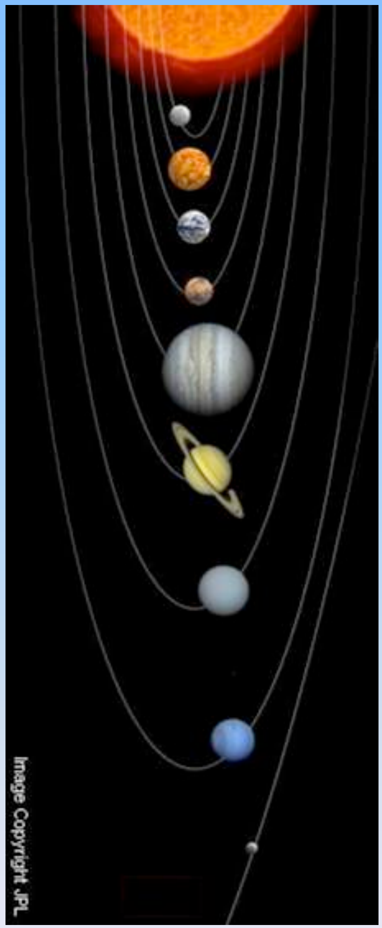
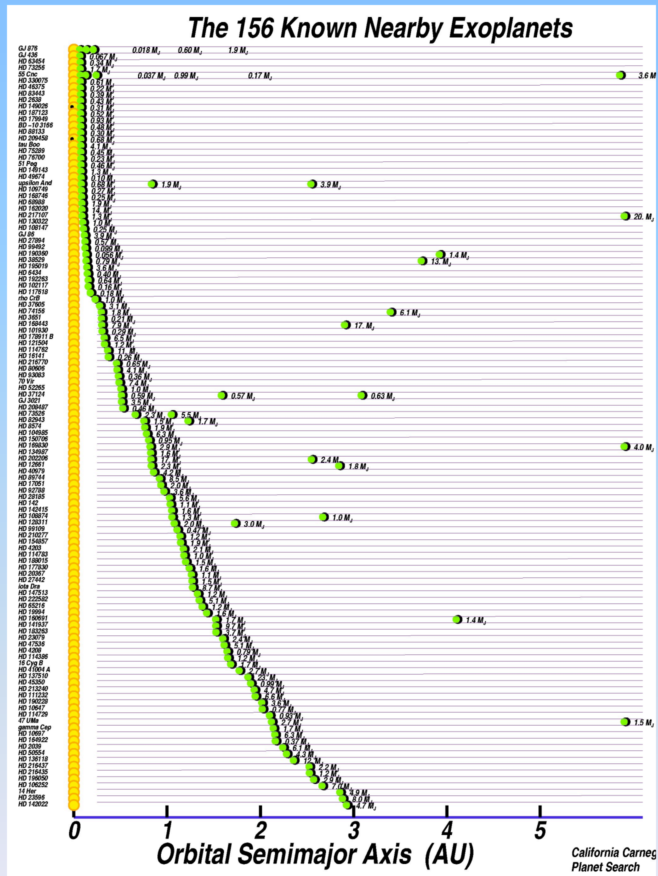
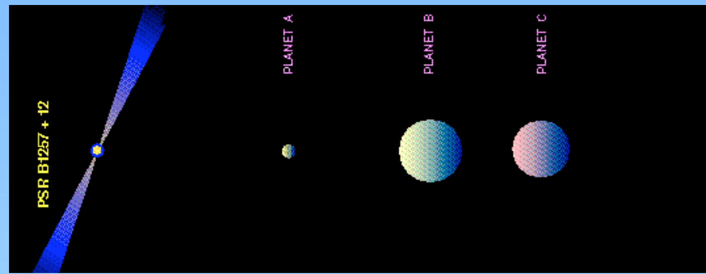
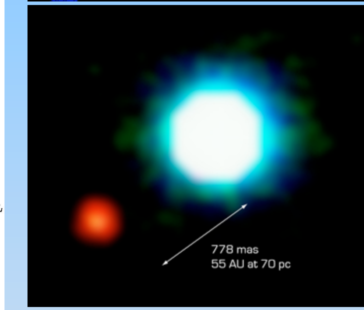
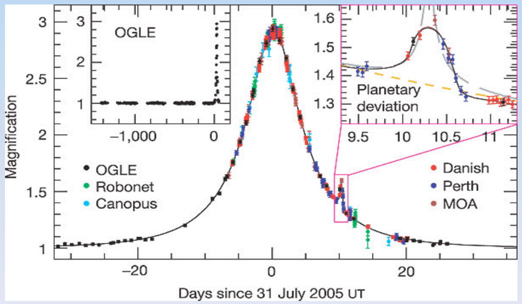
4. Planet formation











So, where did all these planetary systems come from?

Planets form in disks

Reviews of how solar system formed: Lissauer (1993)
Recent reviews of planet formation: Papaloizou & Terquem (2006); also Lissauer+, Durisen+, Nagasawa+, Dominik+ in Protostars and Planets V

There are 14 observations formation models have to explain (Lissauer 1994), two of which are:

- planets orbits are circular, coplanar, and in same direction
- formation took less than a few Myr

Planets form in circumstellar disks in a few Myr

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

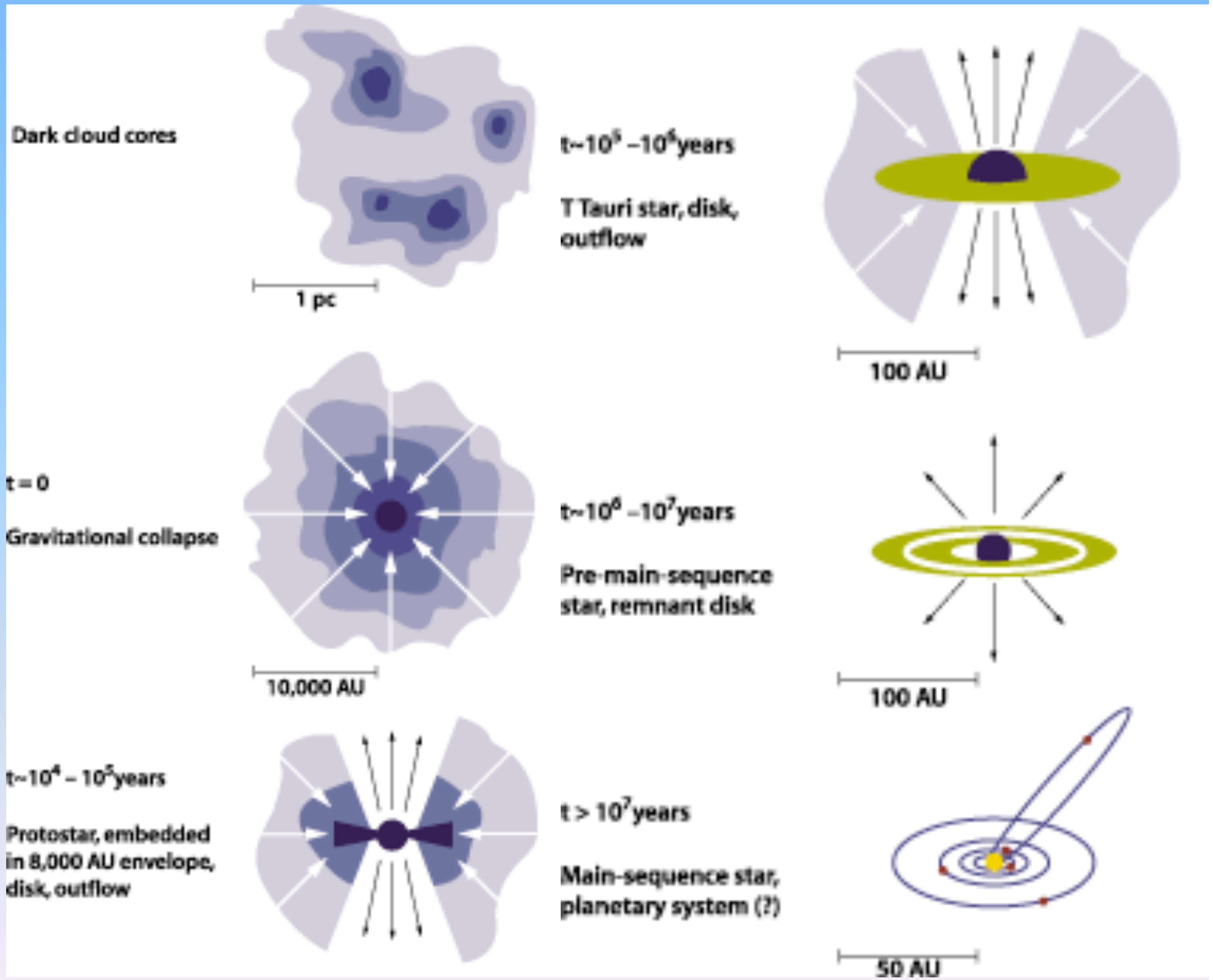
Star Formation

Basic picture (Shu et al. 1987):

Stars form from the collapse of clouds of gas and sub-micron sized dust in the interstellar medium

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, although models not without problems

The origin of the giant planets, and of extrasolar planets, is still debated, but core accretion models reproduce most observations

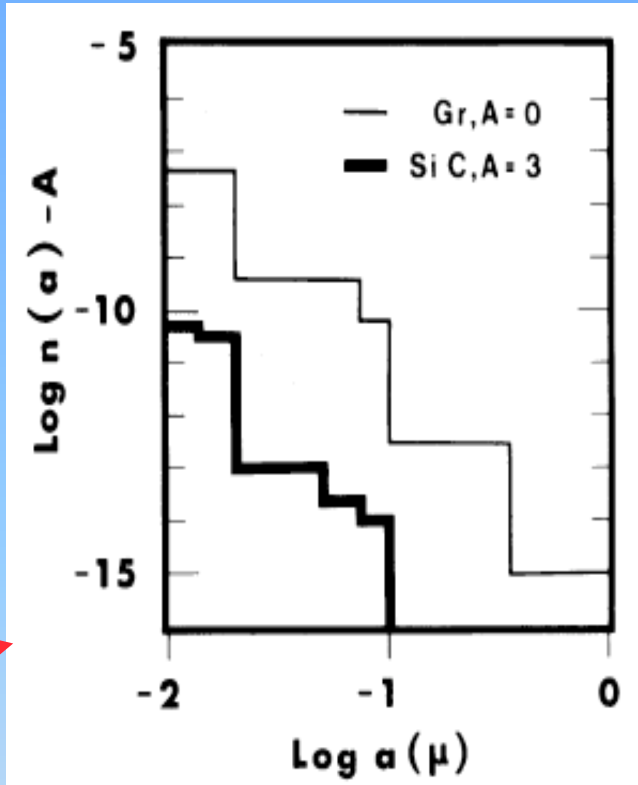
0. Starting conditions

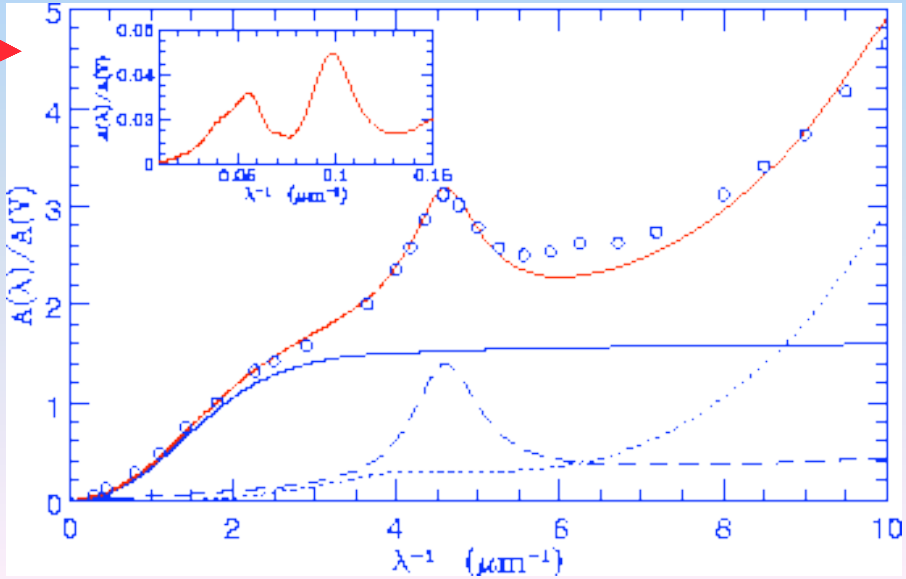
Proto-stellar disk is composed of same material as star, since meteorites have same composition as Sun

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

[see also Dorschner & Henning 1995]





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):

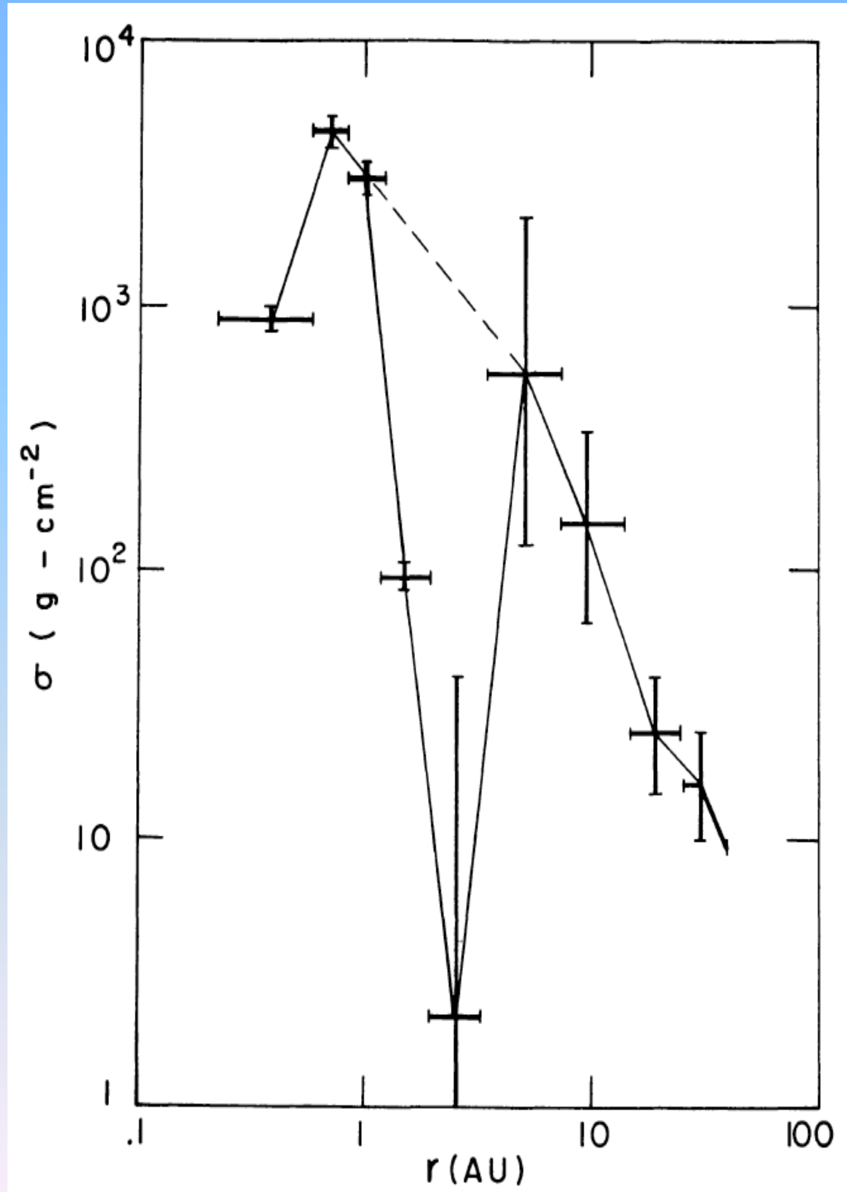
$$\Sigma_{total} = 3\text{-}6\times10^{4}\ r^{-1.5}\ \mathrm{kg/m^2}$$
$$\Sigma_{solid} \approx 0.01\Sigma_{gas}$$

