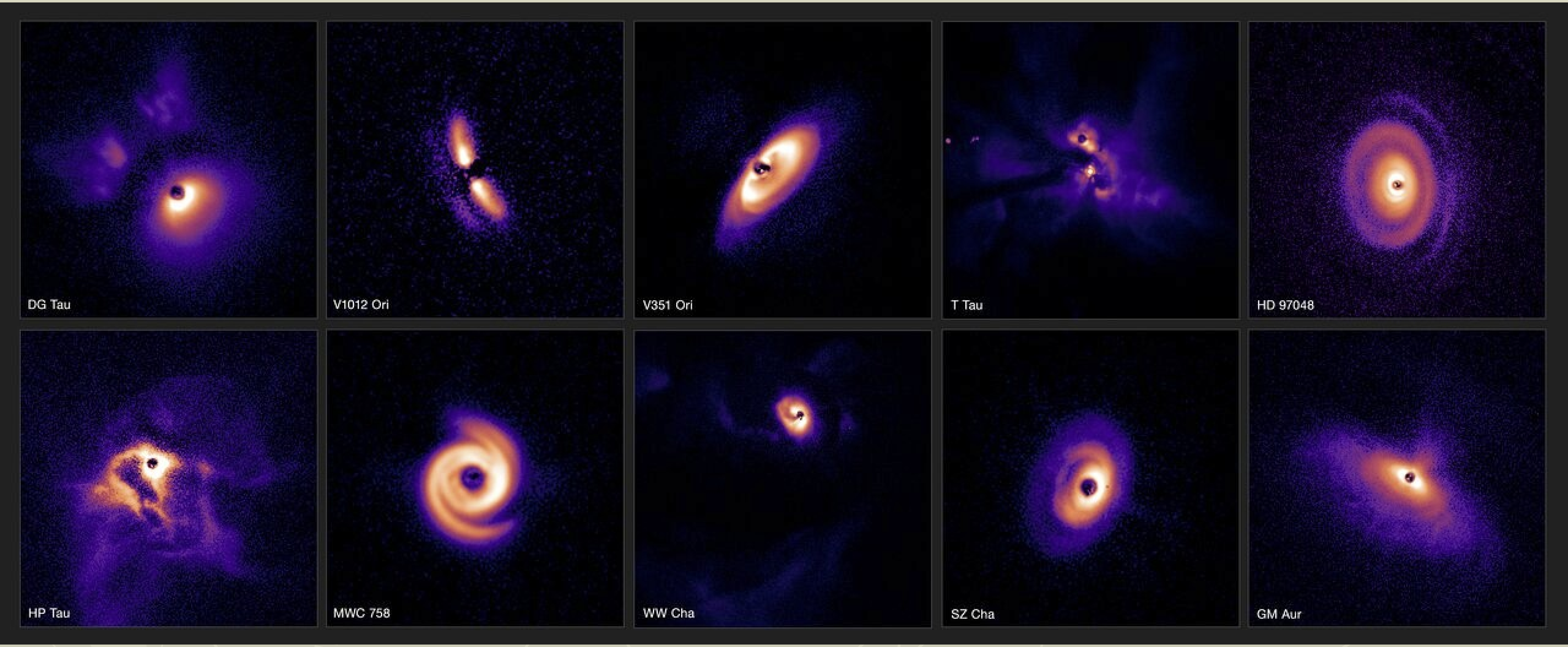
σ Orionis cluster (~ 2-3 Myr).



Lupus low-mass star forming region (~ 1-2 Myr)



...and SPHERE (Spectro-Polarimetric High-contrast Exoplanet REsearch) at VLT



extreme adaptive optics system in a narrow field of view

Solar nebula models

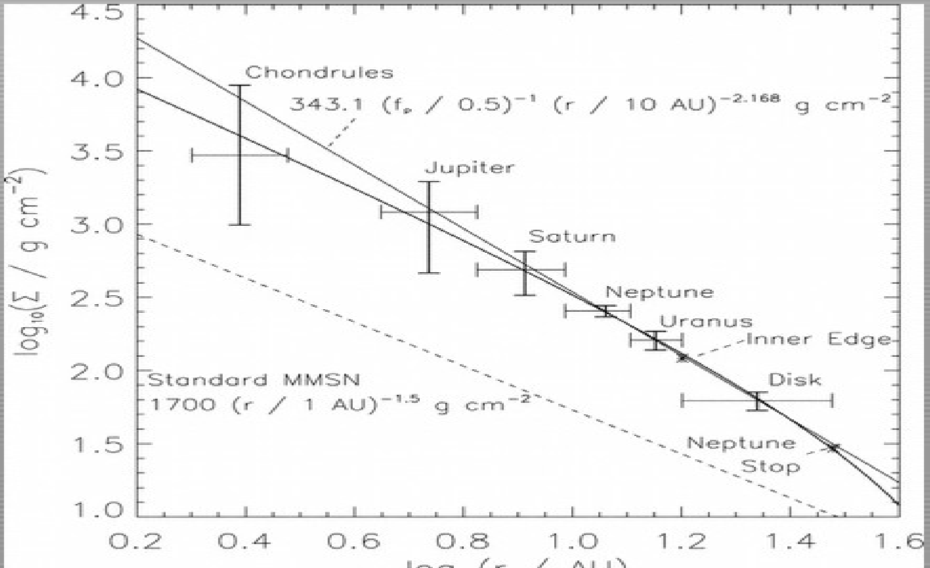
$$\Sigma_g(r) = 1700 \frac{\mathrm{g}}{\mathrm{cm}^2} \left(\frac{r}{\mathrm{AU}} \right)^{-1.5}$$

MMSN

$$\Sigma_g(r) = 5.1 \times 10^4 \frac{\mathrm{g}}{\mathrm{cm}^2} \left(\frac{r}{\mathrm{AU}} \right)^{-2.2}$$

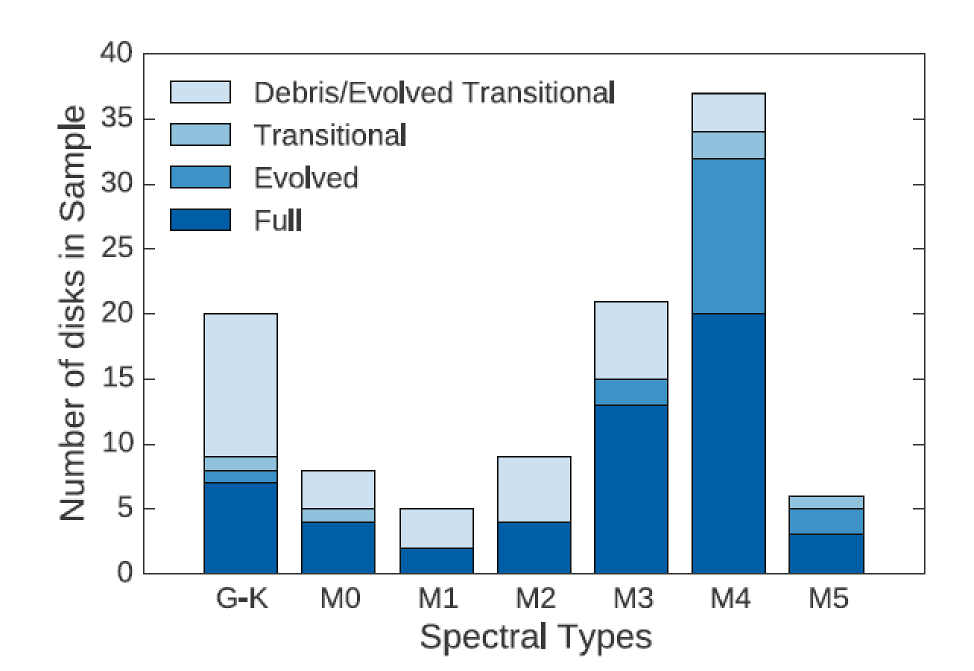
Nice model: the disk is 4 times more massive than the MMSN

Deusch 2007



Recently more massive disks are assumed (5-20 MMSN) to explain STIPS (Tightly Packed Inner Planets like Trappist o Hot/Warm Jupiters…).

More general disk models…..

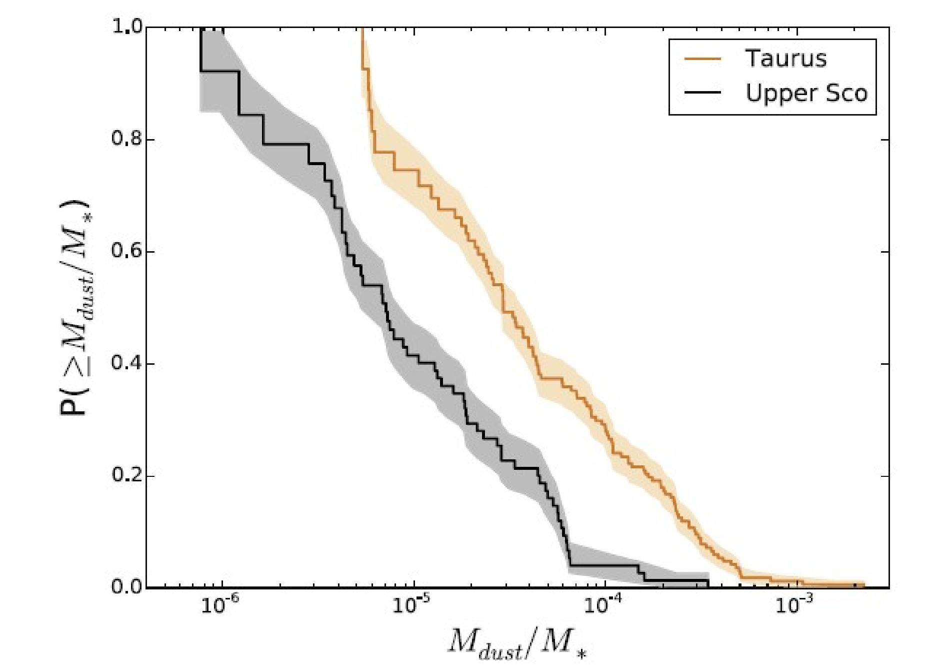


Upper Scorpius OB association (~ 5-11 Myr)

Superficial density distribution

$$\Sigma(r) = \Sigma_c \left(\frac{r}{r_c}\right)^{-\gamma} \exp\left[-\left(\frac{r}{r_c}\right)^{(2-\gamma)}\right]$$

Comparison with the Taurus region (1-2 Myr). Younger region has more massive disks… (Barenfeld et al. 2016)

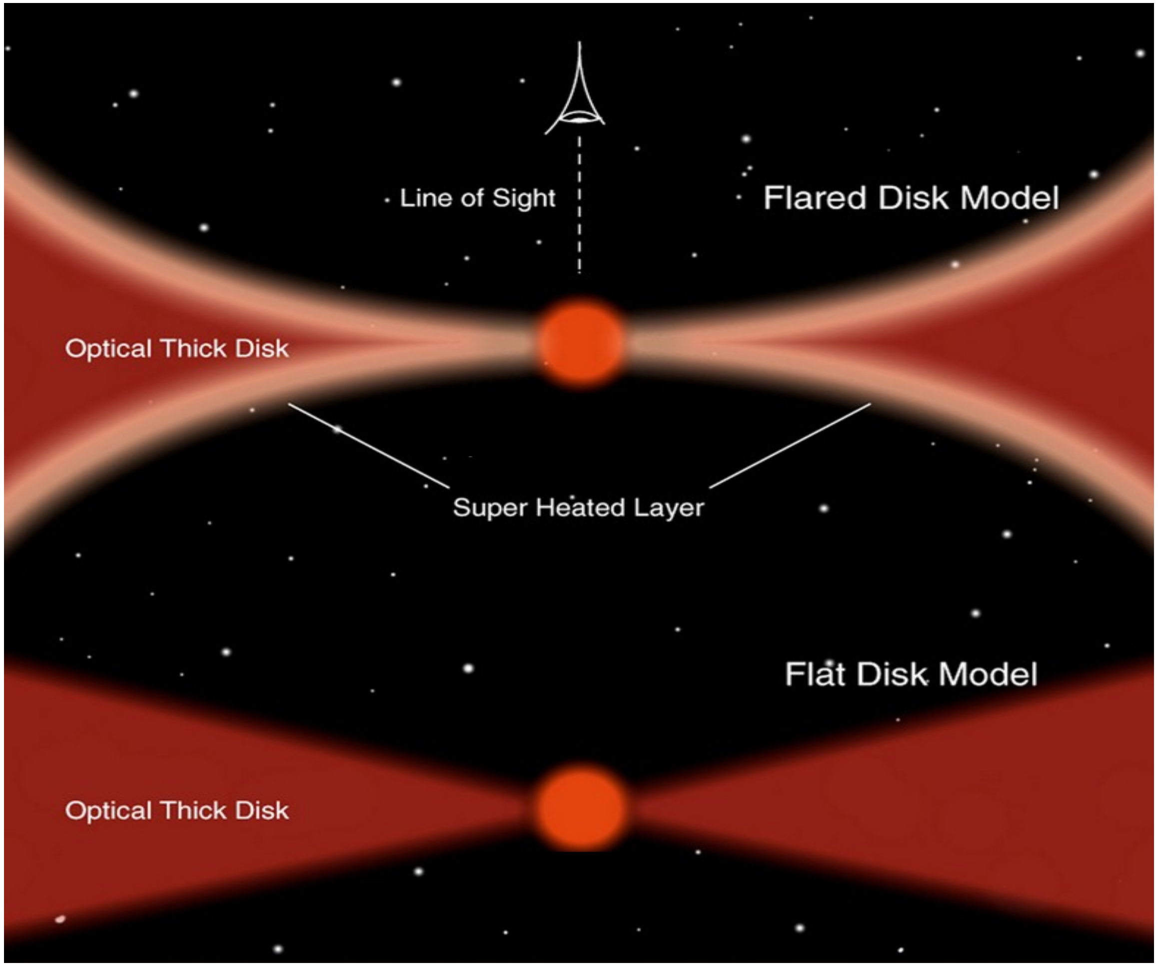


Vertical structure:

$$\rho(z) = e^{-\frac{1}{2}\frac{\Omega^2}{c_s^2}z^2} = e^{-\frac{1}{2}\frac{z^2}{H^2}}$$

h = scale height = H/R usually is around 0.05 for flat disks but the disk can be flared and h can grow going outwards

$$H = \frac{c_s}{\Omega}$$



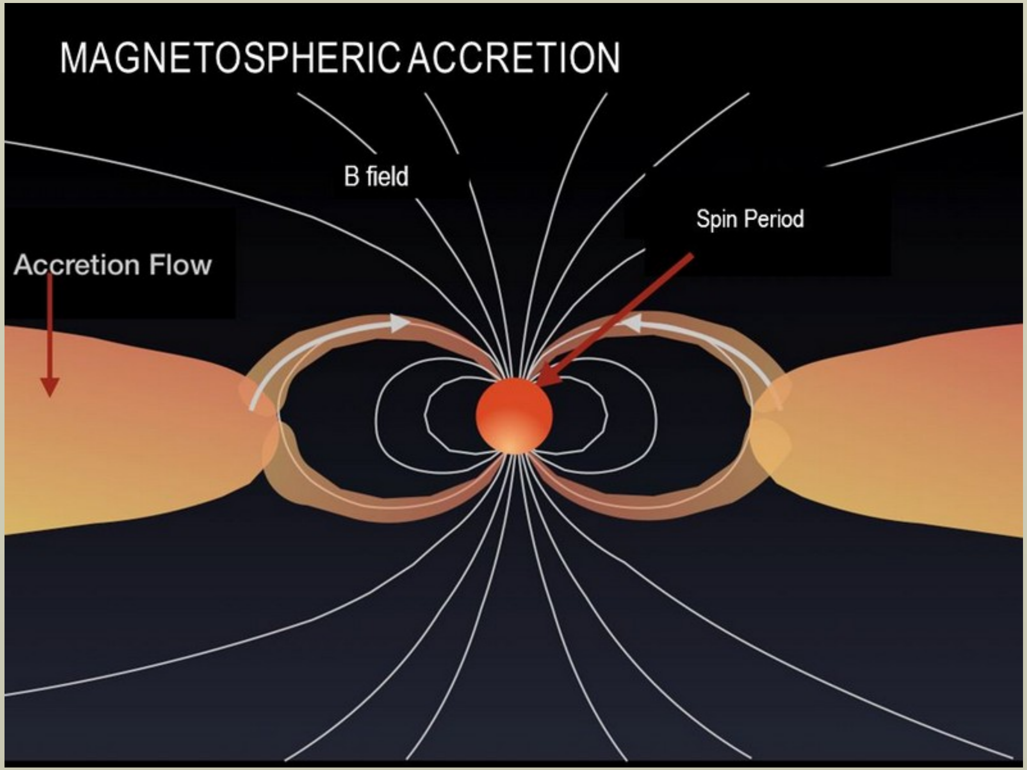
Thermal equilibrium:

- ***Viscous heating***
- ***Star irradiation***
- ***Radiative cooling***

Viscous evolution: dM/dt ~ 10^{-8} M_s /year

Mass accretion onto the star driven by the disk viscosity

Magnetospheric accretion model: disk material impacts the star following the magnetic disk lines (dipole). If the accretion rate is about 10^{-7} M_s/yr, the internal disk radius is about 0.1 AU before the onset of dipolar accretion.



Accretion onto the star is deduced from the ultraviolet and optical excess at the shock where the disk material impacts the star surface at a temperature of about 10^4 K.

