Theory and physics of planet formation

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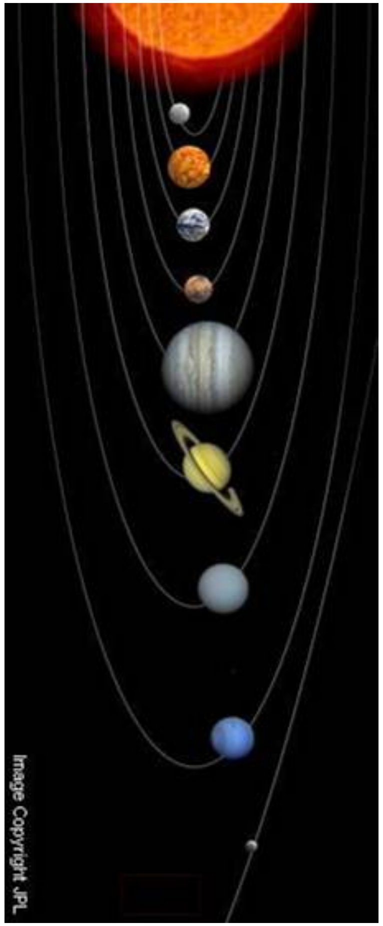
Solar System planets



	Mass	Distance
Mercury	$0.06\ M_{earth}$	0.39 AU
Venus	$0.82\ M_{earth}$	0.72 AU
Earth	$1.0\ M_{earth}$	1.0 AU
Mars	$0.11\ M_{earth}$	1.5 AU
Jupiter	$318\ M_{earth}$	5.2 AU
Saturn	$98\ M_{earth}$	9.5 AU
Uranus	$15\ M_{earth}$	19.2 AU
Neptune	$17\ M_{earth}$	30.1 AU
Pluto	$0.002\ M_{earth}$	39.5 AU

Terrestrial planets

Jovian planets

Dwarf planet

$1\ M_{earth} = 6 \times 10^{24}$ kg $= 3\times10^{-6}\ M_{sun}$, 1 AU $= 1.5 \times 10^{11}$ m

How to detect planets around other stars?

Direct detection
- Direct imaging

Effect on motion of parent star
- Astrometric wobble
- Timing shifts
- Radial velocity method

Effect on flux from stars
- Planetary transits
- Gravitational microlensing

Other indirect techniques
- Disc structures

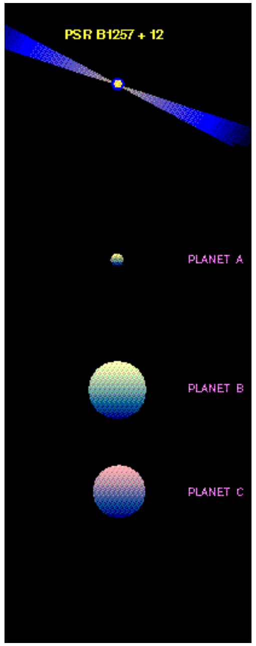
Pulsar Planets

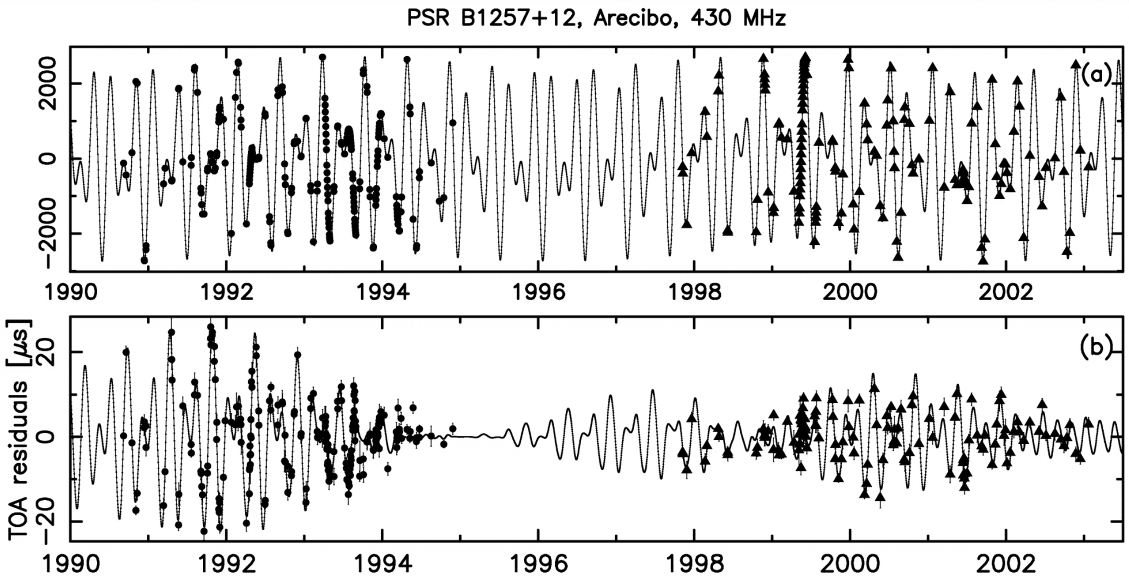
First extrasolar planets detected around 6.2 ms pulsar PSRB1257+12
(Wolszczan & Frail 1992)

The planets are small, coplanar, low eccentricity, and gravitationally interact with each other through 3:2 resonance (Malhotra 1992; Konacki & Wolszczan 2003)

Possible fourth planet or asteroid belt beyond C
(Wolszczan et al. 2000)

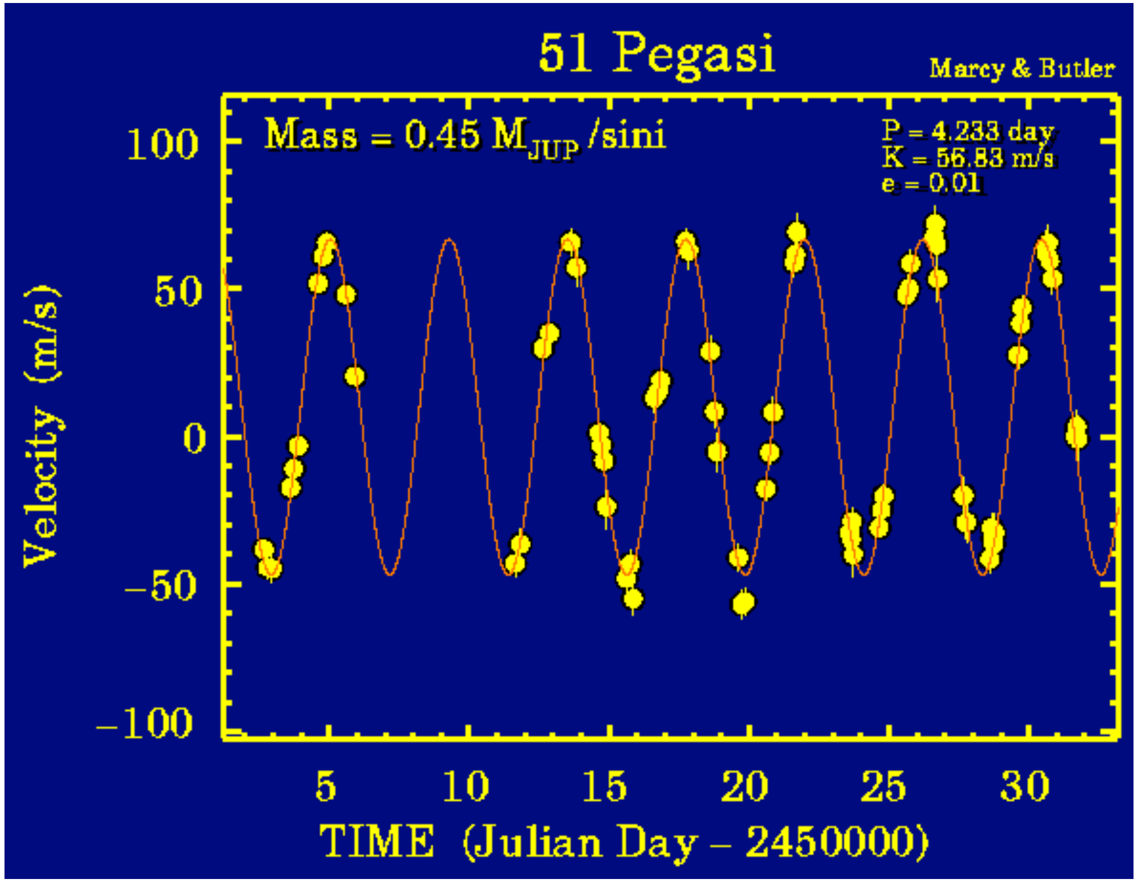
	a, AU	M, M_{earth}	I	e
A	0.19	0.02	-	0
B	0.36	4.3	53^0	0.019
C	0.47	3.9	47^0	0.025





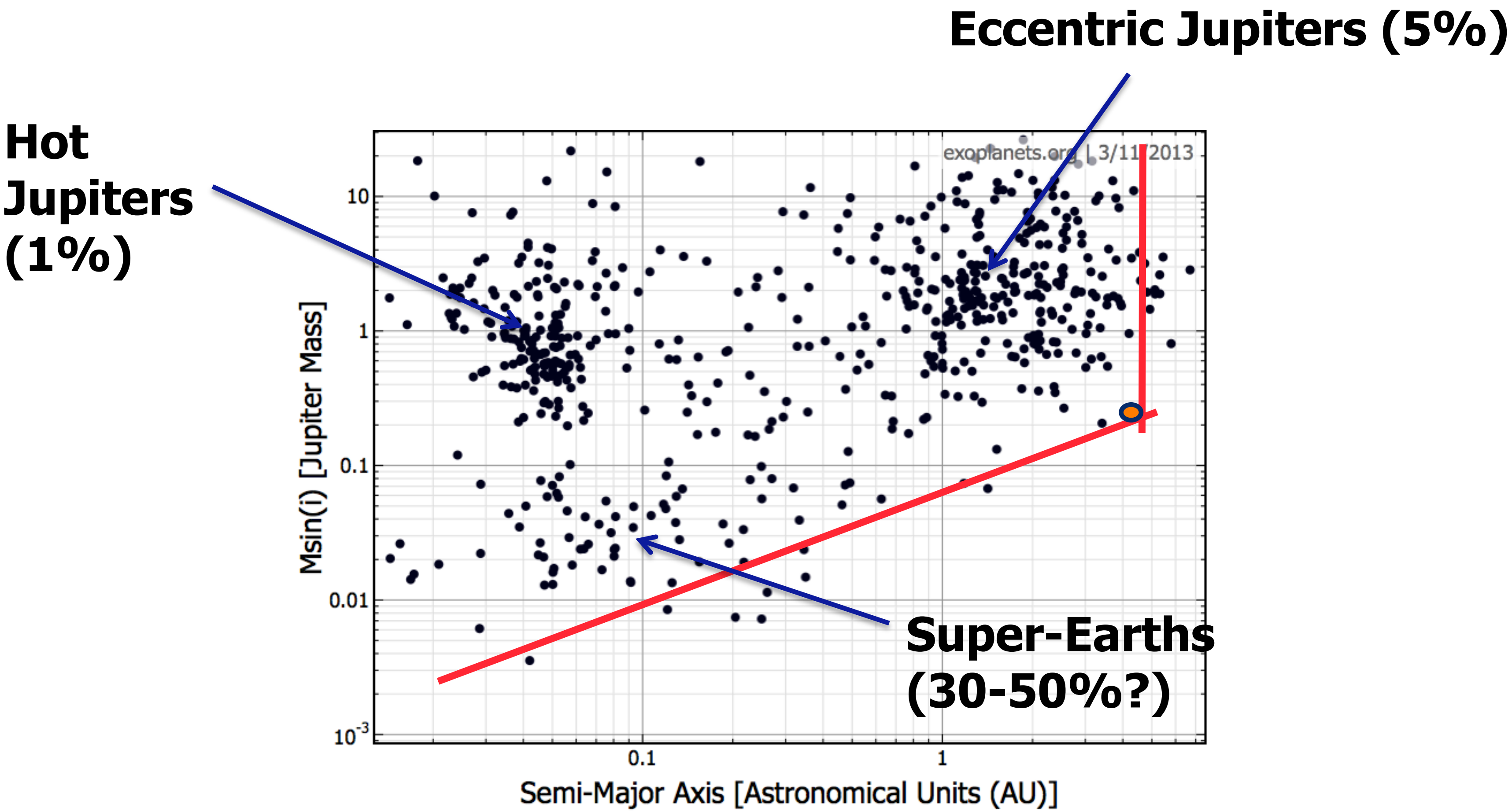
Radial Velocity Planets

First extrasolar planet around main sequence star 51 Peg (G2 at 15pc) used radial velocity method to detect >$0.45M_{jupiter}$ planet at 0.05AU near circular orbit (Mayor & Queloz 1995; Marcy & Butler 1997) = **HOT JUPITER**



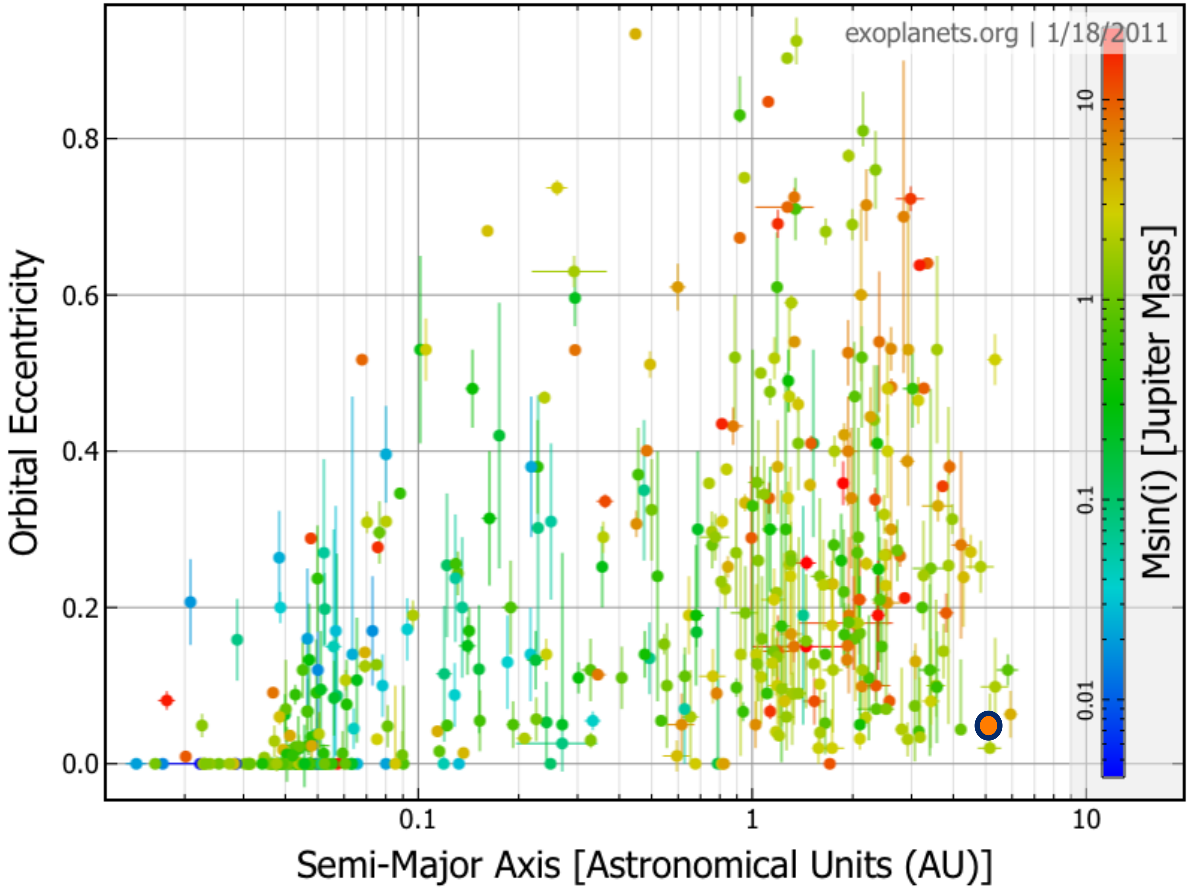
Now >500 planets discovered using this method, >5% of stars have planets (see http://exoplanet.eu or http://exoplanets.org)

Planet discovery space



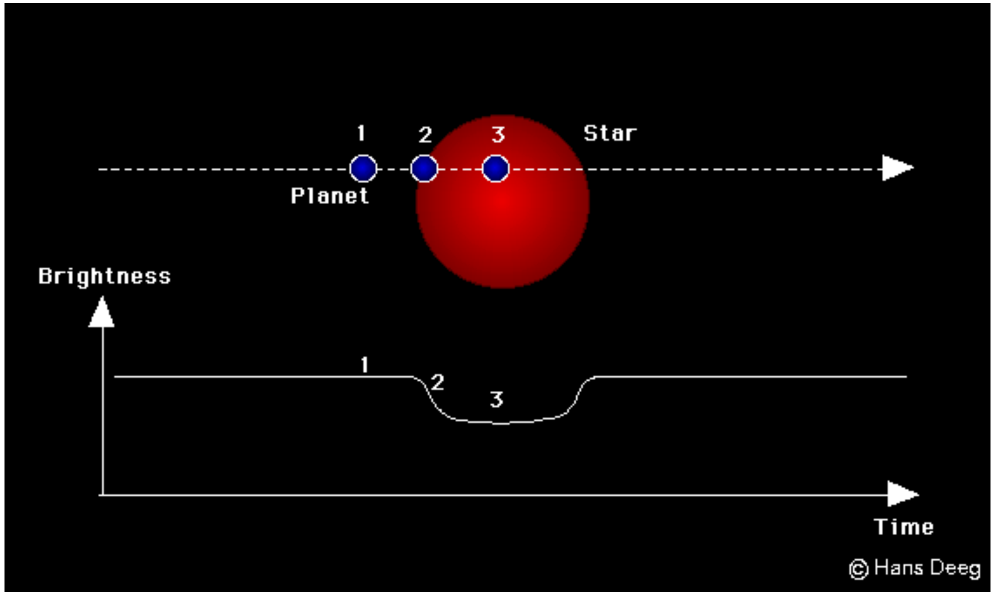
Planet eccentricity distribution

Giant planets at few AU have eccentric orbits: mean 0.32, up to 0.92 (compared with <0.05 for the Solar System)

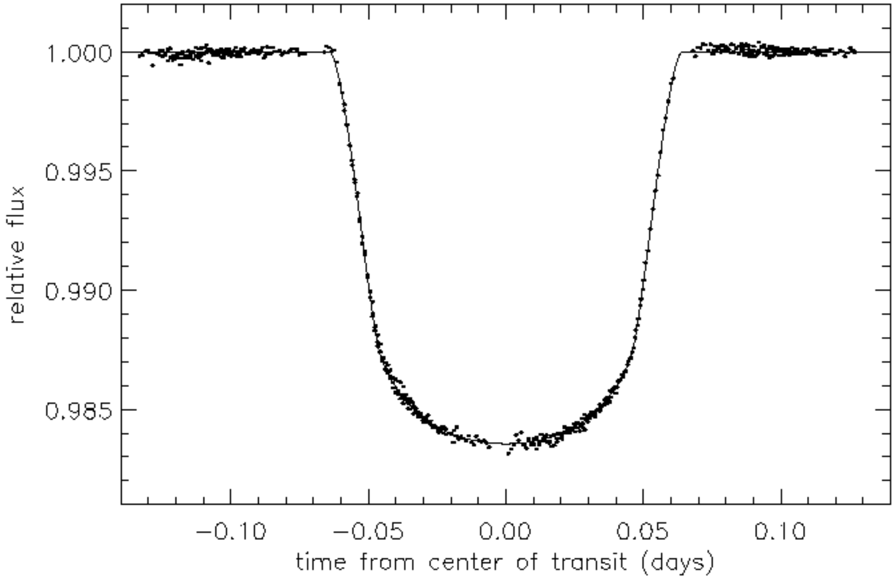


Transit detection method

If orientation just right, star gets fainter when the planet passes in front of it