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

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

Possible jump x4.2 at 2.8AU in density of solids where temperature was low so water ice condenses

But primordial nebula may have had different mass distribution (Desch 2007)

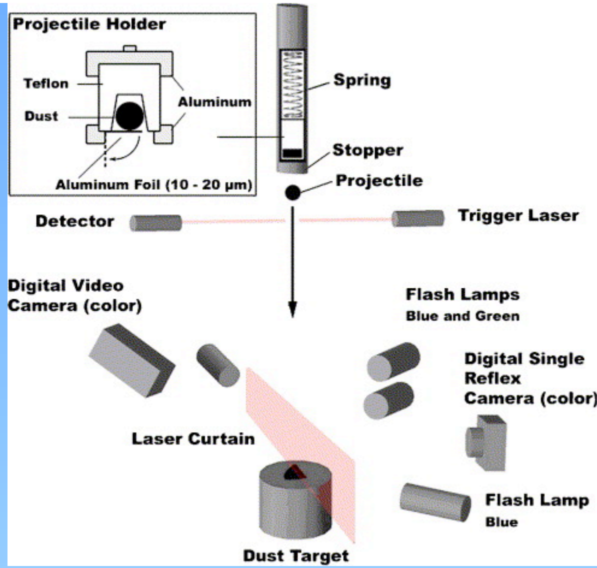


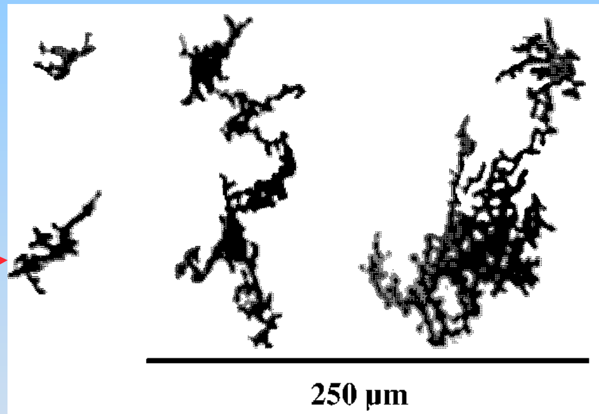
1. Grain growth: 1μm-1m

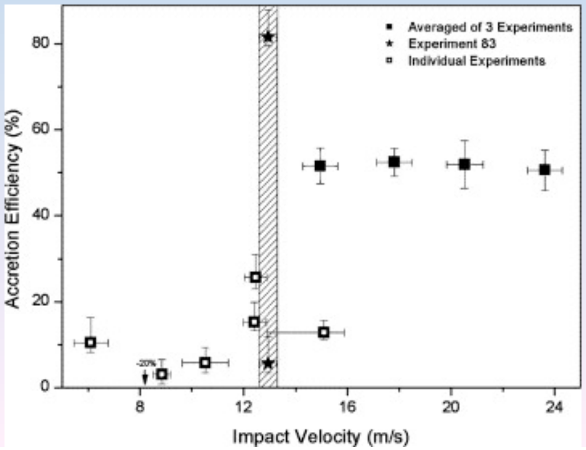
In disks IS grains (and condensates) collide

Outcome depends on collision velocity and sticking properties of grains which are studied both experimentally and theoretically (Heim et al. 1999; Dominik & Tielens 1997; Poppe et al. 2000; Konchi et al. 2002; Wang et al. 2005):

- small grains grow fractally in 0.01m/s collisions, $m \propto D^{1.9}$ (Wurm & Blum 1998, 2000), with porosity 0.67-0.93 (Blum et al. 2007)
- D>1cm collisions compact grains (Blum & Wurm 2000) giving high velocities of ~10m/s, $m \propto D^3$ (Sekiya & Takeda 2003)
- high velocity collisions result fragmentation, but also net accretion (Wurm et al. 2001; 2005)







Gas drag

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

Drag force depends on ratio of:
relative velocity of gas and dust, $\mathbf{\Delta v} = \mathbf{v_g} - \mathbf{v_d}$,
mean thermal velocity of the gas, $v_t=(4/3)[8kT/\pi\mu_g m_H]^{0.5}$

Two regimes:
SUBSONIC ($|\mathbf{\Delta v}|<v_t$) is **Epstein drag law**: $\mathbf{F_g} = -0.25\pi\rho_g D^2 v_t \mathbf{\Delta v}$
SUPERSONIC ($|\mathbf{\Delta v}|>v_t$): $\mathbf{F_g} = -0.25\pi\rho_g D^2 |\mathbf{\Delta v}| \mathbf{\Delta v}$

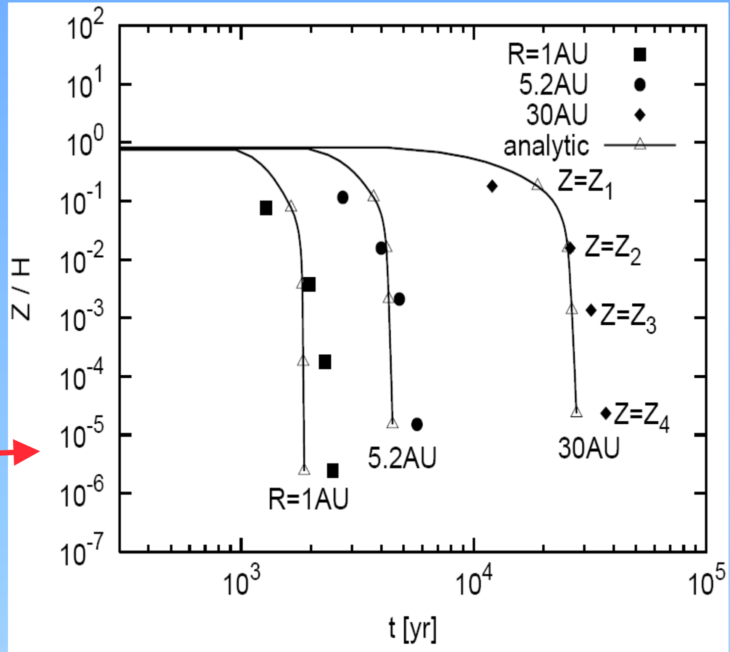
Stopping time is that to cause $|\Delta v|=0$, $t_s=m|\Delta v|/|F_g|$, which compared with orbital velocity, $v_k=\Omega_k r$ gives the ratio

$$T_{ss} = t_s\Omega_k = 2\rho_d D v_k/3\rho_g r v_t = \Sigma_{1p}/\Sigma_g$$

- **DECOUPLED** if $T_{ss}>>1$ (large grains close to star)
- **STRONGLY COUPLED** if $T_{ss}<<1$ (small grains far from star)

Settling to mid-plane

Gas drag causes dust to settle to mid-plane as inclined orbits oscillate vertically, and gas drag damps oscillation



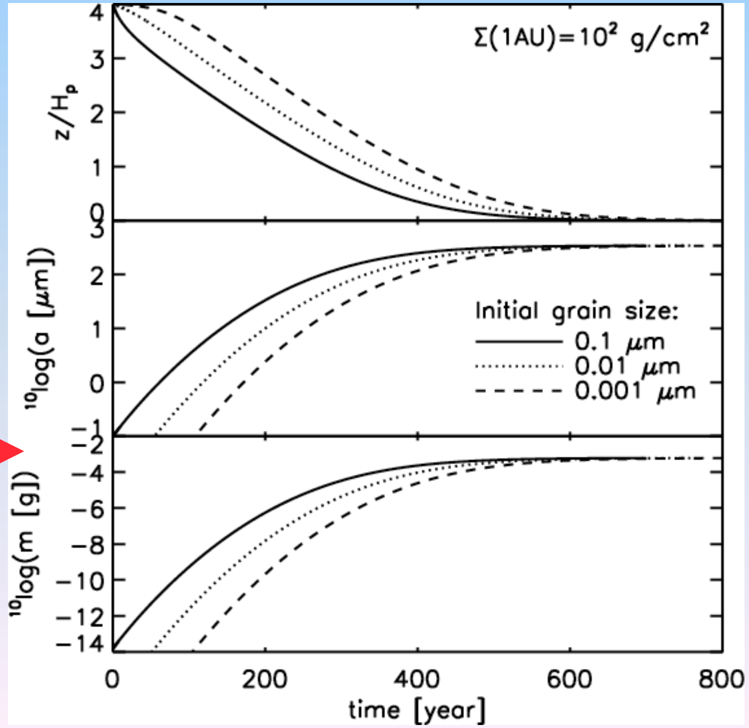
Sedimentation time:

$$1/\Omega_k^2 t_s = 3\rho_g v_t / 2\Omega_k^2 \rho_d D$$

though slower for porous dust

(Ormell et al. 2007)

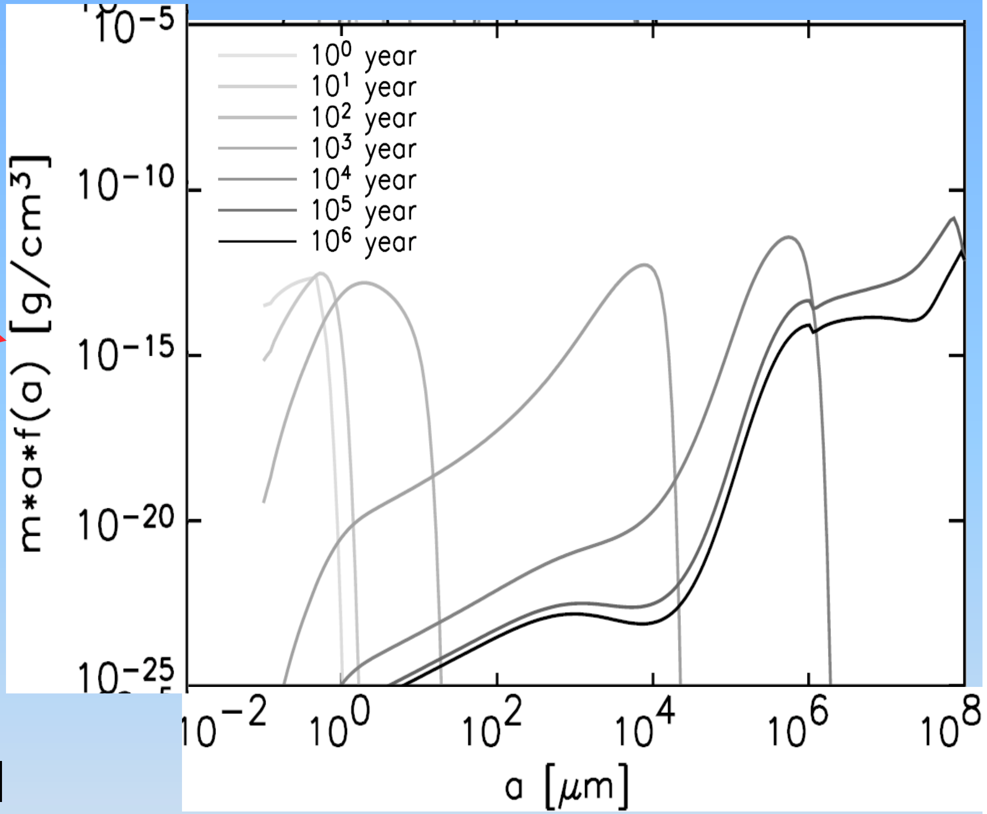
Timescale long for small grains, but these collide during settling speeding process up (Weidenschilling 1980; Nakagawa et al. 1981; Dullemond & Dominik 2005)



Radial migration is then important

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

- creation of small grains in collisions important [absent in models but seen in proto-planetary disks, van Boekel et al. 2004]

- small grains on surface dominate optical depth

While details of turbulence not well understood (Voelk et al. 1980), and are studied using MHD models (e.g., Carbillado, Fromang & Papaloizou 2006), this is not thought to prevent settling (Youdin & lithwick 2007)

2. Grain growth: 1m-10km

Proceeds by collisions between planetesimals?