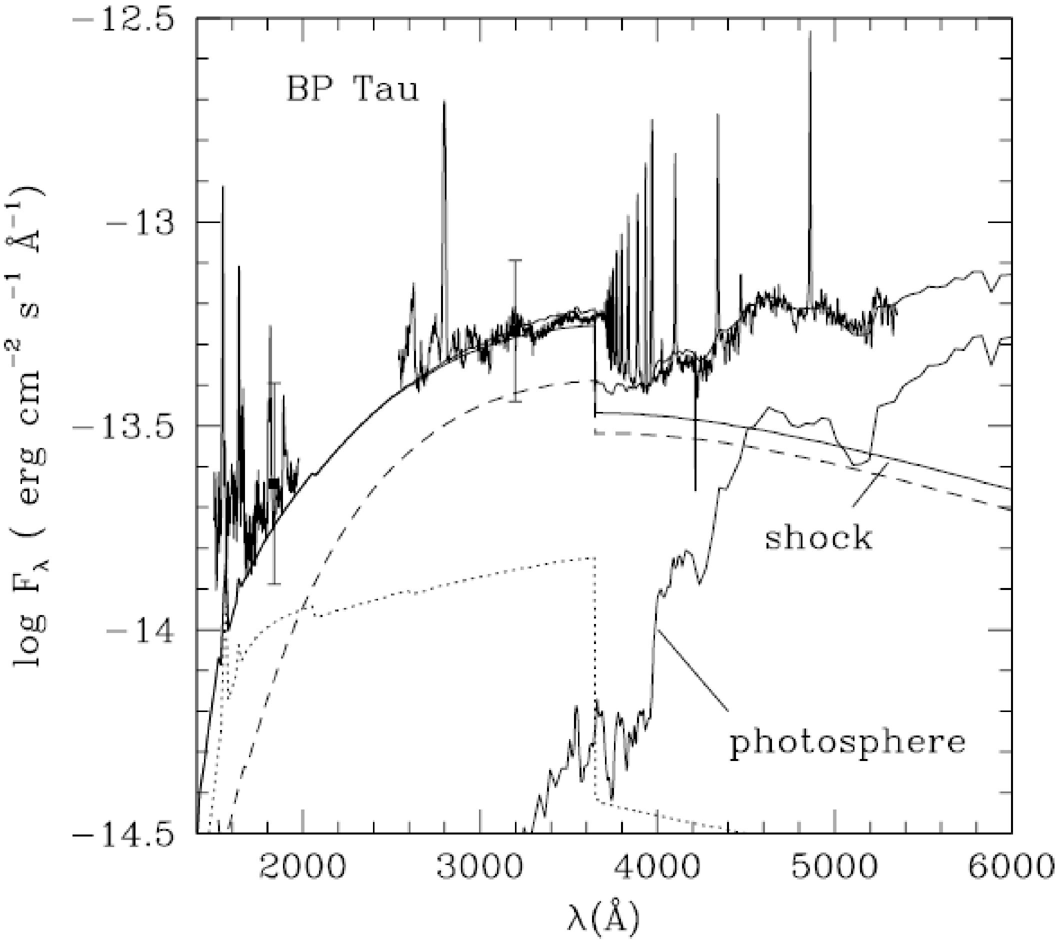


FIG. 1.—Comparison between observed fluxes and shock-model emission for BP Tau. Observations (*thick solid line*) are optical spectrum from GHBC, mean fluxes, and LW and SW *IUE* fluxes, defined by eq. (2). *IUE* fluxes have been smoothened by 5 pixels for better visualization. The mean flux level at 2700 and 1800 Å and range of variability are indicated by the squares and error bars. The theoretical model (*thin solid line*) is composed of the emission from the shock and from the stellar photosphere. The shock emission, in turn, is the sum of the emission from attenuated post-shock and diffuse preshock (*dotted line*) and the heated atmosphere (*dashed line*) (Paper I). The shock parameters used are log $\mathscr{F}$ = 11.5, f = 0.007, and A_V = 0.51.

In the UV, emission from the shock of disk gas impacting the photosphere of the star dominates at short wavelength and gives the mass accretion rate onto the star.

The contribution from the shock is the continous line while the stellar photosphere emission dominates beyond 4000 A.

$$\log\left(\dot{M}_{acc}\right) = -7.9 + 2.1\log\left(M_s\right)$$

Correlation between the mass of the star and the disk accretion rate (Hartmann et al., 2016)

Hydrodynamical modeling of disks

$$\frac{\partial \rho}{\partial t}+\nabla\cdot(\rho \boldsymbol{u})=0$$

$$\rho\frac{D\boldsymbol{u}}{Dt}=-\nabla P-\rho\nabla V$$

Navier Stokes equations of fluid dynamics + energy equation.

$$\rho\left(\frac{\partial u_r}{\partial t}+u_r\frac{\partial u_r}{\partial r}+\frac{u_\phi}{r}\frac{\partial u_r}{\partial \phi}+u_z\frac{\partial u_r}{\partial z}-\frac{u_\phi^2}{r}\right)=-\frac{\partial P}{\partial r}+\rho g_r$$

$$\rho\left(\frac{\partial u_\phi}{\partial t}+u_r\frac{\partial u_\phi}{\partial r}+\frac{u_\phi}{r}\frac{\partial u_\phi}{\partial \phi}+u_z\frac{\partial u_\phi}{\partial z}+\frac{u_r u_\phi}{r}\right)=-\frac{1}{r}\frac{\partial P}{\partial \phi}+\rho g_\phi$$

$$\rho\left(\frac{\partial u_z}{\partial t}+u_r\frac{\partial u_z}{\partial r}+\frac{u_\phi}{r}\frac{\partial u_z}{\partial \phi}+u_z\frac{\partial u_z}{\partial z}\right)=-\frac{\partial P}{\partial z}+\rho g_z$$

$$\frac{D}{Dt}U=\frac{1}{\rho}P\nabla\cdot\boldsymbol{u}+\epsilon-\frac{1}{\rho}\nabla\cdot\boldsymbol{F}_h$$

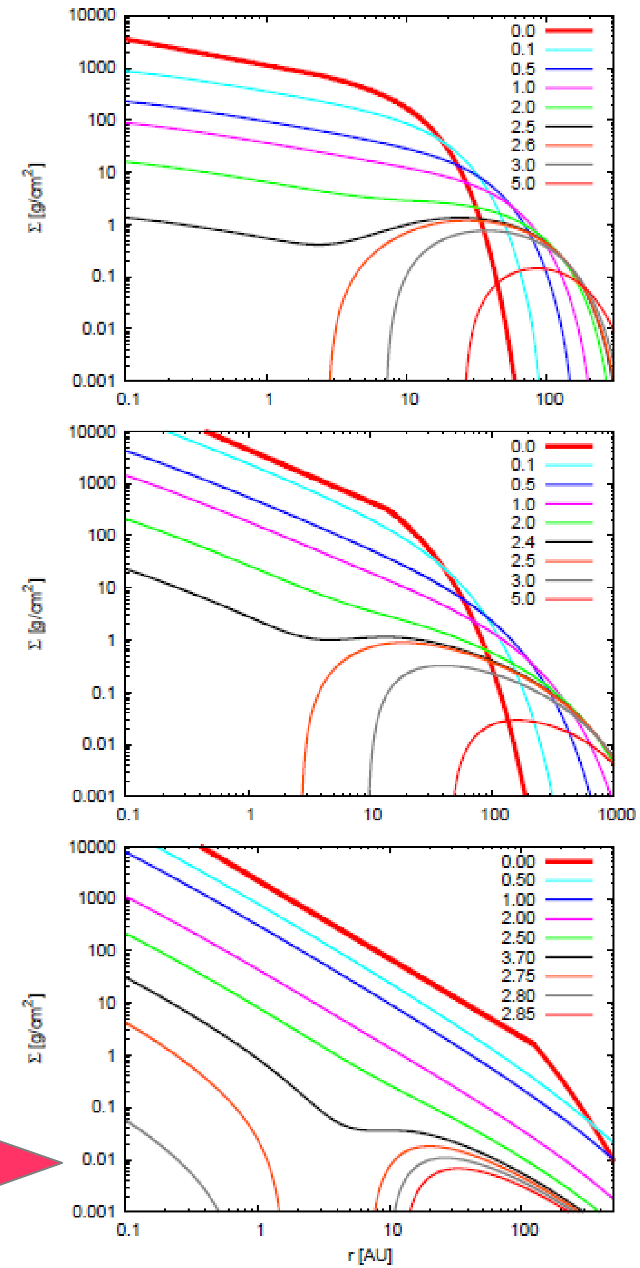
U internal gas energy
F_h heat flux
ε heat source
u velocity

Long term disk evolution: viscous evolution and photoevaporation.

One-dimensional approximated continuity equation (Thin, axis-symmetric disk, isothermal approximation, vertical equilibrium, radial velocity = Keplerian velocity) with viscosity (ν) and photo-evaporation term due to FUV, EUV and X-rays.

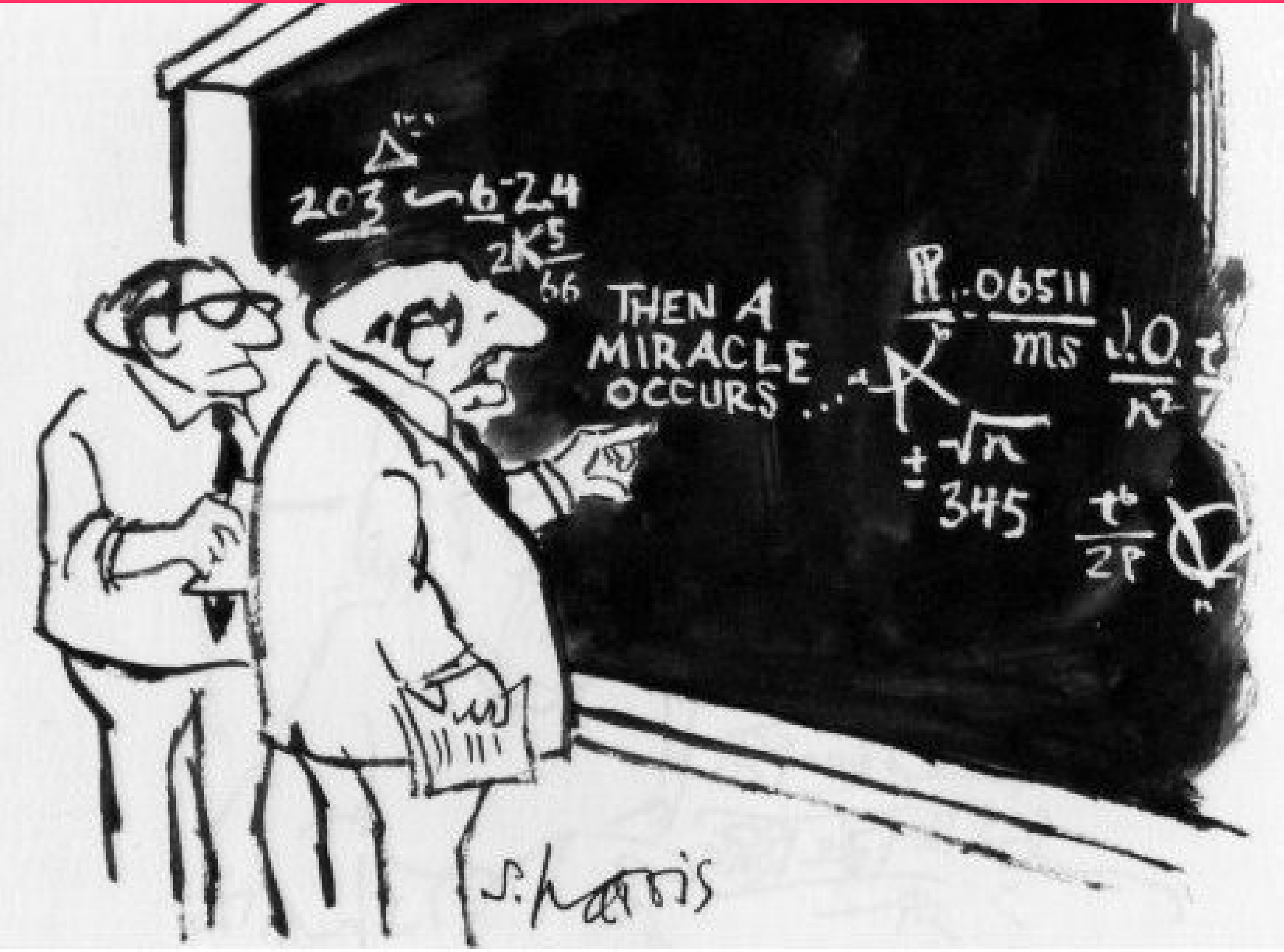
$$\frac{\partial \Sigma}{\partial t} - \frac{3}{r}\frac{\partial}{\partial r}\left(r^{1/2}\frac{\partial}{\partial r}(\Sigma \nu)\right) = -\dot{\Sigma}_{PE}$$

LIFETIME: 3-10 Myr



Inner hole formation like in GM Aur

PLANETESIMAL FORMATION

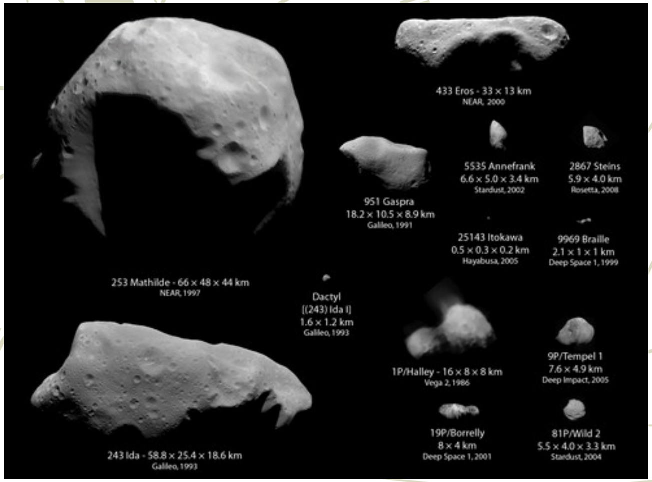


"I think you should be more explicit here in step two."

from *What's so Funny about Science?* by Sidney Harris (1977)

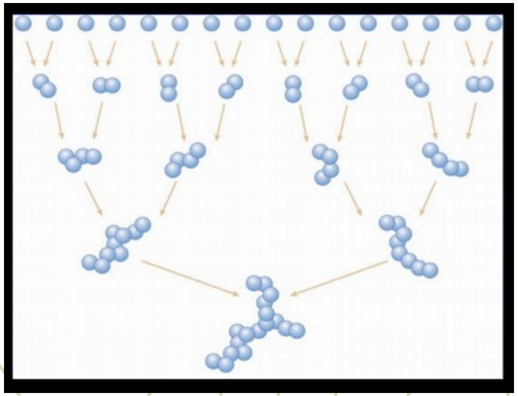
Why the planetesimal theory?

- ***Dust (zodiacal light, circumstellar disks)***
- ***Asteroids***
- ***Comets***



10^{-4} - 10^{-5} cm

$$F_D = -\frac{4}{3}\pi s^2 v_{th} \rho_g v$$

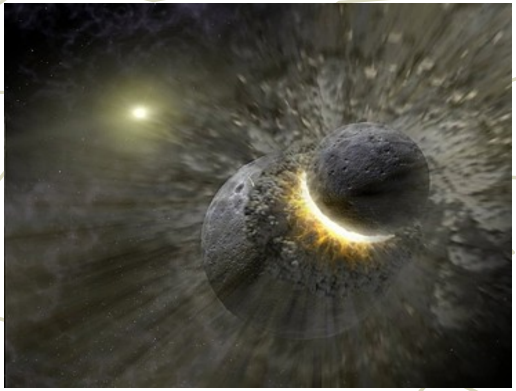


Planetesimals (1-10 km): gravity dominates

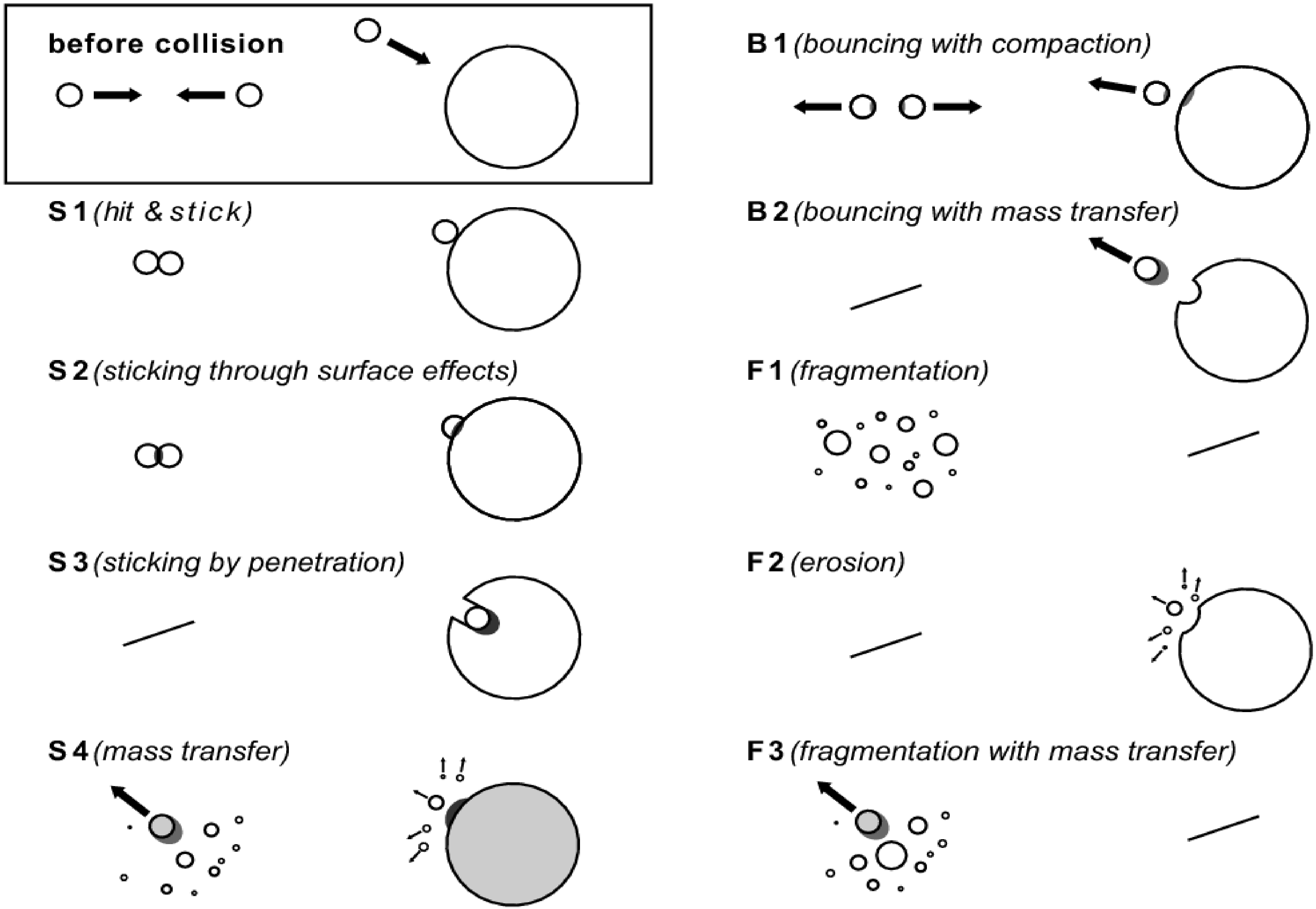
$$F_D = -\frac{4}{3} C_D \pi s^2 \rho_g v^2$$

10^5 – 10^8 cm

Planets



Possible outcomes of dust particle collisions (Guttler et al. 2010)



Relative velocity between dust particles is given by the following contributes:

1) Brownian motion:

$$v_{th} = \sqrt{\frac{8 K_B T}{\pi \mu m_H}}$$

2) Differential vertical settling

$$F_D = -\frac{4\pi}{3} \rho s^2 v_{th} v \qquad F_D = F_G = m \Omega^2 z$$

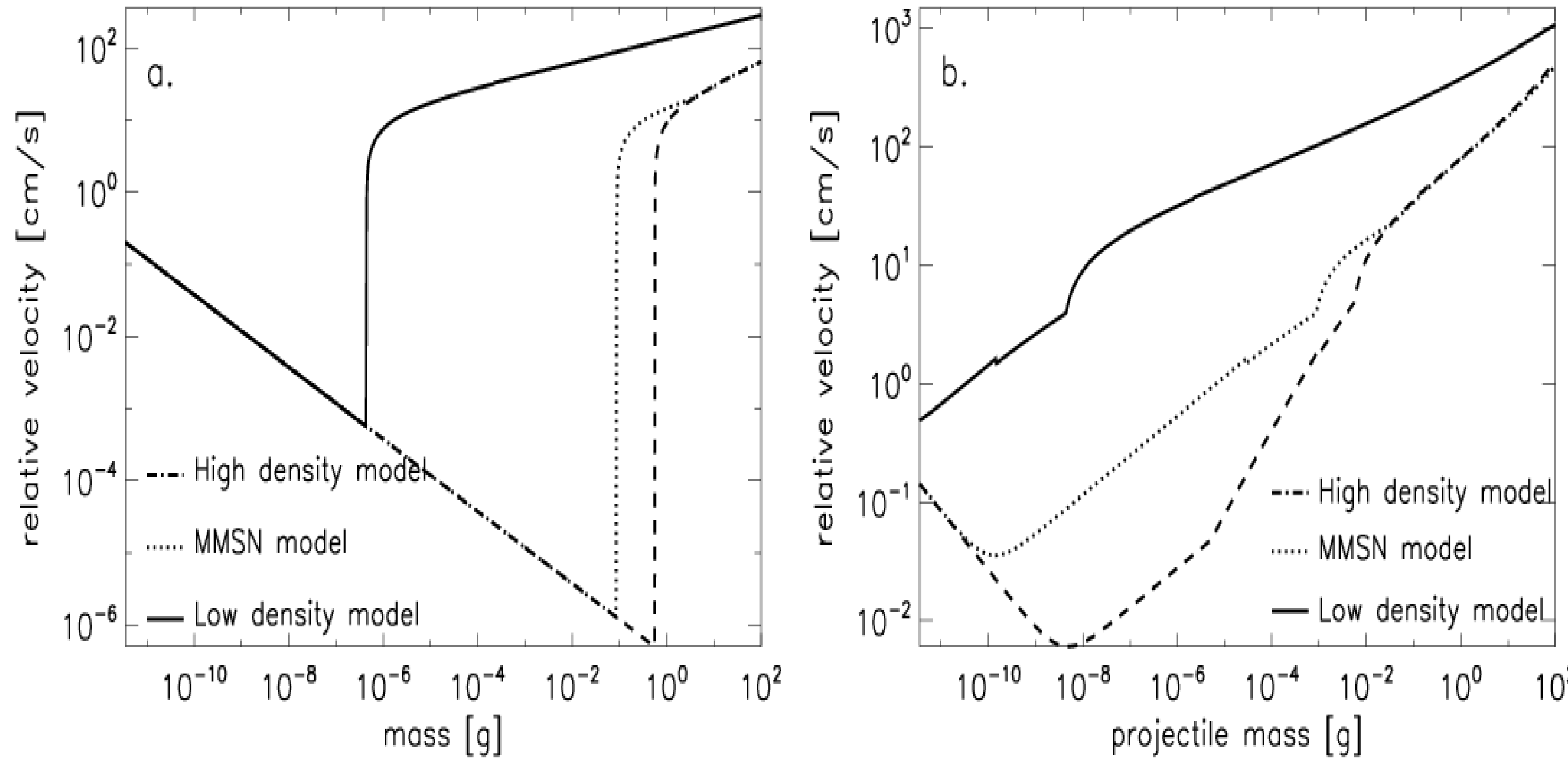
$$v_{settle} = \frac{\rho_d}{\rho_g} \frac{s}{v_{th}} \Omega^2 z$$

3) Differential radial drift

$$v_{radial} = -\eta \frac{\rho_d}{\rho_g} \frac{s}{v_{th}} \Omega^2 R$$

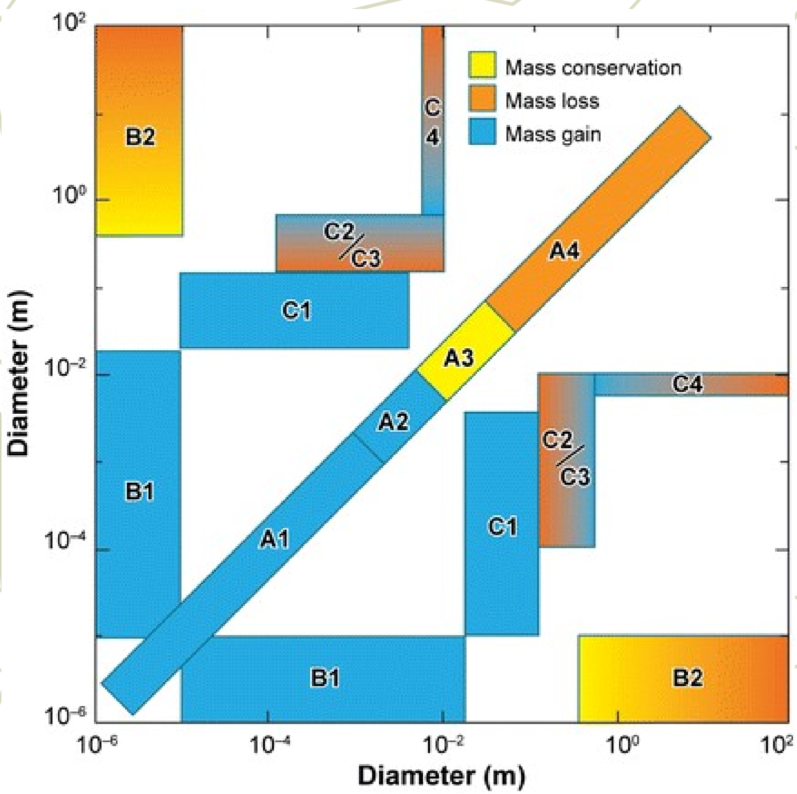
4) Turbulence

Relative velocity for equal size particles (left) and different size particles (1/100 in diameter, right) from Zsom et al. 2010.



Accretion appears not to be possible beyond about 10 cm. The relative velocity grows too much and bouncing and fragmentation dominate.

Wurm & Blum 2008



Blum J, Wurm G. 2008.
Annu. Rev. Astron. Astrophys. 46:21–56

Possible enhancing factors:

- Electric runaway growth: charged particles attract (dipole charging)
- Aerodynamic reaccretion: fragments reaccreted because they feel more the gas friction
- Magnetic forces
- Organic material has stronger sticking properties

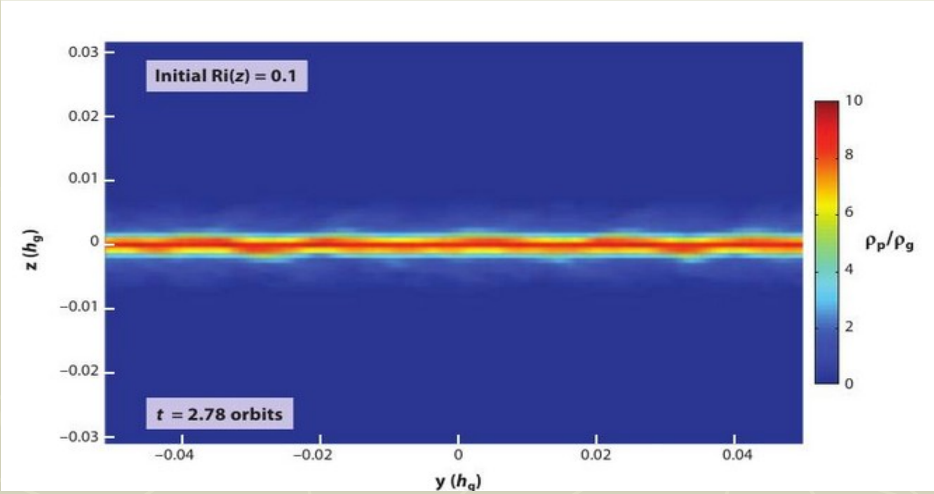
The meter-size barrier

$$v_{radial} = -0.5\,\eta V_K$$

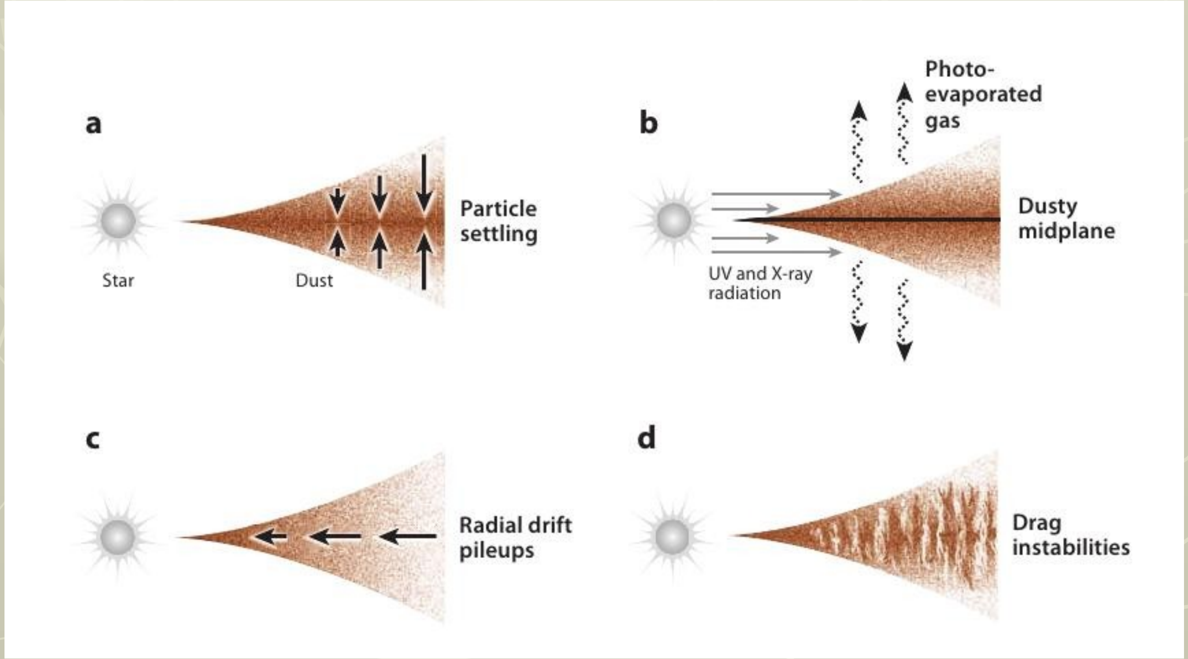
The drift velocity becomes very large when s > 10 cm. Timscale to fall on the star of the order of 10^3 yrs.

- **Planetesimal formation must be rapid**
- **Radial redistribution of material occurs**

Shortcut: gravitational instability ….. but there is the KH instability.

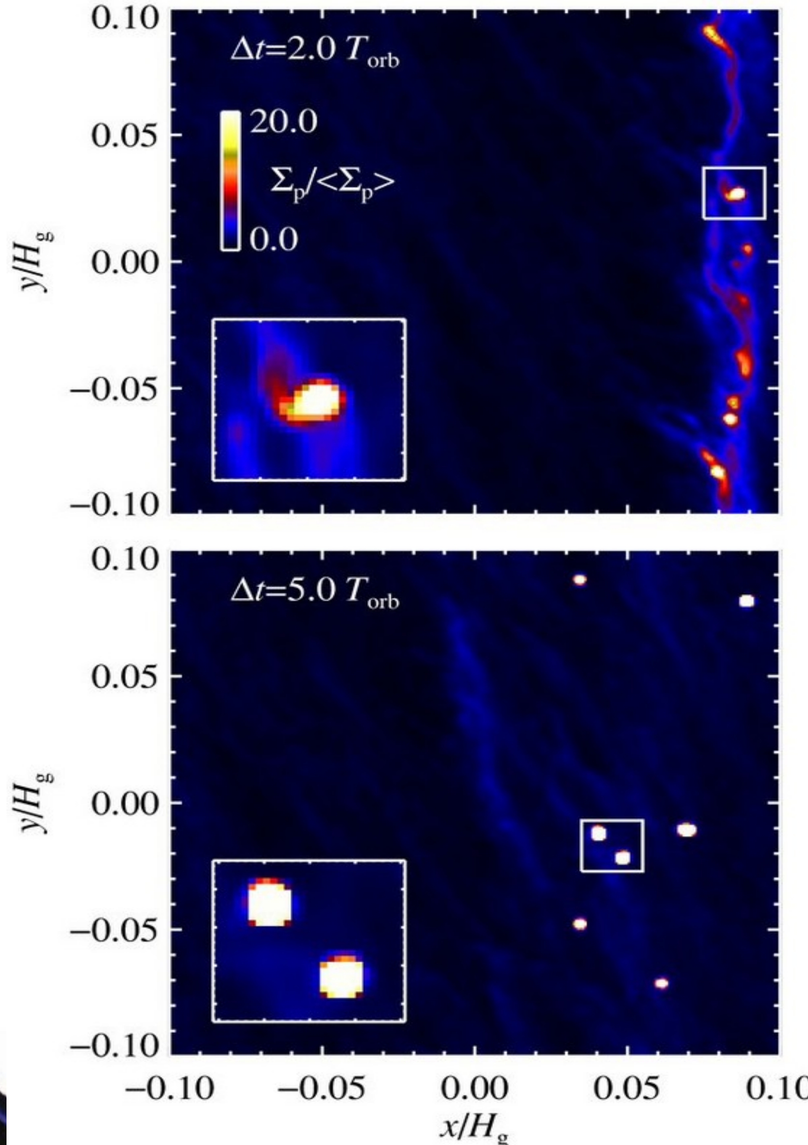


Dust and gas interact and the gas is dragged by the dust -> instability

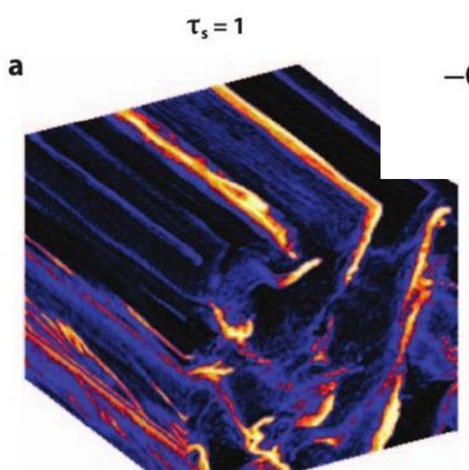


However, if the density of the dust is high enough, GI can occur anyway....

Streaming instability by Johansen & Youdin.
Coupling term in N-S.
Clumping of particles lead to the formation of large bodies.
From cm-size pebbles to bodies 100-200 km in size.
Instability due to back-reaction of dust on gas.
Outcome of pencil code.



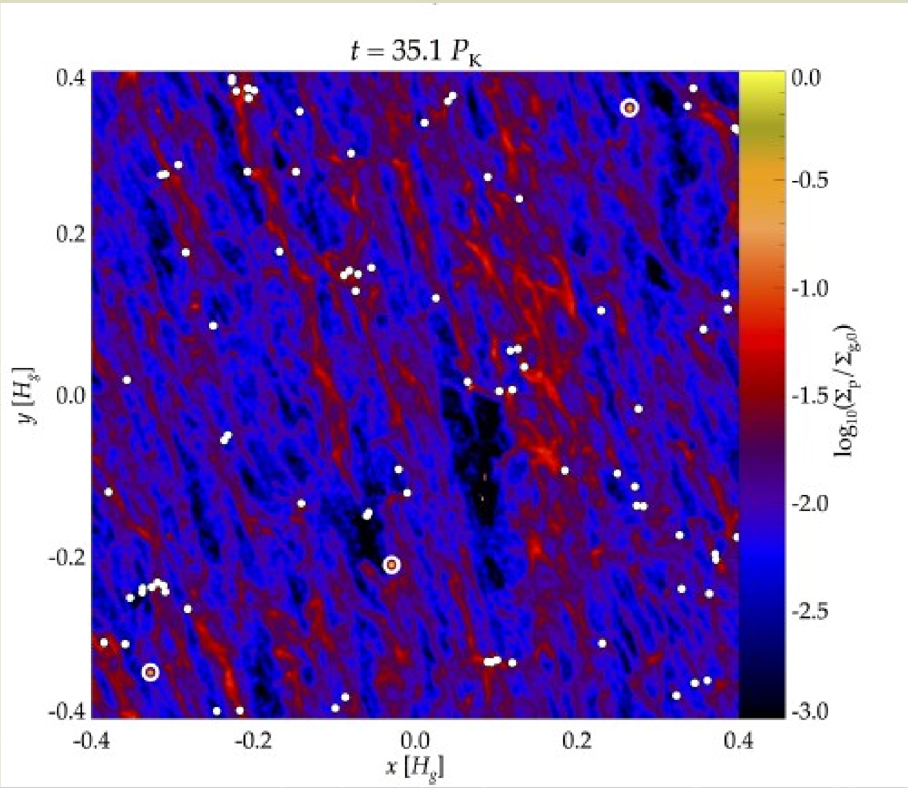
Direct formation of large bodies skipping the 1-10 km size phase



Streaming instability (Johansen...):

$$\frac{\partial \rho_g}{\partial t} + \boldsymbol{u} \cdot \nabla \rho_g - \frac{3}{2}\Omega x \frac{\partial \rho_g}{\partial y} = -\rho_g \nabla \cdot \boldsymbol{u}, \tag{3}$$

$$\frac{\partial \boldsymbol{u}}{\partial t} + (\boldsymbol{u} \cdot \nabla)\boldsymbol{u} - \frac{3}{2}\Omega x \frac{\partial \boldsymbol{u}}{\partial y} = 2\Omega u_y \hat{\boldsymbol{x}} - \frac{1}{2}\Omega u_x \hat{\boldsymbol{y}} - c_s^2 \nabla \ln \rho_g + 2\eta\Omega^2 r \hat{\boldsymbol{x}} - \frac{\epsilon}{\tau_f}(\boldsymbol{u} - \boldsymbol{w}). \tag{4}$$

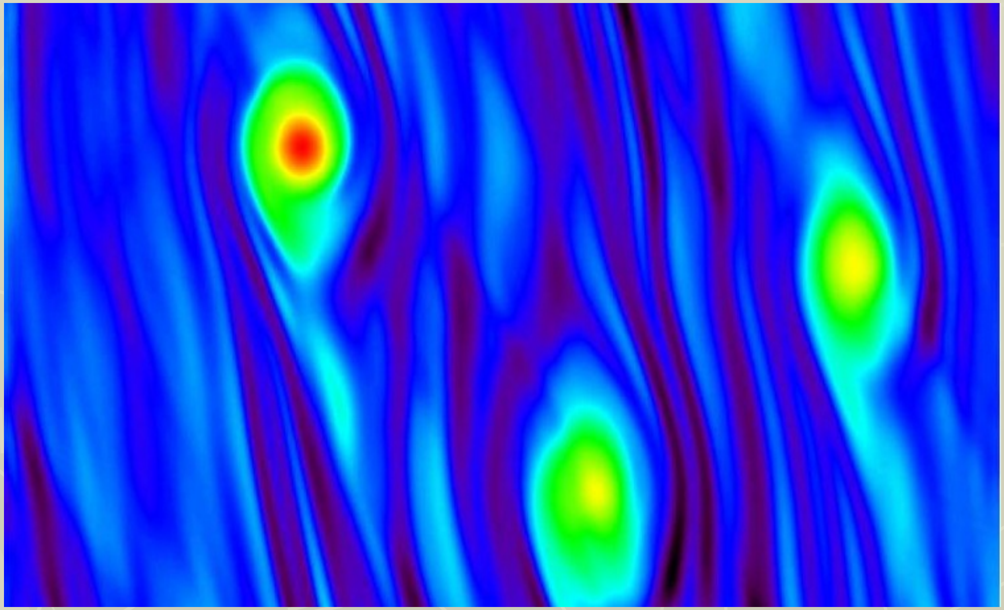


Formation of dense filamentes that become unstable to self gravity and create a planetesimal size distribution ranging from 60 to hundreds of km.