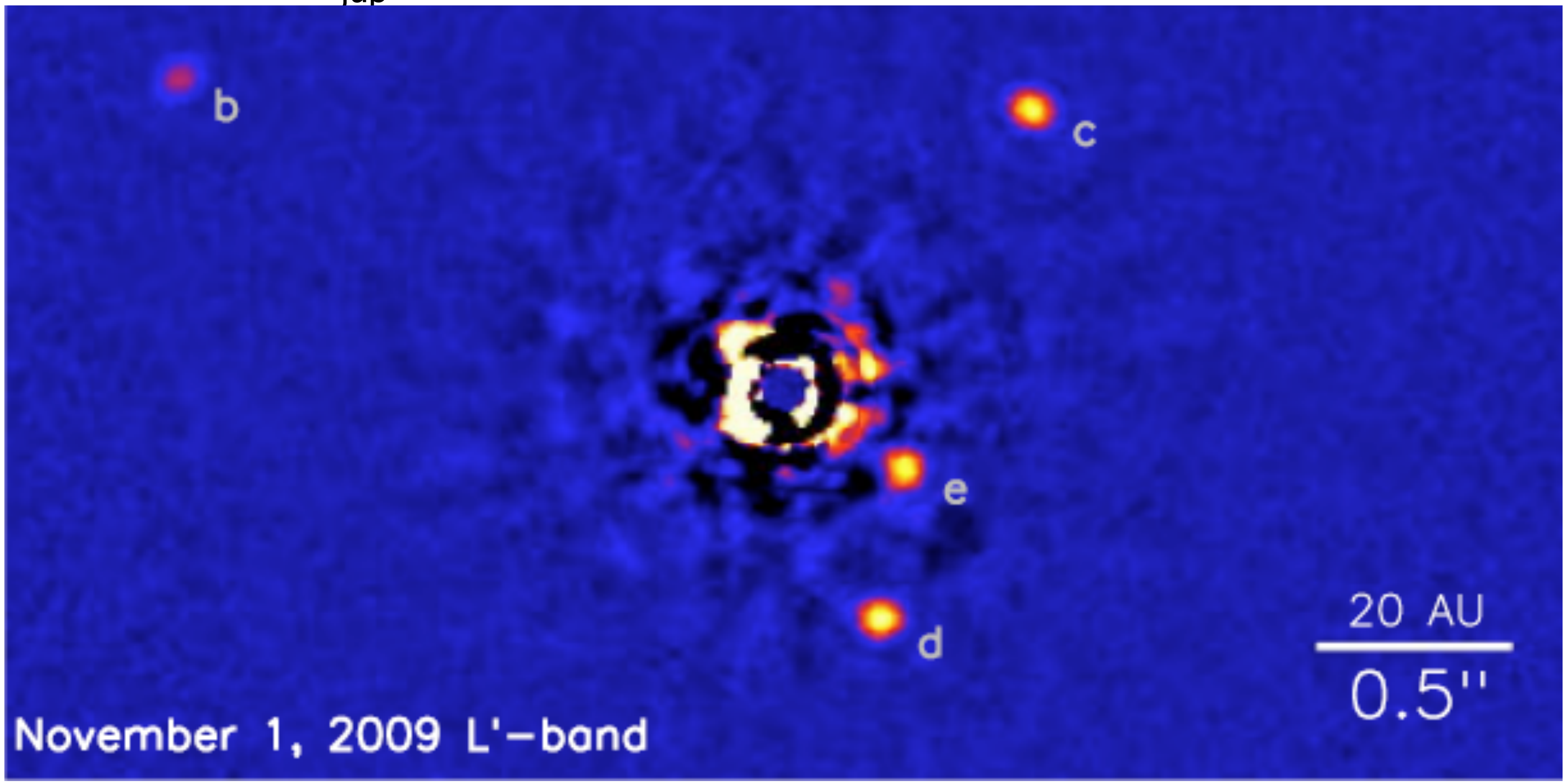
e.g., HD209458b discovered by radial velocity; transit lasts 3hrs every 3.5days



Space missions Kepler and CoRoT can detect few M_{earth} planets (Queloz et al. 2009; Batalha et al. 2011)

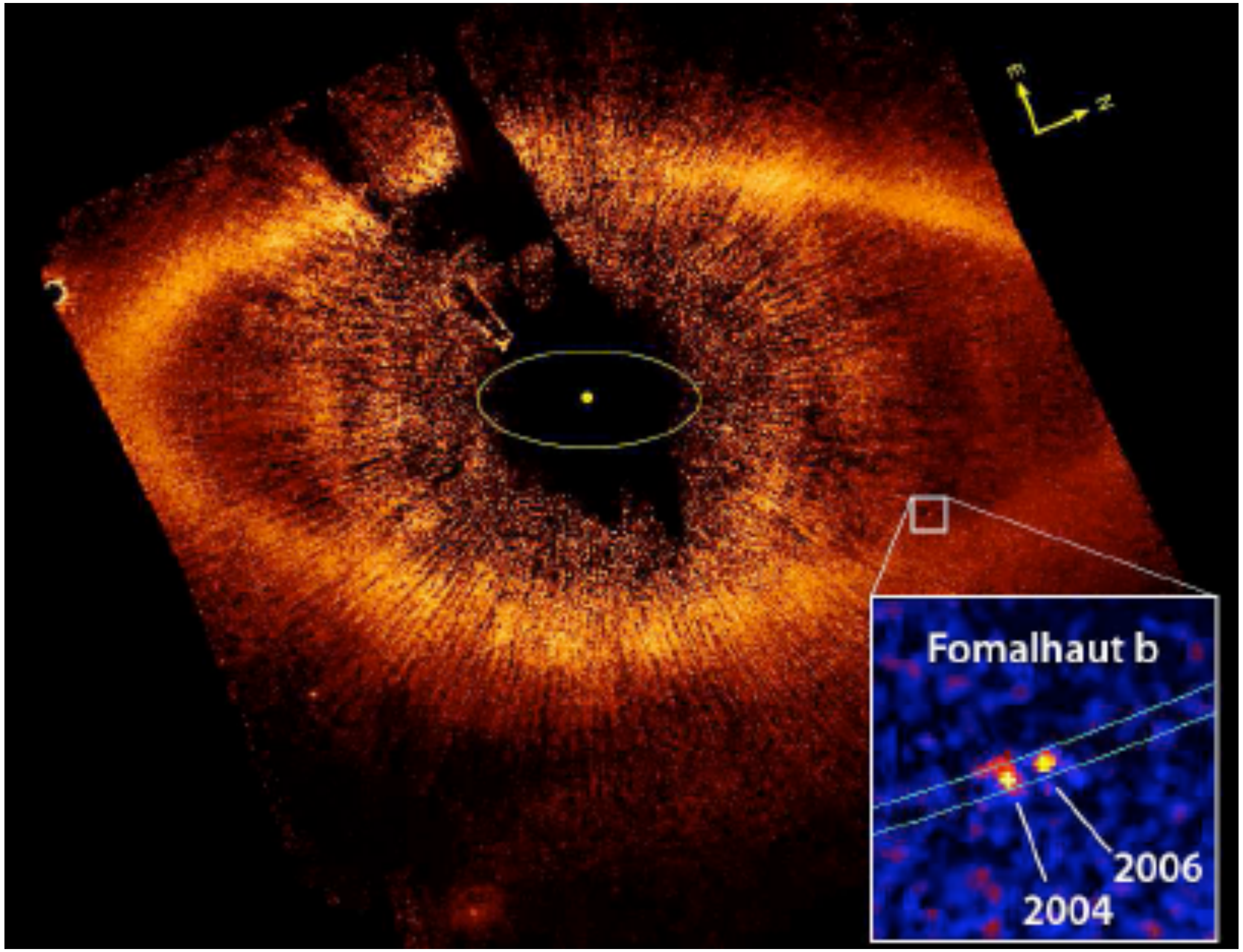
Direct imaging: outer planets

Four planets imaged around 60Myr A star HR8799 with masses 5-13M_{jup} at 14-68AU (Marois et al. 2010)

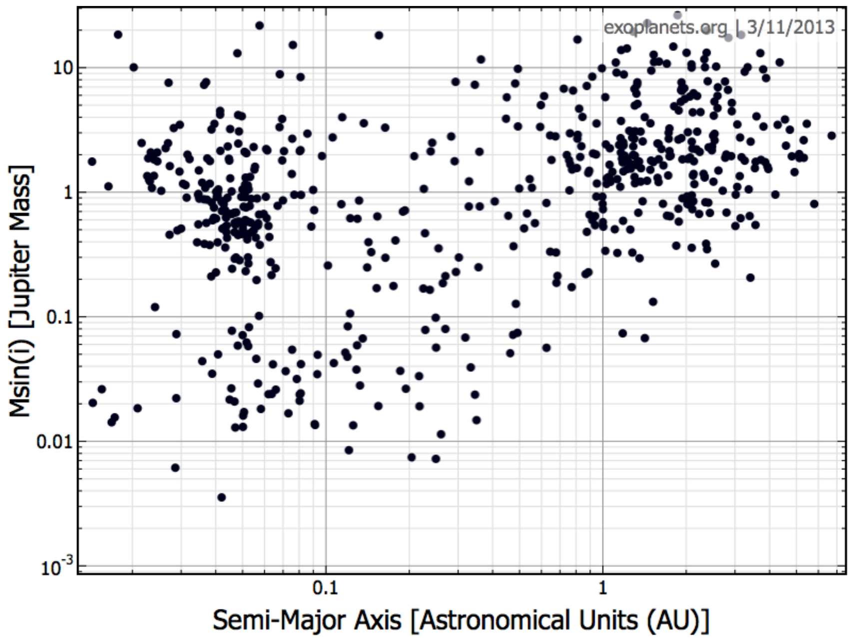


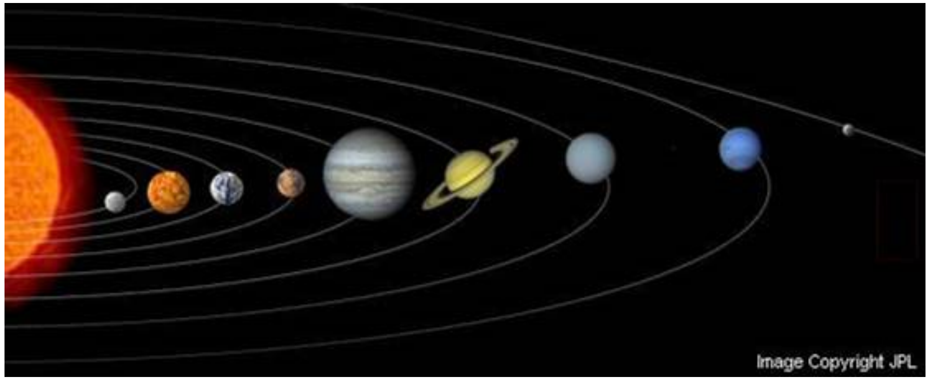
Planet interacting with debris disc

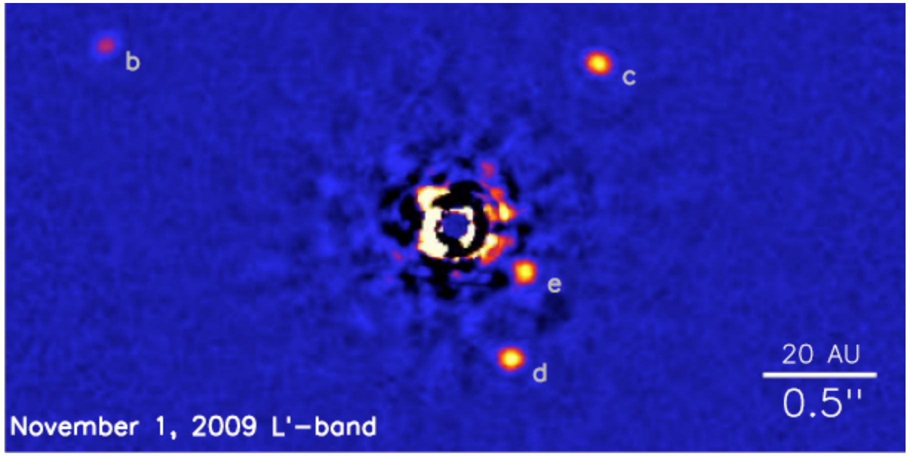
$<2M_{jup}$ planet imaged at inner edge of debris disc around 200Myr A5V star Fomalhaut (Kalas et al. 2008)



So, where did all these planetary systems come from?







Planets form in disks

Two observations of the Solar System that help understand how planets form (Lissauer 1993):

(1) planets orbits are circular, coplanar, and in same direction

Planets form in massive (>>1$M_{jupiter}$) circumstellar disks

NB the idea that planets form in circumstellar disks (the **solar nebula**) goes back to Swedenborg (1734), Kant (1755) and Laplace (1796)

(2) formation took less than a few Myr

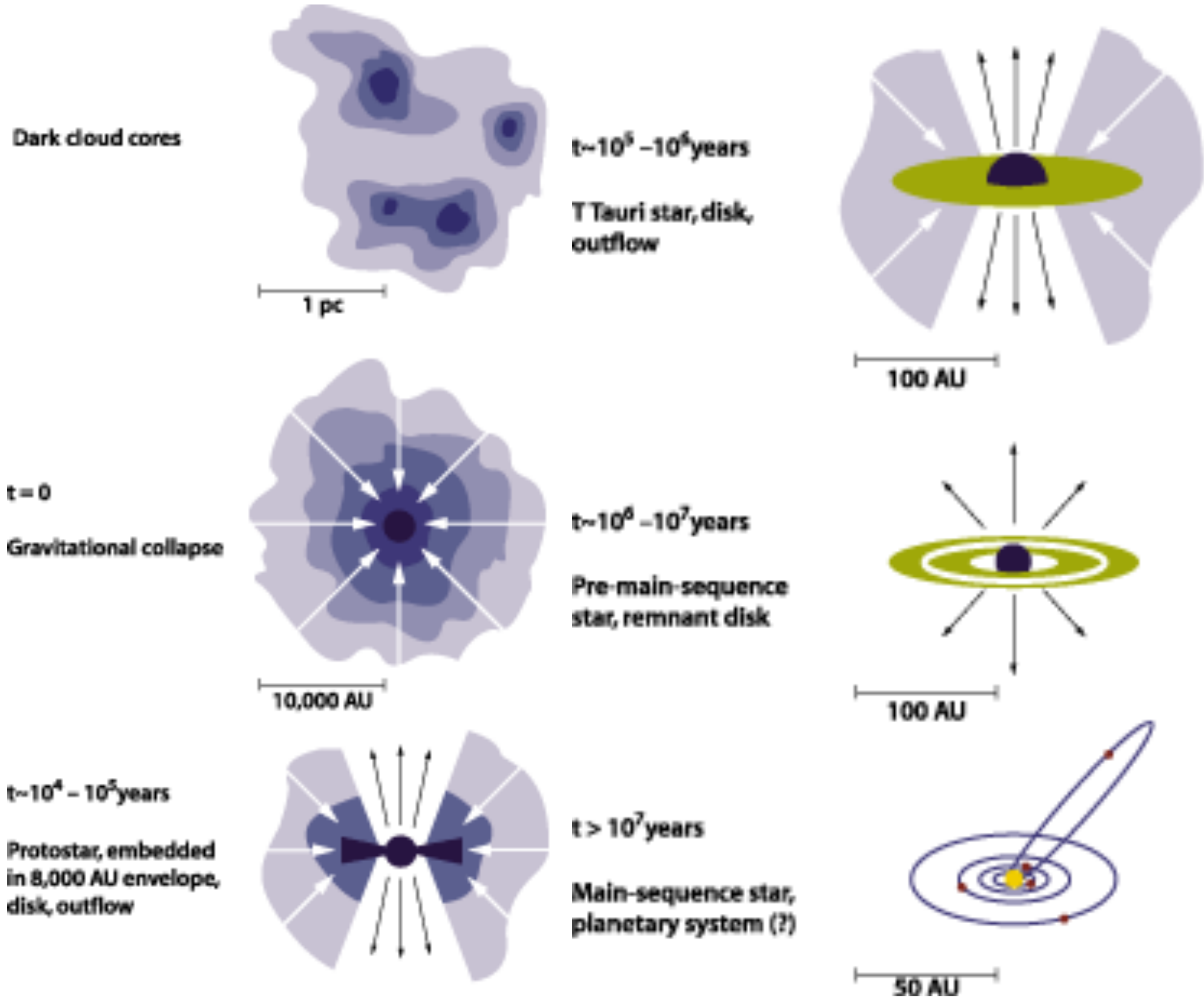
Planets form in circumstellar disks in a few Myr

Star Formation

Stars form from the collapse of clouds of gas and sub-micron sized dust in the interstellar medium (Shu et al. 1987)

After ~1 Myr end up with a star and protoplanetary disk extending ~100 AU

This disk disappears in ~10 Myr and is the site of planet formation



Planet formation models

There are two main competing theories for how planets form:

- **Core accretion** (Safronov 1969; Lissauer 1993; Wetherill, Weidenschilling, Kenyon,…)

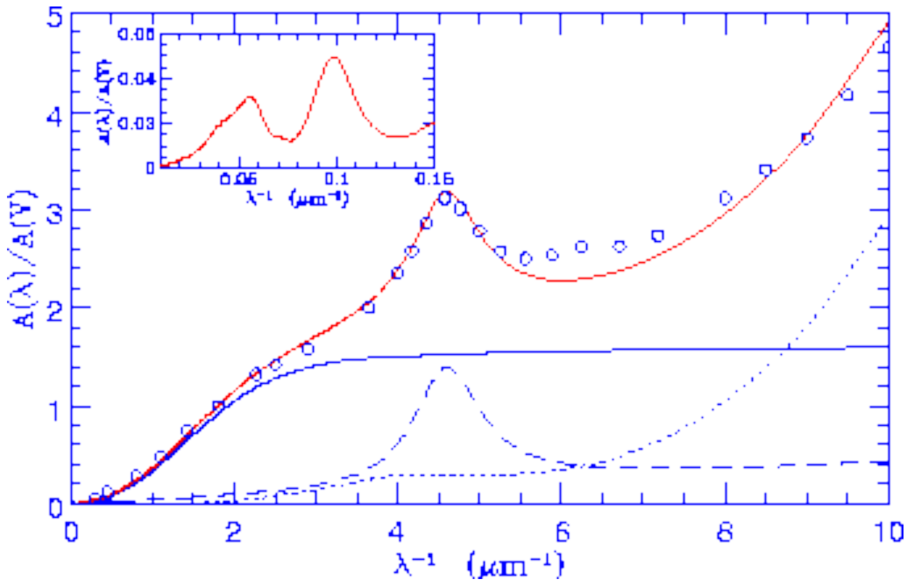
- **Gravitational instability** (Kuiper 1951; Cameron 1962; Boss, Durisen,…)

The core accretion models are more advanced, and this is how terrestrial planets formed, but the origin of the giant planets, and of extrasolar planets, is still debated