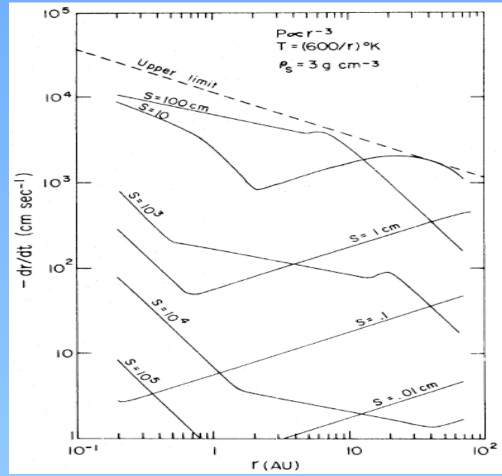
Timescale problem: metre-sized objects migrate in due to gas drag in 100 years, much faster than collisional growth times

Resolution: slow down migration or speed up growth

- gravitational instability
- turbulence/vortices
- spiral structure

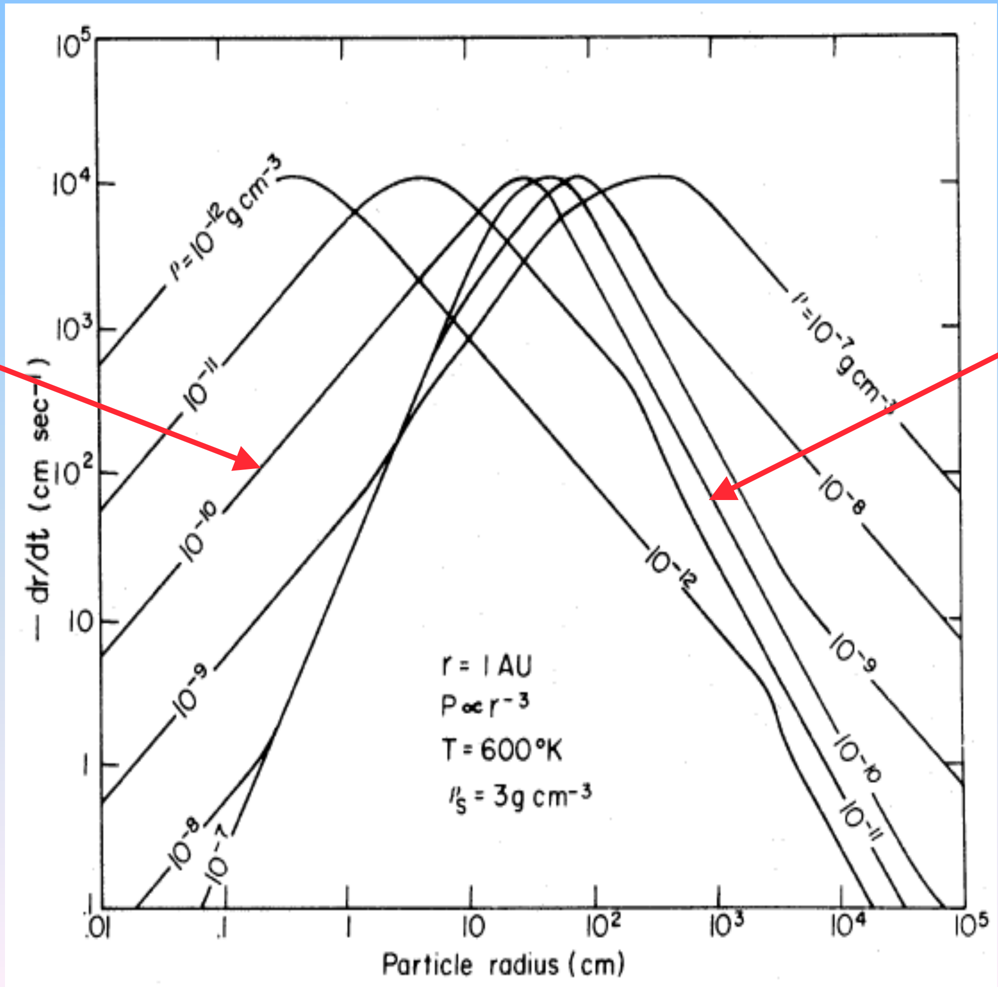
Radial migration

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



1cm/s = 2.1AU/Myr

Grains **coupled** to gas orbit at sub-keplerian gas velocity giving an extra acceleration toward the star = drift in at terminal velocity



Grains **decoupled** from gas spiral in due to the headwind, which means smaller grains migrate faster (due to larger area/mass)

Gas disk structure

(1) 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 gradient decreases with r, so gas velocity is sub-keplerian, dust sees headwind and migrates in

But,

- pressure reverses at disk gap/jump
- radiation pressure gives dust sub-keplerian velocity (Takeuchi & Artymowicz 2001)
- instabilities to changes in radial density distribution (Klahr & Lin 2005)
- turbulence changes pressure gradient (e.g., vortices, Klahr & Bodenheimer 2006)

(2) Vertical component of momentum equation:

$$H = rv_t/v_k \propto r^{1.5}T^{0.5}$$

so as long as $T \propto r^{>-1}$ then disk is flared ($T \propto r^{-0.5}$ for black body dust)

Gravitational instability (GI)

Speed up growth by GI if dust concentrated in mid-plane, since this makes km-sized planetesimals on orbital timescales (Safronov 1969; Goldreich & Ward 1973)

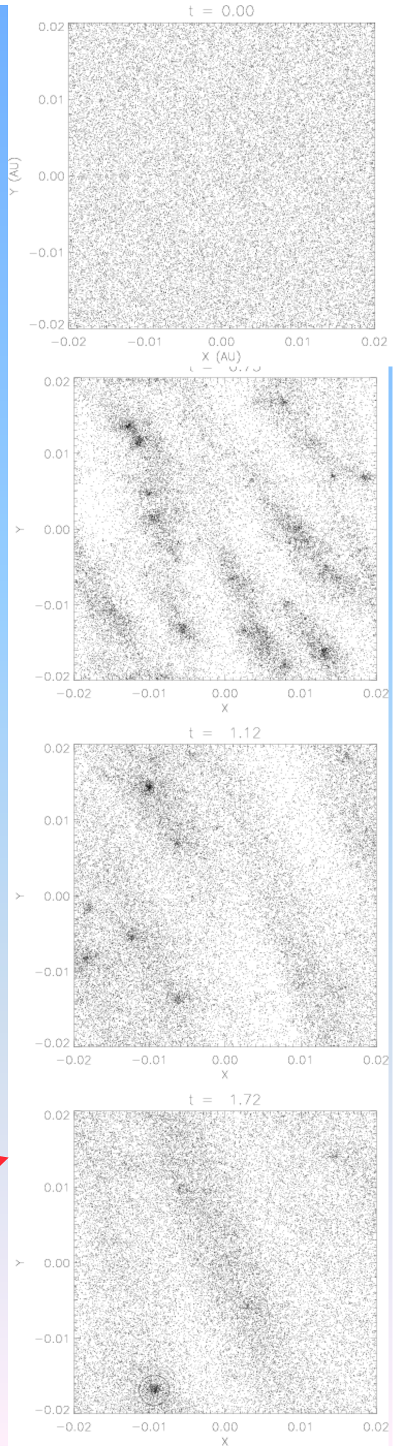
Requires Toomre parameter Q<1

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

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)
- velocity shear doesn't lift all dust (Sekiya 1998; Youdin & Shu 2002)
- inhibited by turbulent stress on particle layer (Weidenschilling 2003)
- helped by size dependent drift rates (Youdin & Chiang 2005)
- N-body simulations of instability process (Tanga et al. 2004)



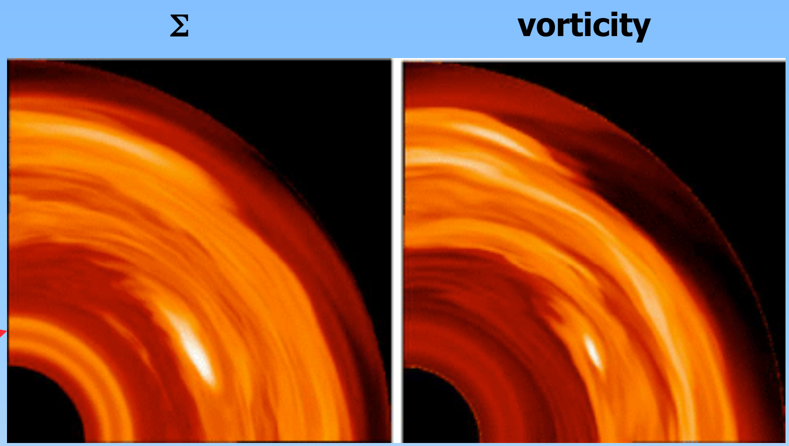
Vortices in proto-planetary disks

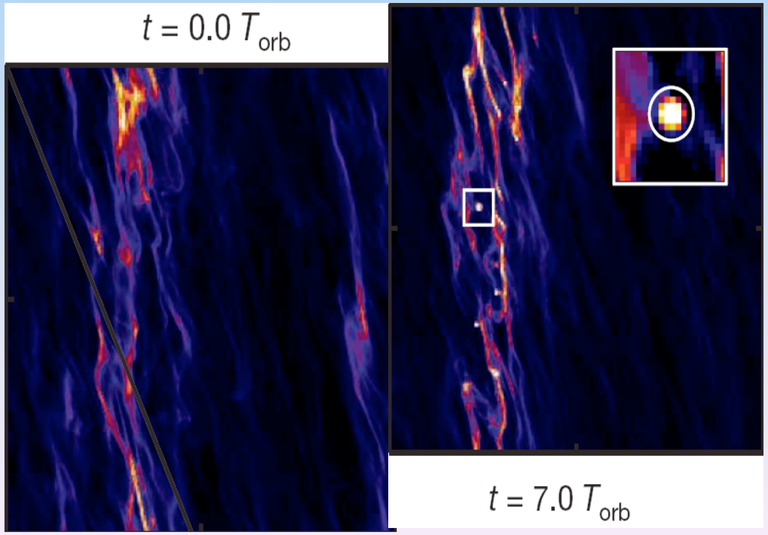
Planetesimal can become trapped in vortices aiding growth (Barge & Sommeria 1995; Tanga et al. 1996; Klahr & Henning 1997; Klahr & Bodenheimer 2003, 2006; Inaba & Barge 2006; Lithwick 2007)

Vortices seen in MHD simulations of dust interacting with turbulent disks, concentrating particles 5cm-10m (Fromang & Nelson 2005; Johansen, Klahr & Henning 2006; Johansen et al. 2007)

Concentrations may be gravitationally unstable, but not clear if vortices last long enough, or if only relevant to specific particle sizes (Godon & Livio 1999; Cuzzi et al. 2001)

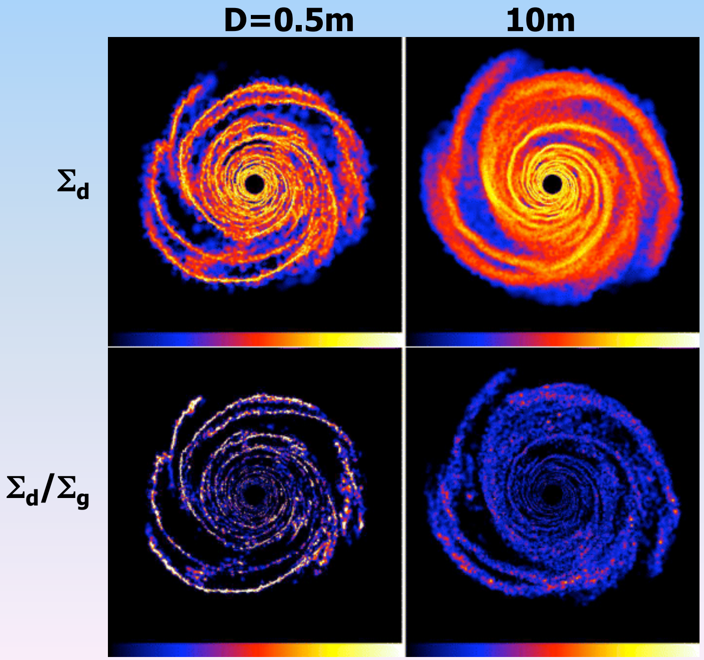
They do reduce drift rates by 40% for D=1m





Spiral and rings in proto-planetary disks

Gas drag also concentrates 1-10m objects in spirals of marginally stable self-gravitating disk (Rice et al. 2004) or of a disk perturbed by a passing star (Theis, Kroupa & Theis 2005; Lodato et al. 2007), although high collision velocities may prevent growth (Britsch et al. 2008)



And in rings:

- Drift rate in turbulent disk $\propto \Sigma^{-1}$, leading to secular instability and dense rings (as if annulus density increases, drag rate decreases) (Goodman & Pindor 2000)

- Clumping instability in optically thin gas disks (Klahr & Lin 2005)

- Photophoresis force (temperature gradient on particle surface) can put up 1μm-10m dust grains at same radius (Krauss & Wurm 2005, Herrmann & Krivov 2008)

3. Runaway growth: 10km-100km

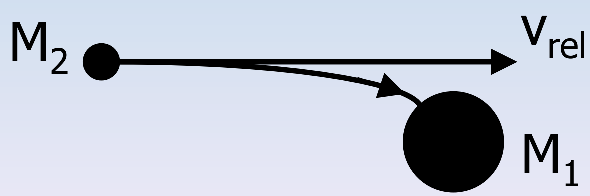
Planetesimals: >km-sized objects compacted by own gravity

Orderly growth: Time to make objects of size m_α:

$t_{acc} = m_\propto/(dm_\propto/dt) = 2r^{1.5}(D_\alpha/1km)(\Sigma/10kgm^{-2})^{-1}$ Myr

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

Runaway growth: Additional factor due to gravitational focussing of $(1+v_{esc}^2/v_{rel}^2)^{-1}$, where $v_{esc}^2=0.25Gm_\alpha/D$. Runaway occurs when $v_{rel}<<v_{esc}$ as $dm_\alpha/dt \propto m_\alpha^{4/3}$ and so large proto-planets decouple from size distribution



Velocity dispersion, v_{rel}, is very important

Modelling methods

Models are either:

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

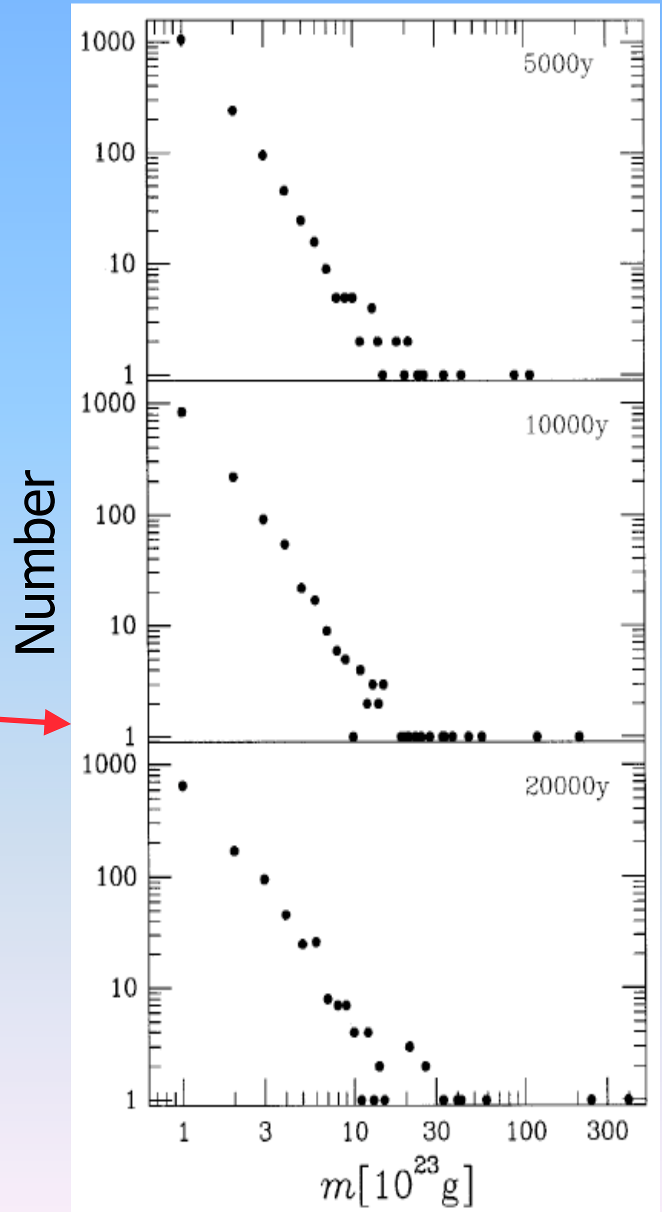
• **Direct:** N-body simulations of gravitational interaction of fewer particles (e.g., Aarseth et al. 1993; Kokubo & Ida 1996)

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 and have low velocity dispersion



Evolution of velocity dispersion

The velocity dispersion is balance of:

- **Gravitational scattering** (increases v_{rel})
  - **Runaway phase:** scattering among planetesimals (m_α) keeps v_{rel} const
  - **Dynamical friction:** scattering m_α by m_β causes v_{rel} of m_β to decrease
  - **Oligarchic phase:** sufficiently massive cores ($3m_\beta\Sigma_\beta > m_\alpha\Sigma_\alpha$), mean scattering amongst cores (m_β) and planetesimals (m_α) increases v_{rel} with m_β

- **Gas drag** (decreases v_{rel})
  - Inclination reduced (settling to mid-plane)
  - Eccentricity reduced (oscillation about r=a also damped)
  - More efficient for small mass particles

- **Disk tides** (decrease v_{rel})
  - Important when $m_\beta > 10^{-2}\text{-}10^{-4} M_{earth}$

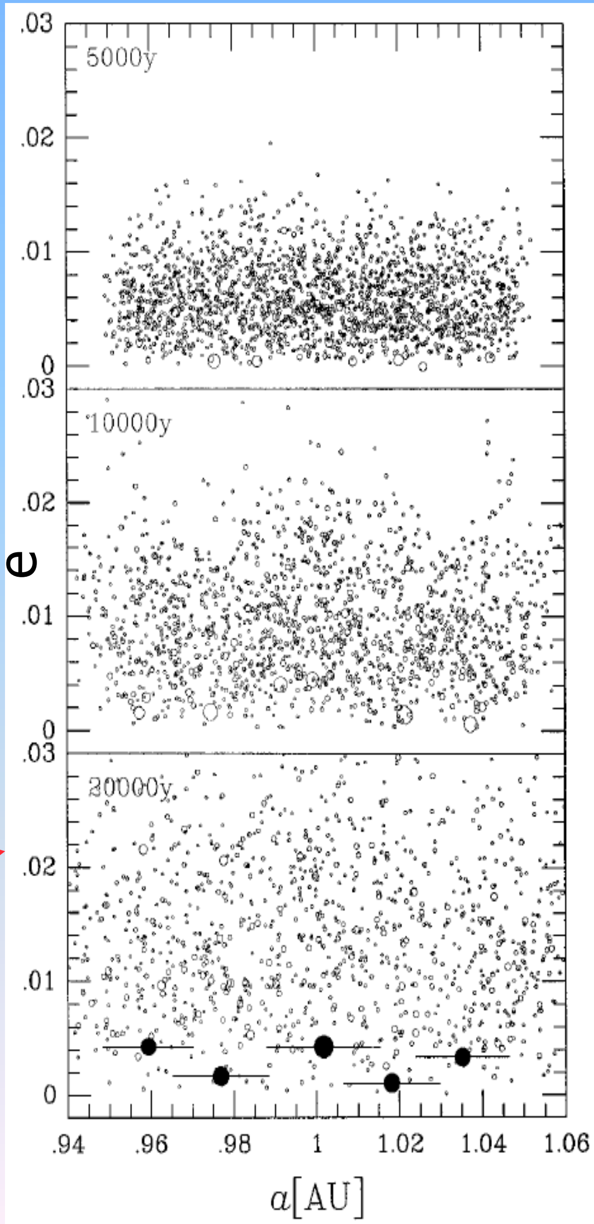
4. Oligarchic growth: 1000km-10,000km

Runaway phase ends when core mass dominates velocity dispersion of planetesimals:

$m_\beta > 2.2\times10^{-7} f^{0.6} r^{6/5} (\Sigma_\alpha m_\alpha / 10^{17} \mathrm{kg}^2 \mathrm{m}^{-2})^{0.6} M_{earth}$

Gravitational focussing strong allowing cores (**oligarchs**) to reach M_{earth} quickly (although slowed if disk is turbulent, Ogihara et al. 2007), 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

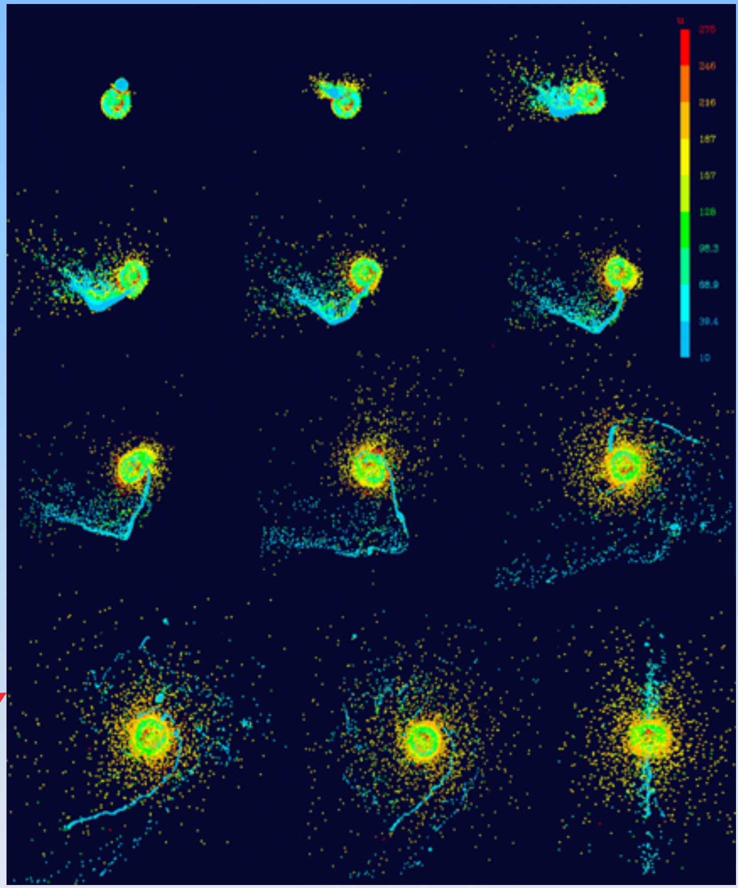
Proto-planet: Massive oligarchs clear **feeding zone** of planetesimals

Isolation mass: (assuming separation of fr_H where f=10) (e.g. Lissauer 1987):

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

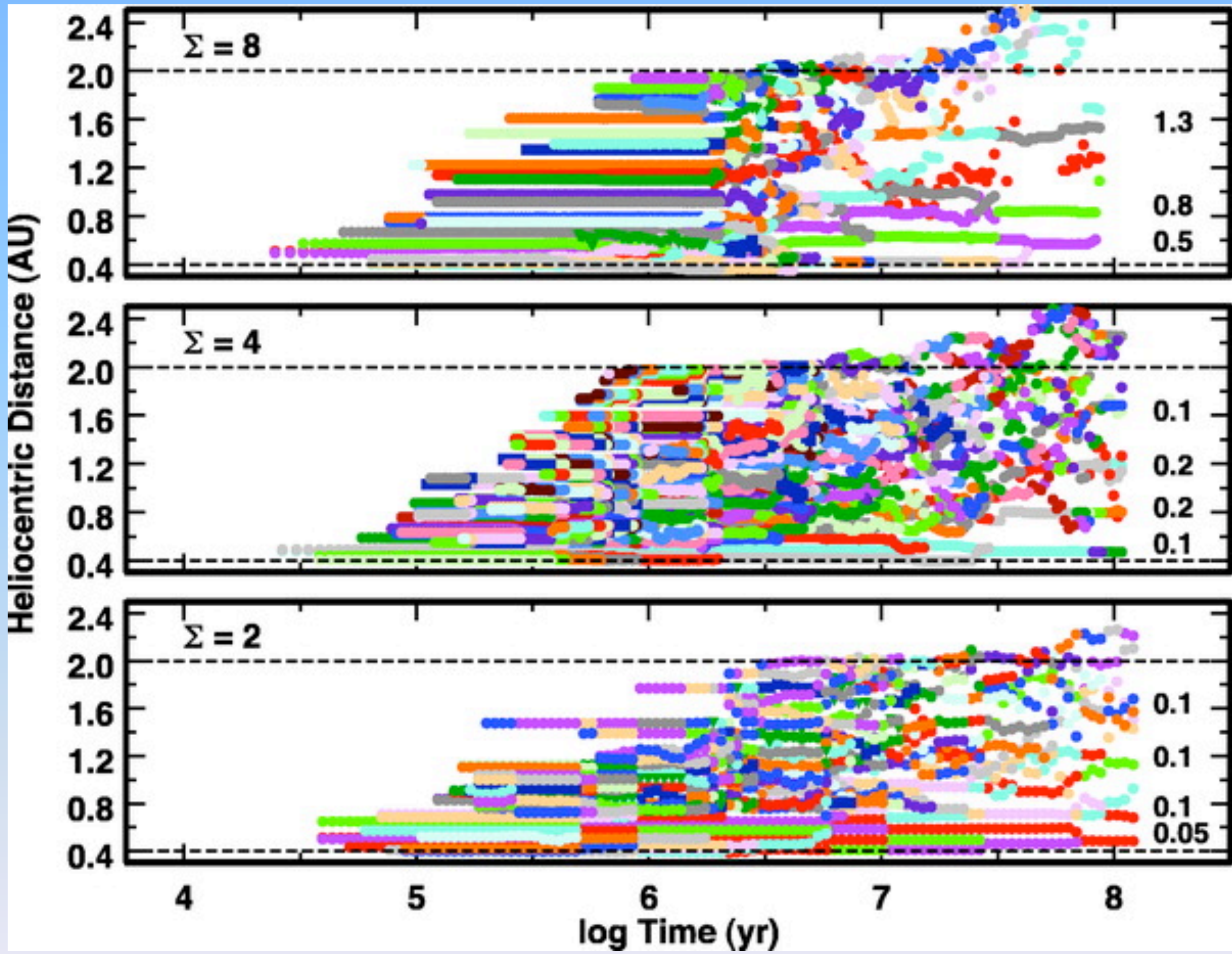
Increase in proto-planet eccentricity, then causes proto-planets to interact (Chambers & Wetherill 1998)

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



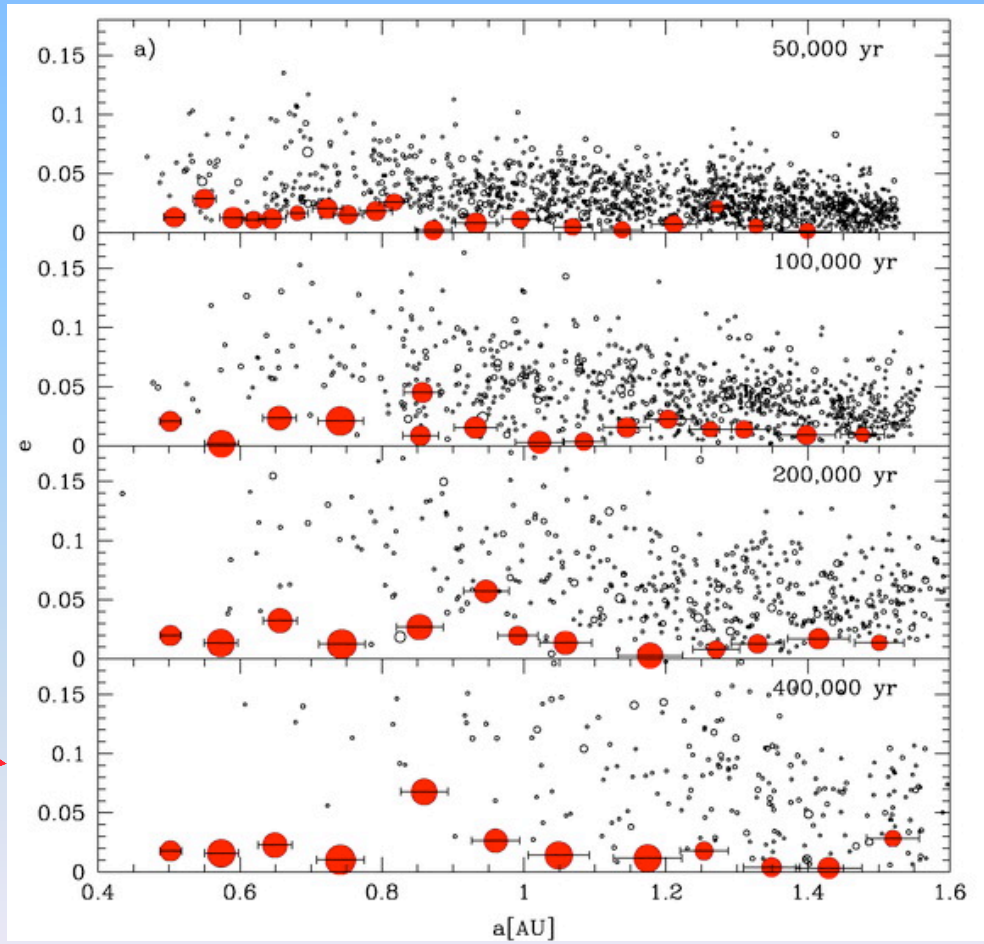
Chaotic growth leads to more mergers resulting in more massive planets; lower density disks form lower mass planets (Kenyon & Bromley 2006)

Role of destructive collisions

Role of small debris created in destructive collisions unclear:

• Analytical arguments of velocity dispersion evolution suggest that small body population significantly damps eccentricities (Goldreich, Lithwick & Sari 2004)

• N-body simulations get not much debris after oligarchic growth both with (Leinhardt & Richardson 2005) and without fragmentation (Kokubo & Ida 2002)



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

Critical core mass:

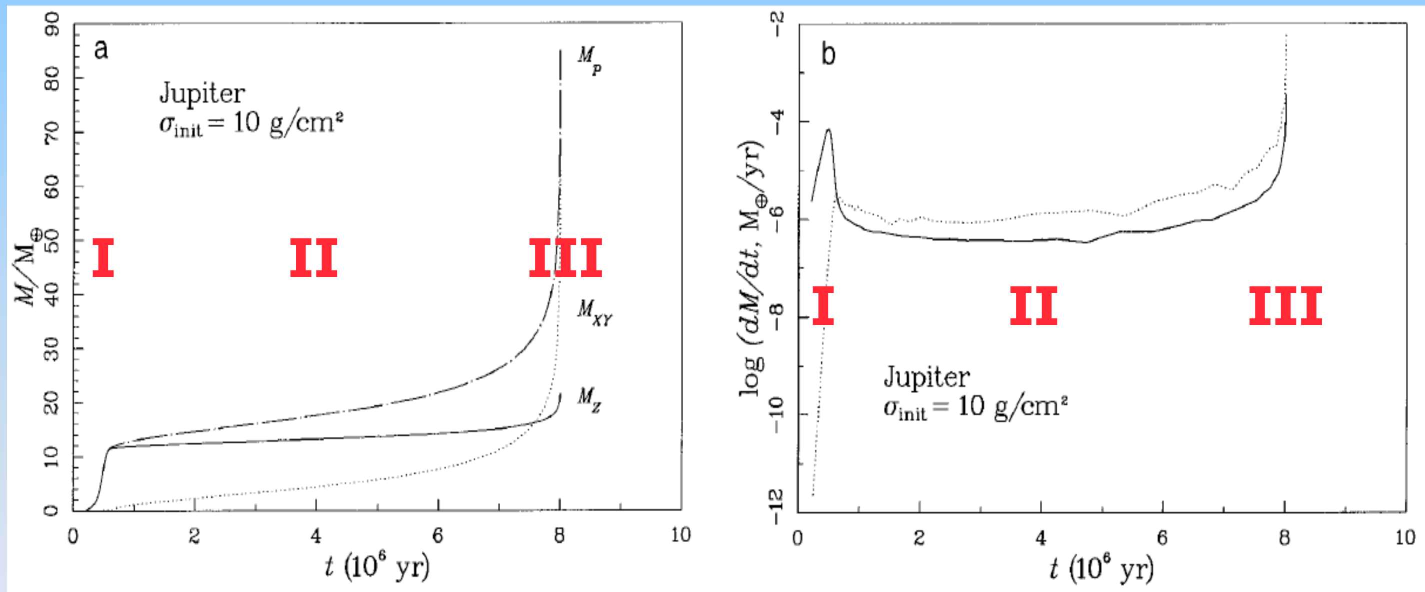
- core grows with atmosphere in quasi-static thermal equilibrium until at critical mass (~10 M_{earth}) when it rapidly accretes gaseous envelope
- final mass determined by available gas and how fast it can be accreted