Problems: 1) it needs a high density in the midplane for the back-reaction term to be important. This is possible within a dead zone where there is no turbulent diffusion. 2) It needs large dust grains con St ~ 10^{-2}-10^{-1} corresponding to size of about 1mm to 1 cm i.e. pebbles 3) numerical models use superparticles and not a size distribution of grain sizes, 4) sensitivity to the resolution and initial conditions

Any kind of gas instability may work if it forms pressure bumps!

MRI, baroclinic instability, gravitational instability, VSI (Vertical Shear Instability) …..



Is it possible to discriminate? Maybe different initial size distributions are created depending on the instability responsible for planetesimal formation.

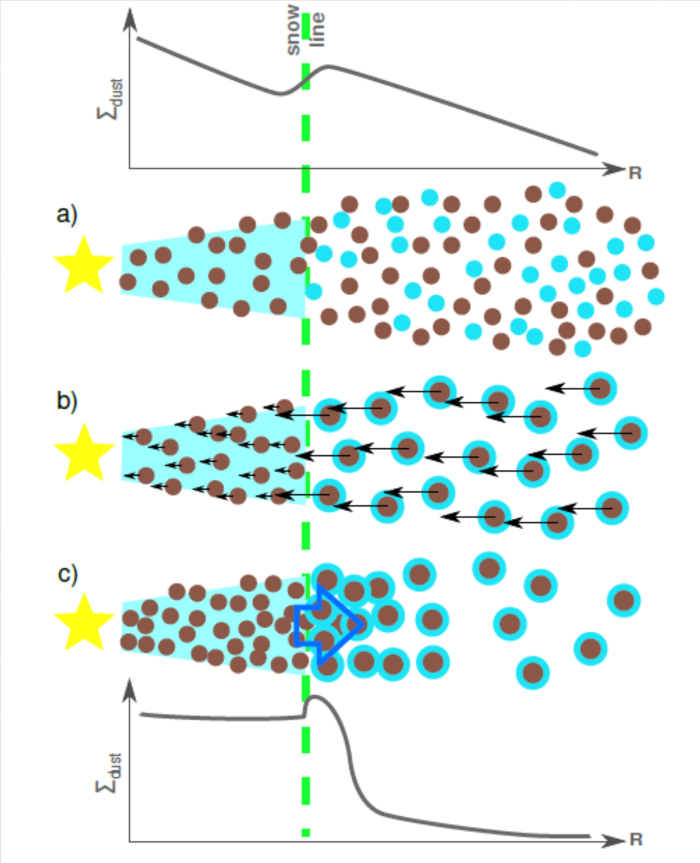
….other sources of pressure bumps...

1) Edge of a dead zone of low ionization. (Dzyurkevich et al. 2010)

2) Spiral arms due to a planet, binary companion or gravitational instability

3) Edges of planet carved gaps (Pinilla et al. 2012)

4) Trapping at the snow line (Drazkovska & Alibert 2017)



Large icy grains from outside migrate faster, reach the snow-line, evaporate and slow down creating a density bump.

Collisional growth: coagulation models based on Smoluchowski equation:

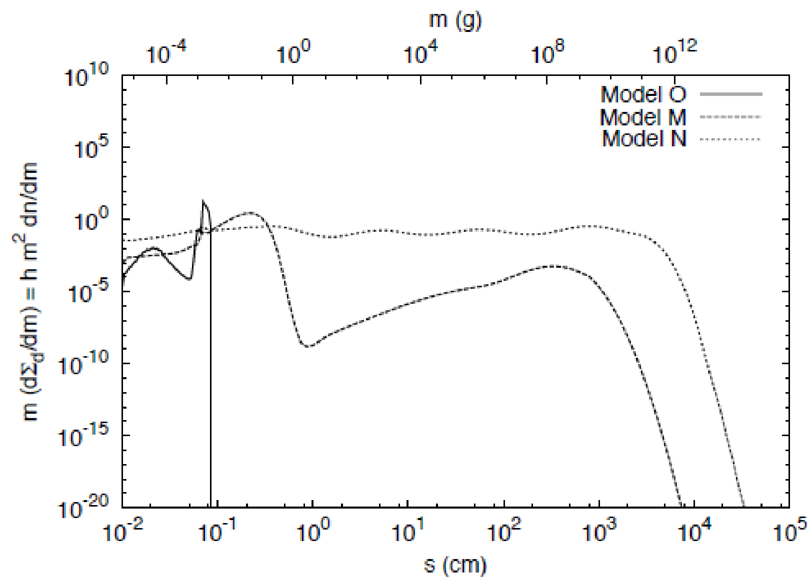
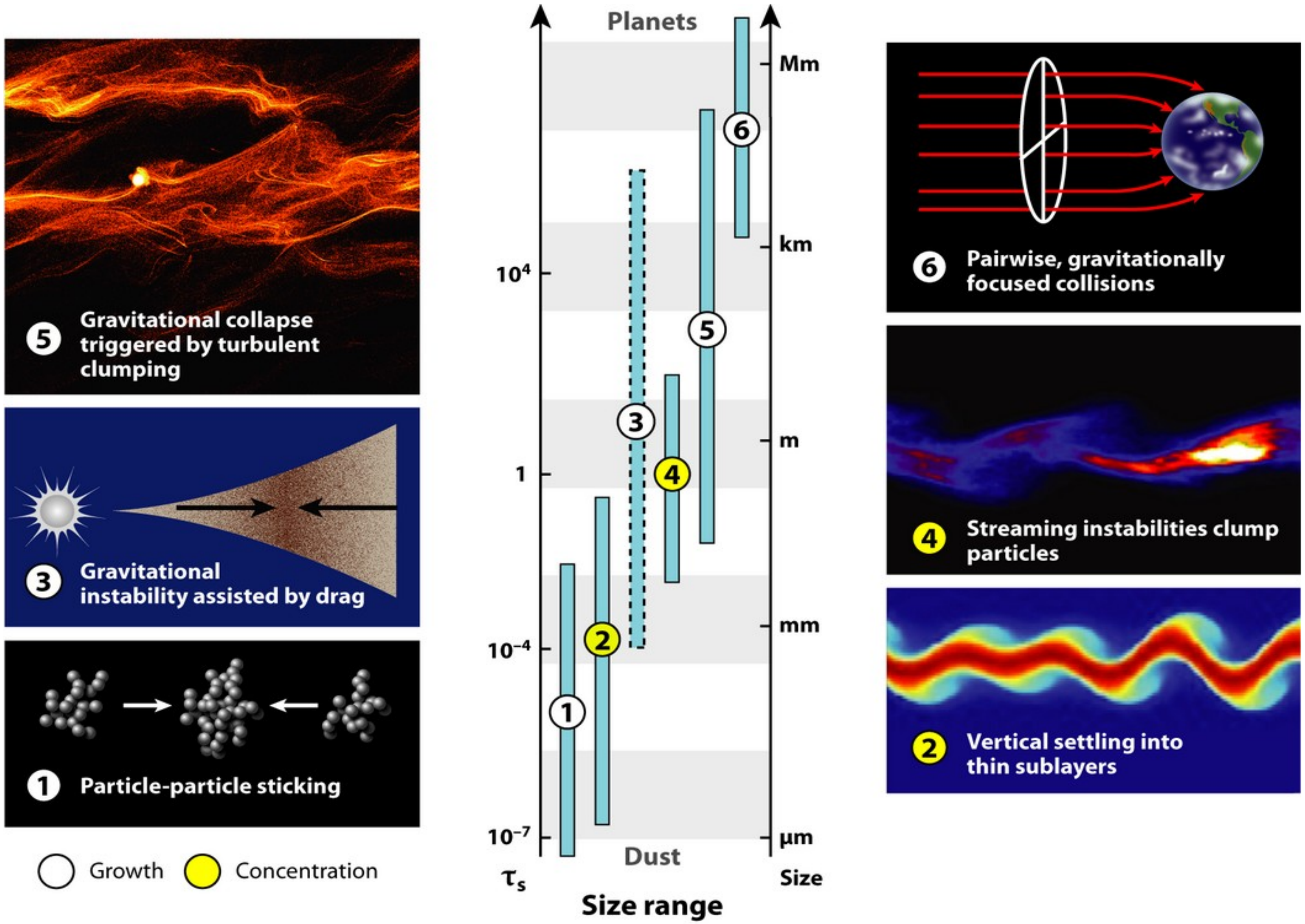


Fig. 6.— Particle evolution at 1AU after 10,000 yrs, with bouncing and mass transfer, in models where the bouncing velocity is $v_b = 5\mathrm{cm/s}$, and where the mass ratio above which every collision (that would otherwise lead to fragmentation) sticks is $\phi = 50$. The size range and resolution for Model O is the same as in Figure 5. The size range for Models M and N was increased to $s \in [10^{-2}, 10^{7}]\mathrm{cm}$, with a corresponding increase in the number of mass-points to keep the same resolution.

Garaud et al, ApJ 2013: adopt frequency distribution of relative velocities among planetesimals (M is for Maxellian, N more complex including effects of turbulence etc..). Bouncing, fragmentation barriers overcome…..



Chiang E, Youdin AN. 2010.
Annu. Rev. Earth Planet. Sci. 38:493–522

Possible different mechanisms for planetesimal formation.

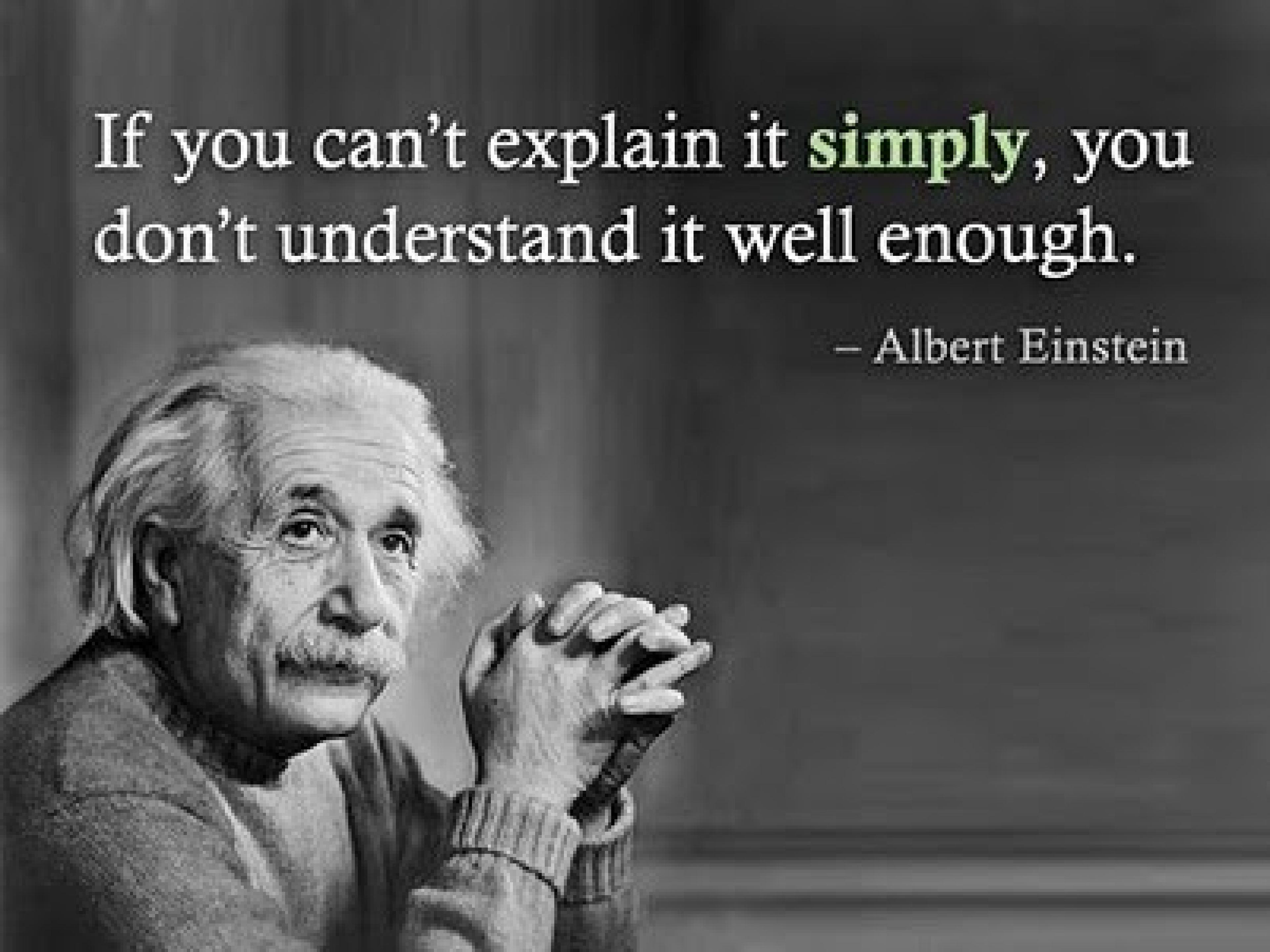
Blum 2018 (astroph)

Table 1 Comparison between the three formation scenarios of planetesimals described in Sect. 3.

	Gravitational collapse (Sect. 3.1)	Mass transfer (Sect. 3.2)	Icy agglomerates (Sect. 3.3)
Size of planetesimals [km]	$\lesssim 1000$ [1]	$\lesssim 1$ [2-4]	~ 10 [5]
Volume filling factor	$0.36 \times 0.6 \approx 0.2$ [6-7] ~ 0.4 *	~ 0.4 [8]	~ 0.1 [5]
Tensile strength of interior [Pa]	$\sim 1 - 10$ [9-10]	$\sim 10^3 - 10^4$ [8,11]	$\sim 10^3 - 10^4$ (guess)
Critical fragmentation energy for 1 m-sized body [J kg^{-1}]	$\sim 10^{-5}$ [12]	$\sim 10^2$ [12]	$\sim 10^2$ [12]
Normalised Knudsen diffusivity	$\equiv 1$	$\sim 10^{-4} \ldots 10^{-3}$ [13]	$\sim 10^{-5} \ldots 10^{-4}$ [13]
Thermal conductivity [Wm^{-1}K^{-1}]	$10^{-3} - 1$ [14] (conduction/radiation)	$10^{-2} - 10^{-1}$ [14] (conduction)	$10^{-2} - 10^{-1}$ [14] (conduction)

References:
[1] Schäfer et al. (2017), [2] Windmark et al. (2012b), [3] Windmark et al. (2012a),
[4] Garaud et al. (2013), [5] Kataoka et al. (2013), [6] Weidling et al. (2009),
[7] Zsom et al. (2010), [8] Kothe et al. (2010), [9] Skorov and Blum (2012),
[10] Blum et al. (2014), [11] Blum et al. (2006), [12] Krivov et al. (2018),
[13] Gundlach et al. (2011), [14] Gundlach and Blum (2012)

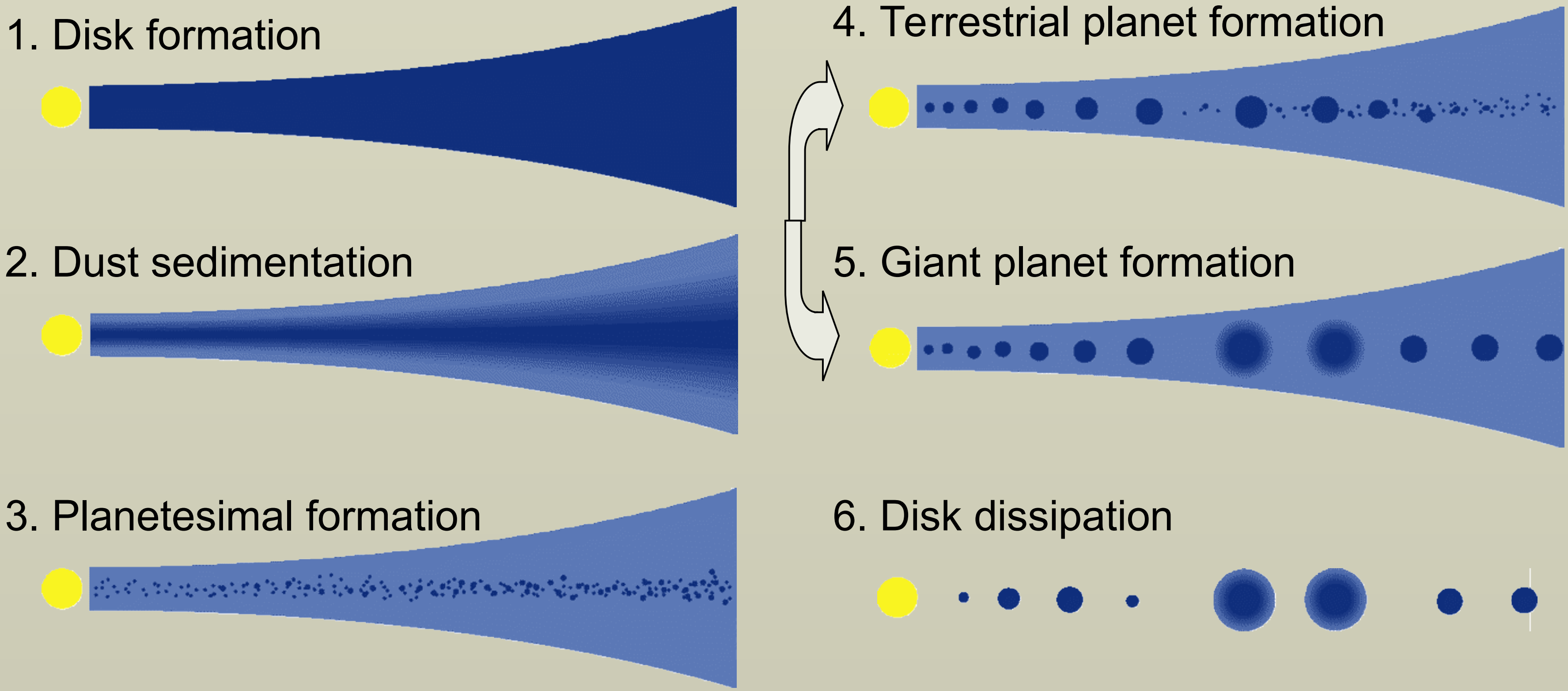
* For planetesimals with $R \gtrsim 10 - 50$ km


If you can't explain it simply, you don't understand it well enough.
– Albert Einstein

...concluding....