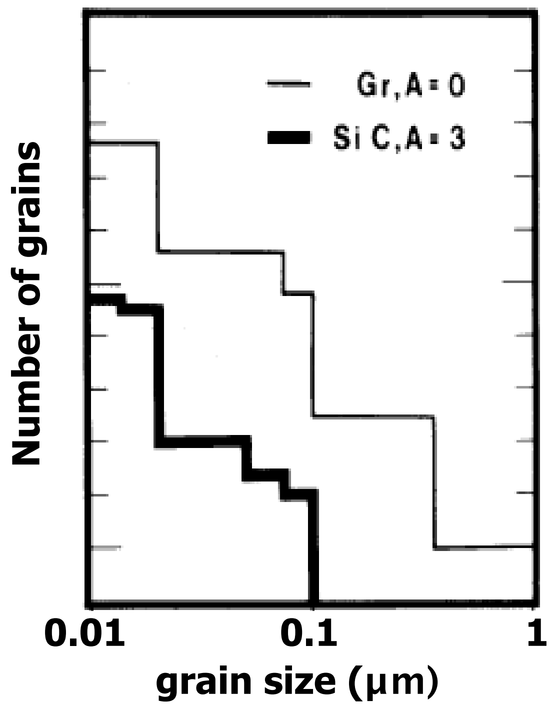
0. Starting conditions

Proto-stellar disk is composed of same material as star (e.g., meteorites have same composition as Sun)

ISM dust distribution determined from modelling extinction and polarization curves (Mathis, Rumpl & Nordsiek 1977, Li & Greenberg 1997):



Size distribution n(a) $\propto a^{-3.5}$ from 0.005 to 1μm including silicate/ organic refractory and graphite (carbonaceous) grains and PAHs



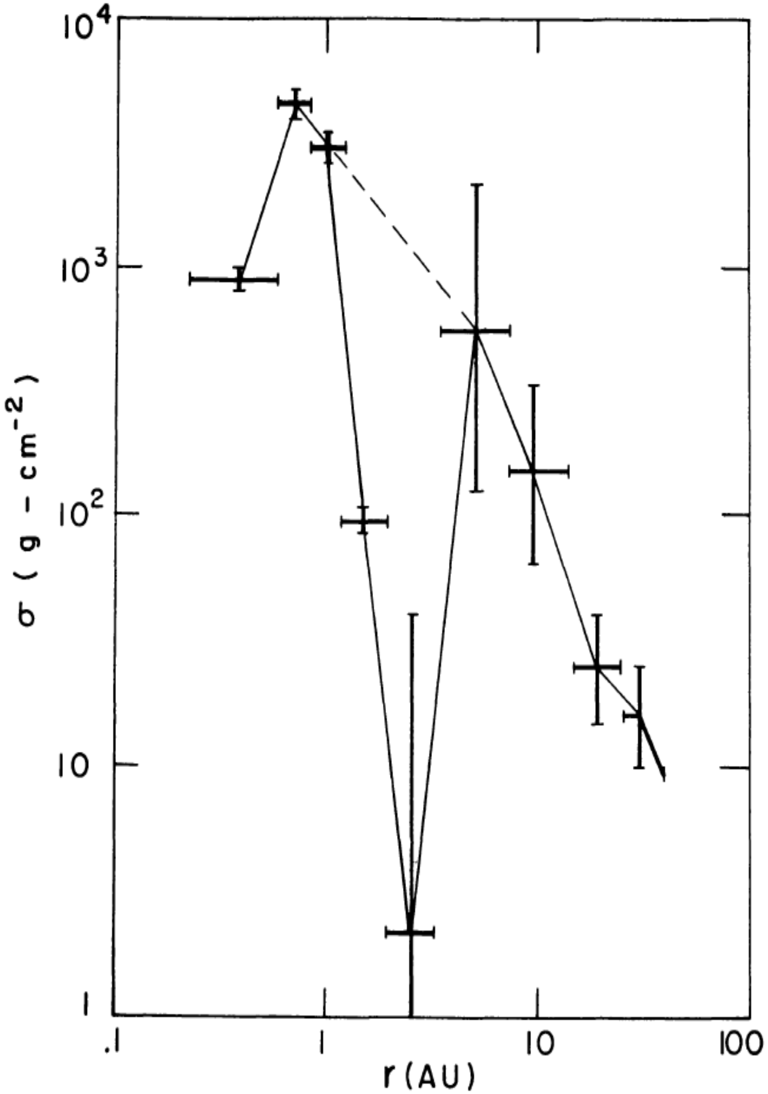
Minimum mass solar nebula

A common concept in planet formation is the **minimum mass solar nebula**, the current distribution of mass (solid and gas) restored to solar composition, which is the minimum the Sun's proto-planetary disk must have had (Weidenschilling 1977; Hayashi 1981; Desch 2007):

$$\Sigma_{total} = 3\text{-}6\text{x}10^4\ r^{-1.5}\ \text{kg/m}^2$$
$$\Sigma_{solid} \approx 0.01\Sigma_{gas}$$

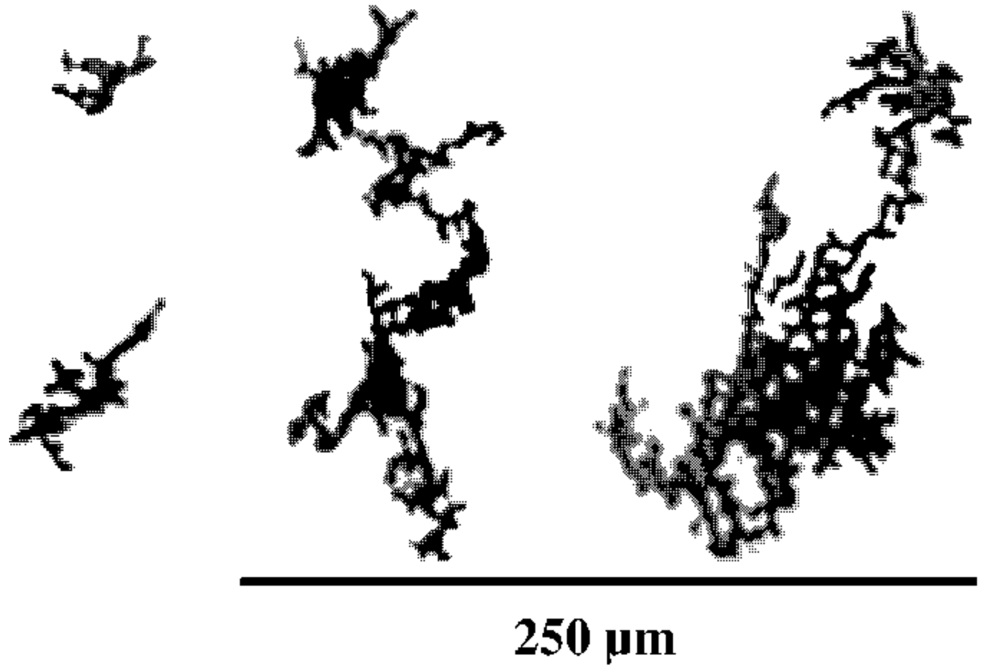
with total mass of $0.01\text{-}0.1M_{sun}$

$$M_{solid}(r_1\text{-}r_2) = 14\text{-}28M_{earth}[r_2^{0.5}\text{-}r_1^{0.5}]$$



1. Grain growth: 1μm-1m

In disks dust grains (and condensates) collide, the outcome of which depends on collision velocity and sticking properties of grains (Heim et al. 1999; Dominik & Tielens 1997; Poppe et al. 2000; Konchi et al. 2002; Wang et al. 2005):



Micron-sized grains grow fractally ($m \propto D^{1.9}$) with porosity ~0.8 in 0.01m/s collisions (Wurm & Blum 1998, 2000; Blum et al. 2007)

D>1cm collisions compact grains and give higher velocities ~10m/s (Blum & Wurm 2000; Sekiya & Takeda 2003)

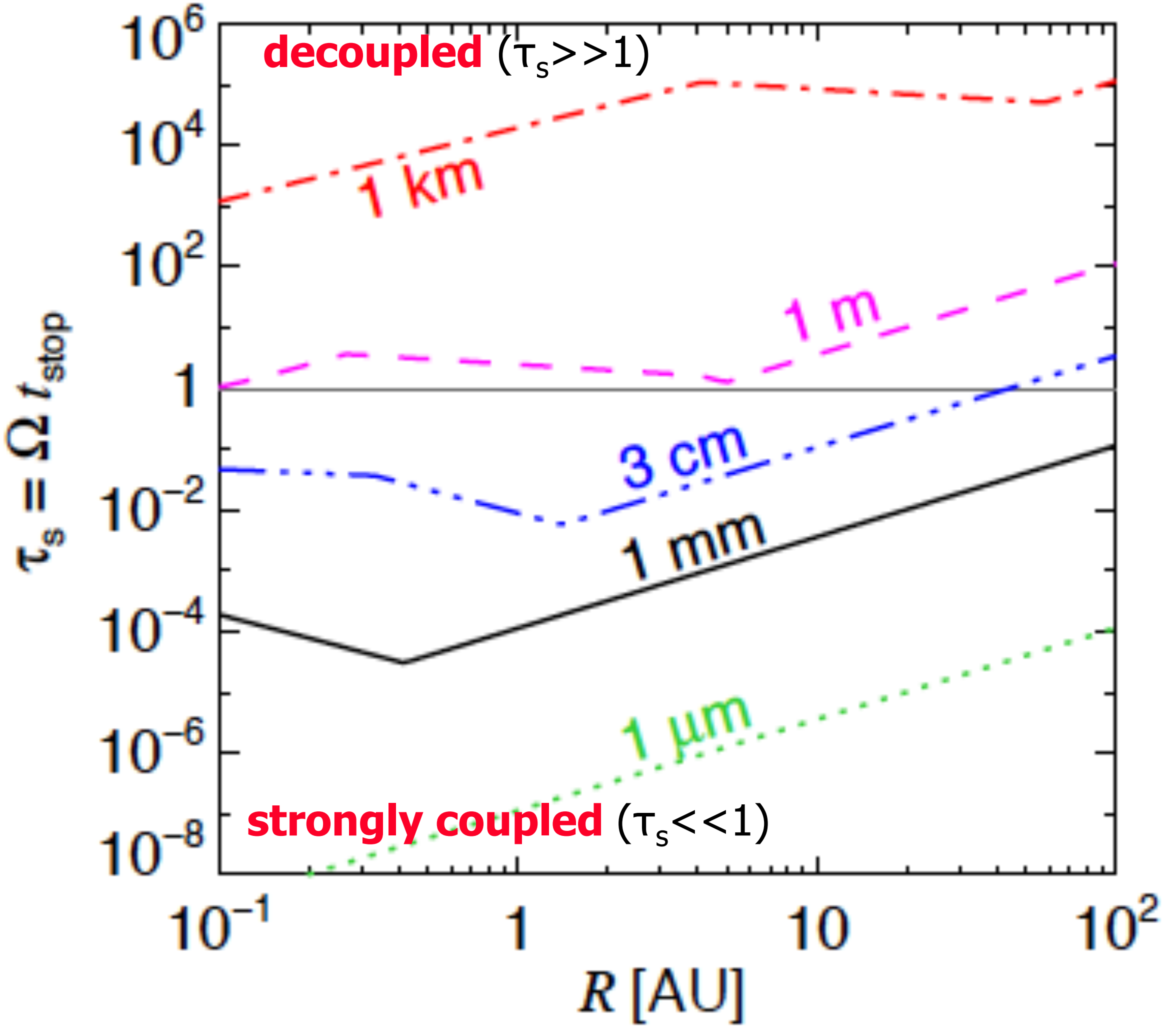
High velocity collisions result in fragmentation (Wurm et al. 2001; 2005)

Gas drag: effect on different size grains

Dust orbits the star, but motion dominated by gas drag (e.g., Weidenschilling et al. 1977)

Drag force depends on relative velocity of gas and dust $\mathbf{F_g} = -0.25\pi\rho_g D^2 v_t \Delta\mathbf{v}$

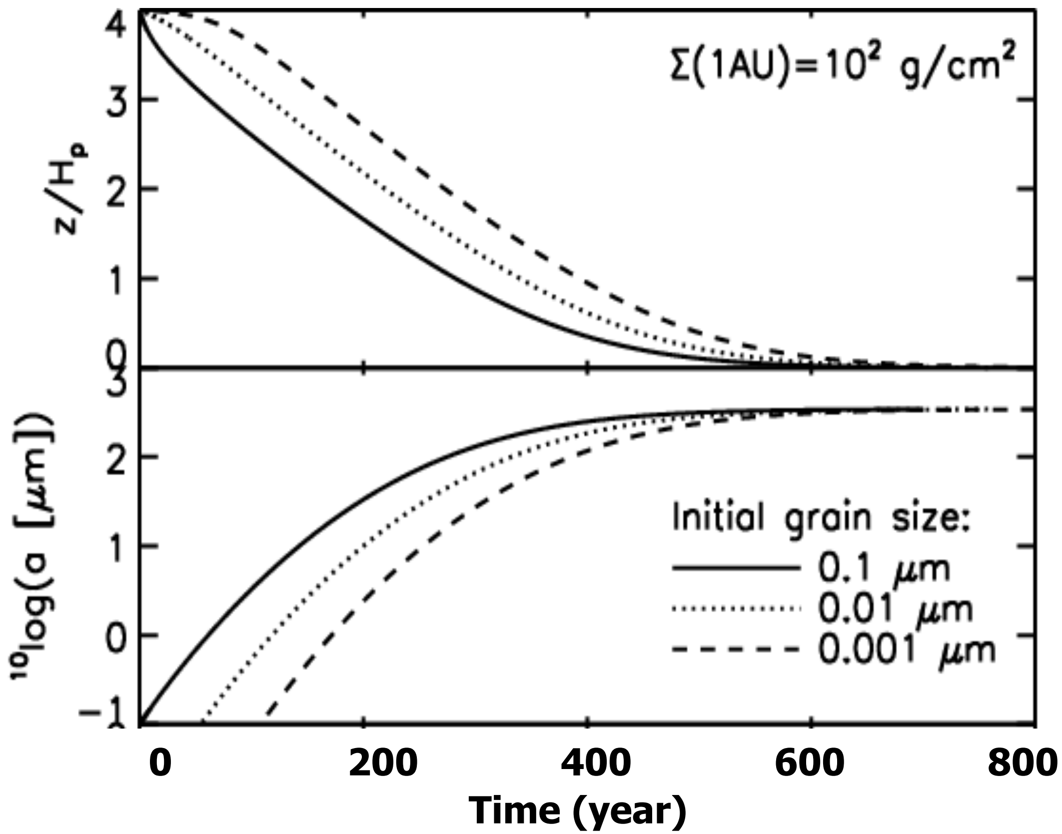
Stopping time is that to cause $\Delta v=0$, which compared with orbital velocity, $v_k=\Omega_k r$ gives the ratio $\tau_s = t_s\Omega_k = \Sigma_{1p}/\Sigma_g$



Settling to mid-plane

Gas drag causes dust to settle to mid-plane (e.g., inclined orbits oscillate vertically, and gas drag damps oscillation)

Timescale long for small grains, but these collide during settling speeding process up (Weidenschilling 1980; Nakagawa et al. 1981; Dullemond & Dominik 2005)

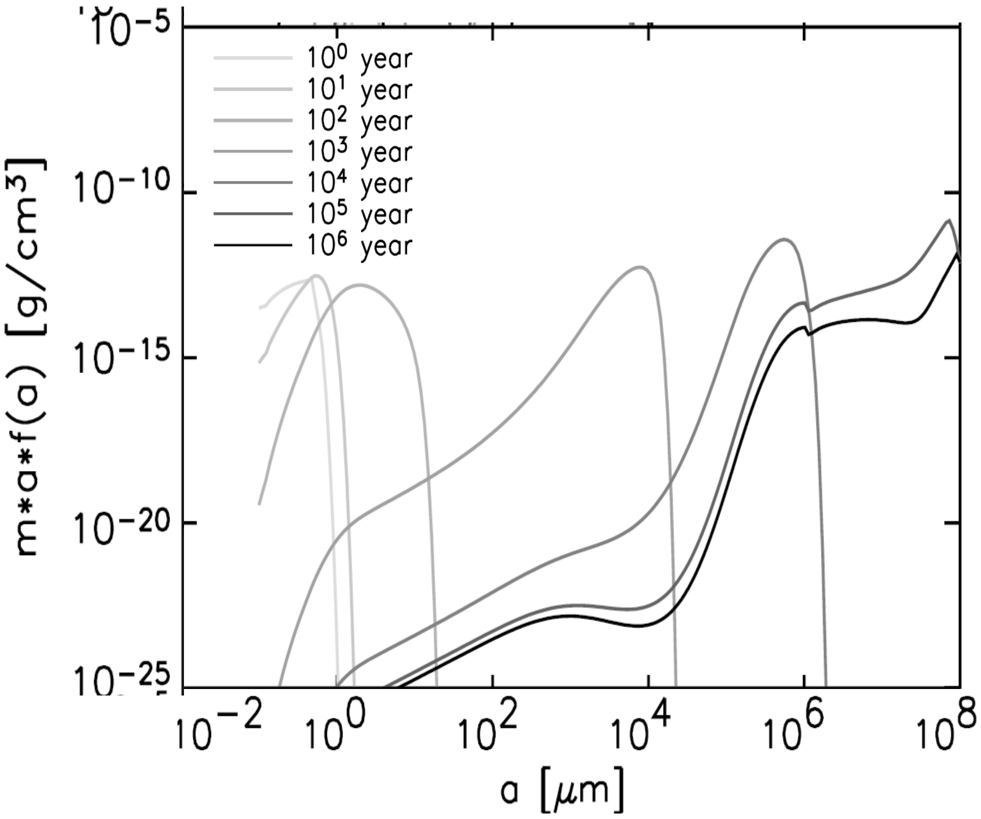


Coagulation models

Models solve coagulation equation with dust settling, turbulent mixing, brownian motion (e.g., Dullemond & Dominik 2005; Tanaka et al. 2005; Nomura & Nakagawa et al. 2006):

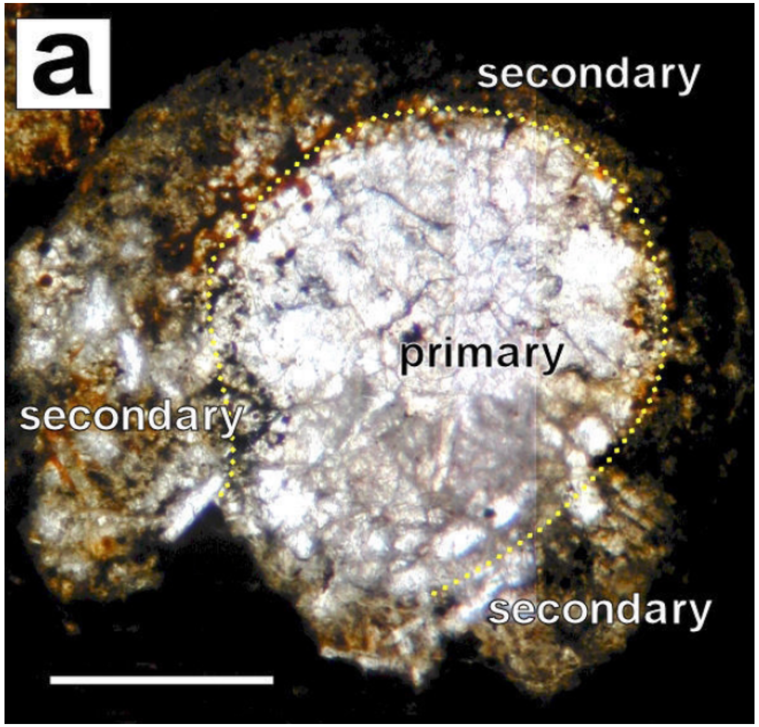
- growth to ~1m easy in 1Myr

- more recent models include creation of small grains in collisions (boosting micron-sized grain population at late times), and also find a bouncing barrier (i.e., no growth beyond 1m)



Chondrule formation

Chondritic meteorites are largely composed of 0.1-10mm previously molten silicate particles (chondrules) with inclusions of older refractory elements (CAIs) and ~1Myr older chondrules (Akaki et al. 2007; Moynier et al. 2007), implying repetitive flash heating and cooling on 1hr timescales