Three main stages (Pollack et al. 1996):

(I) runaway growth to isolation

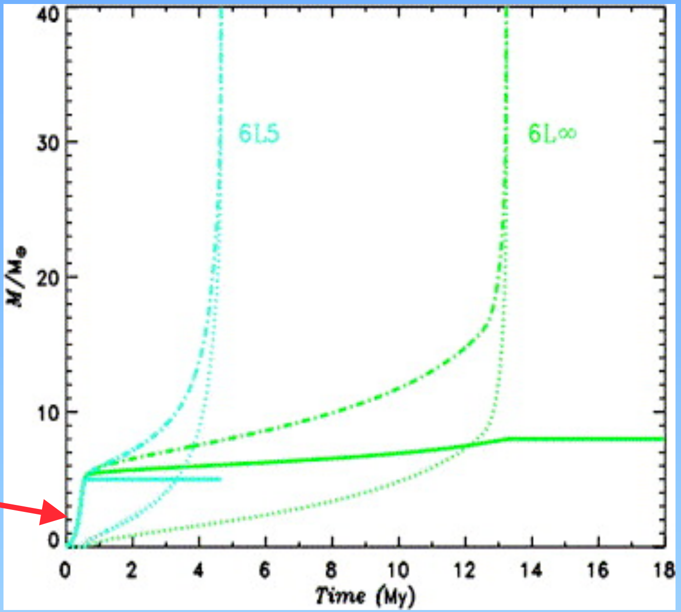
(II) small time independent accretion rates

(III) rapid accretion, when $M_{solid}=M_{gas}$ envelope contracts, outer boundary expands



Jupiter can form in 10Myr with core of 15M_{earth} if proto-solar nebula was a few times MMSN

Modifications to gas accretion

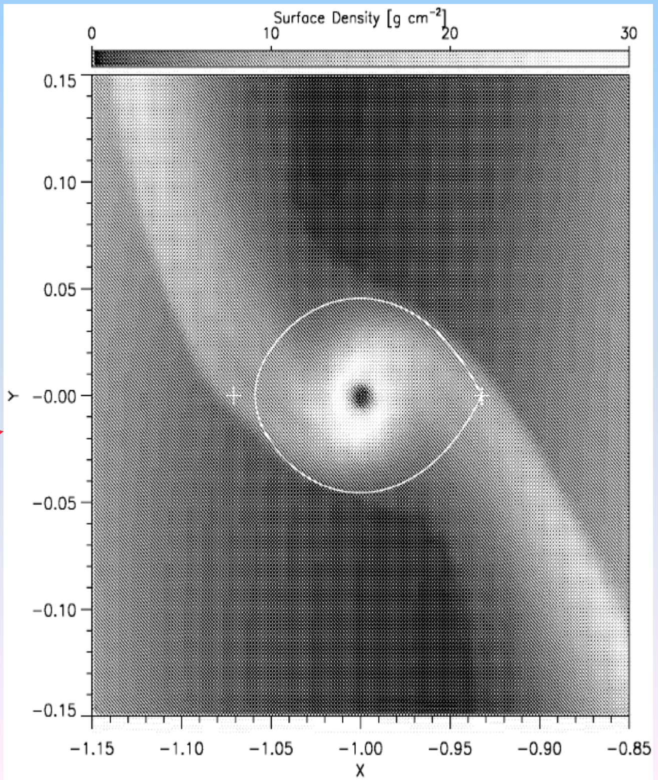


Motivation: low core mass of Jupiter, timescales longer than gas disk lifetimes

Opacity: reducing opacities to 2% ISM halves formation times (Hubickyj, Bodenheimer & Lissauer 2005; Papaloizou & Nelson 2005)

Stopping planetesimal accretion: helps runaway envelope (if core mass already large)

Now easy to form Jupiter in 5Myr with $5M_{earth}$ core



However, planet-disk interactions important:

- Non-axisymmetric, shocked flows (Lubow et al. 1999) and circumplanetary disk (Bate et al. 2003; Machida et al. 2008)
- Flow through disk gap (Lubow & D'Angelo 2005)
- Thermodynamics (Klahr & Kley 2006)
- Dust accretion (Paardekooper & Mellema 2006)

Planet migration

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

- can form in situ (Bodenheimer et al. 2000)
- and in scattering between planets (Weidenschilling & Marzari 1996)

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)

Planet migration: type I

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

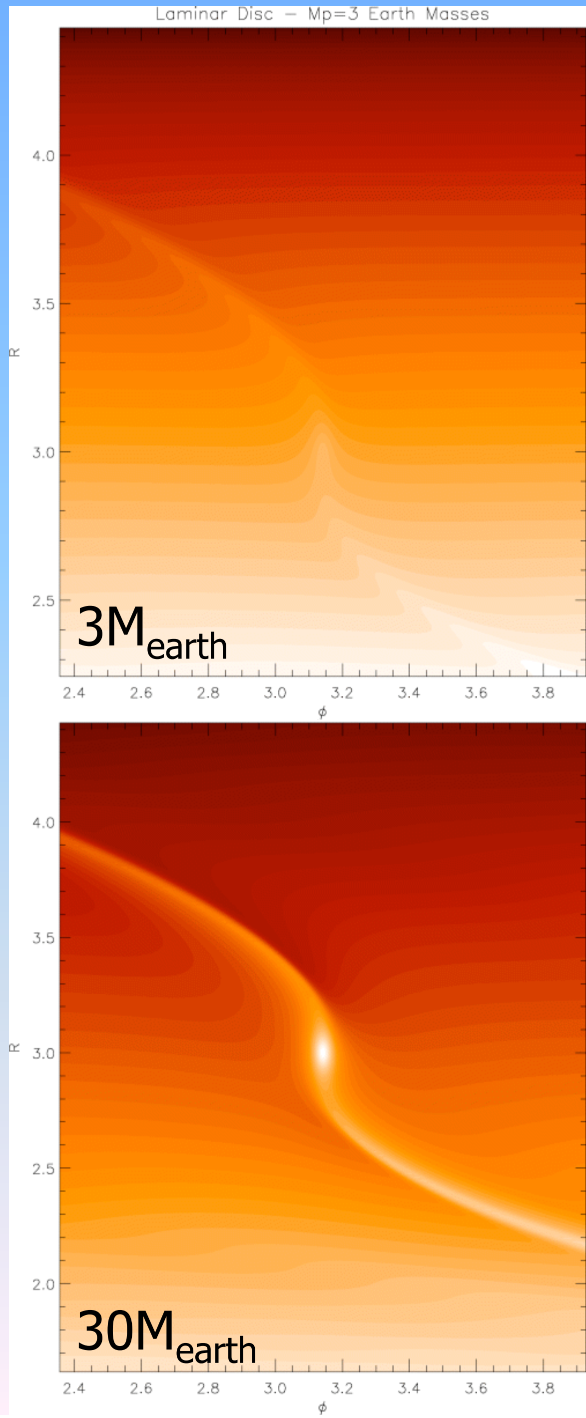
- waves interior to the planet exert positive torques
- exterior waves exert negative torques

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):

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

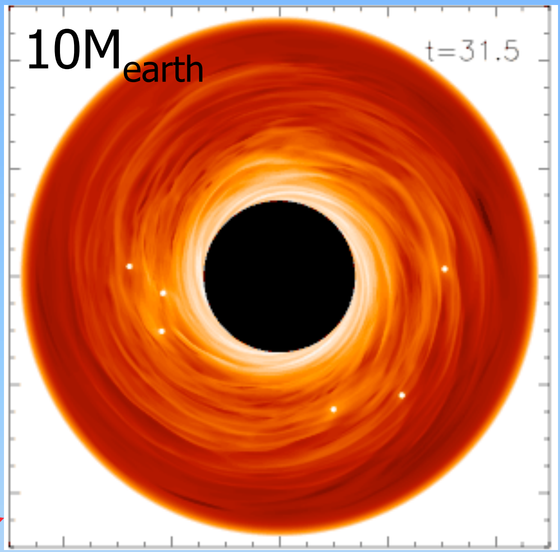


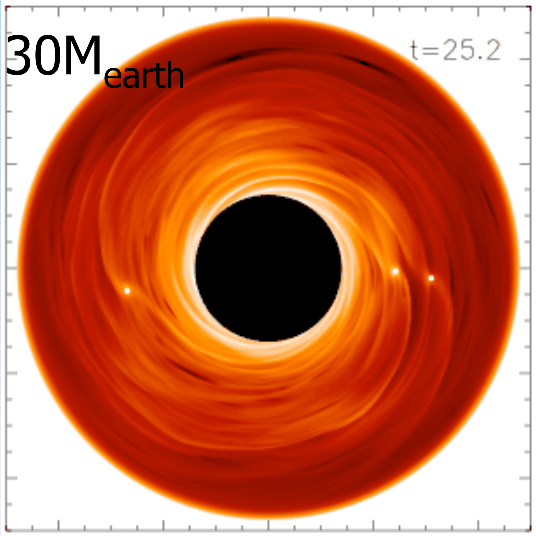
Why don't all planets migrate in?

Short migration times pose question: why don't all planets migrate in before they can accrete gas?

Several solutions to this problem:

- migration aids growth (Tanaka & Ida 1999; Alibert et al. 2005)
- turbulence slows migration (Nelson et al. 2005)
- planetesimal disk torque no help (Kominami, Tanaka & Ida 2005)
- magnetic fields stop migration (Fromang, Terquem & Nelson 2005)
- jump in surface density halts migration (Masset et al. 2006)





Planet migration: gap opening and type II

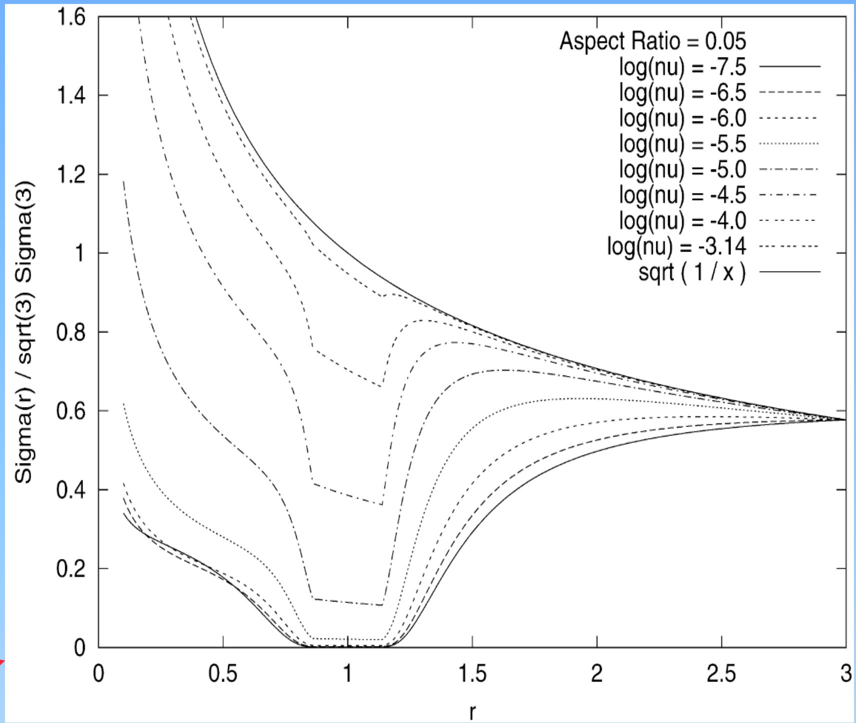
Linearity breaks down when $M_{pl}/M_{*}>(H/r)^3$ which is $\sim 30M_{earth}$

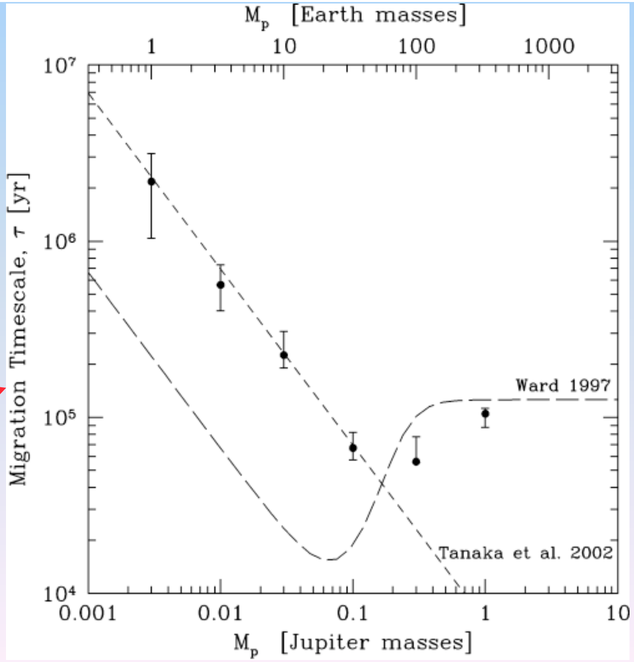
Gap opening and gap structure depends on: planet mass, disk height, and viscosity (Crida, Morbidelli & Masset 2006; Rafikov 2002; Edgar & Quillen 2007)