Planetesimal formation is still an open problem, in particular concerning the mechanisms that built them up and their initial size

Planet formation: short summary



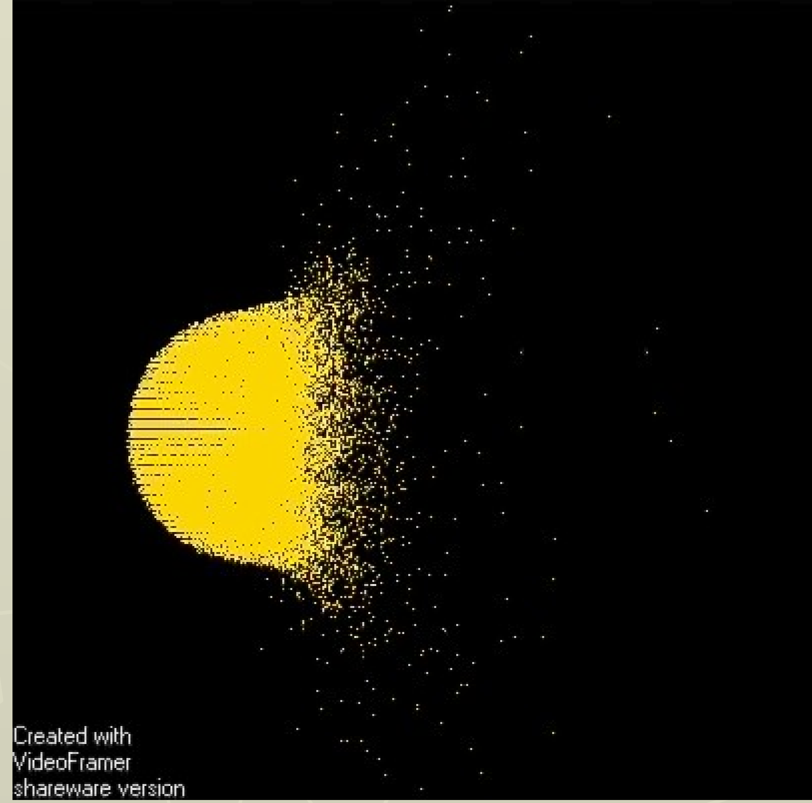
Planetesimal gravitational accumulation via mutual collisions

Billions of planetesimals around the protostar...

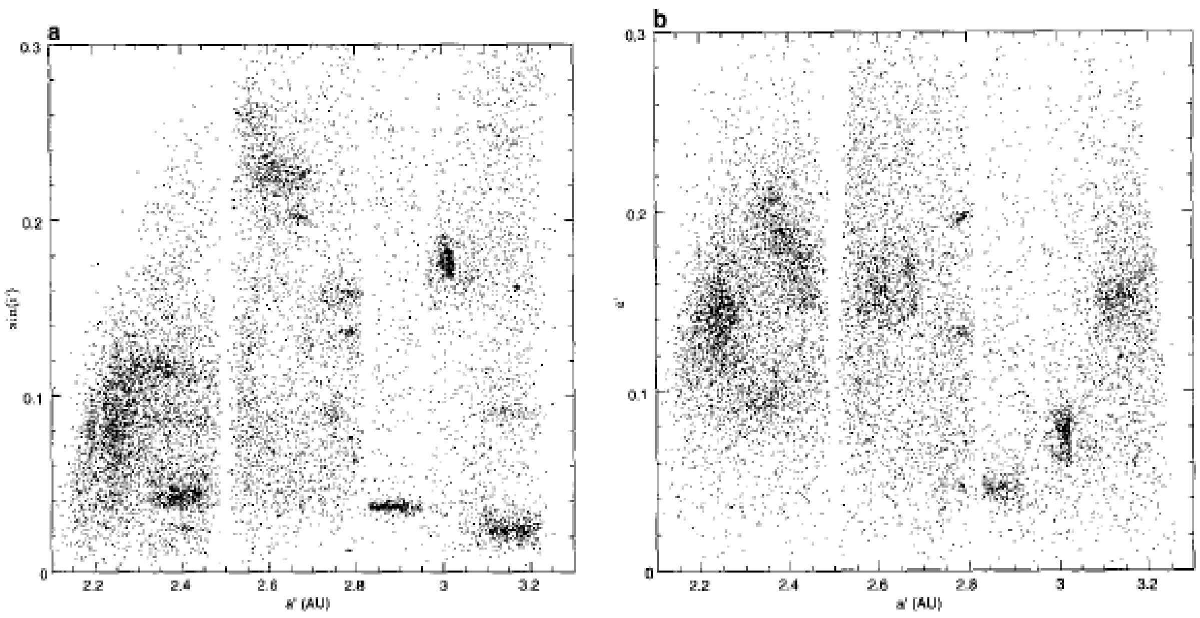
..they colide forming bigger bodies (planets). Leftovers of this process are asteroids and comets.

Planet building code: numerical model to simulate planetesimal evolution and estimate 1) the timescale of planetesimal accumulation 2) the different type of growth (runaway or oligarchic) 3) the final masses of the protoplanets.

Today collisions between asteroids in the Asteroid Belt are similar to planetesimal collisions, but are more energetic because their orbits are more eccentric and inclined .

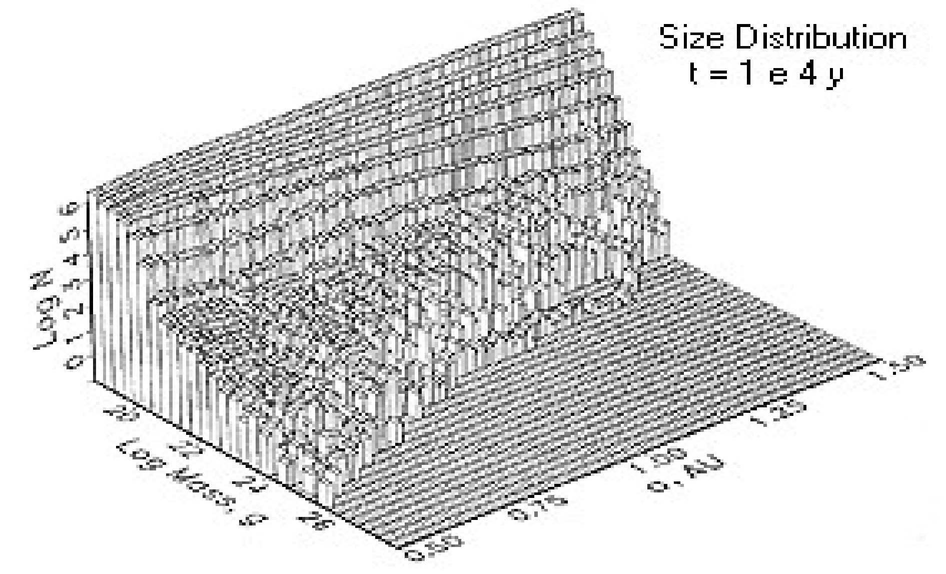


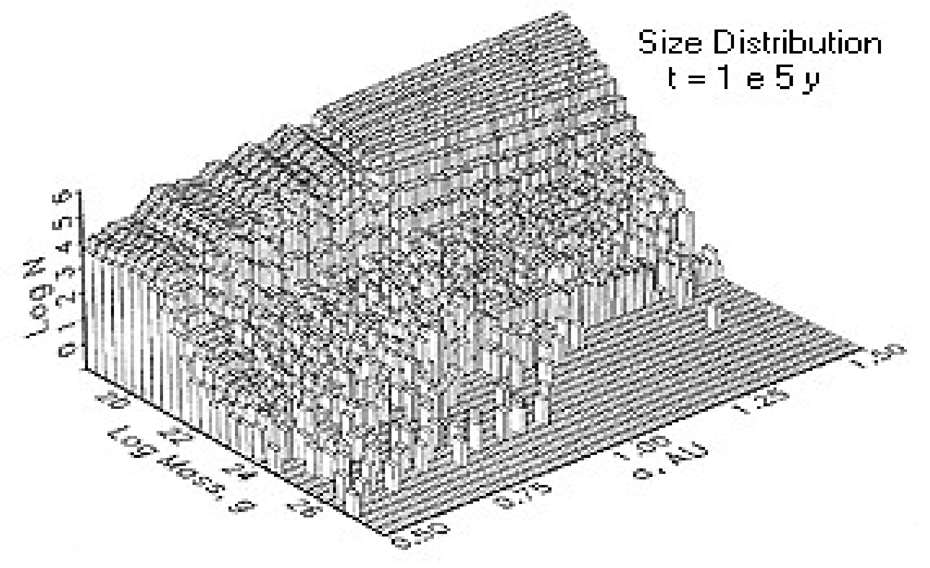
Collisions produce Hirayama families.

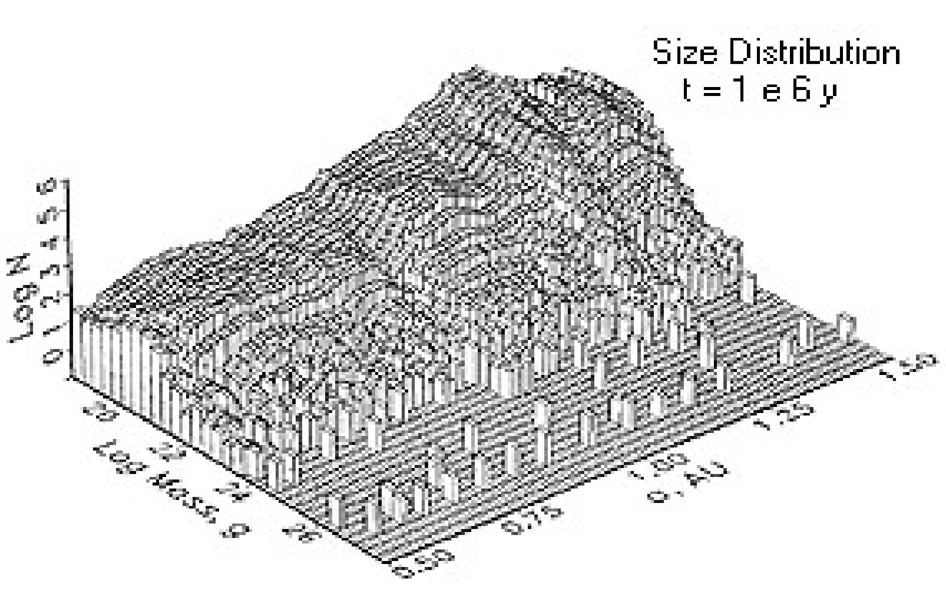


Hydrodynamical models to study the outcome of high energy impacts.

Planet building code outputs: the size distribution of planetesimals at different radial distances



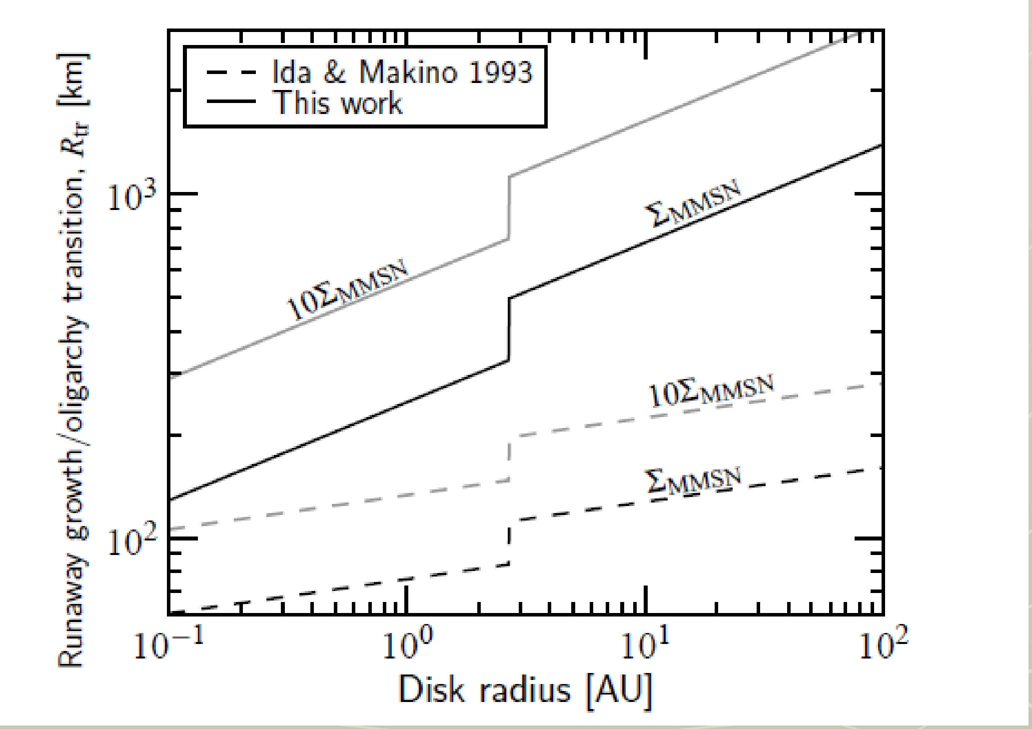


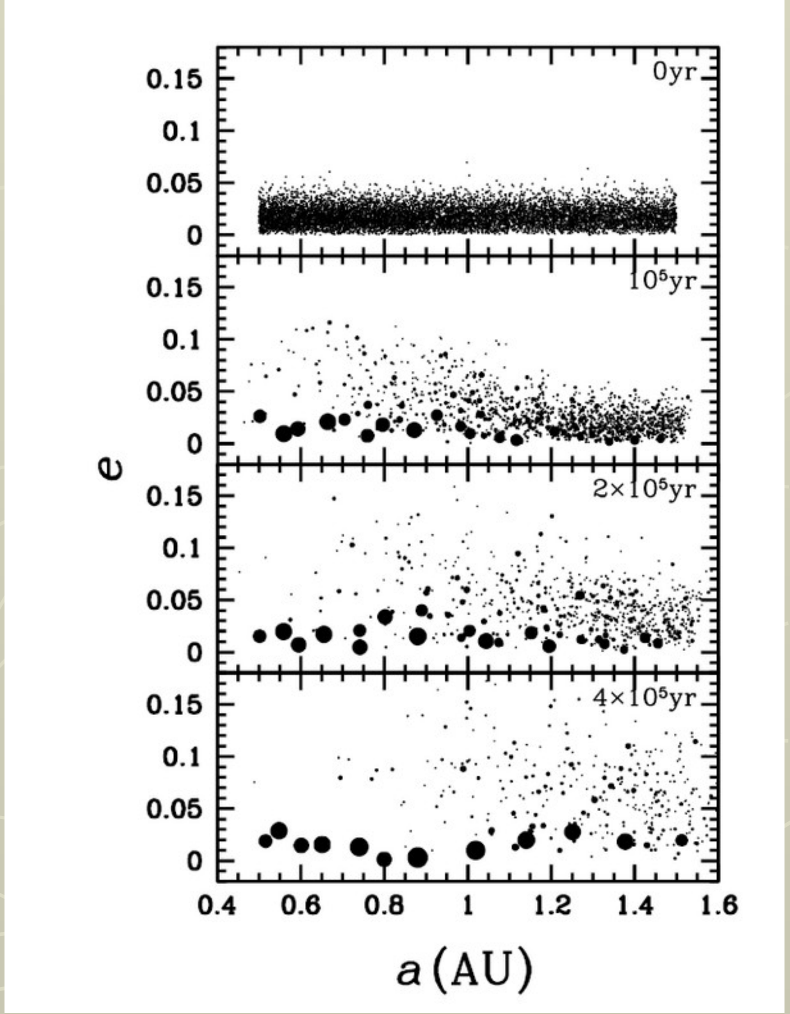


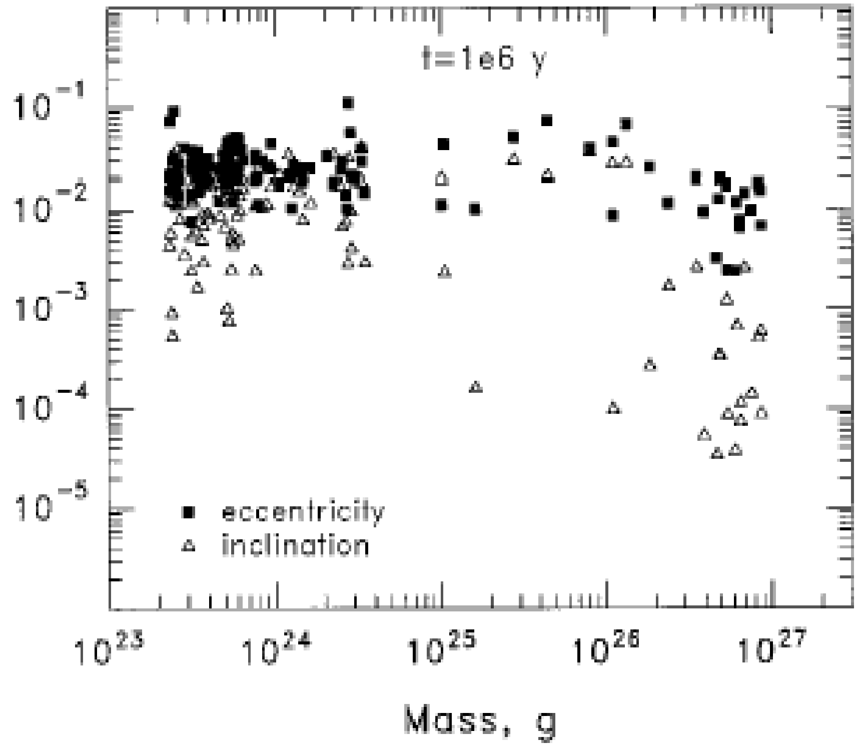
Terrestrial planet formation: 1- the old way

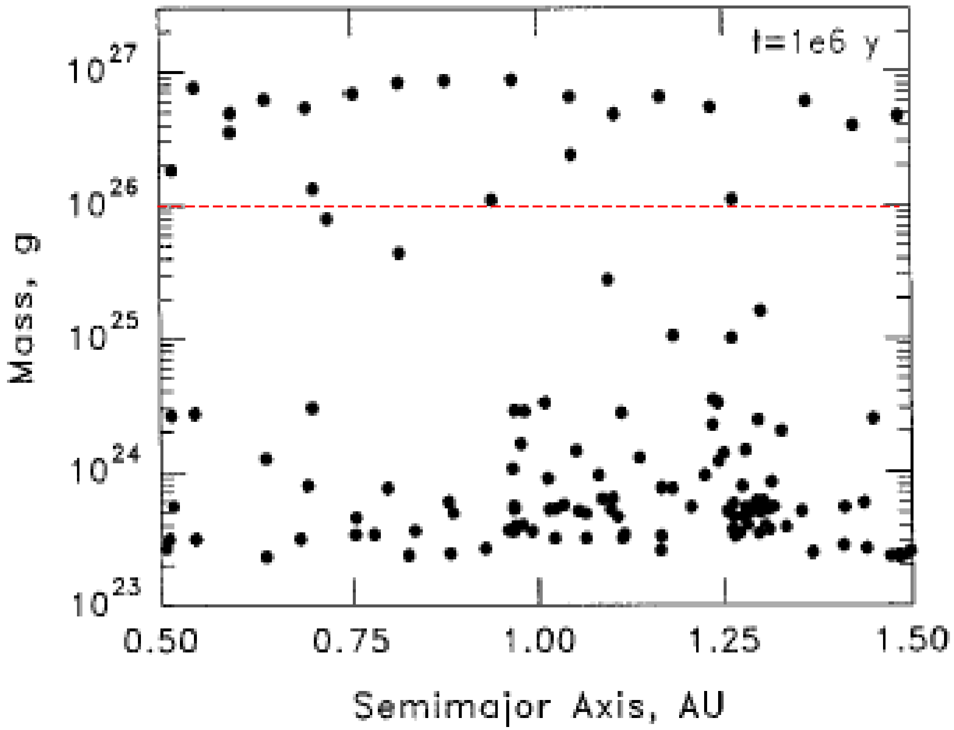
Runaway growth → Oligarchic growth → Giant impacts (chaotic growth)

Transition from runaway and oligarchic growth.