Proposed heating mechanisms include:

- gamma ray burst (Duggan et al. 2003),
- lightning in PPN (Desch & Cuzzi 2000, MacBreen et al. 2005),
- passage through shocks (Ciesla & Hood 2002, Miura & Nakamoto 2007),
- giant (Krot et al. 2005) or small (Miura et al. 2007) imapcts

Radial migration

Gas disk structure from radial component of momentum equation:

$$GM_*/r^2 = \omega^2 r + (1/\rho_g)dP_g/dr$$

gives

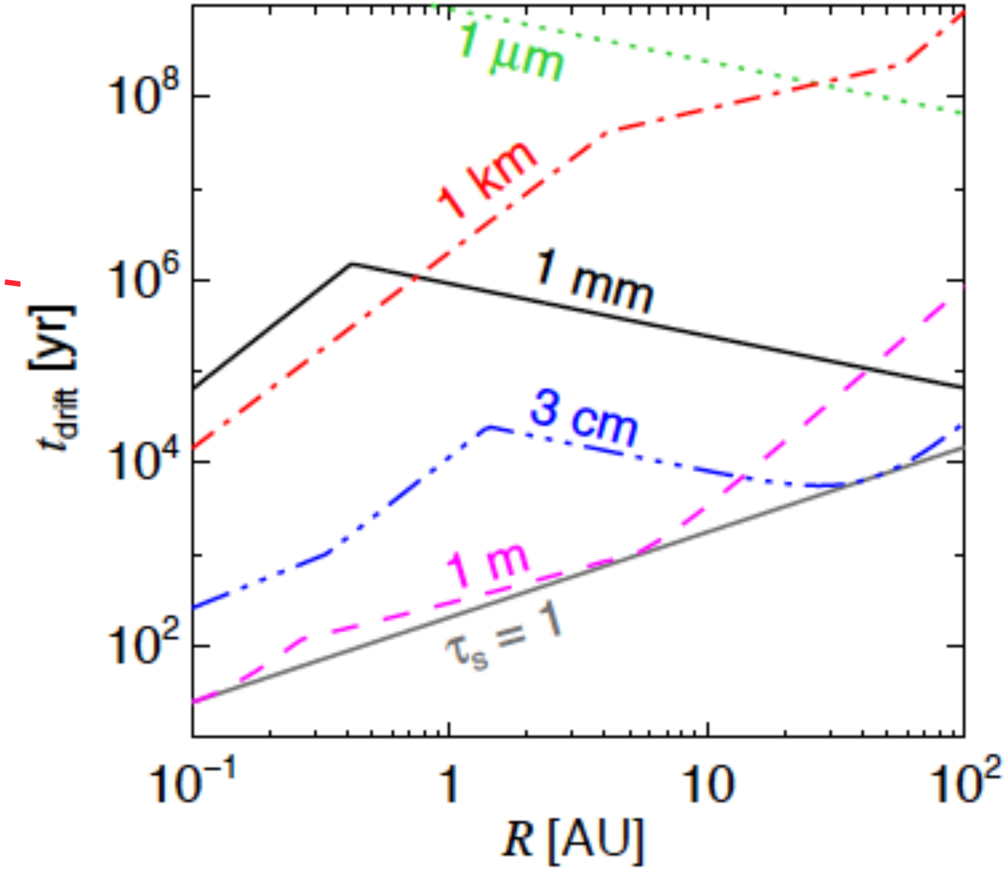
$$v_g = v_k\,(1 - \eta)^{0.5}, \text{ where } \eta = -\,(r\Omega_k{}^2\rho_g)^{-1}\, dP_g/dr$$

Generally pressure decreases with r, so gas velocity is sub-keplerian, so dust migrates in (as it sees a headwind or extra acceleration toward star)

Radial migration

Gas drag on metre-sized objects causes them to fall onto star in 100 yr
(Weidenschilling et al. 1977)

Small grains immune as they **couple** to gas, and terminal velocity due to extra acceleration toward star (due to their sub-keplerian velocity) is small



Large grains are also immune as they **decouple** from gas, and their large area/ mass means headwind is slow to move grains

2. Grain growth: 1m-10km

Proceeds by collisions between planetesimals?

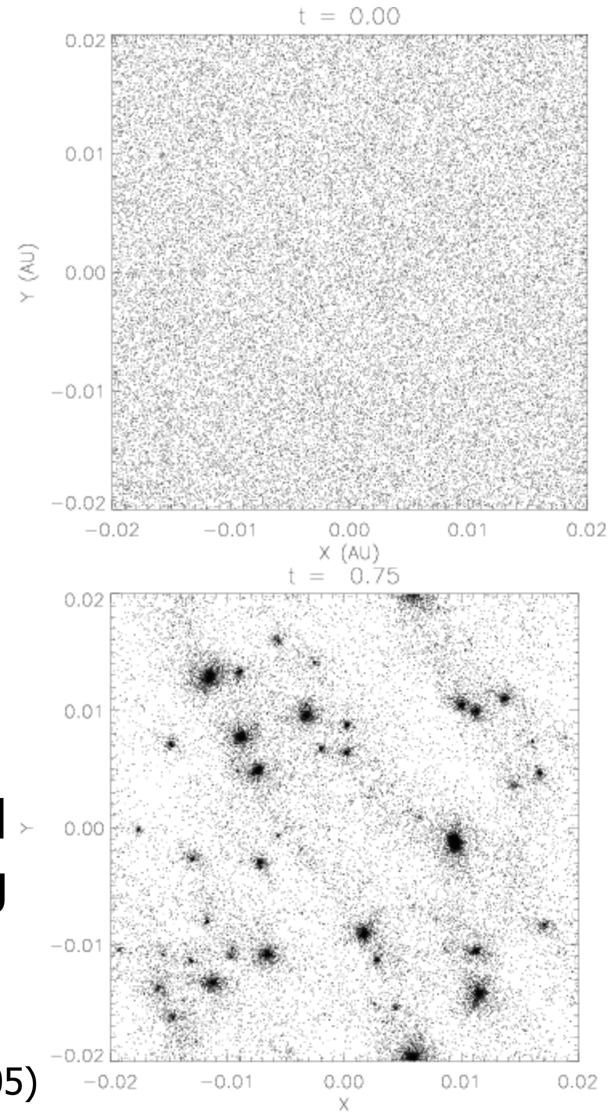
Timescale problem: metre-sized objects migrate in due to gas drag in 100 years, much faster than collisional growth times

Resolution:

- speed up growth
- slow down migration

Gravitational instability (GI)

Speed up growth by GI if dust concentrated in mid-plane, making km-sized planetesimals on orbital timescales (Safronov 1969; Goldreich & Ward 1973; Tanga et al. 2004)



Requires Toomre parameter Q<1

$$Q = \Omega_k c_d/(\pi G \Sigma_d)$$

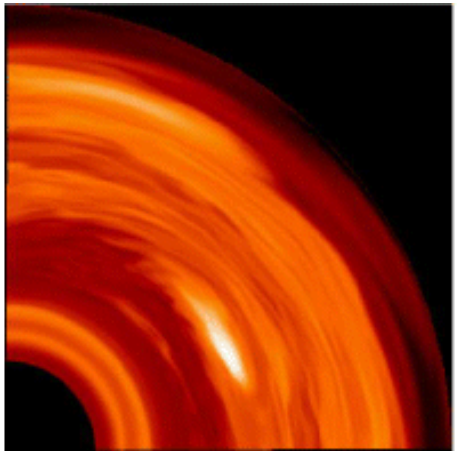
Typically, dust mass densities $>10^{-7}$ g/cm^3

Ongoing debate:

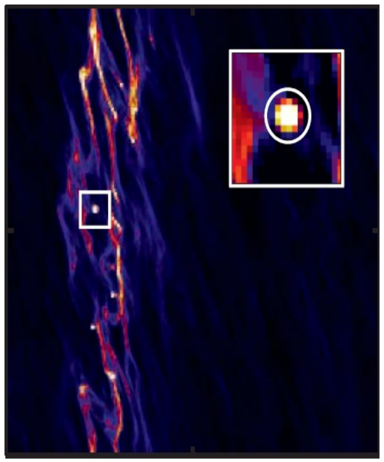
- dust entrains gas causing vertical velocity shear and Kelvin-Helmholtz instability thus turbulence increasing velocity dispersion (Weidenschilling 1980; 2003), but velocity shear doesn't lift all dust (Sekiya 1998; Youdin & Shu 2002)
- helped by size dependent drift rates (Youdin & Chiang 2005)

Vortices / spirals in proto-planetary disks

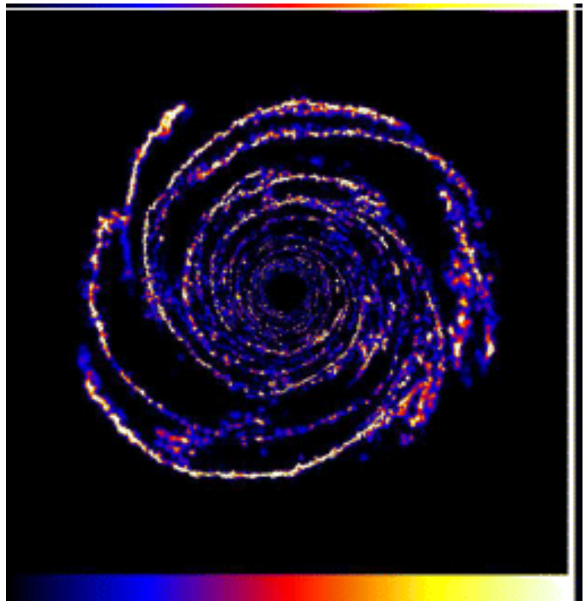
Particles 0.1-1m can become trapped in vortices (Barge & Sommeria 1995; Tanga et al. 1996; Klahr & Henning 1997; Klahr & Bodenheimer 2003, 2006; Inaba & Barge 2006; Lithwick 2007; Johansen et al. 2007; Youdin et al. 2012)



Concentrations may be gravitationally unstable, if they last long enough (Godon & Livio 1999; Cuzzi et al. 2001)



1-10m particles also concentrated in spirals of marginally stable self-gravitating disk (Rice et al. 2004) or if perturbed by a passing star (Theis, Kroupa & Theis 2005; Lodato et al. 2007)



Slowing down / stopping migration

Radial component of momentum equation:

$$GM_*/r^2 = \omega^2 r + (1/\rho_g)dP_g/dr$$

giving

$$v_g = v_k (1 - \eta)^{0.5}, \text{ where } \eta = - (r\Omega_k^2\rho_g)^{-1} dP_g/dr$$

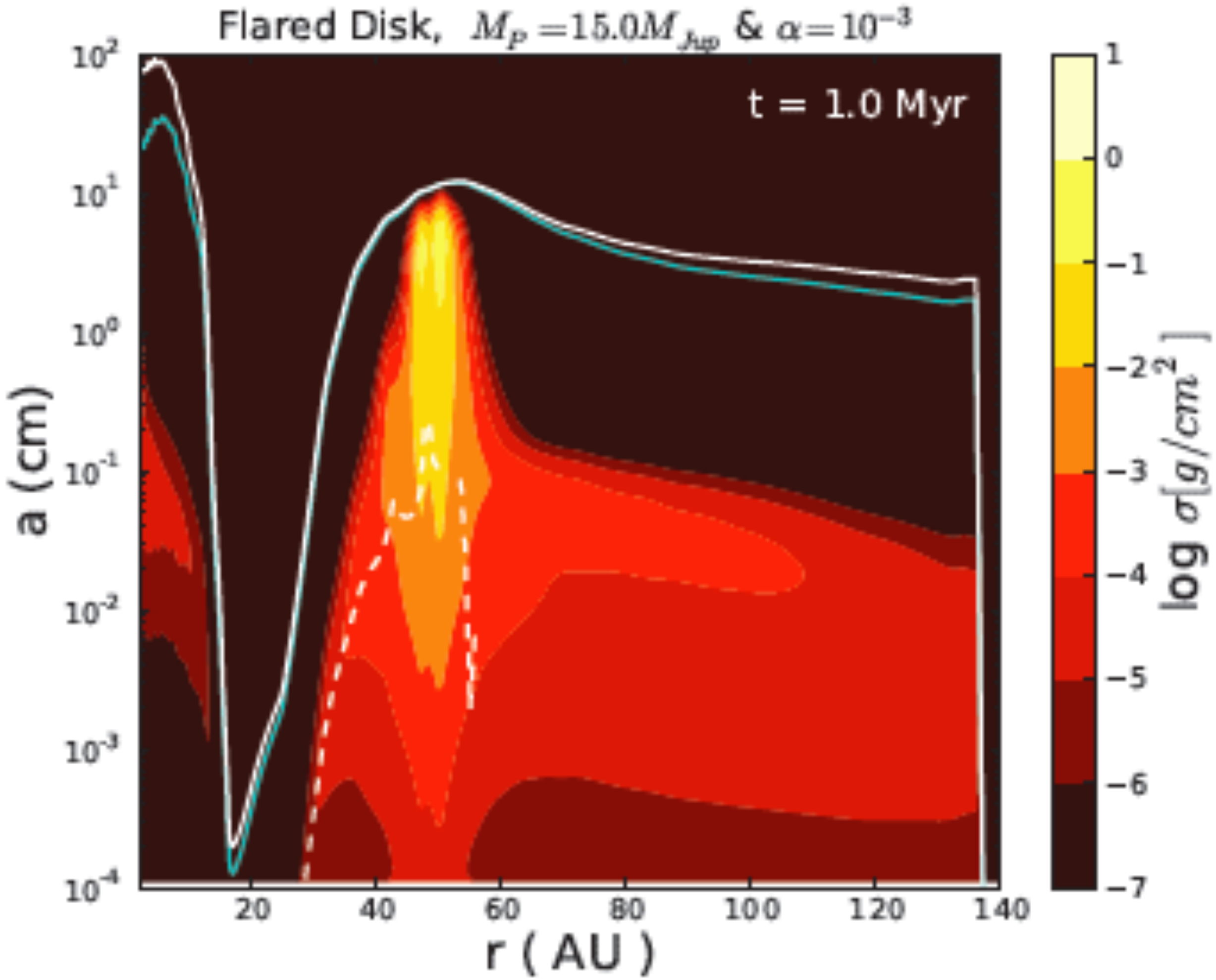
Generally pressure decreases with r, so gas velocity is sub-keplerian, so dust migrates in (as it sees a headwind or extra acceleration toward star)

But,

- pressure reverses at disk gap/jump
- turbulence changes pressure gradient (e.g., vortices, Klahr & Bodenheimer 2006)

Dust coagulation in disk with gap

Coagulation models encounter problems with radial drift, but if the disk has a gap (e.g., due to planet), then dust is all pushed to just outside the gap where growth to metre-size is possible
(Pinilla et al. 2012)



Summary of lecture 1

- Solar System and extrasolar planet summary
- Planets form in disks of gas and dust
- Dust inevitably settles to the mid-plane and grows to ~1m in size
- However, 1m-sized boulders have a short lifetime, so to make km-sized planetesimals we need to invoke as yet poorly understood mechanism like gravitational instability

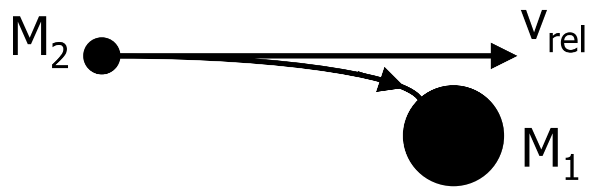
3. Runaway growth: 10km-1000km

Orderly growth: If collision cross-section is ~ $\pi D^2/4$, time to make objects of mass "m" is:

$t_{acc} = m/(dm/dt) = 2r^{1.5}(D/1km)(\Sigma/10kgm^{-2})^{-1}$ Myr

i.e., 10-100km objects take 0.6-6Myr to grow in a MMSN at 5AU

Runaway growth: Gravitational focussing enhances collision cross-section by $(1+v_{esc}^2/v_{rel}^2)$, where $v_{esc}^2=0.25Gm/D$. Runaway occurs when $v_{rel} \ll v_{esc}$ as $dm/dt \propto m^{4/3}$ and so large proto-planets decouple from size distribution



Velocity dispersion, v_{rel}, is very important

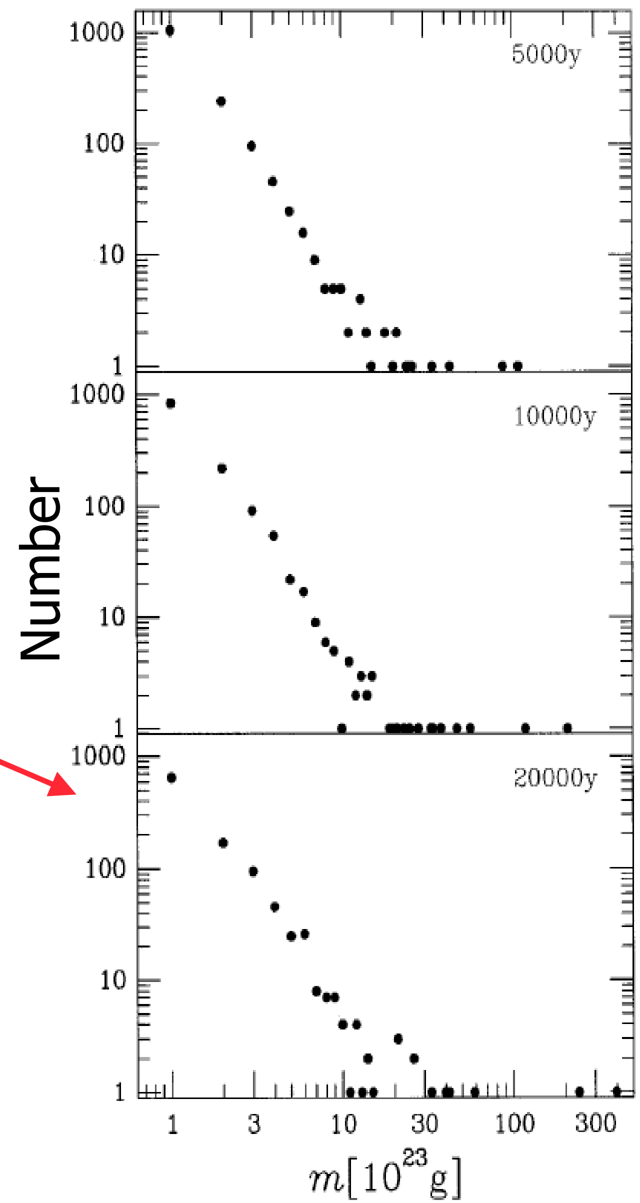
Modelling methods

• **Statistical:** particle-in-a-box with the Fokker-Planck equation follows distributions of orbital elements of many particles (e.g., Wetherill & Stewart 1989)

• **Direct:** N-body simulations of gravitational interaction of fewer particles (e.g., Aarseth et al. 1993; Kokubo & Ida 1996)

Runaway seen using both methods

Two particle approximation:
• disk made up of planetesimals $m_{\alpha} = 10^{15}$kg (D_{α}=10km) which do not grow with time
• and cores of size m_{β} which do grow



Evolution of velocity dispersion

The velocity dispersion is balance of:

- **Gravitational scattering** (increases v_{rel})
    - **Runaway phase:**
        - v_{rel} of planetesimals (m_α) is const and dominated by mutual scattering
        - v_{rel} of cores (m_β) small by **dynamical friction** (scattering planetesimals)
    - **Oligarchic phase:**
        - once cores sufficiently massive ($3m_\beta\Sigma_\beta > m_\alpha\Sigma_\alpha$), v_{rel} of planetesimals is dominated by scattering with cores