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

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

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





Planet migration: type II

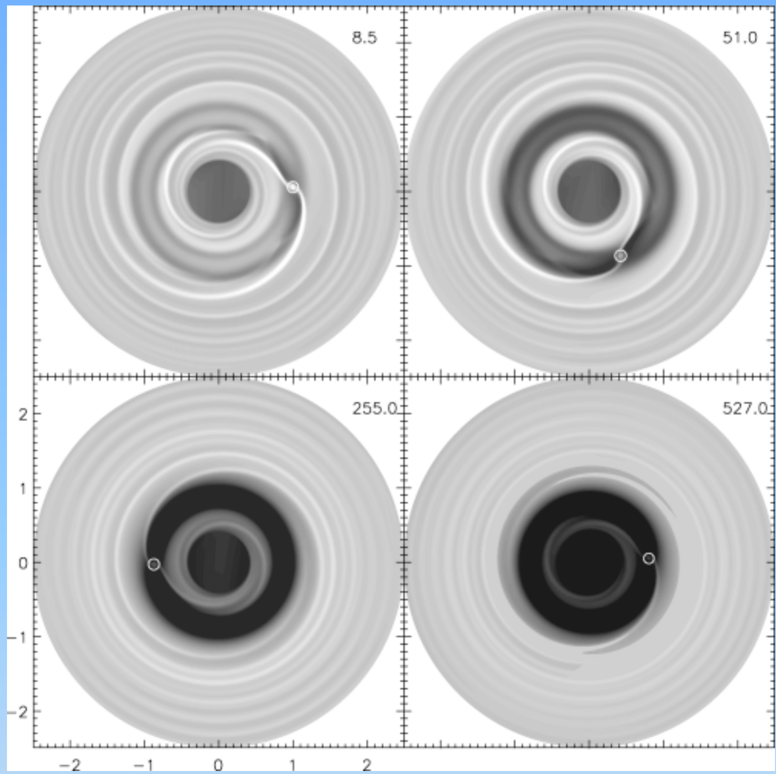
Planet migrates in on viscous timescale regardless of whether it is accreting (10,000 orbital periods, Nelson et al. 2000):

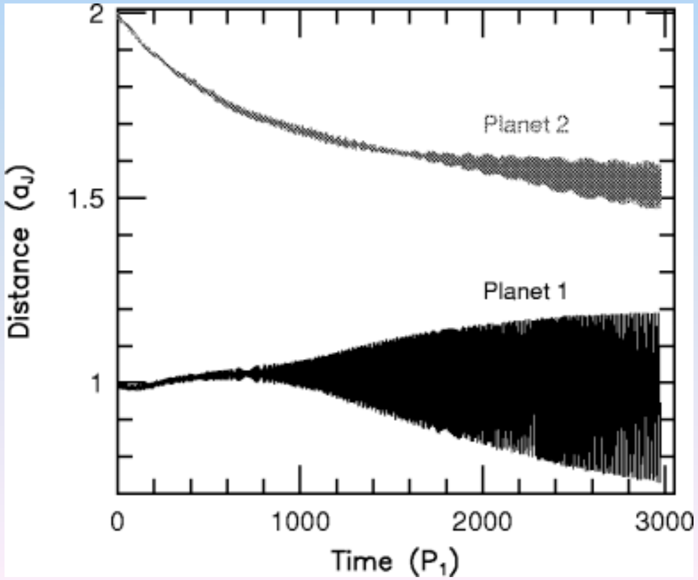
$$dr/dt = -1.5\nu/r$$

although if planet is more massive than the disk its inertia can slow down the migration

This is too fast, so need mechanisms for slowing down and stopping migration (Kuchner & Lecar 2002):

- accreting matter on the way in (Alibert et al. 2005)
- stop in region with low viscosity (with no MRI)
- due to multiple planets clearing (Kley 2000)
- migrate out if e_{pl}>0.2 (D'Angelo, Lubow & Bate 2002)
- trapping in resonance (Morbidelli & Crida 2007)



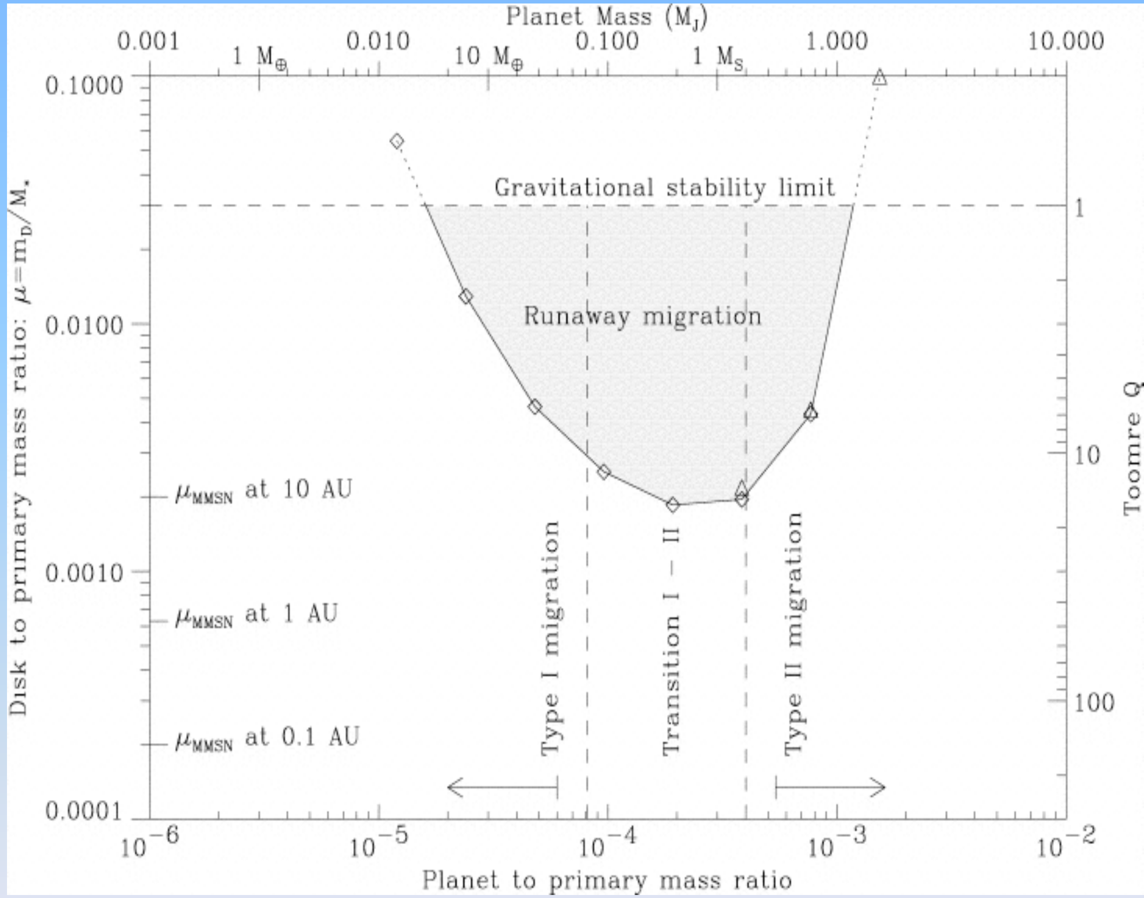


Planet migration: type III

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)

Runaway because magnitude of torque depends on migration rate



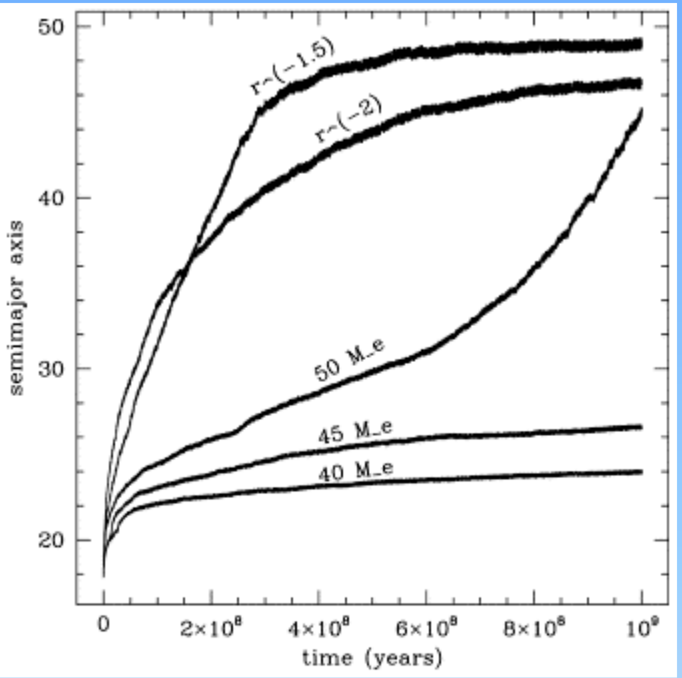
This result may be a numerical effect, since it is not reproduced in higher resolution simulations (D'Angelo, Bate & Lubow 2005) but still discussed (Peplinski et al. 2007)

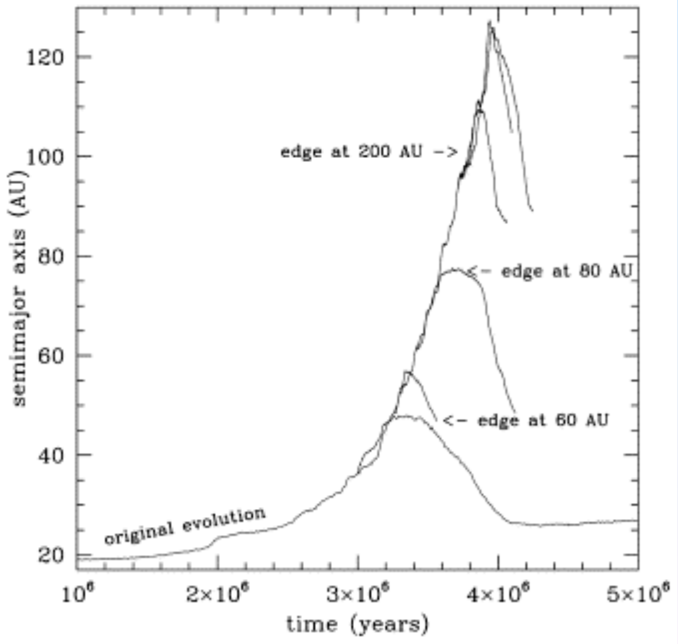
Planet migration: type IV

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 has been studied for Kuiper belt structure (e.g., Hahn & Malhotra 1999)

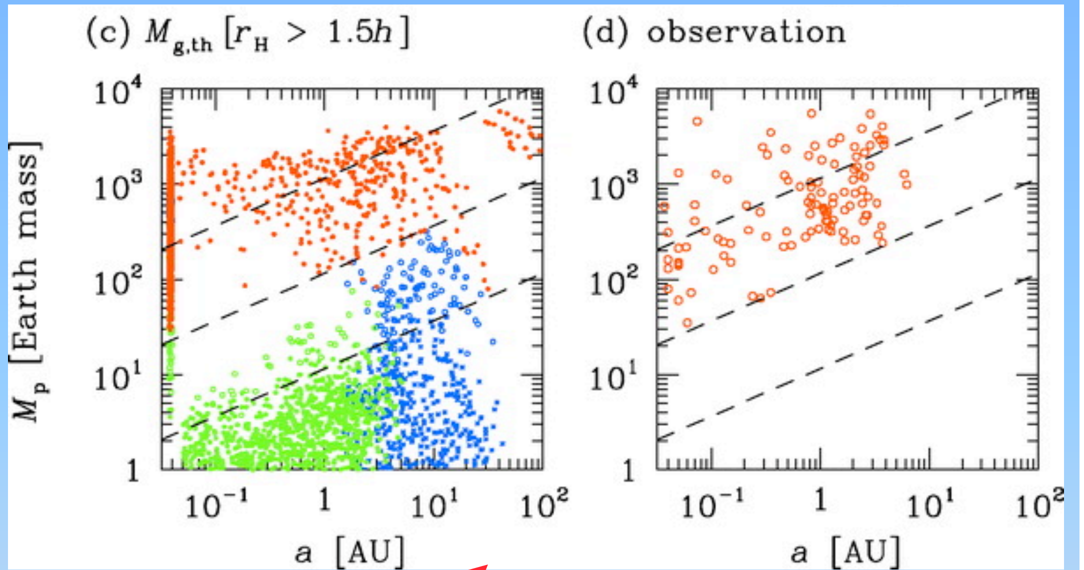
Generalised more recently (Gomes et al. 2004):
• migration speeds up in massive disk
• migration reversed when planet encounters the outer edge of planetesimal disk





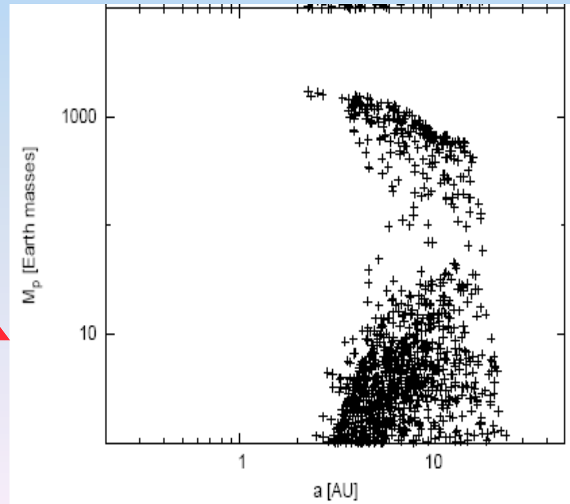
Formation+migration models

Accretion with type I migration: (McNeil, Levison & Duncan 2005) forming Earth possible with enhanced proto-stellar disk, but planet separation/mass are high ($20r_H$, $0.4M_{earth}$) and large planetesimal population (but see Alibert et al. 2005 and Ida & Lin 2008)



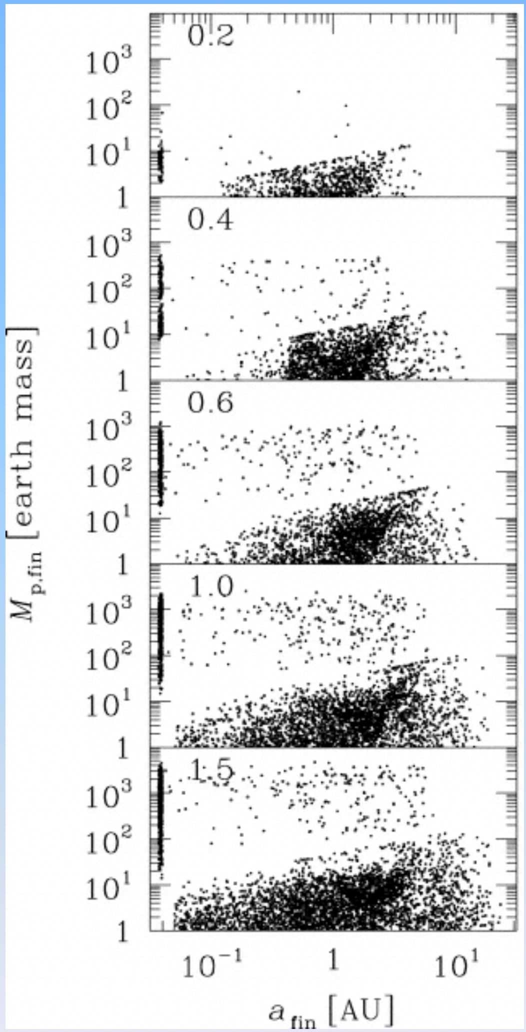
Core growth, envelope accretion and type II migration:

- (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
- Kornet & Wolf (2006) found more massive planets migrate easier, but didn't include disk mass distribution and did include gas accretion after gap opening and different H/r function