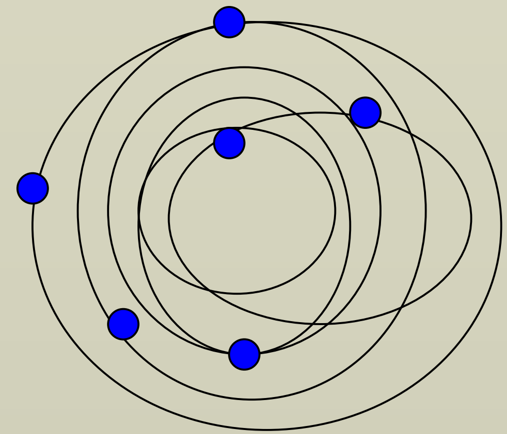


From Runaway growth to the phase of giant impacts (Moon formation…)

Protoplanets, because of their large mass, mutually excite their eccentricities and impact each other.

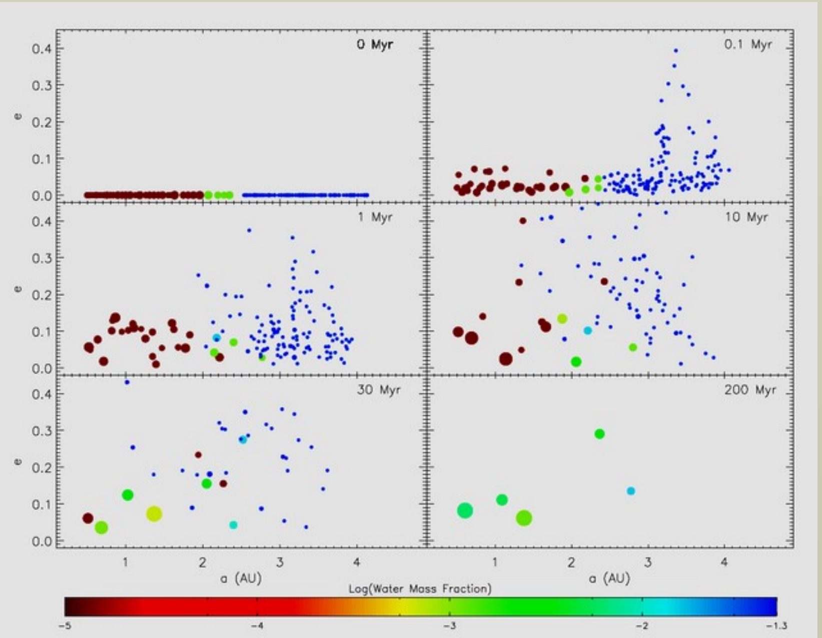
$M_p = (4 \pi B r^2 \Sigma_s / 3 M_\star)^{1/2}$ limiting mass for dynamical isolation. Beyond, impacts!

- $B \approx 4$ (multiple of Hill's sphere) $r_H = r\,(m_p / 3 M_\star)^{1/3}$
- Σ_s superficial density of planetesimals
- $M_\star$ Star mass



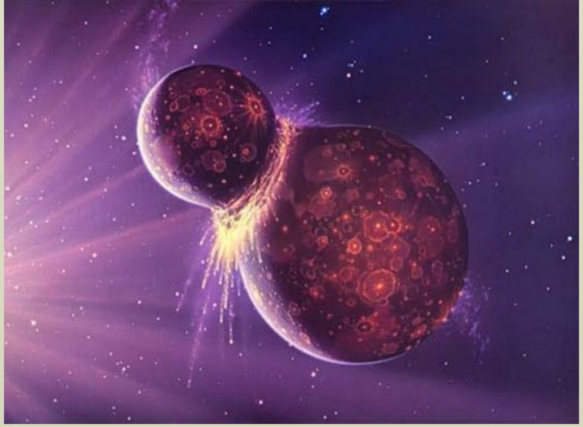
$M_p \approx 0.05\ M_\oplus$ at 1 AU

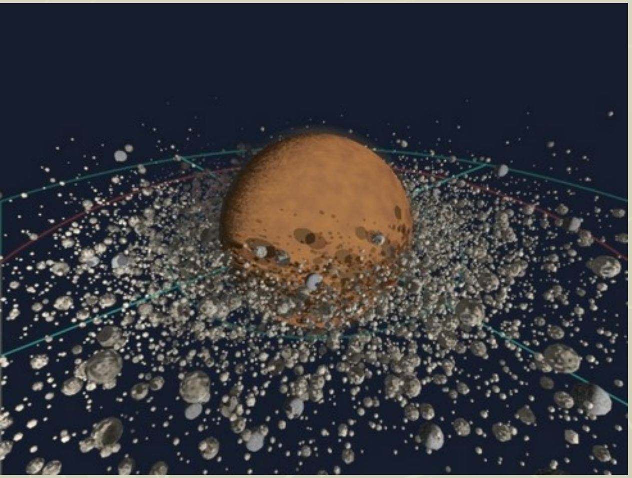
$M_p \approx 1.4\ M_\oplus$ at 4 AU

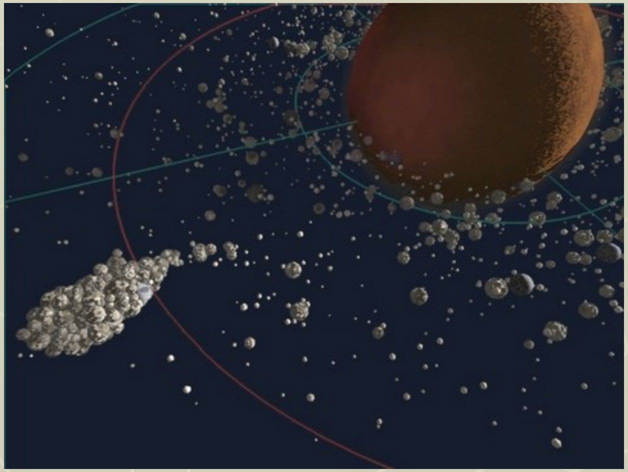


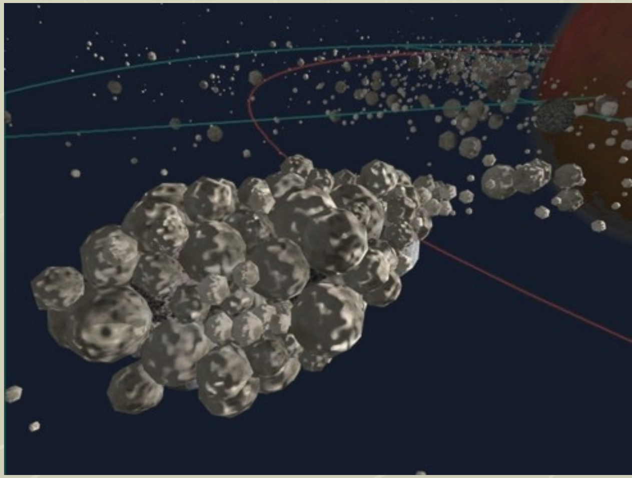
N-body simulation of protoplanets impacts leading to terrestrial planets.

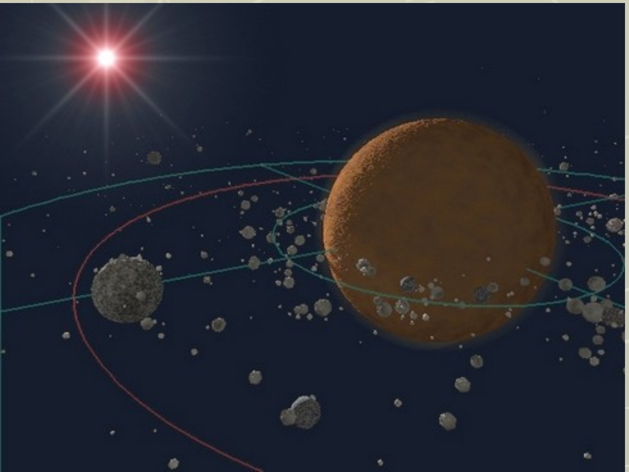
Moon origin by a *giant impact*. The projectile was approximately of Mars size.

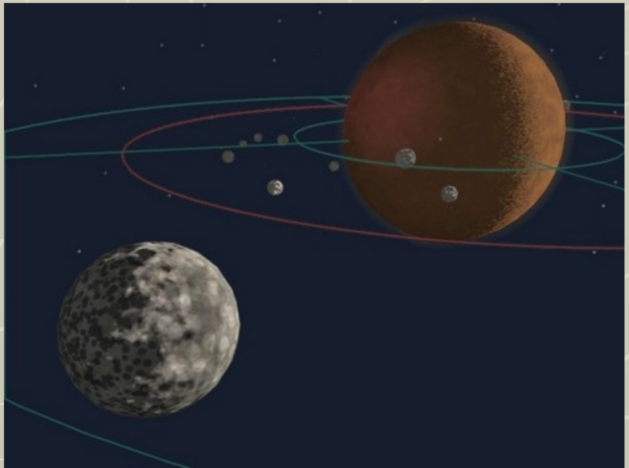






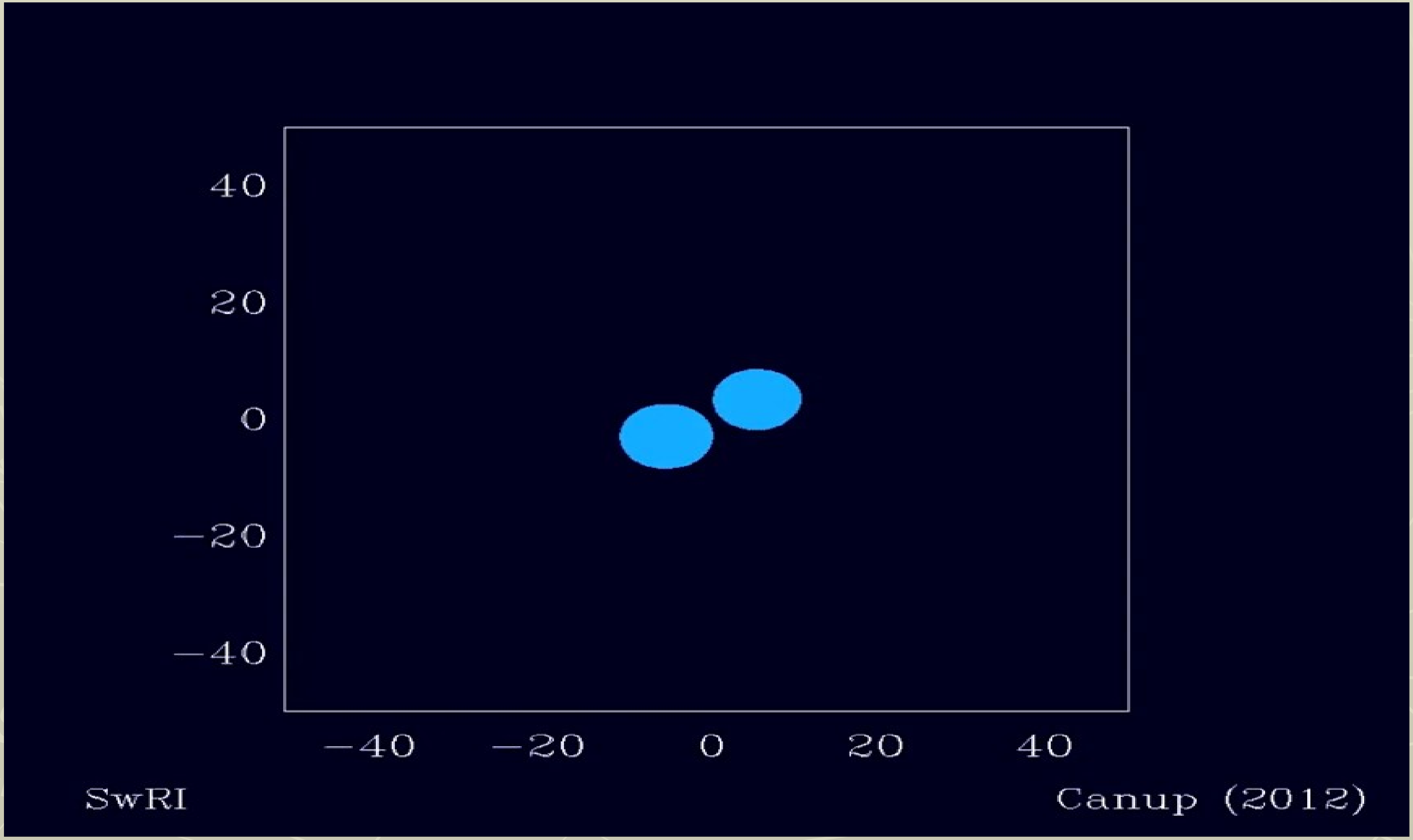






The Moon re-accumulates from the debris disk.

Modeling the 'Giant Impact' which formed the Moon.

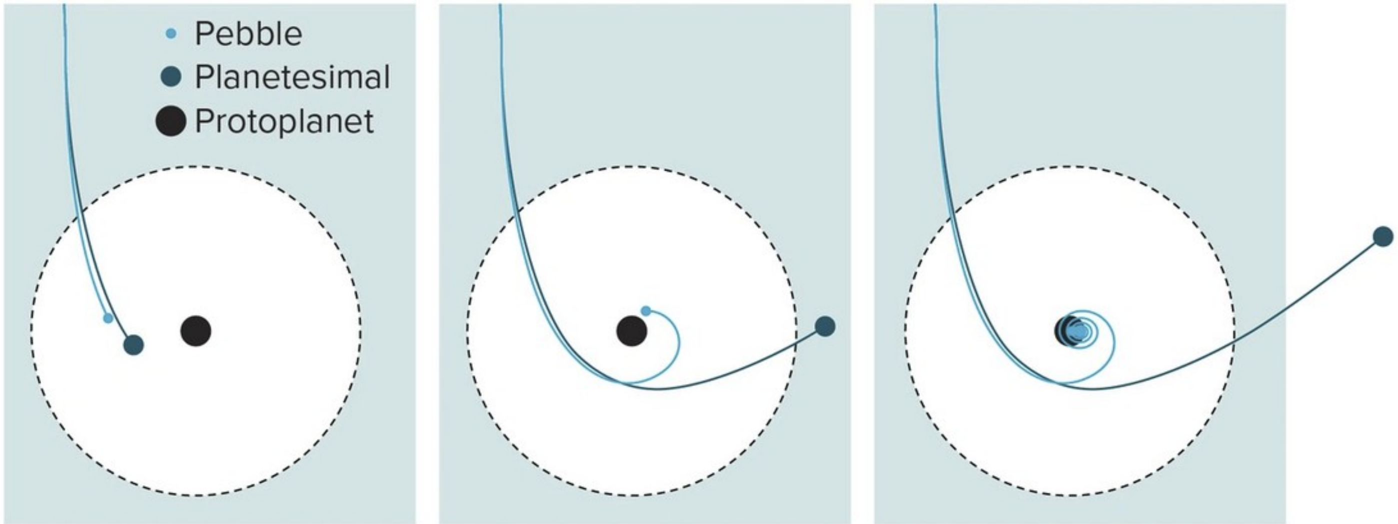


Pebble accretion: not all dust into planetesimals. Pebbles significantly contribute to the planet growth and speed up the planet formation timescale.