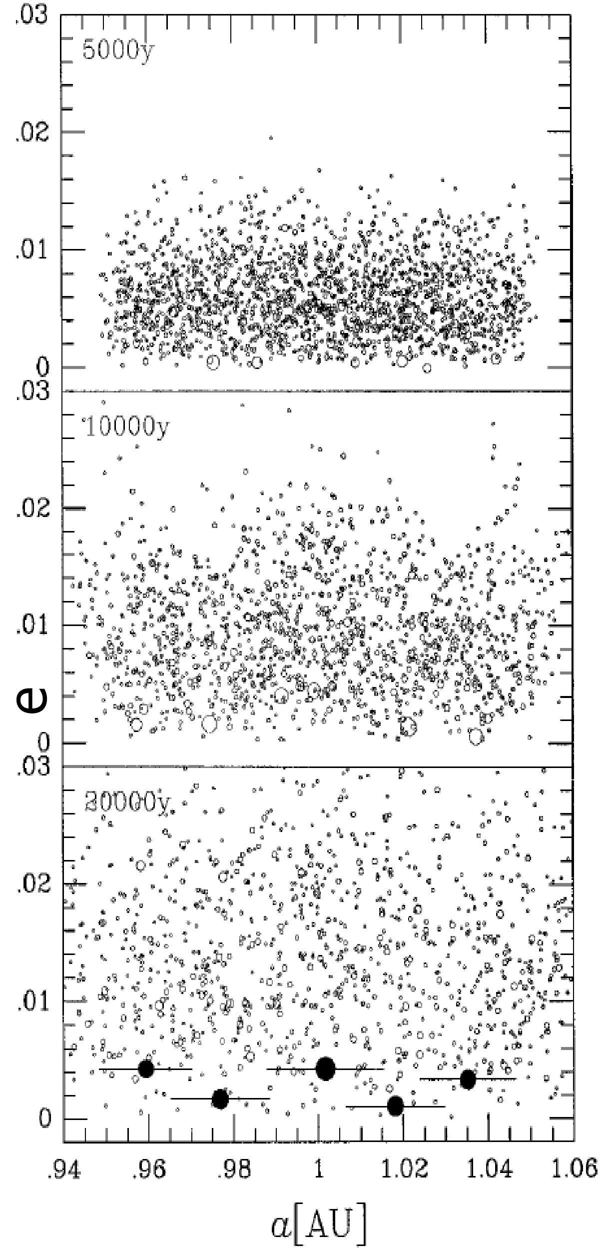
- **Gas drag and disk tides** (decrease v_{rel})
    - Inclination and eccentricity reduced (e.g., settling and circularisation)
    - Gas drag only important for small particles
    - Tides only important when $m_\beta > 10^{-2}$-$10^{-4} M_{earth}$

4. Oligarchic growth: 1000km-10,000km

Runaway phase ends when core mass dominates velocity dispersion of planetesimals:
$m_\beta > 2.2\mathrm{x}10^{-7} f^{0.6} r^{6/5} (\Sigma_\alpha m_\alpha / 10^{17} \mathrm{kg}^2 \mathrm{m}^{-2})^{0.6}\ M_{earth}$

Gravitational focussing is still strong allowing cores (**oligarchs**) to reach M_{earth} quickly, but velocity dispersion increases with m_β meaning large and small planetesimals grow at same rate

Oligarchs grow at 5 Hill's radii separation: as they grow r_H increases, meaning some are squeezed out resulting in collisions and scattering (Kokubo & Ida 1995, 1998)

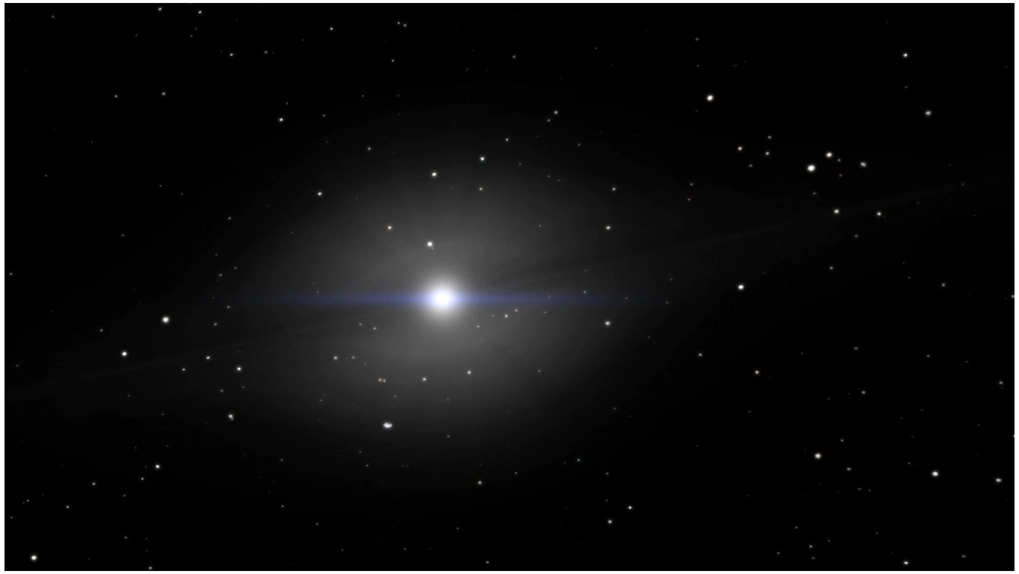


5. Chaotic growth

Massive oligarchs clear **feeding zone** of planetesimals, reaching **isolation mass** of (assuming separation of fr_H where f~10) (e.g. Lissauer 1987):

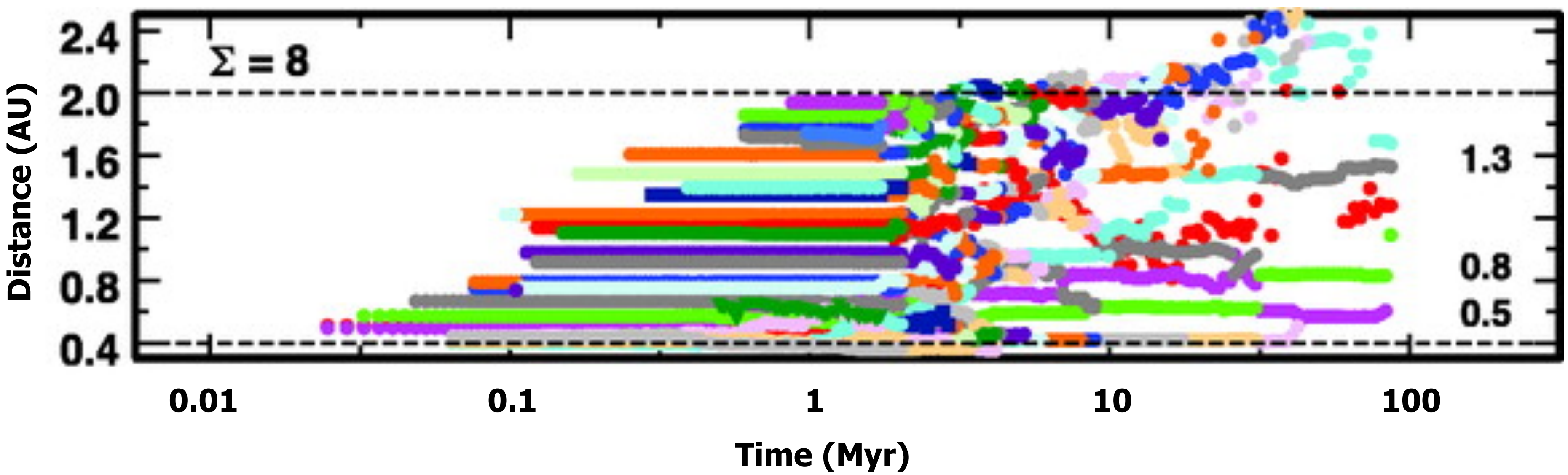
$$m_\beta = 3.3\times10^{-3} f^{1.5} (\Sigma_\beta / 10\,\mathrm{kgm}^{-2})^{1.5} r^3\ M_{earth}$$

After gas dissipates (or before), proto-planet eccentricities increase causing interactions (Chambers & Wetherill 1998), whereby proto-planets grow slowly through massive collisions, although ejection of proto-planets up to $1M_{earth}$ common in outer solar system (Goldreich, Lithwick & Sari 2004)



Transition to chaotic growth

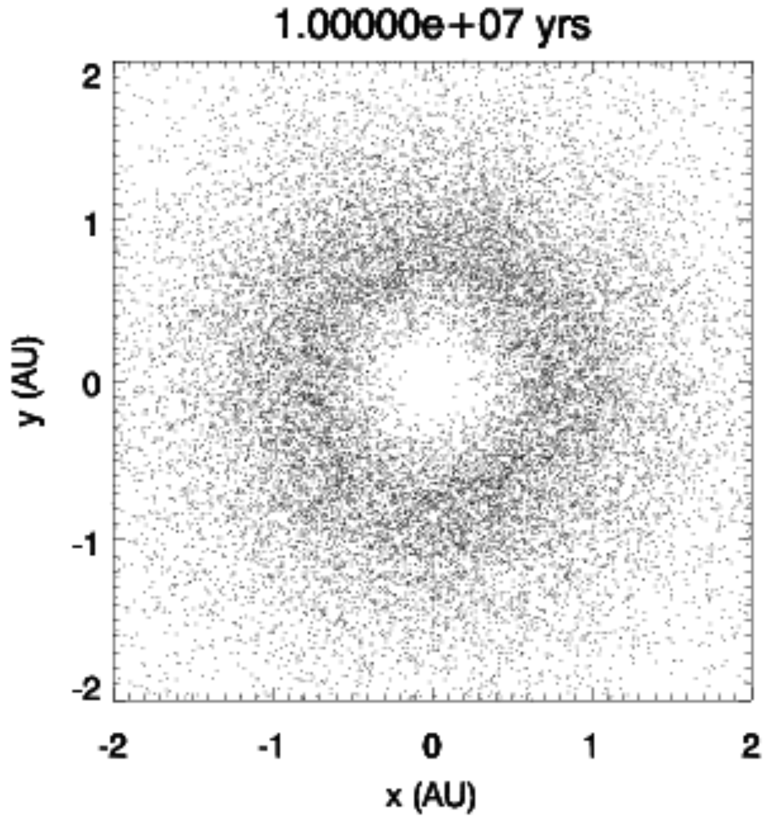
Hybrid simulations which follow oligarchs using N-body and planetesimals using statistics show transition to chaotic growth requires mass in oligarchs to be more than that in planetesimals and for the disk density to be above a threshold (Kenyon & Bromley 2006)



Role of destructive collisions

• Analytical arguments suggest small body population damps eccentricities (Goldreich, Lithwick & Sari 2004), and planets could continue to grow by hoovering up small particles (Schlichting et al.)

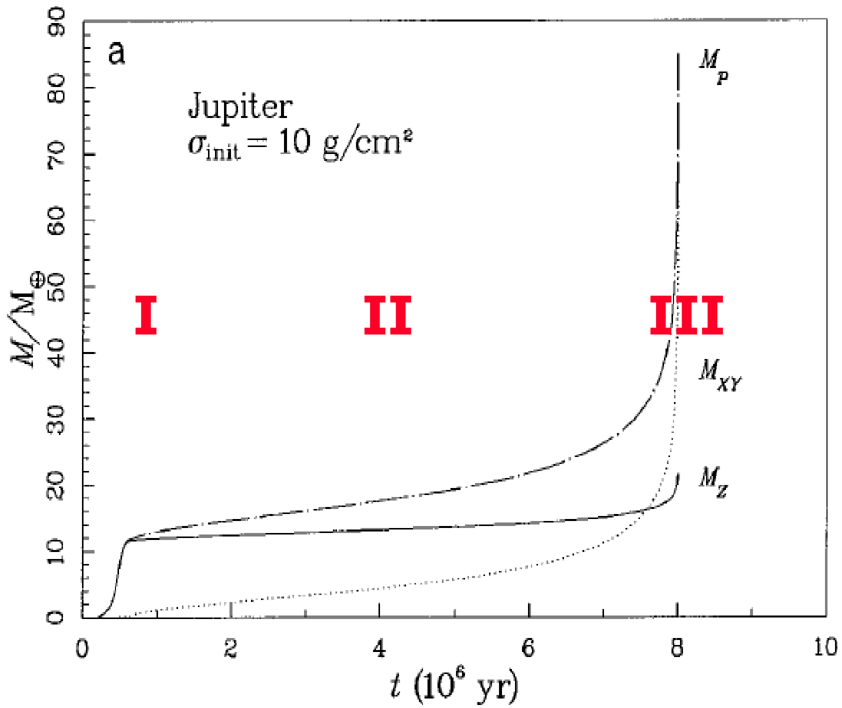
• Though N-body simulations conclude not much debris after oligarchic growth (Leinhardt & Richardson 2005), debris from collisions may be observable and prevalent (Jackson & Wyatt 2012)



6. Gas accretion: M_{earth} to $M_{jupiter}$

Core grows with atmosphere in quasi-static thermal equilibrium until critical mass (~10 M_{earth}) when it rapidly accretes gaseous envelope; final mass determined by available gas and how fast it can be accreted

Three stages (Pollack et al. 1996):
(I) runaway growth to isolation
(II) small time independent accretion rates
(III) rapid accretion, when $M_{solid}=M_{gas}$ envelope contracts, outer boundary expands



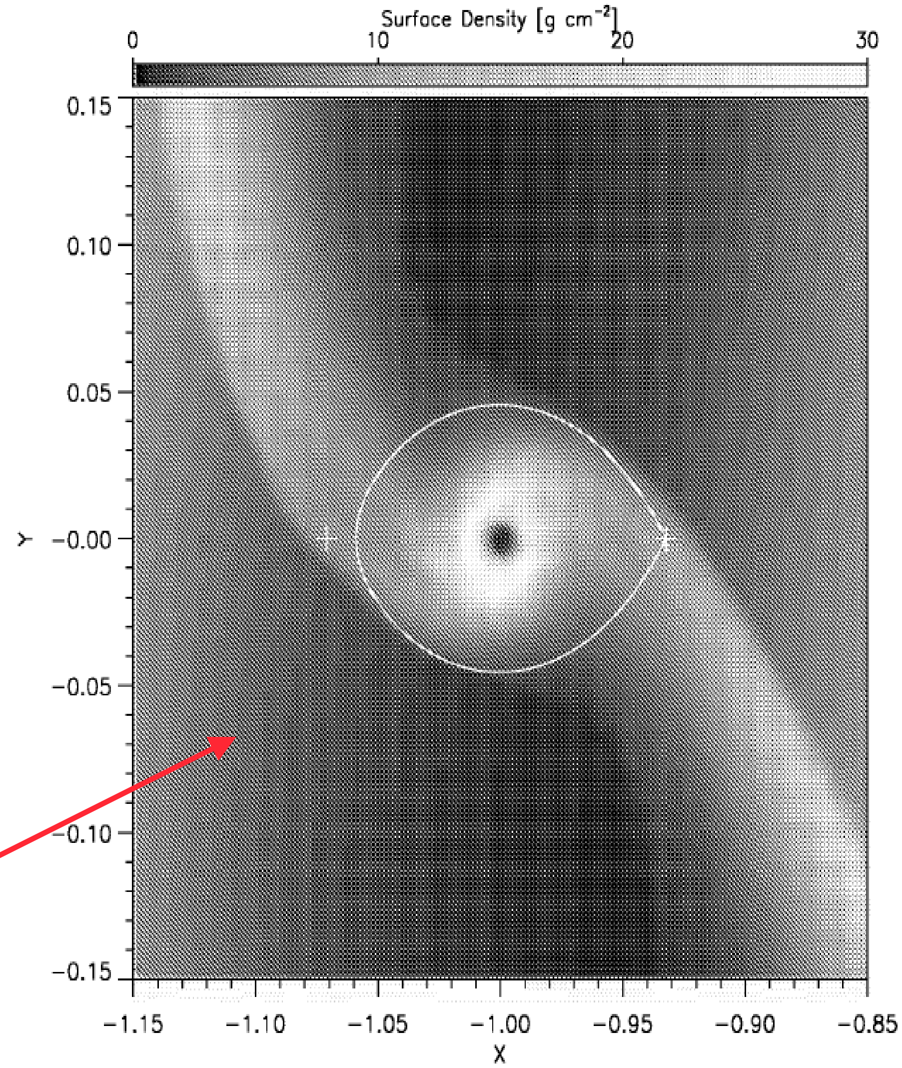
Jupiter forms in <10Myr with 15M_{earth} core from >1MMSN solar nebula

Modifications to gas accretion

Problem: low core mass of Jupiter, timescales longer than gas disk lifetimes

Solution: reducing opacities to 2% ISM and stopping planetesimal accretion (Hubickyj, Bodenheimer & Lissauer 2005; Papaloizou & Nelson 2005), as limitation is loss of energy from the envelope; now easy to form Jupiter in 5Myr with $5M_{earth}$ core

However, 1D models ignore planet-disk interactions that result in non-axisymmetric, shocked flows (Lubow et al. 1999), circumplanetary disk (Bate et al. 2003; Machida et al. 2008) and detailed structure of flow through disk gap (Lubow & D'Angelo 2005)



Planet migration

Hot Jupiters (HJs) were/are widely believed to have formed farther out then migrated in, although

- can form in situ (Bodenheimer et al. 2000)
- and in scattering between planets or Kozai + tidal circularisation (Weidenschilling & Marzari 1996; Naoz et al. 2012)

Proposed migration mechanism is interaction with the proto-planetary disk which results in three types of migration (Papaloizou et al. 2007):

Type I: small mass planets, treated in linear regime (Ward 1997)
Type II: larger mass planets open a gap (non-linear) (Lin & Papaloizou 1984)
Type III: runaway migration from co-orbital torques (Masset & Papaloizou 2003)

See Seba's talk for more details

Planet migration: type I

Acts on small proto-planets which excite density waves at Linblad resonances (Goldreich & Tremaine 1979):

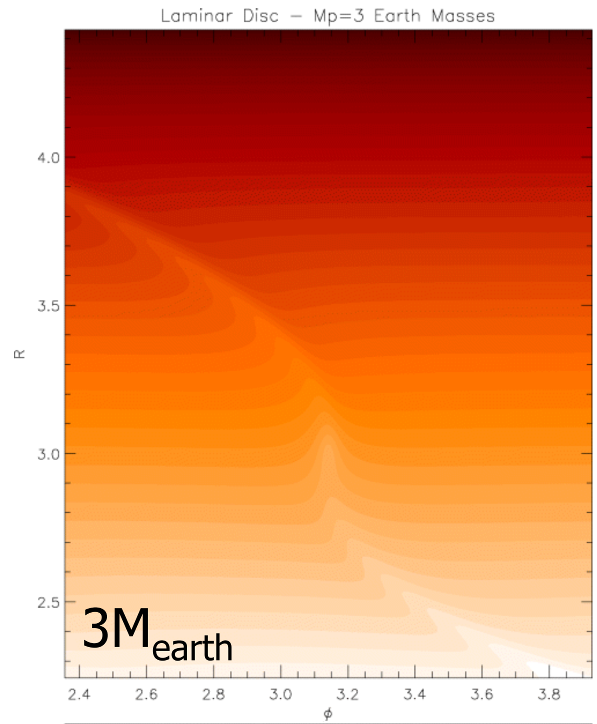
- waves interior to the planet exert positive torques
- exterior waves exert negative torques