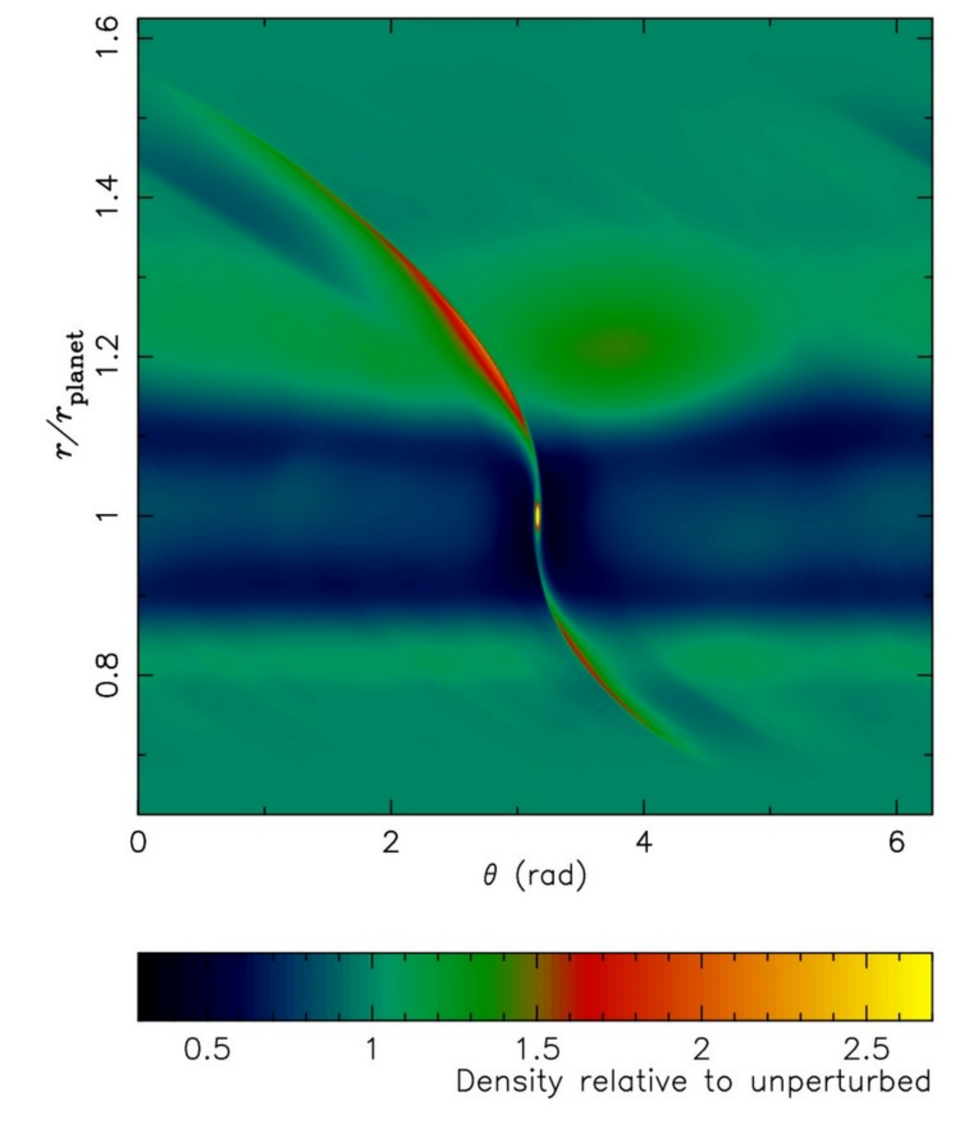
(*Ormel & Klahr 2010; Lambrechts & Johansen, 2012, 2014; Johansen & Lambrechts, 2017*)

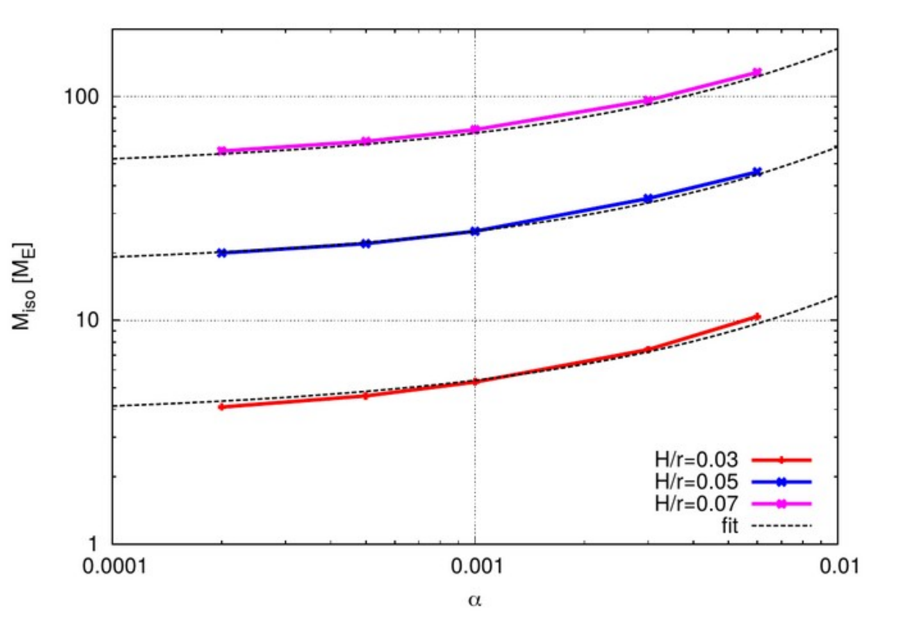
A pull on pebbles

A pebble flying past a protoplanetary body is slowed by friction from surrounding gas as it enters the protoplanet's gravitational influence (dotted line). That slowdown allows the small pebble to be captured by the protoplanet's gravity and spiral in for a smash-up, whereas a larger planetesimal just zips by. Over time, many pebbles will coalesce with the protoplanet, allowing it to grow large quickly.



Pebble isolation mass: accretion of mass stops when the gap carved by the planet in the gaseous disk generates a pressure maximum outside of its orbit, which stops the inward flux of pebbles (pressure trap).

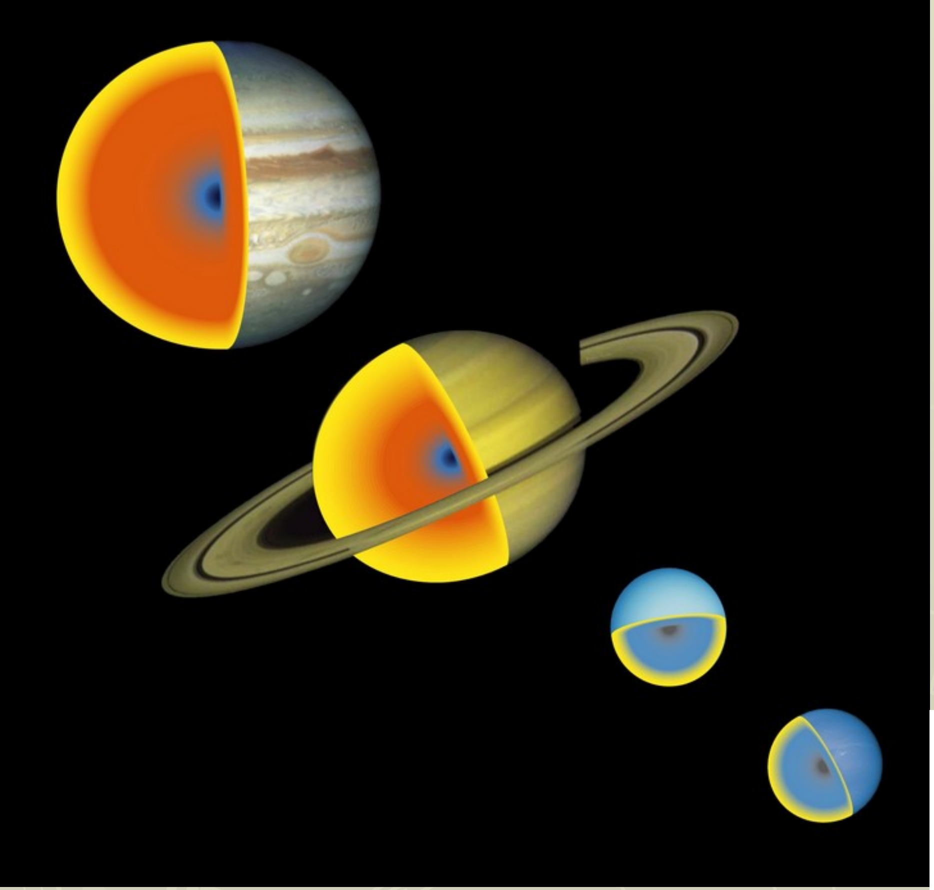




(*Lambrechts et al., 2014; Bitsch et al. (2018)*)

It depends on the viscosity of the disk alpha and on the scale height.

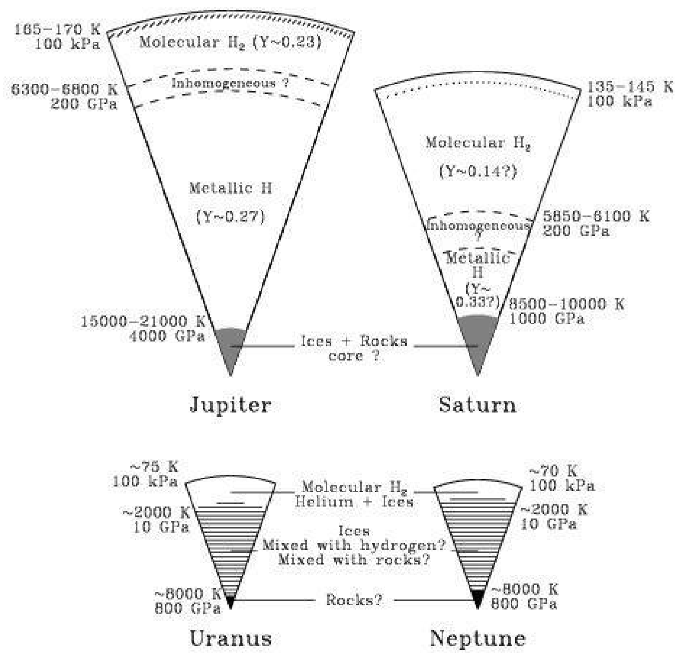
Question: how much mass in the core of giant planets or in terrestrial planets is due to pebble accretion?



Giant planets: structure

- Yellow: molecular hydrogen
- Red: metallic hydrogen
- Blue: ices
- Black: rock

Enrgy source: helium sedimentation.

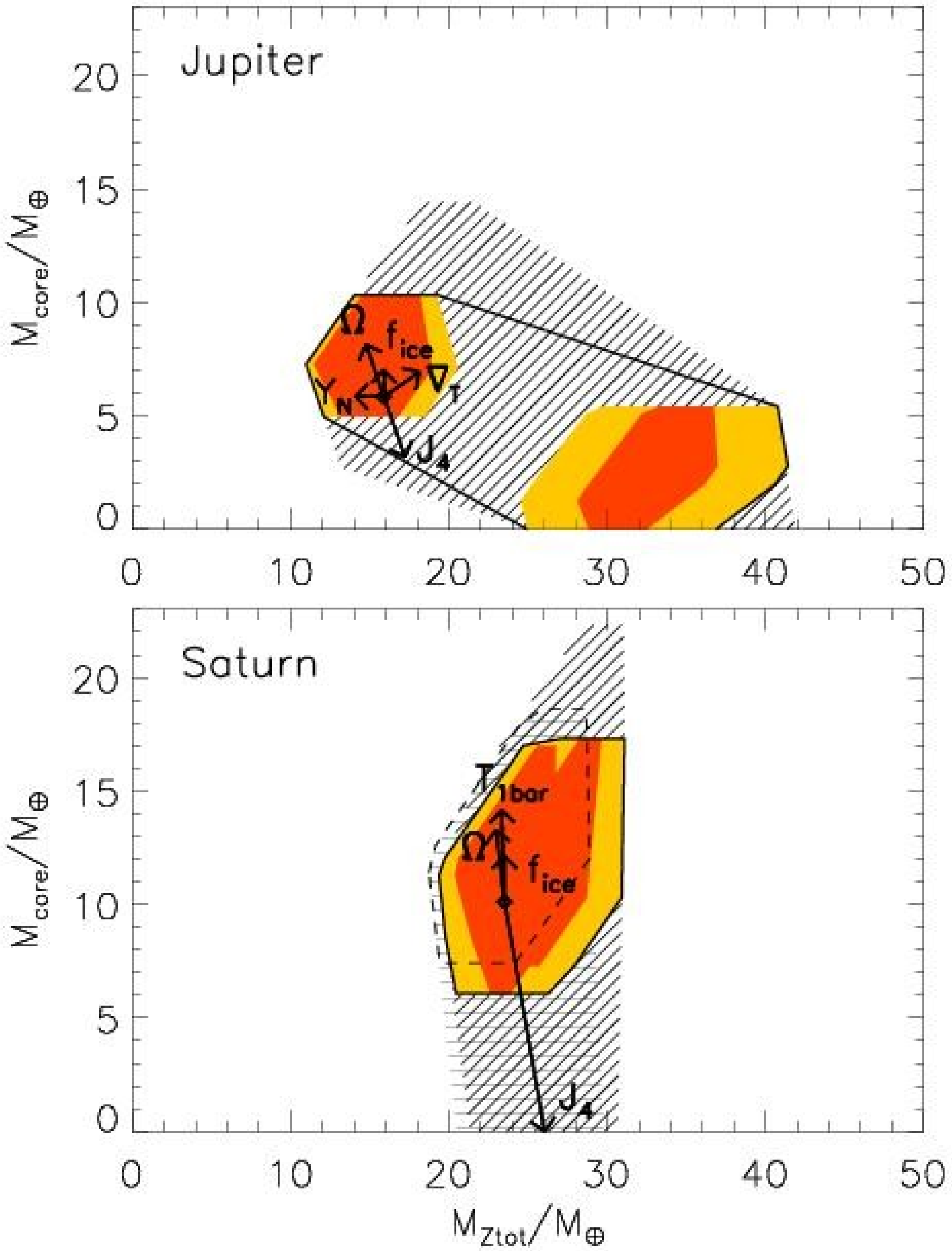


Y = He mass mixing ratio

Jupiter: Y=0.238 ± 0.007

Saturn: Y=0.140 ?

Sun: Y=0.280 ± 0.005



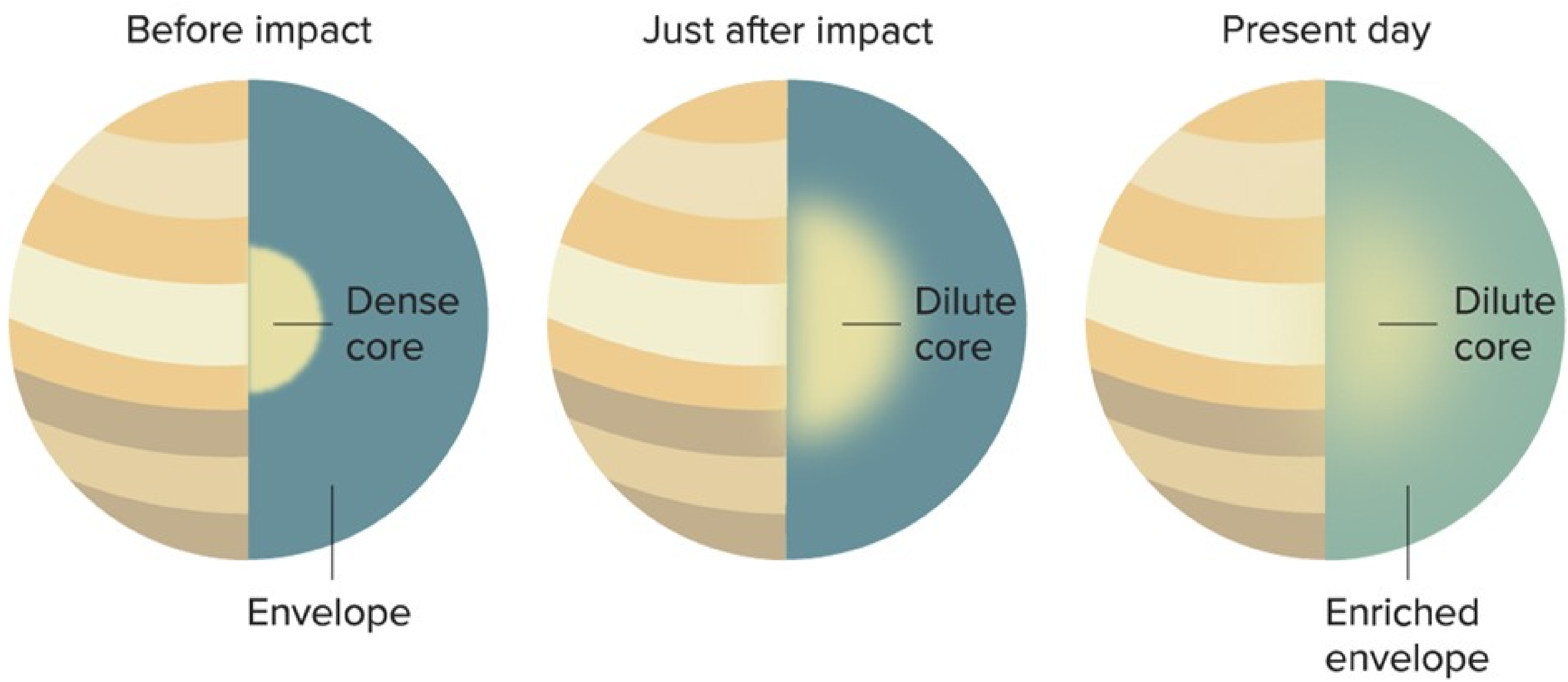
CORE of Giants

Hydrostatic models based on ground and space observations and data:

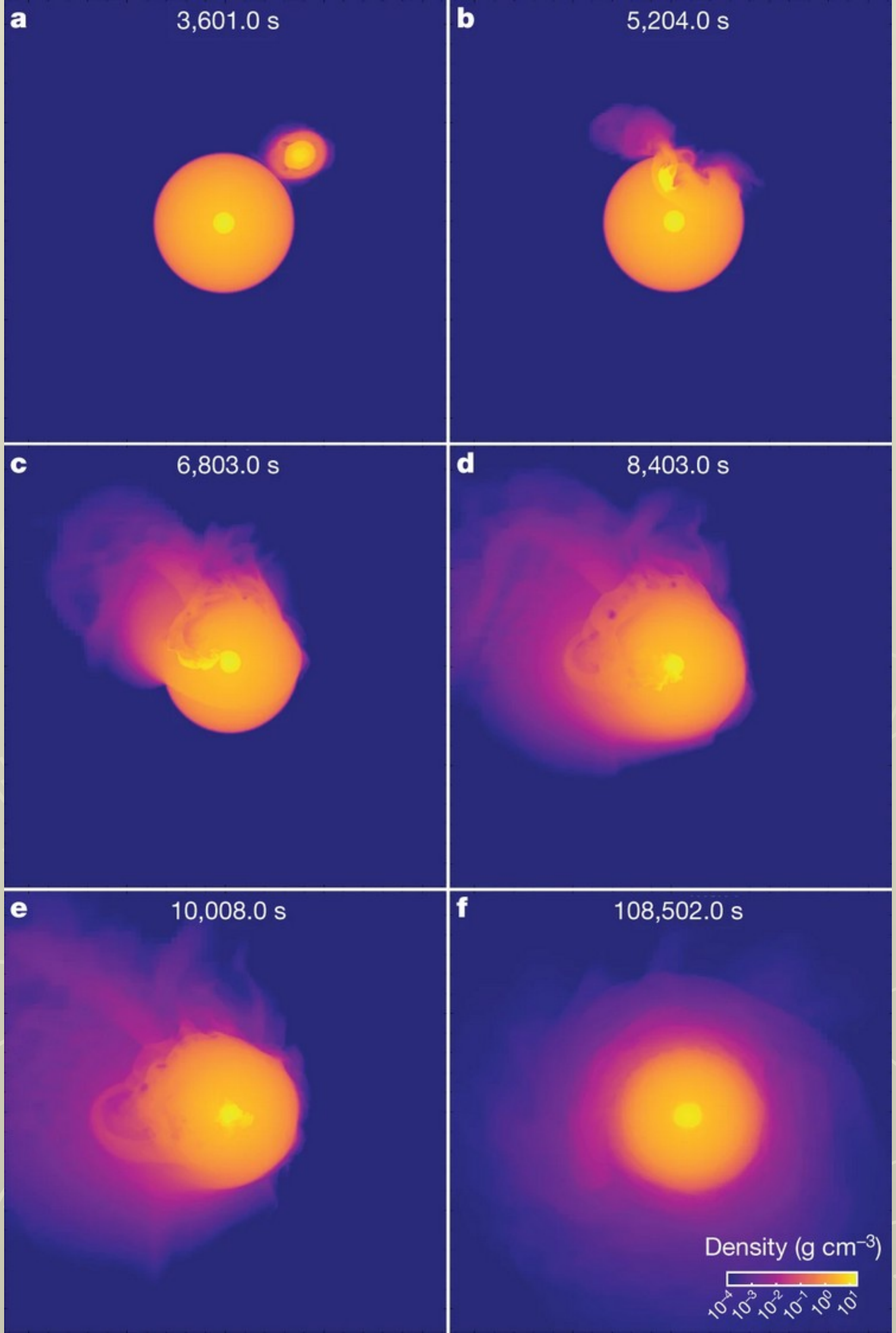
- **Gravitational field of the planet.**
- **Magnetic field**
- **Energy emission**
- **Superficial temperature and pressure**
- **Chemical composition of the atmosphere.**
- **mixing ratios**
- **JUNO mission → more detailed results!**

Recent models from JUNO mission data suggest a diluted core, maybe due to an impact with a large planetary embryo (Liu, Nature 572, 355. 2019) .

A collision may have left Jupiter with a fuzzy core

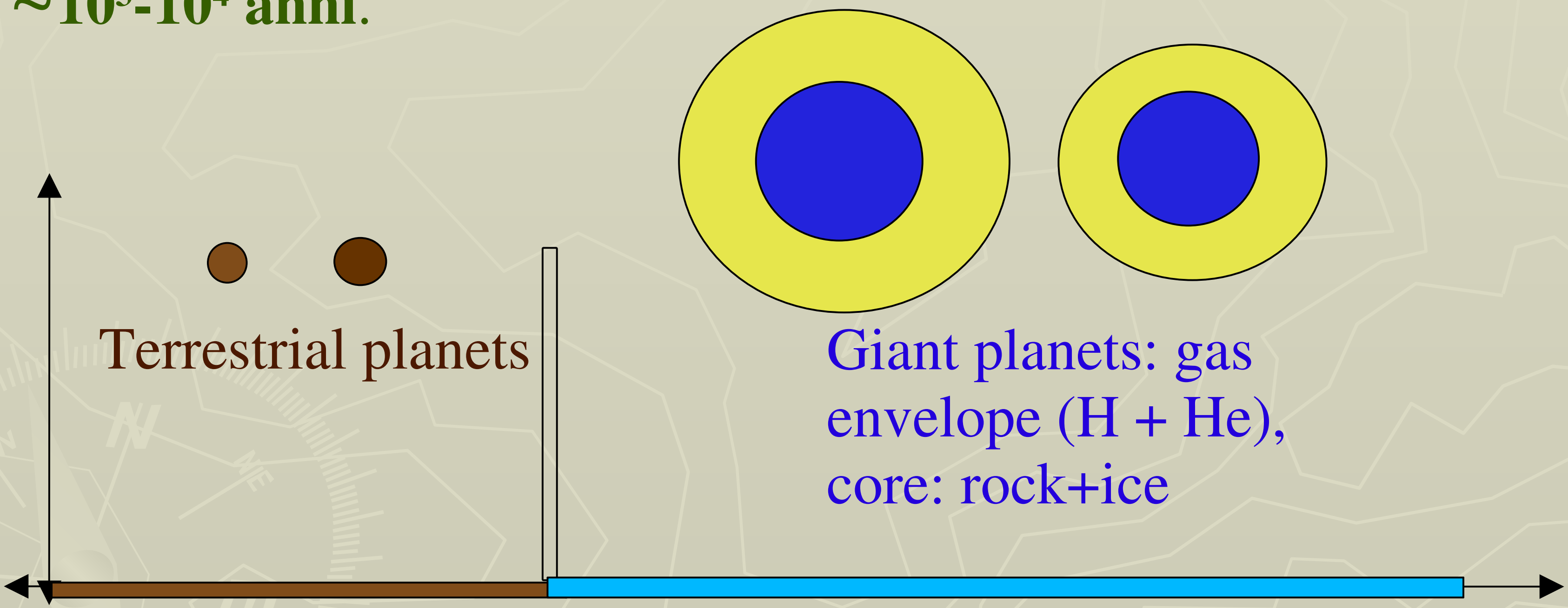


SOURCE: T. GUILLOT / *NATURE NEWS & VIEWS* 2019 KNOWABLE MAGAZINE



Hydro-simulation of the impact of a 10 Earth Mass protoplanet with a proto-Jupiter having a core of 10 Earth masses (Liu, 2019).

- **Planetesimal & pebble accumulation into protoplanets t ~ 1 – 5 Myr (terrestrial planets), ≤ 1 Myr for giant planets?**
- **Giant impacts: t ~ 1-100 Myr in the terrestrial planets region, much faster for giant planets which have to form before in a few Myr**
- **"Gas infall" on giant planets after core formation very fast: t ~10^3-10^4 anni.**

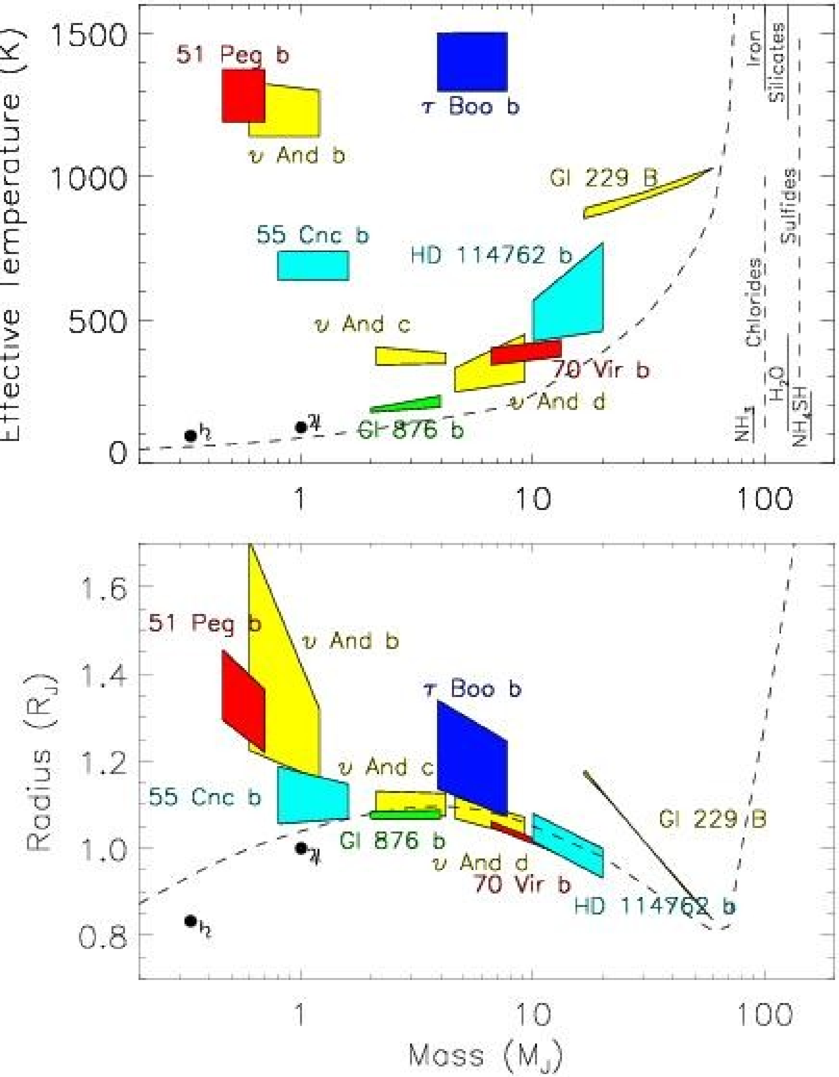


4 UA (Frost line) **T ~ 170 K**

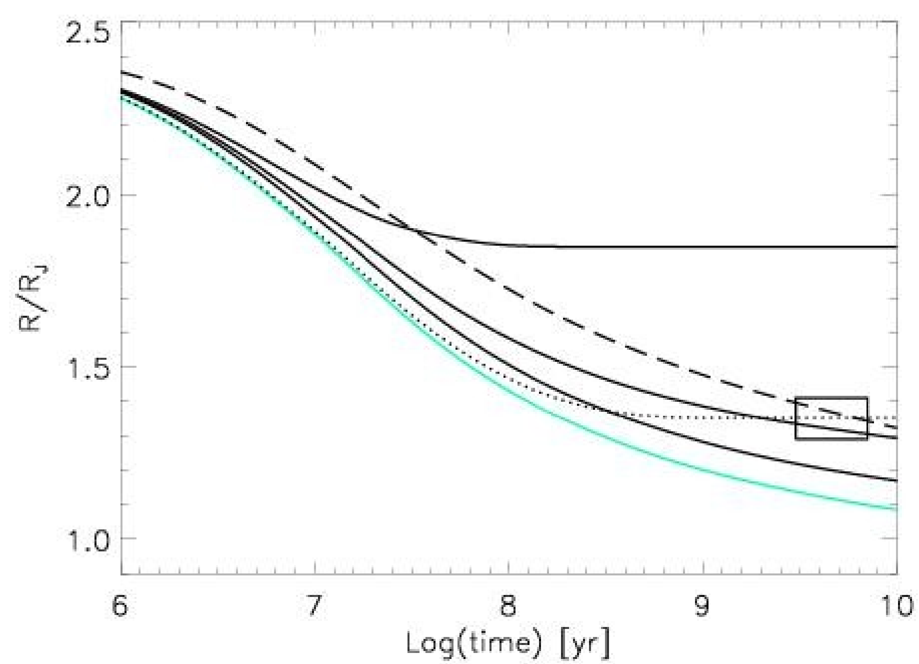
Within the frost line only silicates and methal are solid. M_s / M_g ~ 1/240

Beyond the frost line ices condenstate (H_2O, CH_4, NH_3) M_s / M_g ~ 1/60

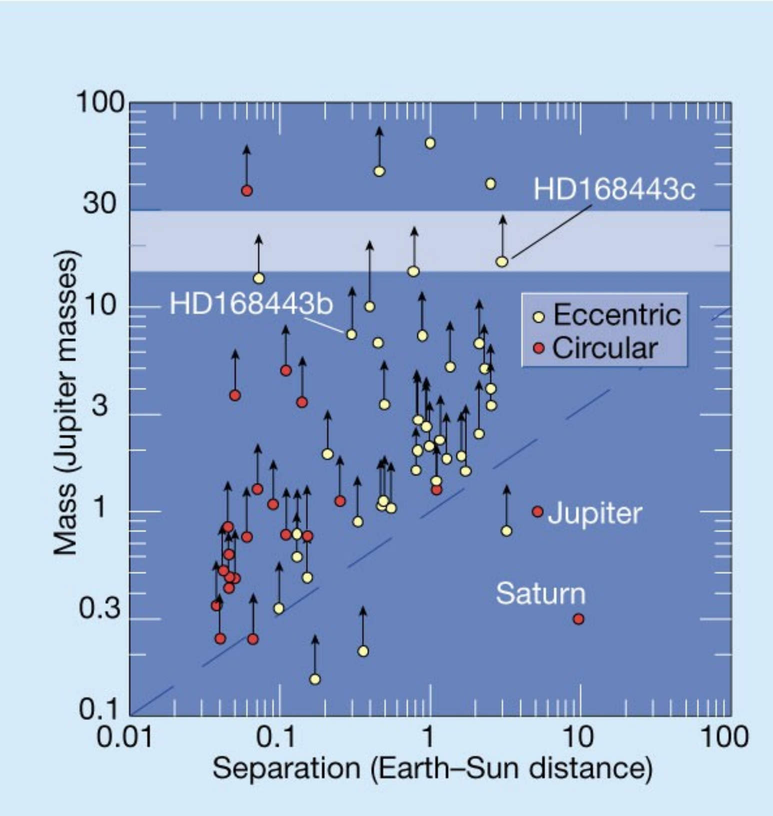
Lo stesso modello puo’ essere applicato ai pianeti extrasolari. Calcolo delle temperature e raggi dei vari pianeti (www.obs-nice.fr/guillot).



Evoluzione nel tempo di un ‘Hot Jupiter’

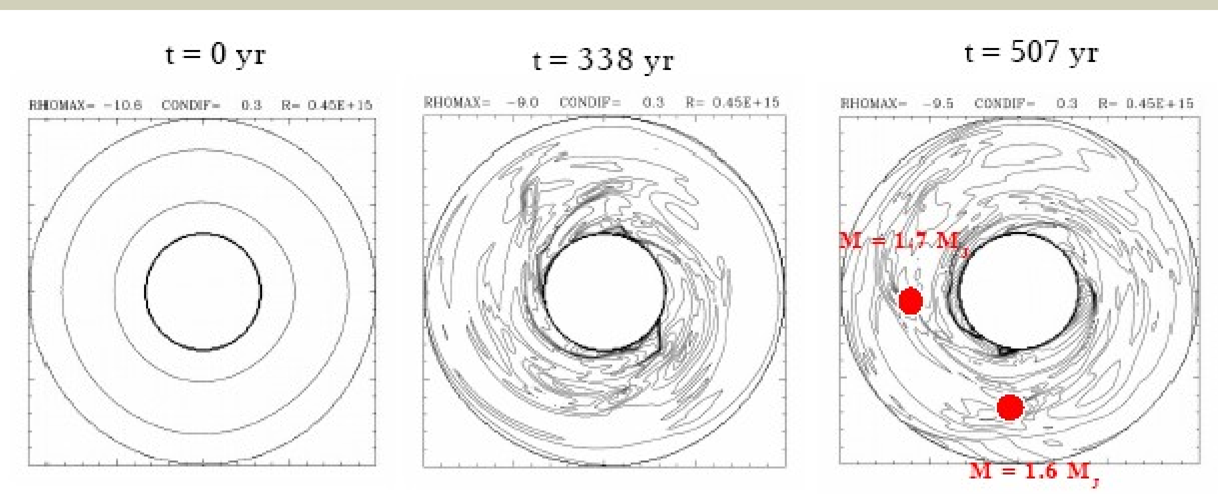


MODELLO DI FORMAZIONE PER INSTABILITA' DEL DISCO



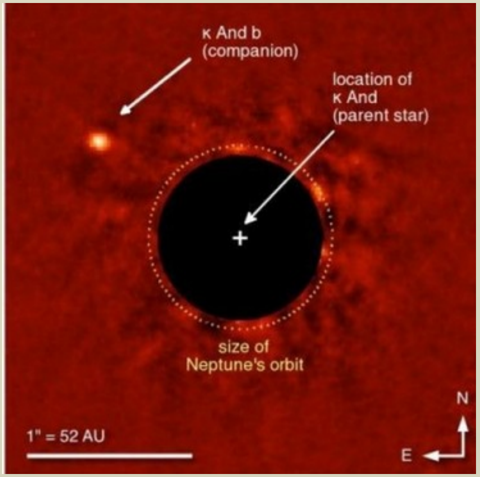
Tempi scala di formazione: 10^2-10^3 anni

Non funziona per il sistema solare (Saturno troppo piccolo, Urano e Nettuno troppi ghiacci)

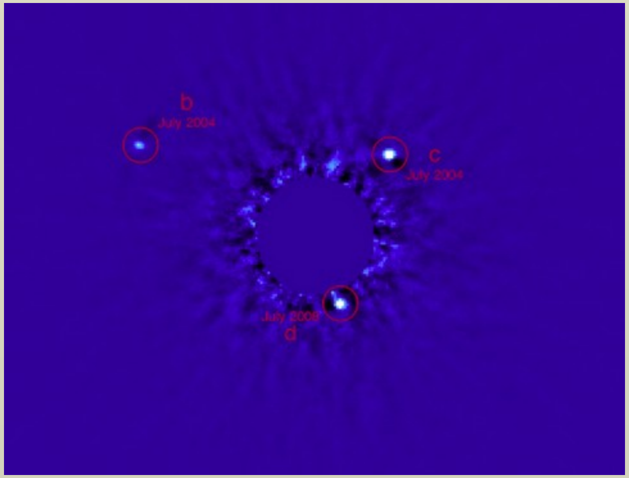


Risultati di calcoli con un modello idrodinamico gravitazionale e radiativo 3D (Boss).

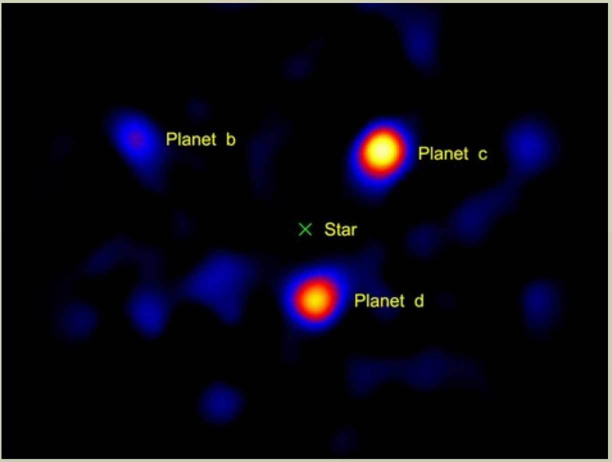
M_{disco} = 0.106 M_{sun}, T(10 AU) = 50 K



IRXS J160929.1-210524
(young star (5Myr)



HR8799



Si sono formati per instabilità gravitazionale?

Pianeta	Massa	Periodo orb.	Sem. maggiore	Scoperta
e	7+3−2 MJ	~ 45 anni	~ 14,5 ± 0,5 UA	2010
d	7+3−2 MJ	~ 100 anni	~ 24 UA	2008
c	7+3−2 MJ	~ 190 anni	~ 38 UA	2008
b	5+2−1 MJ	~ 460 anni	~ 68 UA	2008
Cintura asteroidale	—	—	~75 UA	2008