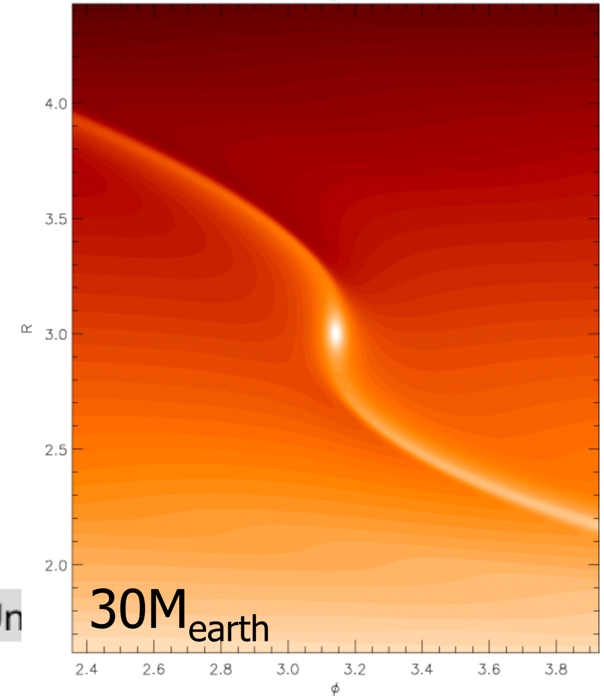
Sum of torques is negative leading to inward migration on timescales of 0.2Myr for $1M_{earth}$ at 5AU (Korycansky & Pollack 1993; Ward 1997; Tanaka, Takeuchi & Ward 2002)

$$dr/dt = -2.7\ (M_{pl}/M_*)\ r\Omega_k\ (\Sigma r^2/M_*)\ (r\Omega_k/v_t)^2$$

Same torques also damp planet eccentricity on timescale (Artymowicz 1993; Tanaka & Ward 2004):

$$t_e = 3.46\ (v_t/r\Omega_k)^2\ (r/|dr/dt|)$$





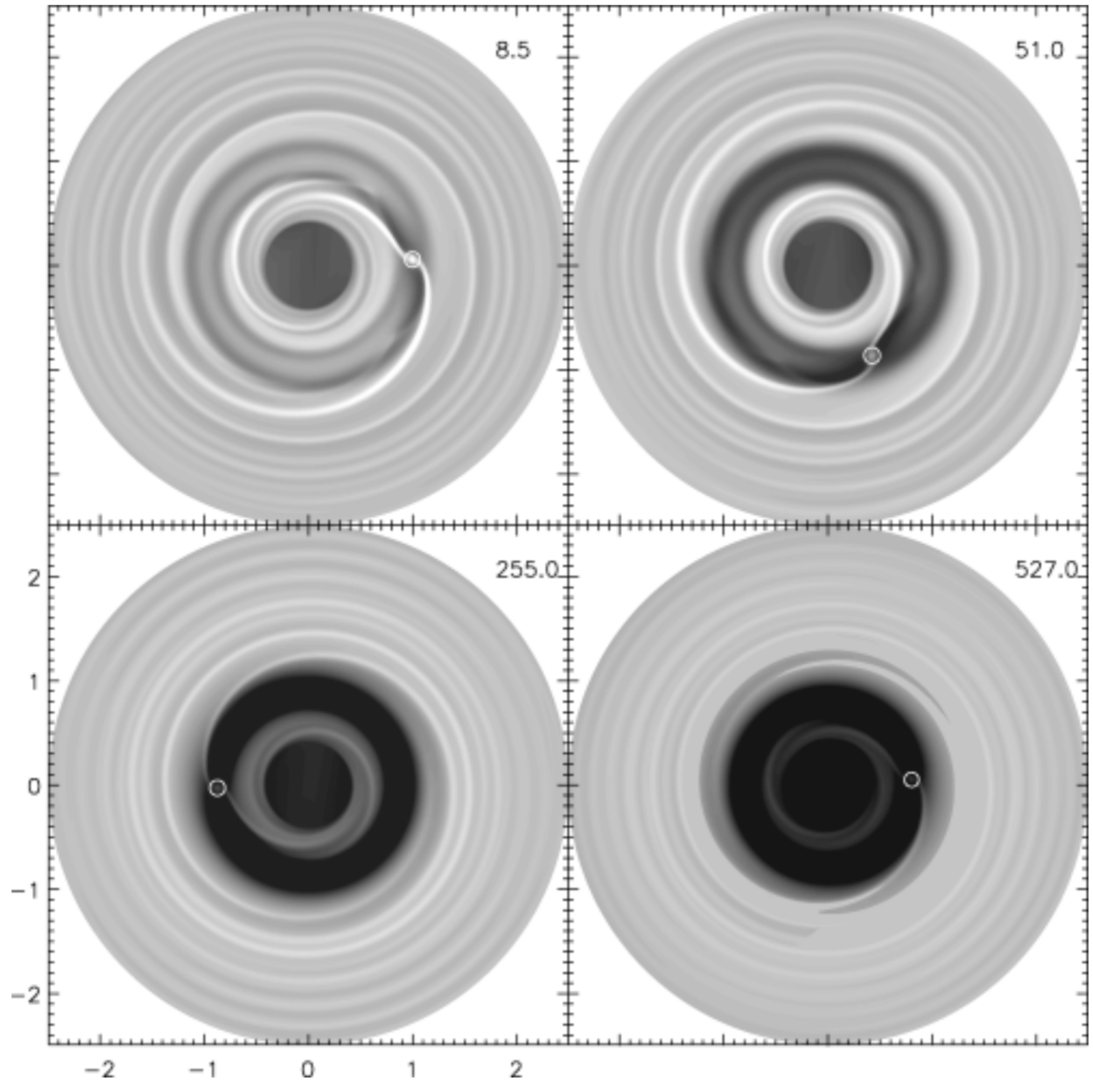
Gap opening: transition to type II

Linearity breaks down when $M_{pl}/M_{*} > (H/r)^3$ which is $\sim 30 M_{earth}$

Gap opening and gap structure depends on: planet mass, disk height, and viscosity (Crida, Morbidelli & Masset 2006; Rafikov 2002; Edgar & Quillen 2007)

$0.75 H/R_H + 50(M_{*}/M_{pl})/Re < 1$

where $Re = r^2 \Omega_k / \nu$

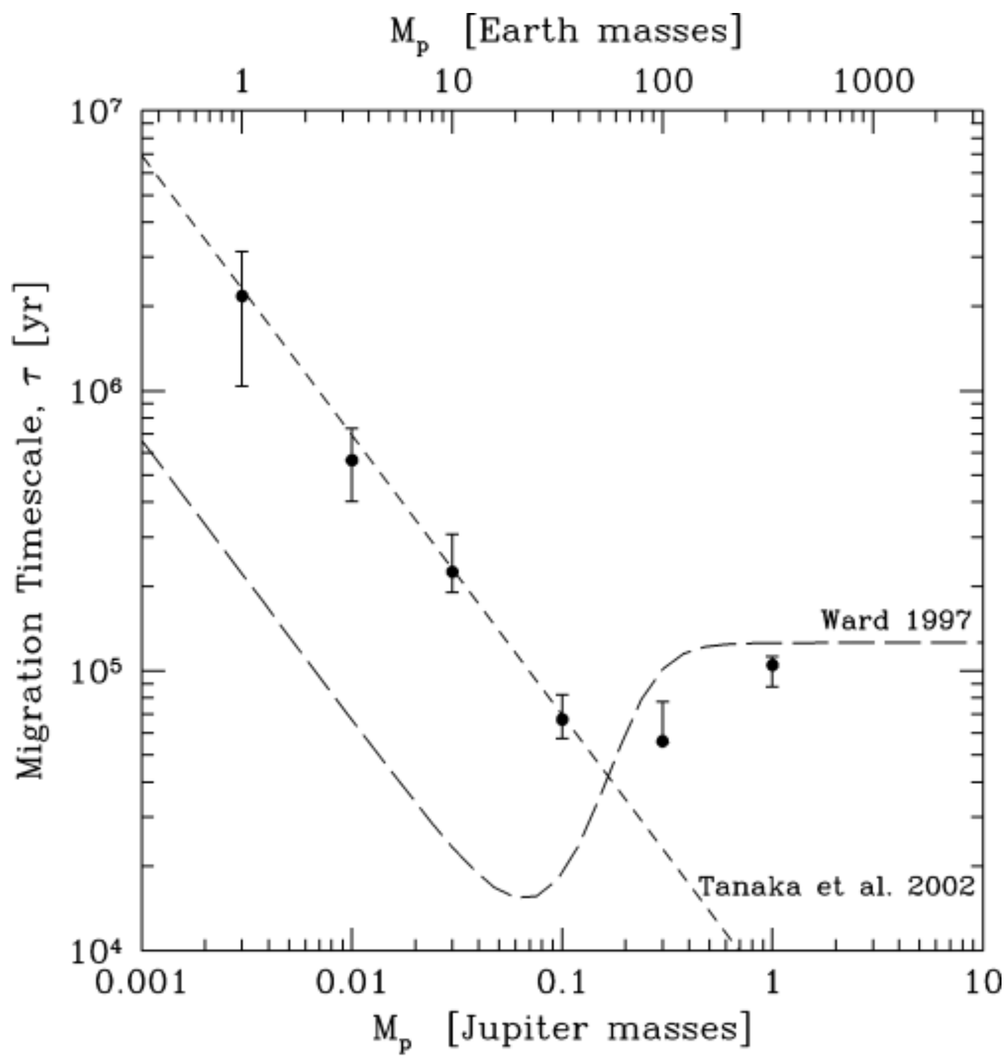


Planet migration: type II

After gap opening, planet migrates in on viscous timescale (10^4 orbital periods, Nelson et al. 2000), though inertia of massive planets can slow migration:

$$dr/dt = -1.5\nu/r$$

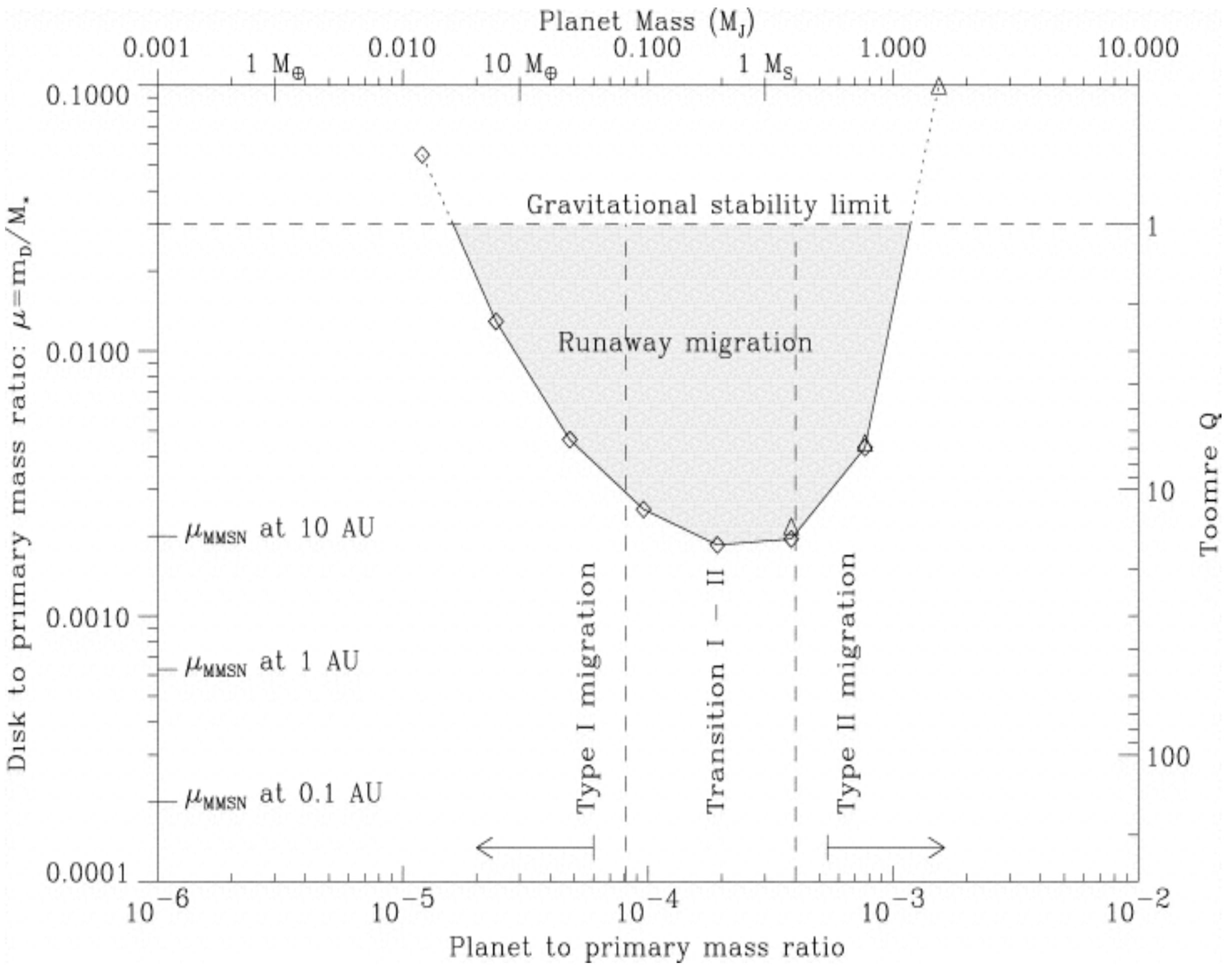
The transition from type I to type II migration is smooth (Bate 2003)



Planet migration: type III

Type III migration is associated with coorbital torques and acts very fast on ~Saturn mass planets massive disks in which there is a partial gap (Masset & Papaloizou 2003)

Radial migration means that torques from co-orbiting material do not average to zero (Ogilvie & Lubow 2003)

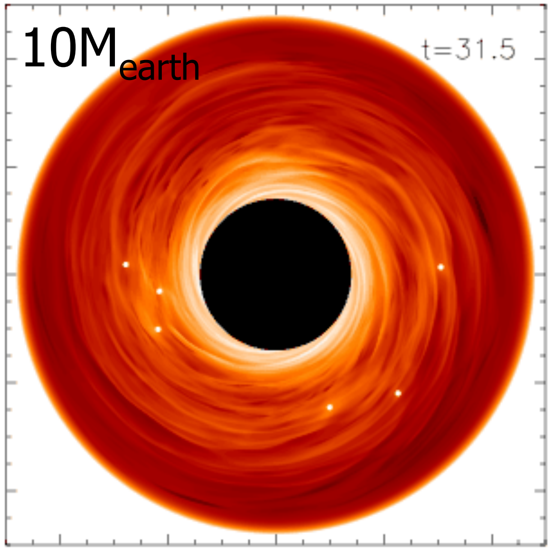


But may be a numerical effect (D'Angelo, Bate & Lubow 2005)?

Why don't all planets migrate in?

Short migration times, so why don't all planets migrate in before they can accrete gas?

Possible solutions:



• migration aids growth (Tanaka & Ida 1999; Alibert et al. 2005)

• turbulence slows migration (Nelson et al. 2005)

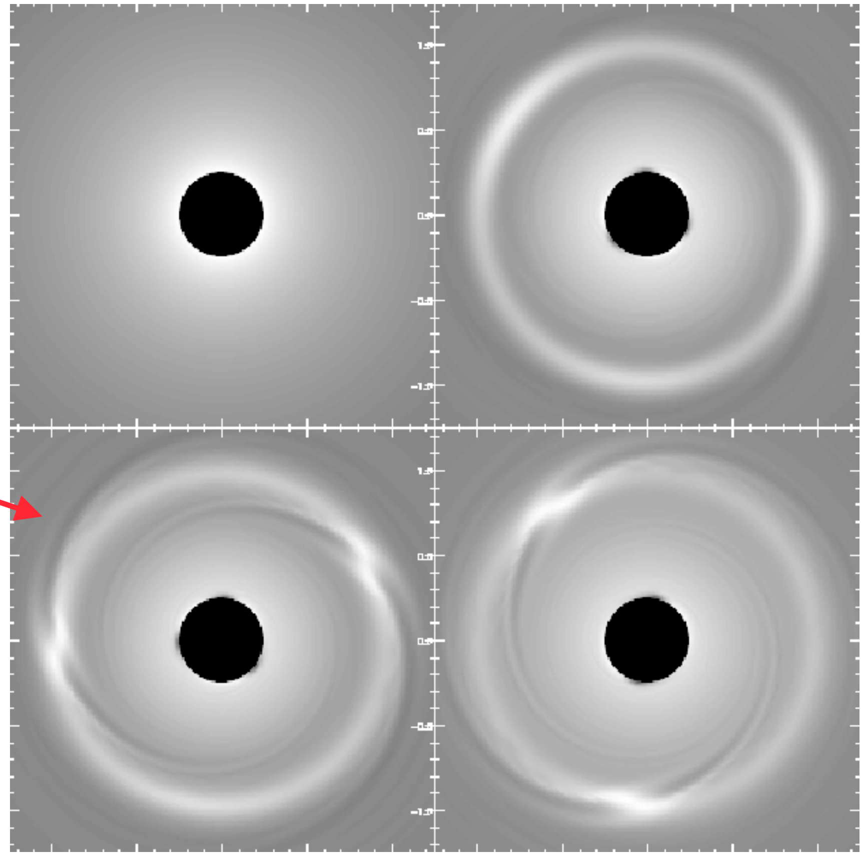
• planet traps = jump in surface density halts migration (Hasegawa & Pudritz 2011; Sandor et al. 2011)

• migration outward for certain conditions (D'Angelo et al. 2002; Crida et al. 2009)

• trapping in resonance (Morbidelli & Crida 2007)

PPD properties: dead zones

Dead zone: disk region (<12AU) is poorly ionised (Turner et al. 2007) and so growth of magneto-rotational instability (MRI, Balbus & Hawley 1991) against ohmic dissipation cannot be sustained (Gammie 1996) leading to low viscosity causing:

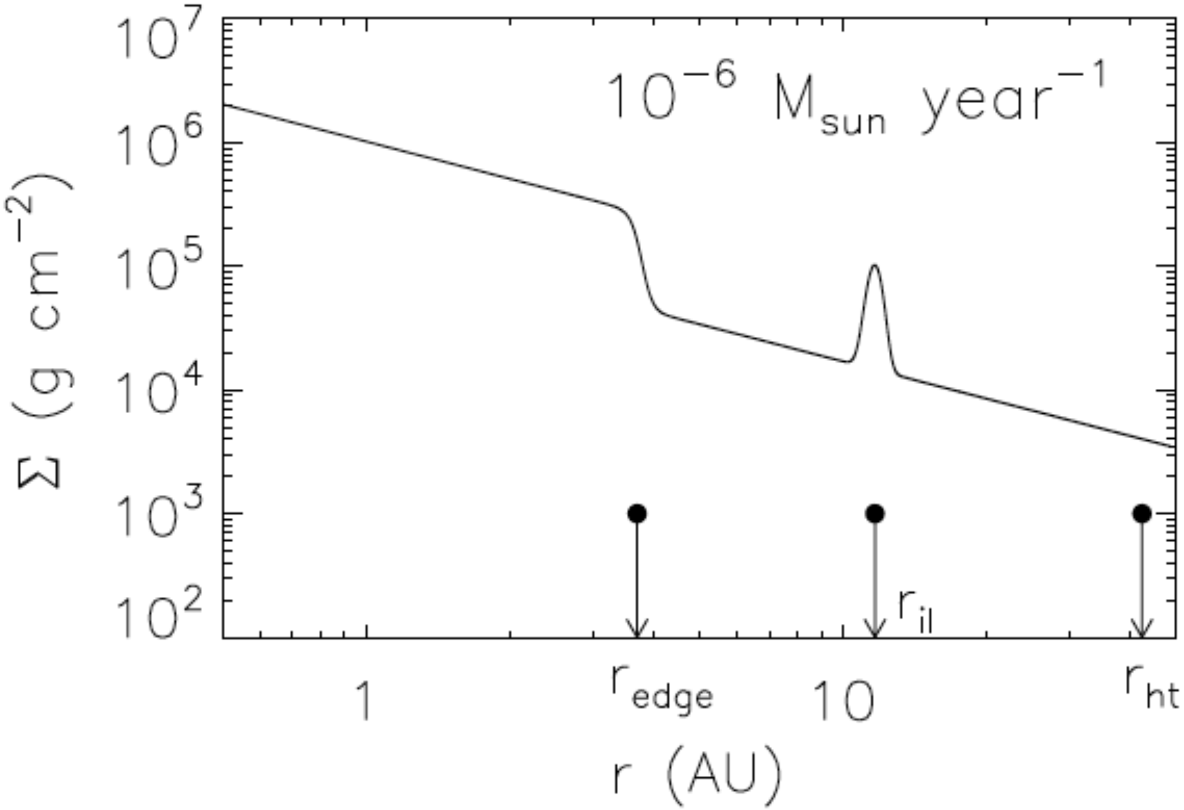


- mass pile-up (Morbidelli et al. 2007) promoting GI or Rossby Wave Instab (Varniere & Tagger 2006)
- long type I & II migration times (Thommes 2005; Chiang, Fischer & Thommes 2002; Matsumura et al. 2006)
- decrease in active layer thickness causes pressure maximum halting type I migration (Ida & Lin 2008b)

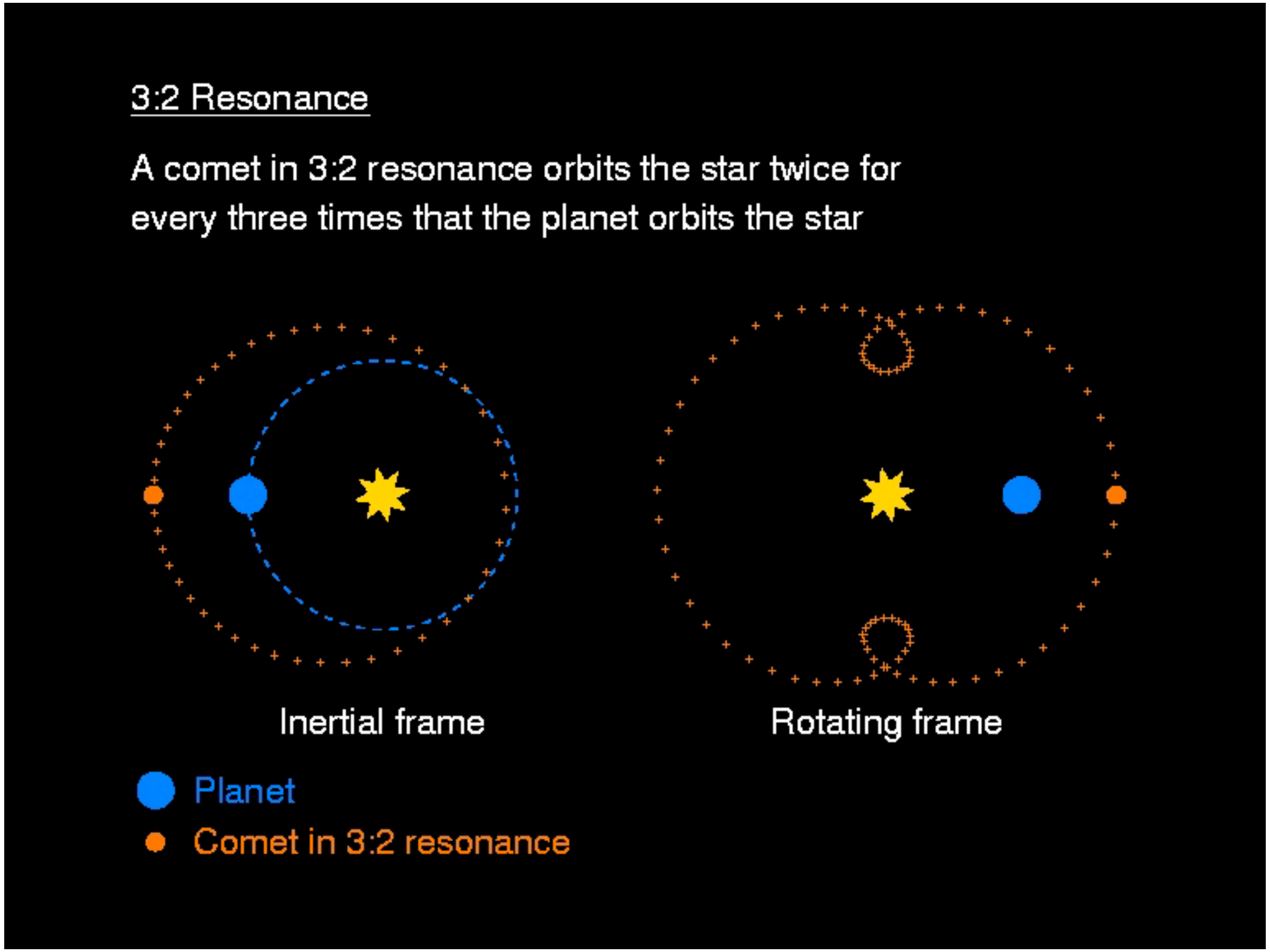
Planet traps

Dead zones cause two planet traps at edge and also a thermal transition, and also a trap from ice-line where solid surface density jumps by x4 due to ice condensation (Hasegawa & Pudritz 2011)

Ice-line is at T<145-170K depending on partial pressure of water vapour (Podolak & Zucker 2004), ~2.7AU in solar system from abundance of icy C-class asteroids (Rivkin et al. 2002; Hsieh & Jewitt 2006) though note the snowline moves during PMS evolution (Kennedy et al. 2006)



Importance of mean motion resonances

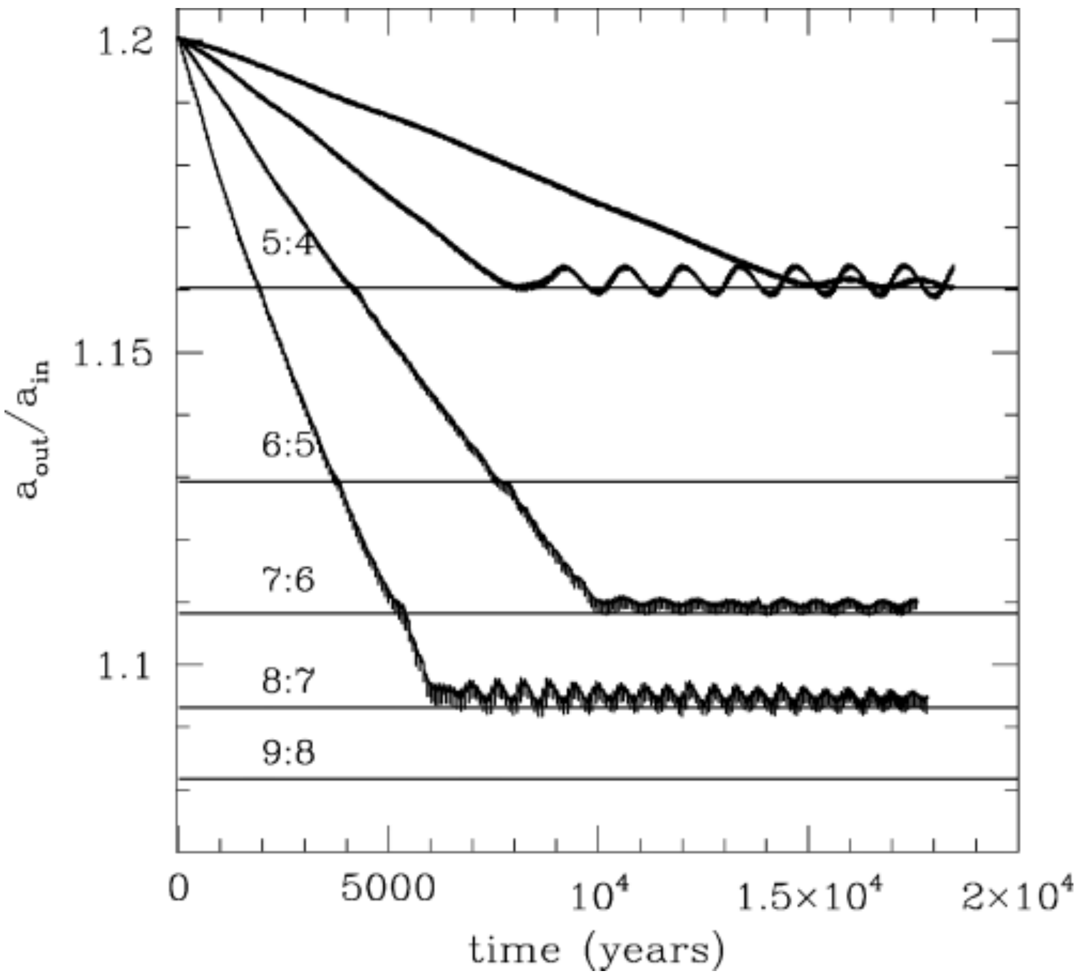


Planet migration with multiple planets: resonances

Trapping only occurs for converging orbits; it is probabilistic but probability of trapping is readily worked out (Mustill & Wyatt 2011)

e.g., Earth-mass planets with type I migration trapped in first order resonances (7:6 etc) (Papaloizou & Szuszkiewicz 2005; Terquem & Papaloizou 2007)

More massive planets get trapped into 2:1 and 3:2 resonances

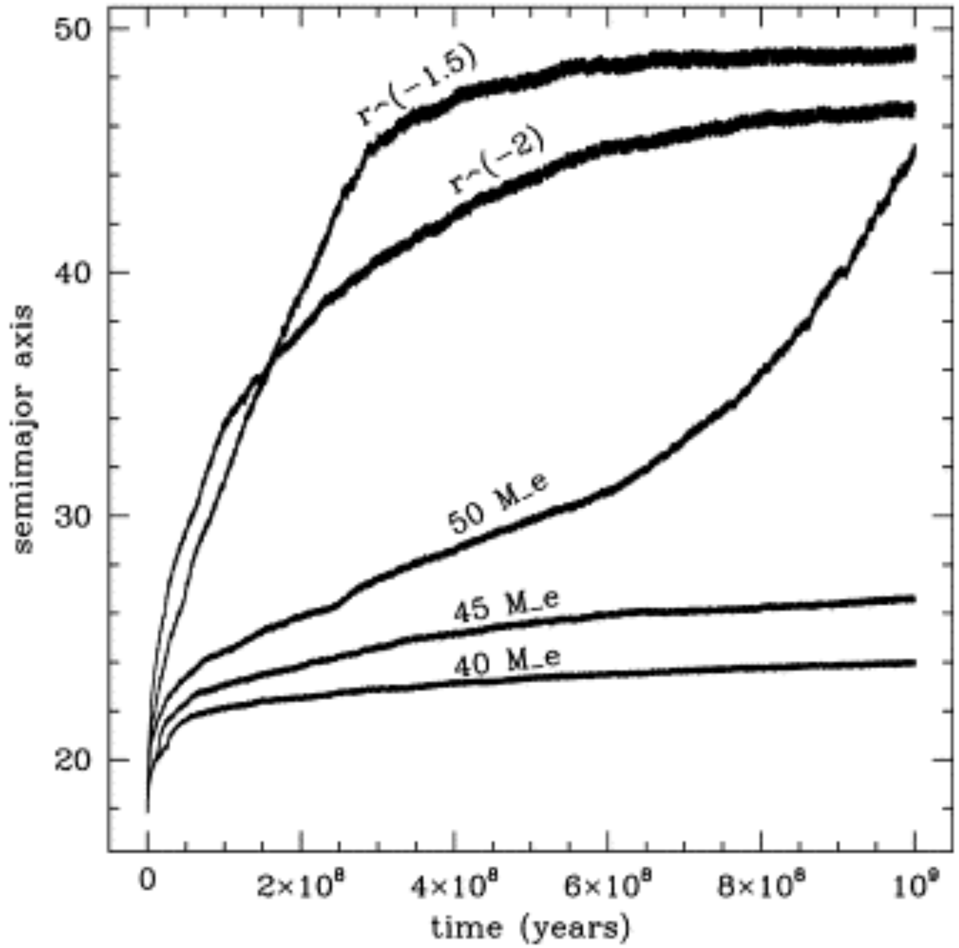


Late stage migration due to planetesimal scattering

During chaotic growth proto-planet and planetesimal scattering results in exchange of angular momentum and so radial migration of planets (Fernandez & Ip 1984)

This type of migration applies to planets less massive than Saturn, and is responsible for Kuiper belt structure (e.g., Hahn & Malhotra 1999)

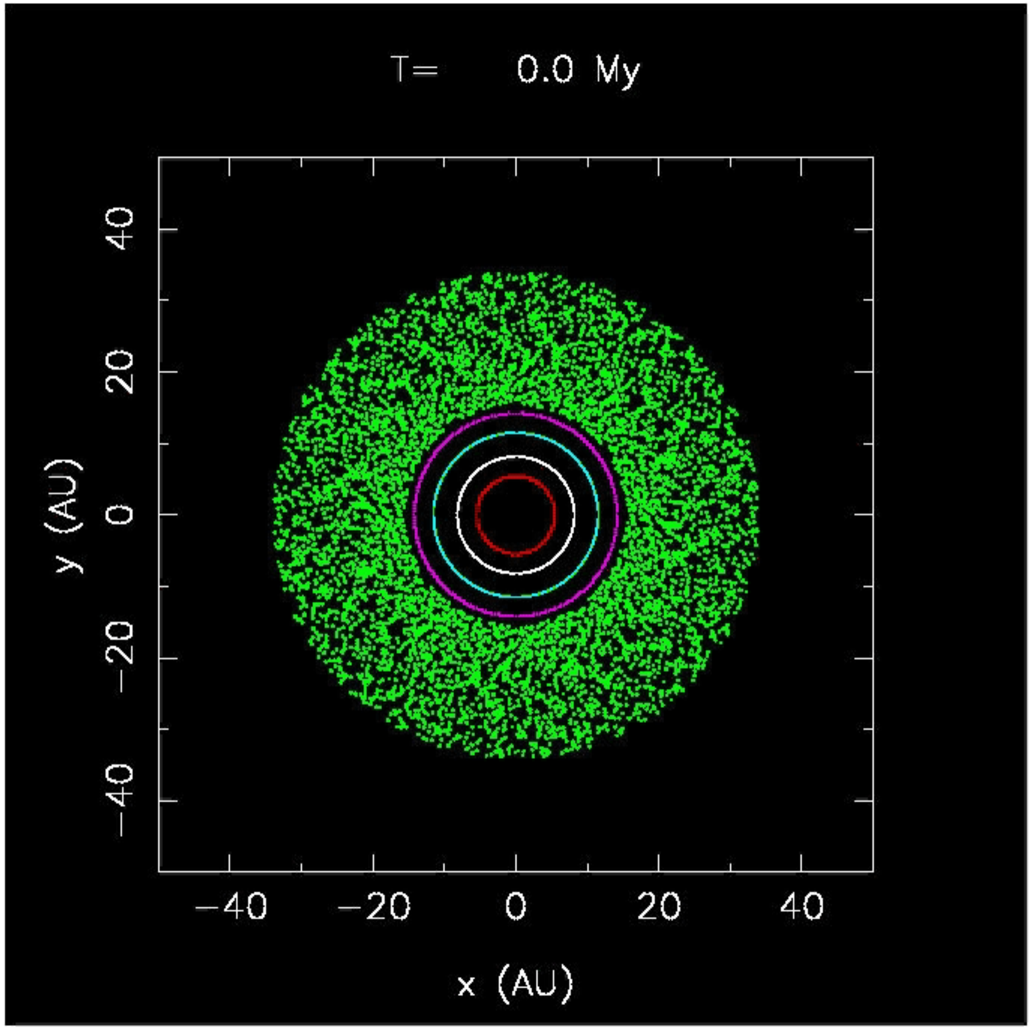
Migration faster in more massive disk (Gomes et al. 2004), and can reverse when planet encounters planetesimal disk outer edge



Chaotic evolution

Multiple planet systems can be chaotic and evolution of outer solar system still mystery:

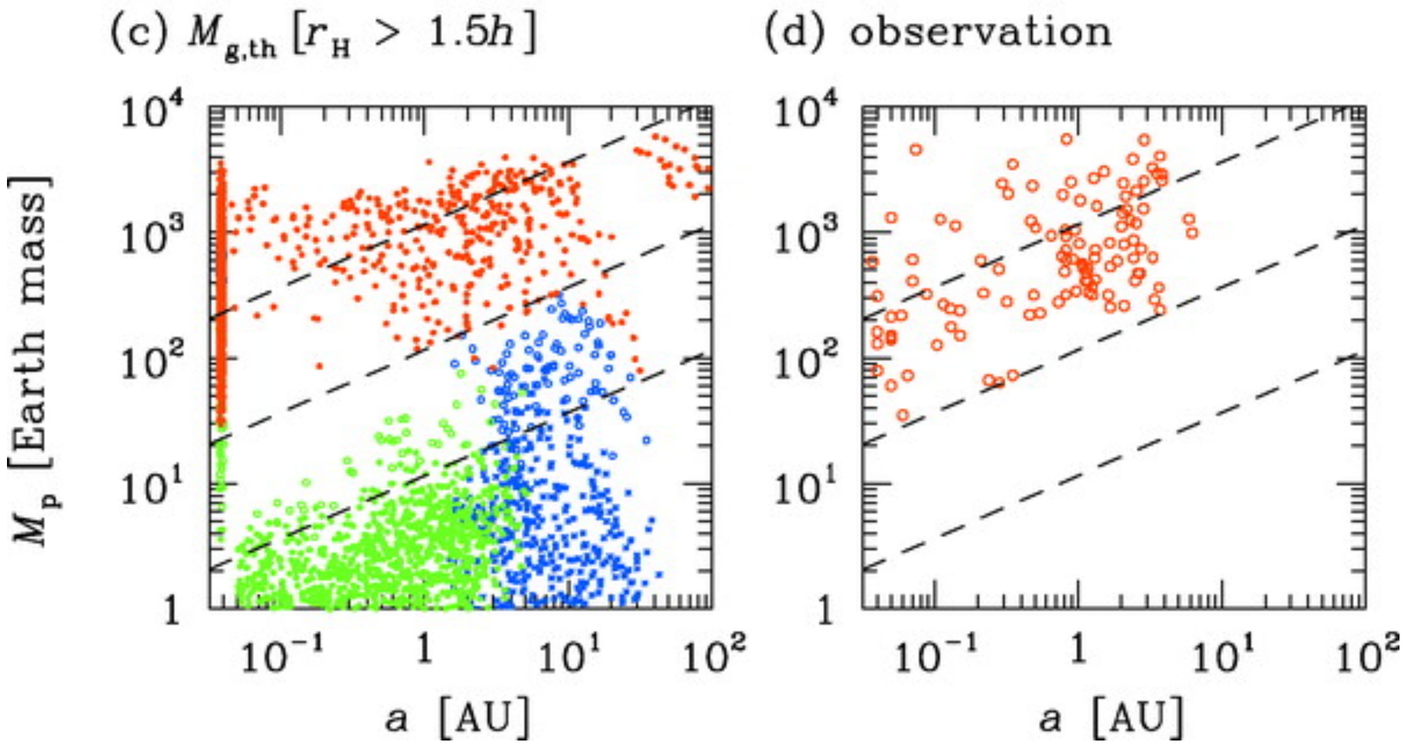
Current model has slow migration causing Jupiter and Saturn to cross 2:1 resonance pumping up eccentricities of UN (Tsiganis et al. 2005)



Population models

Monte Carlo simulation, drawing initial radius, disk mass, and disk lifetime randomly, growing planetesimals into embryos that can accrete gas and migrate (Ida & Lin 2004)

Predicts a desert in mass-semimajor axis distribution caused by rapid growth from a few to >100M_{earth} and slow core growth at >3AU

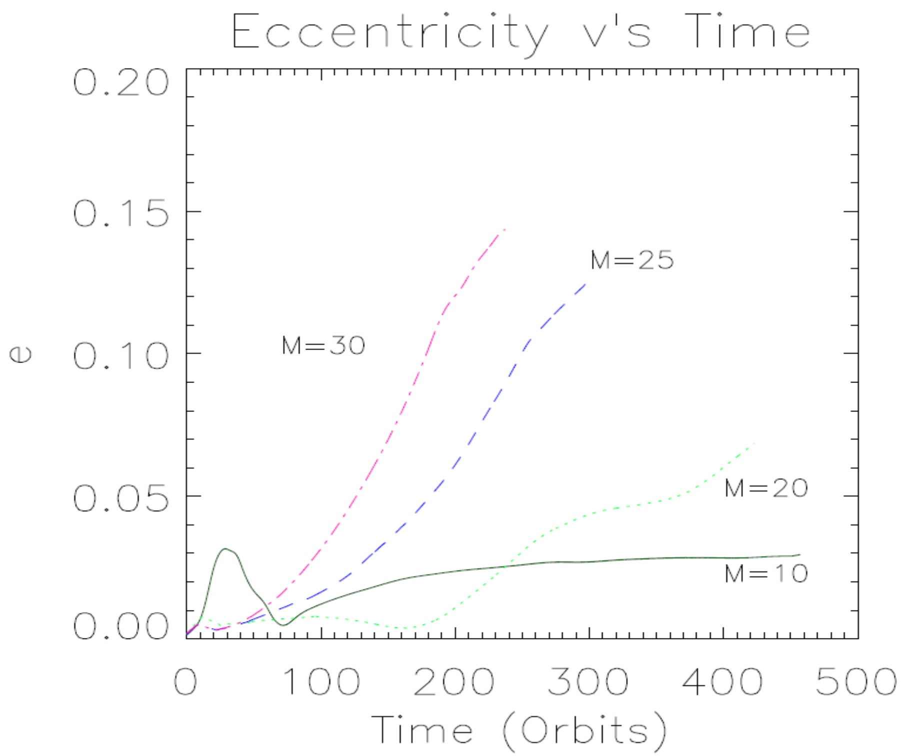


Eccentricity distribution: planet-disk interaction

Outstanding question is the origin of the large eccentricities of planets:

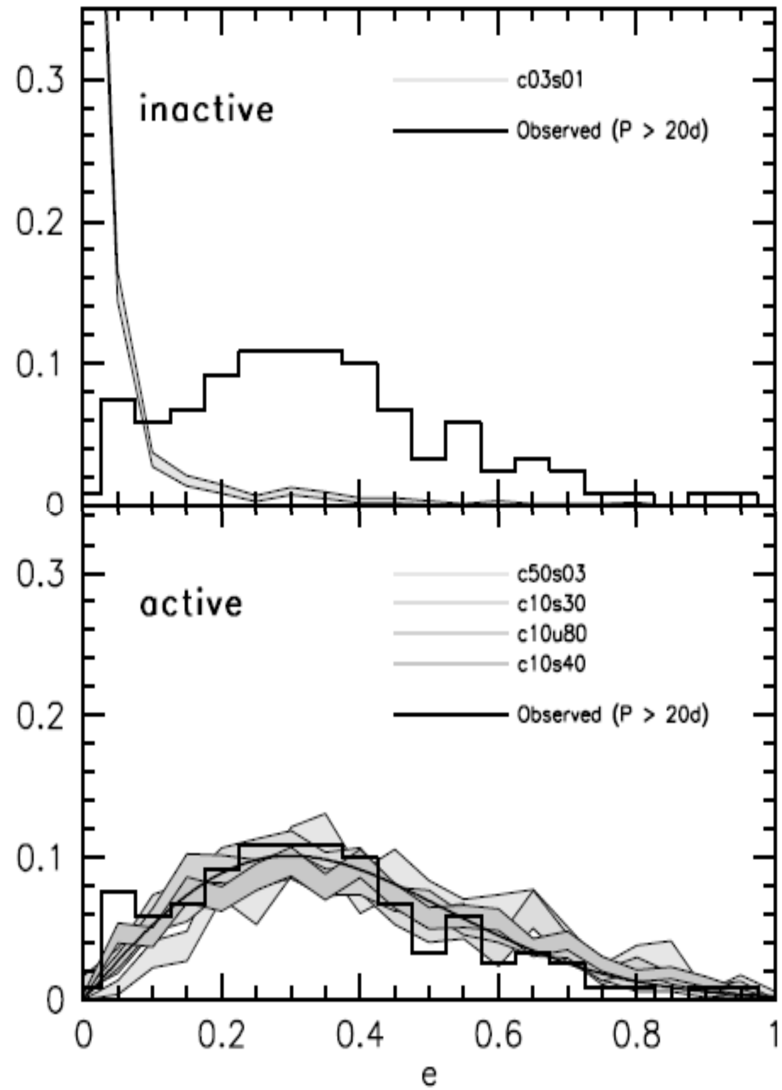
Theory: External first order Linblad resonances pump e_{pl} (Goldreich & Tremaine 1980), but e_{pl} damped by corotational (Artymowicz 1993) and apsidal (Ward & Hahn 2000) resonances

Simulations: Back reaction damps e_{pl} and e_{pl}~0.2 when >10$M_{jupiter}$ (Papaloizou, Nelson & Masset 2001; Kley & Dirksen 2006; Matsumura & Pudritz 2006; **Dunhill** et al. 2013)



Eccentricity distribution: planet-planet interaction

- Dynamical instability of 2 planets ejects outer planet leaving closer-in planet with high e_{pl} (Rasio & Ford 1996; Ford & Rasio 2007)
- **Jumping jupiter** = instability with 3 planets in which one ejected (Weidenschilling & Marzari 2002) predicts high e_{pl} systems have Jupiter on wide orbit
- Multiple planet systems with random parameters relax to observed eccentricity distribution (Juric & Tremaine 2007)

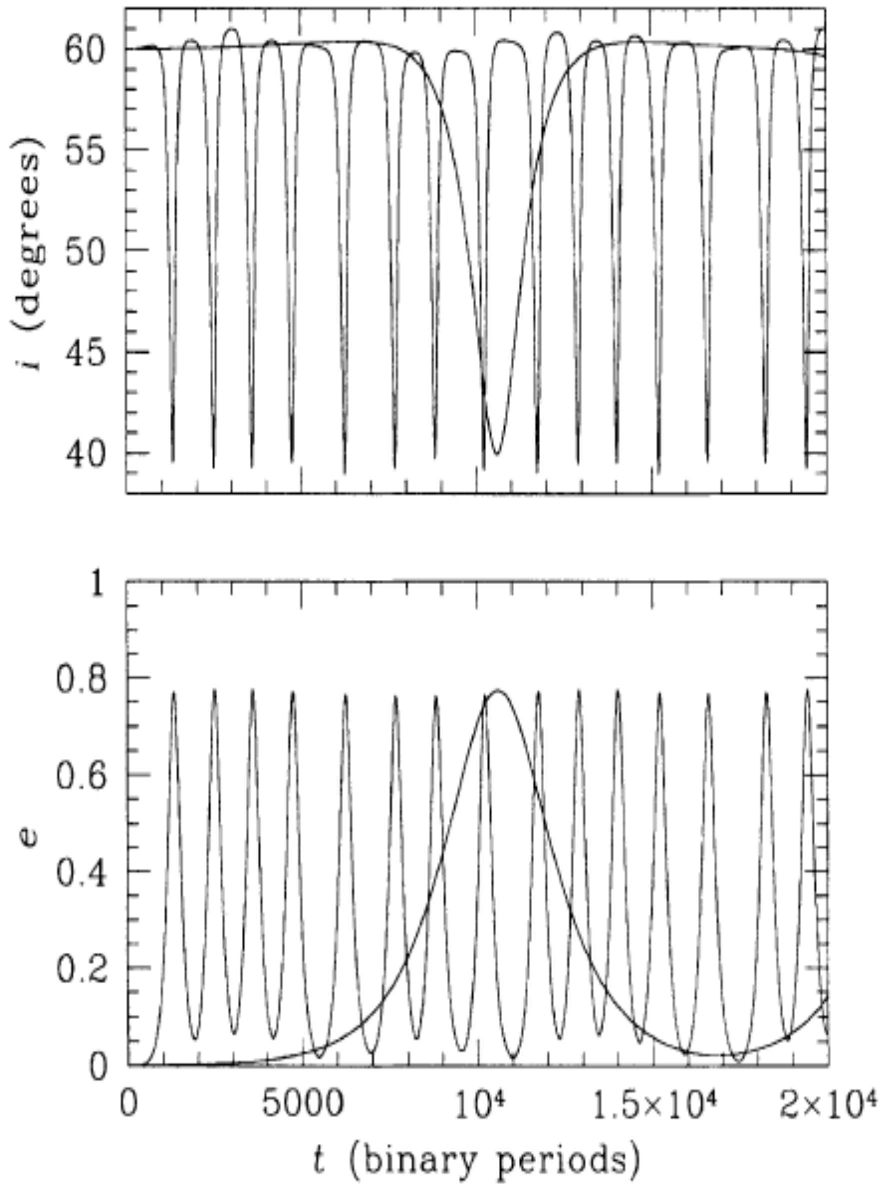


Kozai interactions

- For planets on orbits that are highly inclined to orbit of binary star, secular perturbations called Kozai oscillations cause high e_{pl} (Holman et al. 1997))

- But produces more low e_{pl} than high e_{pl} planets (Takeda & Rasio 2005)

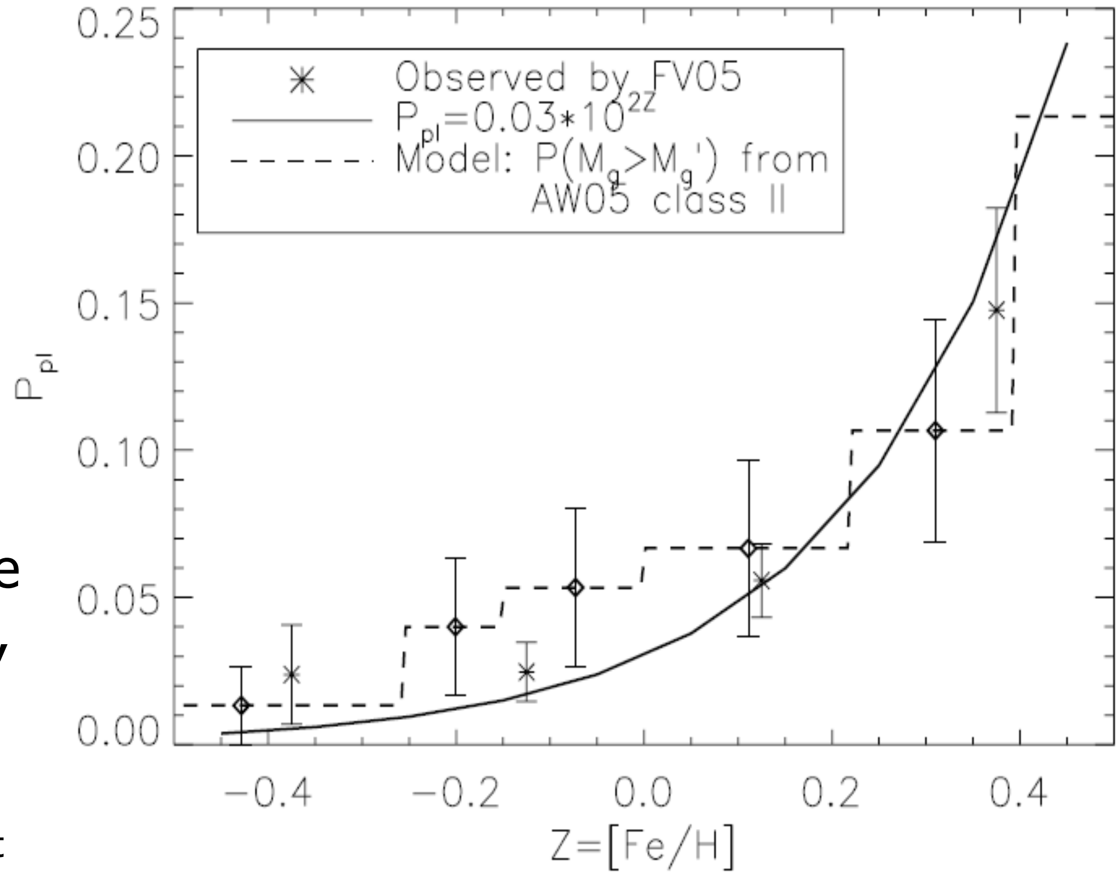
- However, Kozai oscillations in conjunction with stellar tides is one possible origin of Hot Jupiters



Metallicity distribution

• Planet host stars are metal rich, seen as proof of formation by core accretion, since faster growth predicted in higher Z (metallicity) disks because of the higher density of solids ($t_{growth} \propto \Sigma_d^{-1.5}$, Ida and Lin 2004; Kornet et al. 2005)

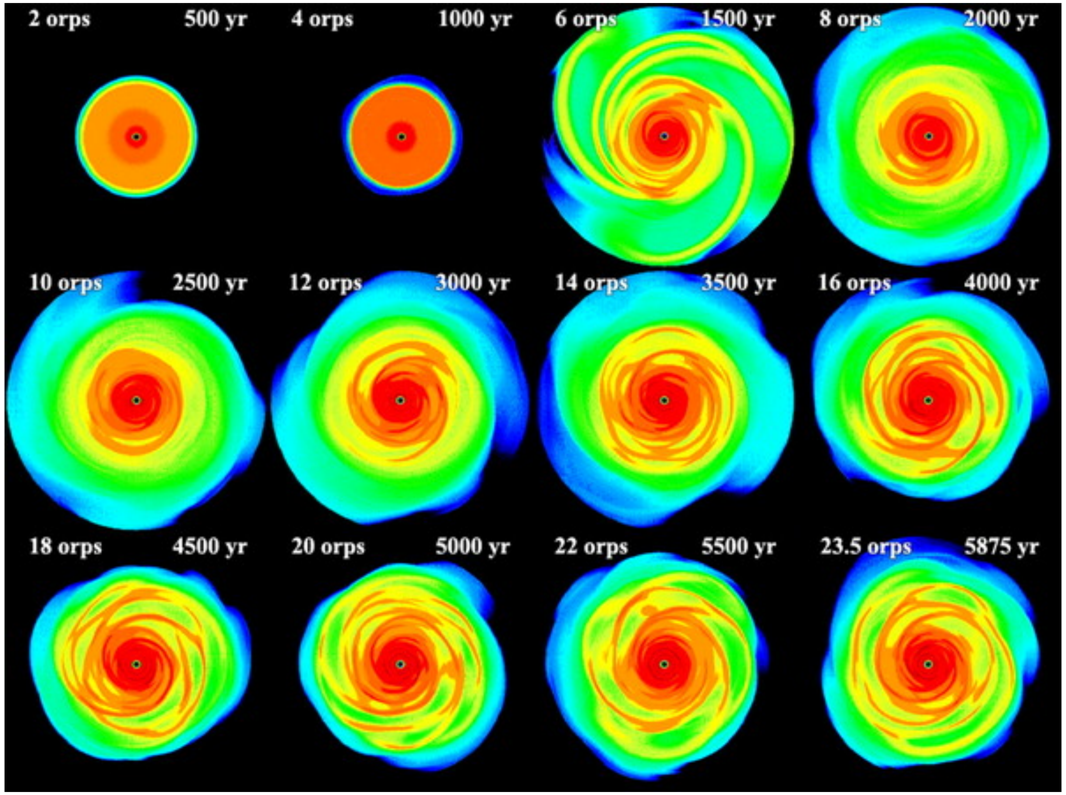
• Form of metallicity dependence from distribution of PPD masses, since if $M_s = 0.01\ M_g\ 10^Z$ and a planet forms when $M_s > M_{s,\,crit}$ then $P_{pl} = P(M_g > 100 M_{s,\,crit}\ 10^{-Z})$ Wyatt, Clarke & Greaves (2007)



Gravitational instability model

Planets form on orbital timescales when part of disk becomes gravitationally unstable (Kuiper 1949, Cameron 1978): $Q \sim (M_{star}/M_{disk})(H/r) < 1$
Characteristic size is H and so mass $\sim M_{jupiter}$ (assuming $H/r \sim 0.1$)

But, as disk builds up mass from envelope (decreasing Q), instabilities as Q approaches 1 (Laughlin & Bodenheimer 1994) lead to angular momentum transport, so Q never reaches 1 (Vorobyov & Basu 2007) unless the disk is cooled (so H/r decreases) or matter added (so M_{disk} increases) quicker than orbital timescales (Gammie 2001)



Gravitational instability model

Cooling and formation location:

• Radiative transfer can't cool mid-plane sufficiently, but convection currents can and GI possible >8AU (Boss 2004; Rafikov 2006)

• Disks forming planets by GI <10AU would be uncommonly luminous (Rafikov 2005)

• Gas giants difficult to form at >100AU due to rapid inward mass transport by spiral arms (Boss 2006)

• Only 1-10$M_{Jupiter}$ planets at 10-100AU by GI are possible

Are clumps long-lived?

• Simulations show clumps may not be long lived (Durisen et al. 2001; Mejia et al. 2005; Pickett & Durisen 2007)

• But survival lifetime in simulations increases with resolution (Boss 2005)

GI: Confrontation with known planets

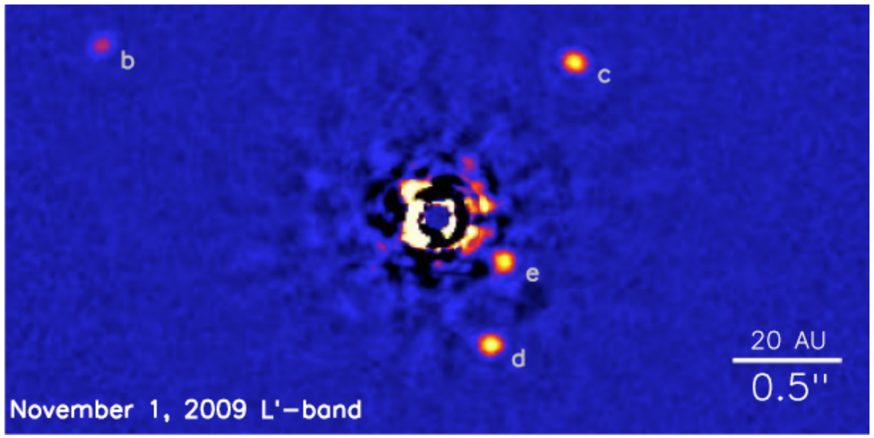
Can it explain origin of cores of Solar System giant planets?

• Yes: as rock and ice cores can form after planet through sedimentation (predicts 6 and 2 M_{earth} cores, mostly Si for Jupiter and Saturn) (Boss 1998; Helled et al. 2008)

Does it have correct dependence on metallicity?

• No: stellar metallicity does not affect planet formation by GI because disk radiative energy loss is controlled by star not disk radiation (Boss 2002; Cai et al. 2006), although lower compressional heating may help (Mayer et al. 2006)

But it's the best explanation for HR8799 (Dodson-Robinson et al. 2010)



Summary of lecture 2

- **Growth 10km – Jupiter through core accretion is relatively easy, though details not finalised (like role of debris, and circumplanetary environment)**

- **Also need to prevent migration, and proto-planetary disk structure may help through planet traps**

- **Most exoplanet observations explained by core accretion using “population synthesis” models, but gravitational instability needed for massive outer planets**