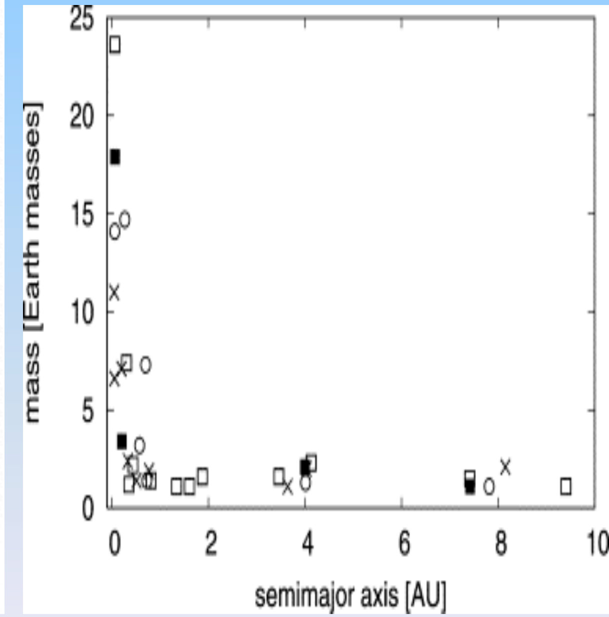


Formation+migration models: Hot Neptunes

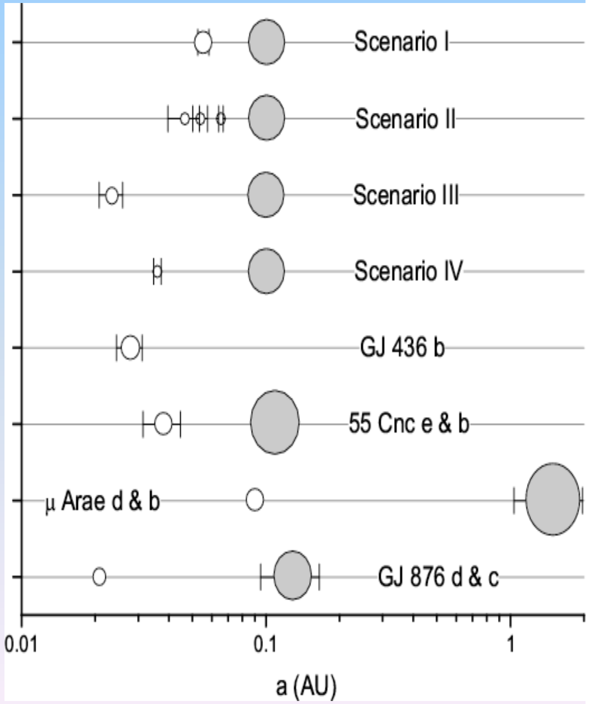
(1) M stars should have hot Neptunes since migration before rapid gas accretion (Ida & Lin 2005)



(2) N-body of core accretion with type I migration predicts rocky/icy hot Neptunes (Brunini & Cionco 2005)



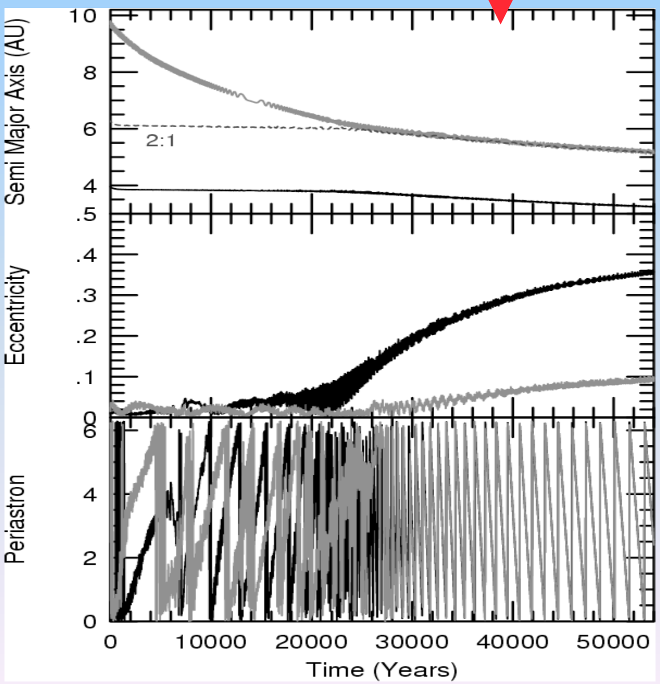
(3) Jupiter migrating by type II shepherds planetesimals interior to the planet which accrete into Hot Neptunes (Fogg & Nelson 2005; Mandell et al. 2007)

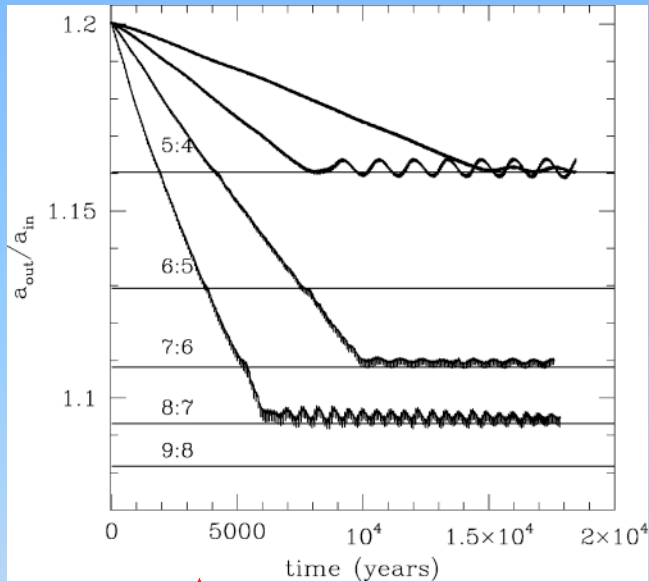


(4) Subsequent evolution of irradiated planet shows Hot Neptunes could be depleted Jupiters (Baraffe et al. 2006)

Planet migration with multiple planets: resonances

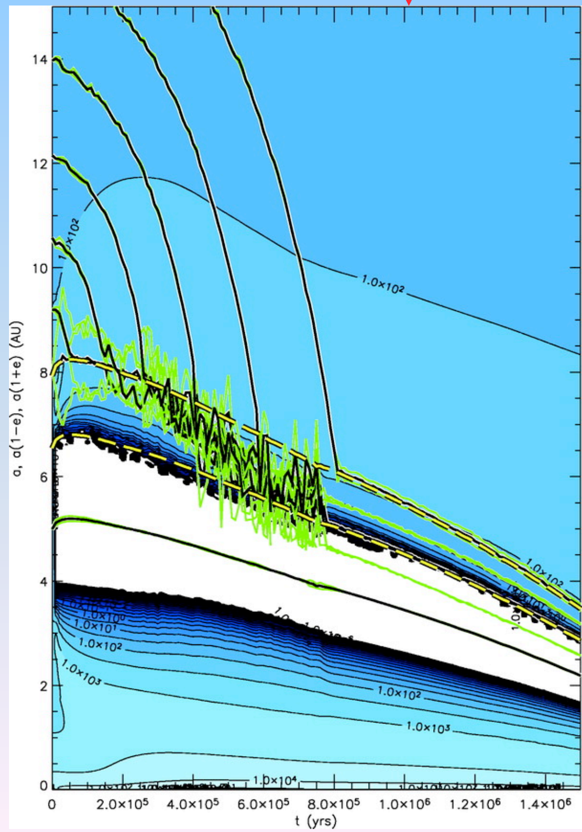
Hydrodynamic+N-body simulations of migration of two Jupiter-mass planet systems give similar results to N-body models with dissipation
(Kley, Peitz & Bryden 2003)





Earth-mass planets with type I migration trapped in first order resonances (7:6 etc) (Papaloizou & Szuszkiewicz 2005) but may be lost following circularisation making hot Neptunes
(Terquem & Papaloizou 2007)

Proto-planets forming outside Jupiter which clears a gap quickly migrate into 3:2 and 2:1 resonances
(Thommes 2005)

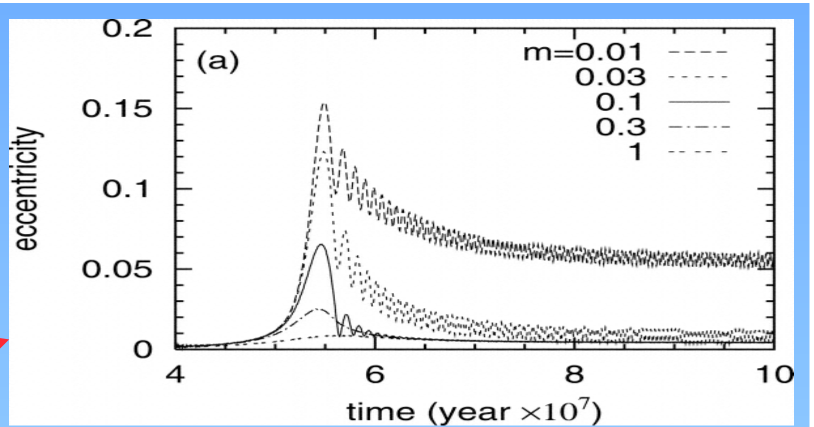


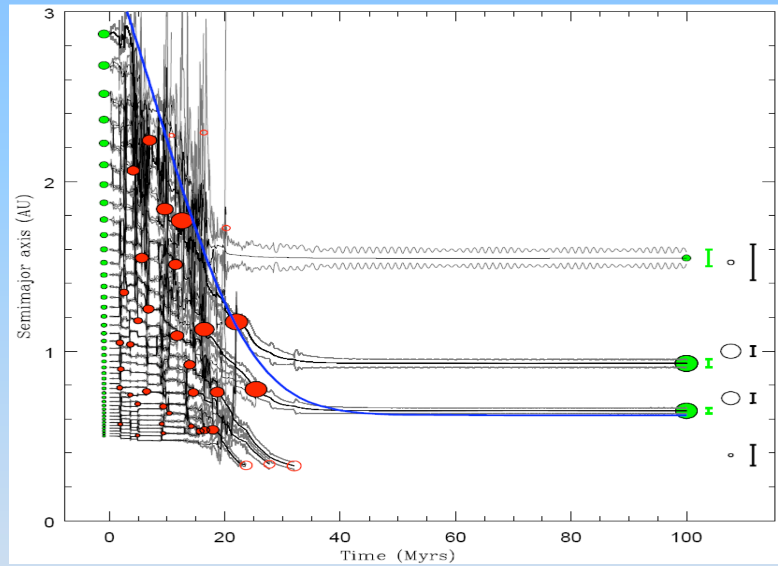
Role of secular resonances

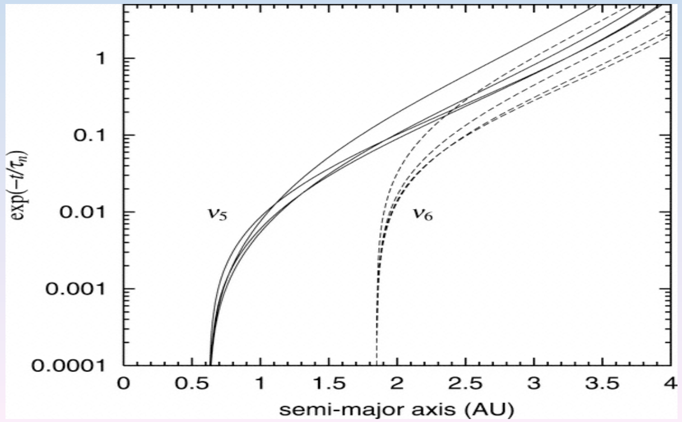
Formation of Jupiters far out affects growth of terrestrial planets without migration

- Secular perturbations excite planetesimal eccentricities while gas drag damps them, balance causes proto-planets to migrate in with secular resonance (Nakagawa et al. 2005) reproducing low e,I of terrestrial planets (Thommes et al. 2008)

- Secular resonances move as the gas disk dissipates = **secular resonance sweeping**, application to solar system sets constraints on nebula removal time (Ward 1981) and may clear asteroid belt (Lecar & Franklin 1997)



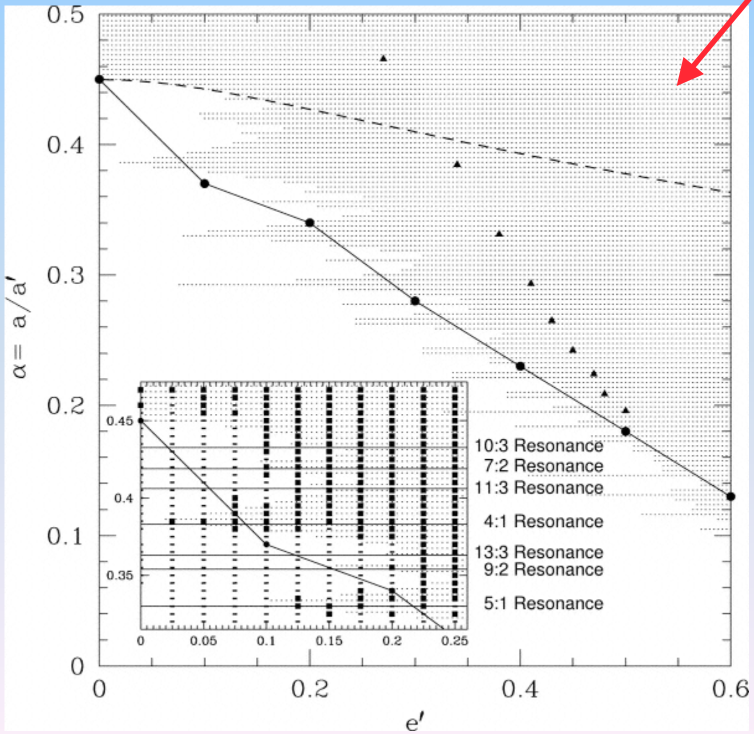


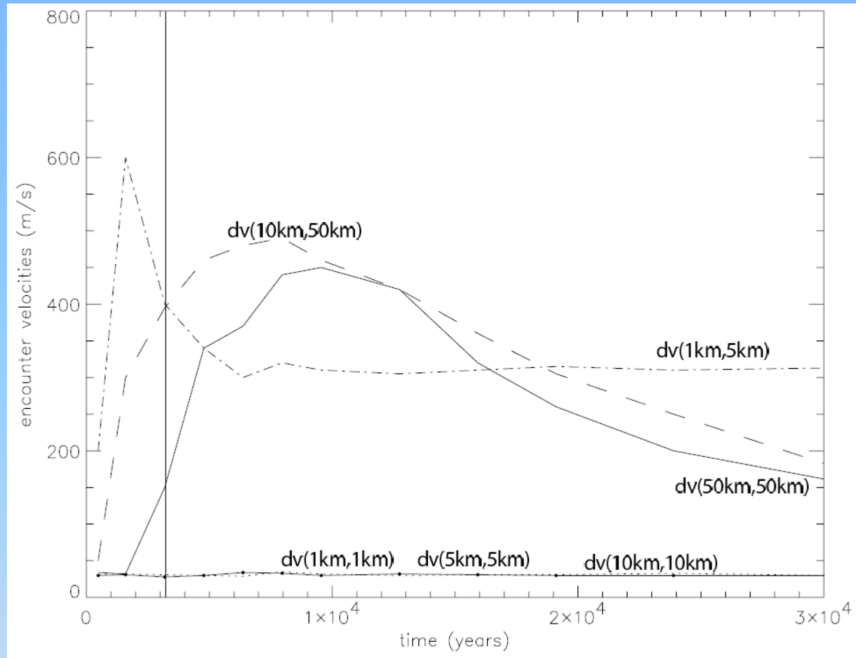


Secular perturbations: core accretion in binary systems

The secular effect of a binary companion affects planet formation:

• Resonance overlap means binary companion clears material close to its orbit (Holman & Wiegert 1999; Mudryk & Wu 2006)



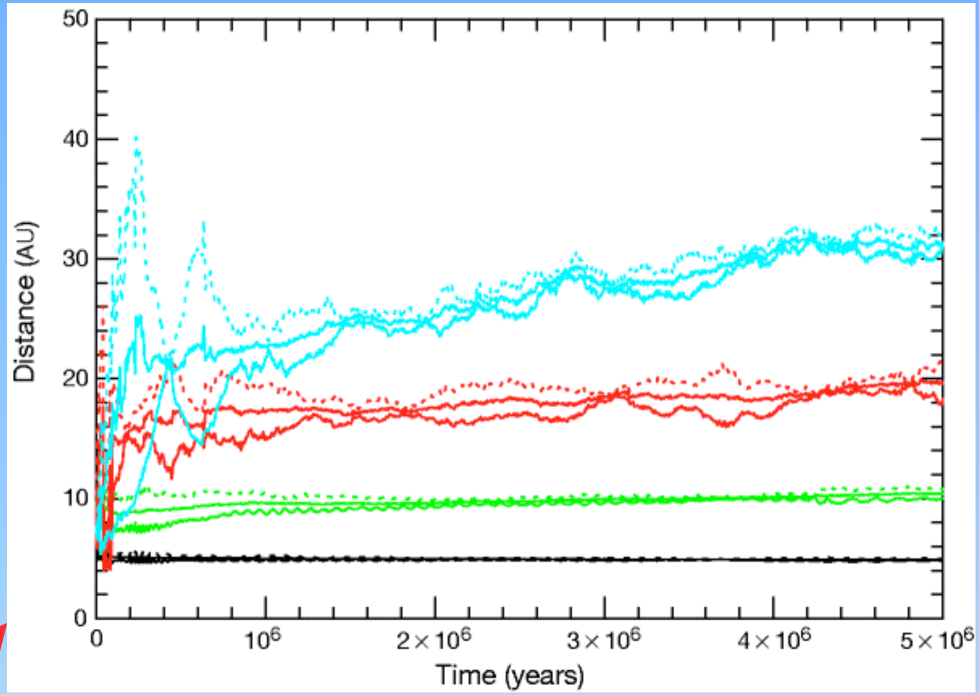


• Close in orbits stable (Quintana et al. 2007), but secular perturbations and gas drag mean collisions between similar size objects have low velocity leading to runaway growth (Kortenkamp, Wetherill & Inaba 2001; Thebault, Marzari & Scholl 2006), although gas disk eccentricity may prevent growth (Paardekooper et al. 2008)

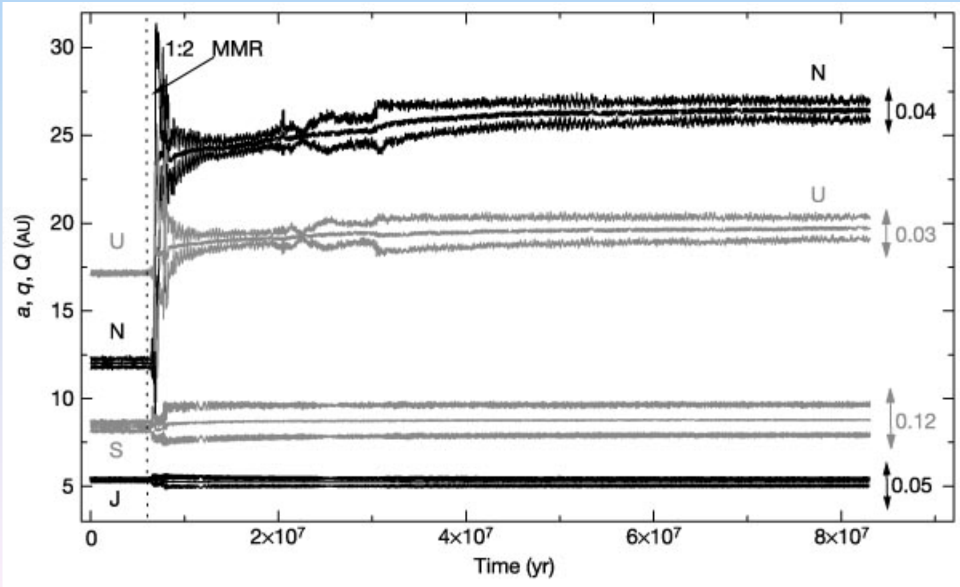
Chaotic evolution

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

• Uranus and Neptune could be cores formed between Jupiter and Saturn, later flung out to interact with the primordial Kuiper belt (Thommes et al. 1999)



• Slow type IV migration could have caused Jupiter and Saturn to cross 2:1 resonance pumping up eccentricities of UN (Tsiganis et al. 2005)



PPD properties: snowline

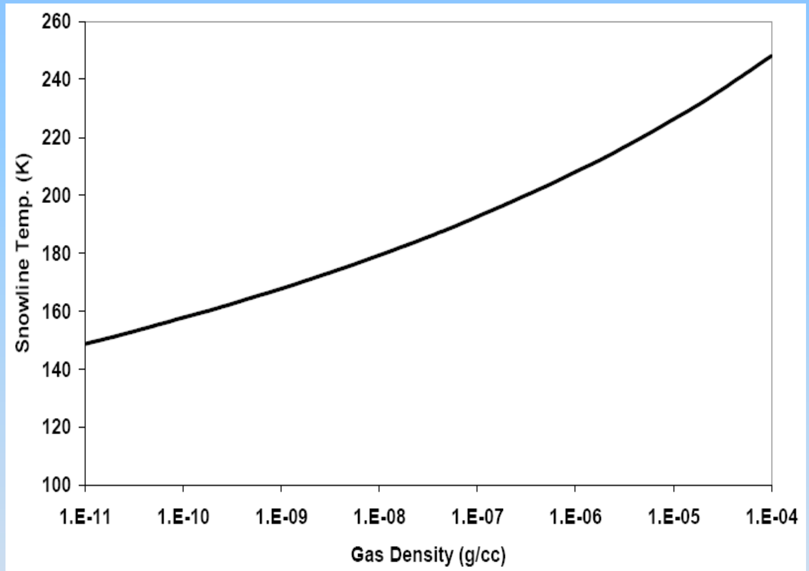
Solid surface density Σ_d jumps by x4 at snowline where ices condense, and since isolation mass $\propto \Sigma_d^{1.5}$ cores of gas giants thought to form there

Solar system:

- snowline at ~2.7AU from abundance of icy C-class asteroids (Rivkin et al. 2002) and presence of water on asteroids (Hsieh & Jewitt 2006)

Theory:

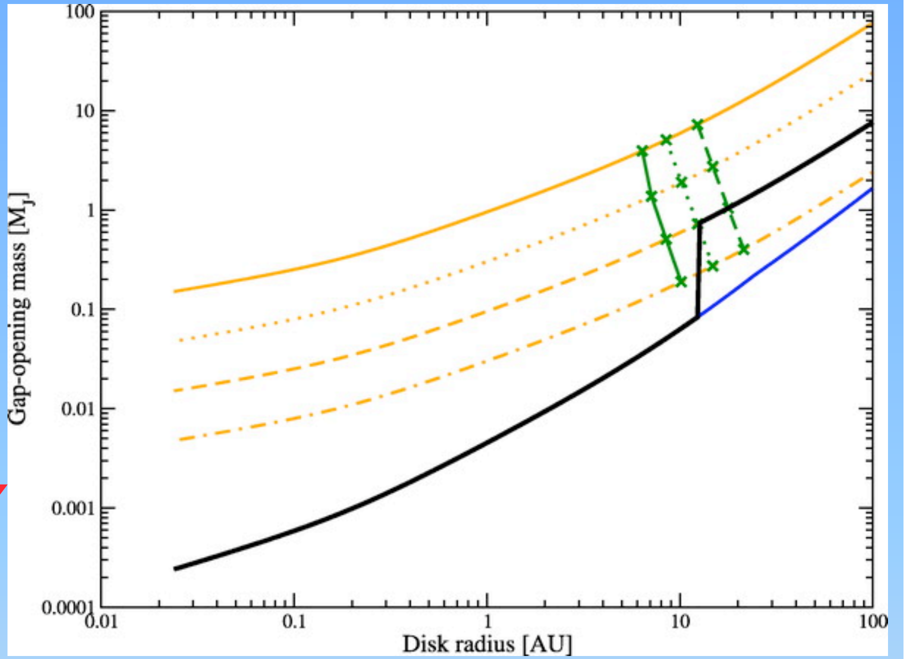
- when T<145-170K depending on partial pressure of water vapour (Podolak & Zucker 2004), putting snowline at 1.6-1.8AU in solar system (Lecar et al. 2006)



- Increasing grain opacities, including heating by ^{26}Al, and full coagulation/ settling models push snowline out (Grimm & MacSween 1993; Kornet, Rozyczka & Stepinski 2004)
- Snowline moves in during PMS evolution aiding planet formation (Kennedy et al. 2006)
- Effect on dead zone may enhance planet formation there (Ida & Lin 2008b)

PPD properties: dead zones

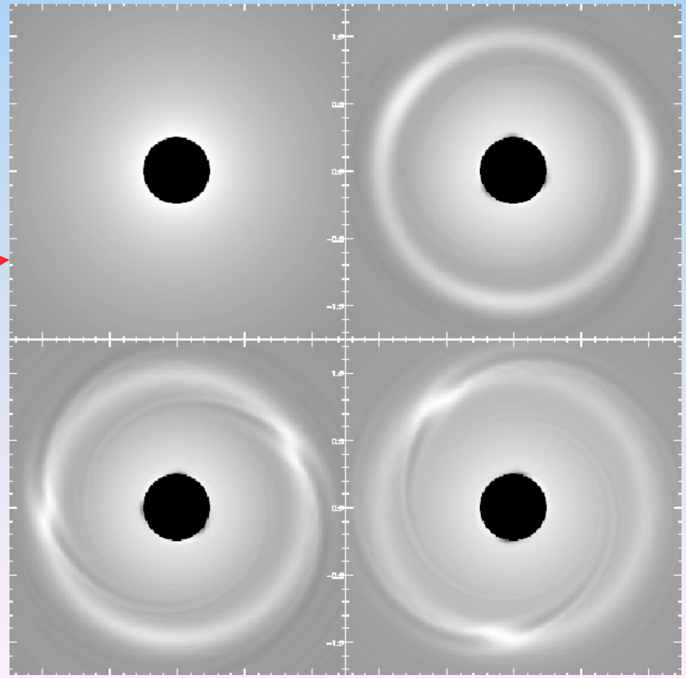
Dead zone: disk region (<12AU) is poorly ionised (Turner et al. 2007) and so growth of magneto-rotational instability (MRI, Balbus & Hawley 1991) against ohmic dissipation cannot be sustained (Gammie 1996) leading to low viscosity causing:



• gap opening at low planet mass (Matsumura & Pudtritz 2005)

• long type I & II migration times (Thommes 2005; Chiang, Fischer & Thommes 2002; Matsumura et al. 2006)

• mass pile-up (Morbidelli et al. 2007) promoting GI or Rossby Wave Instab (Varniere & Tagger 2006)



• high eccentricity from large gap (Matsumura & Pudritz 2006)

• decrease in active layer thickness causes pressure maximum halting type I migration (Ida & Lin 2008b)

Gravitational instability model

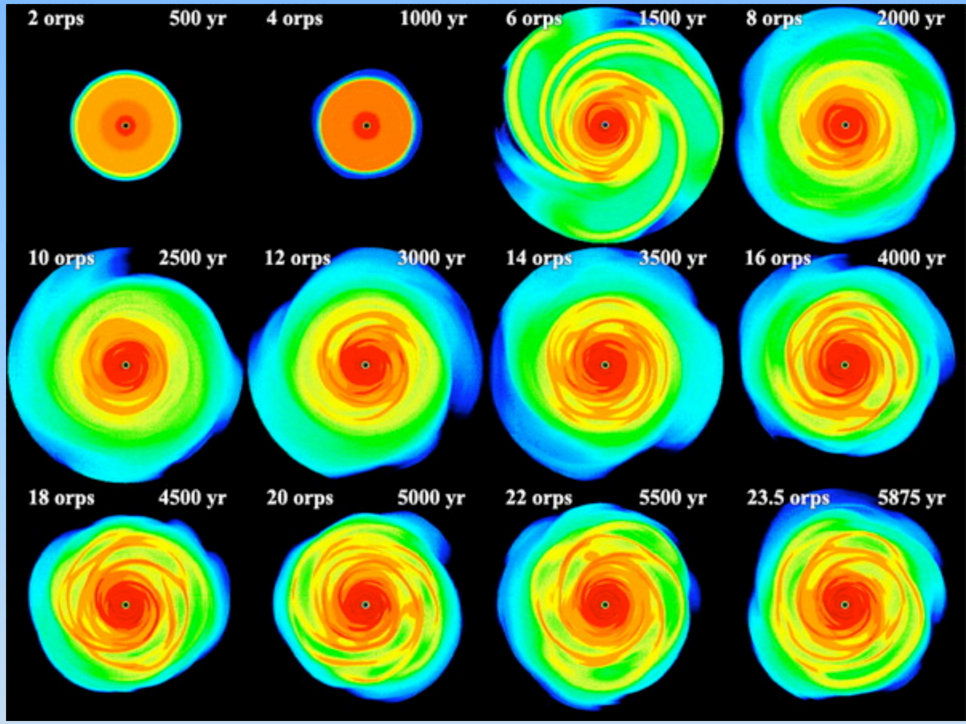
Gravitational instability:
Planets form on orbital timescales when part of disk becomes unstable (Kuiper 1949, Cameron 1978):

$Q \sim M_{star}H/(M_d r) < 1$

Characteristic size is H and so mass $\sim M_{jupiter}$ (assuming $H/r \sim 0.1$)

Could a collapsing cloud result in unstable disk?

- Disk builds up mass from envelope (decreasing Q)
- Non-axisymmetric spiral modes develop when Q approaches 1 (Laughlin & Bodenheimer 1994) leading to angular momentum transport on orbital timescales



Q never reaches 1 (Vorobyov & Basu 2007) unless the disk is cooled (so H/r decreases) or matter added (so M_d increases) quicker than orbital timescales ($\tau_c < 3\Omega_k^{-1}$) (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, although see Cai et al. 2006)

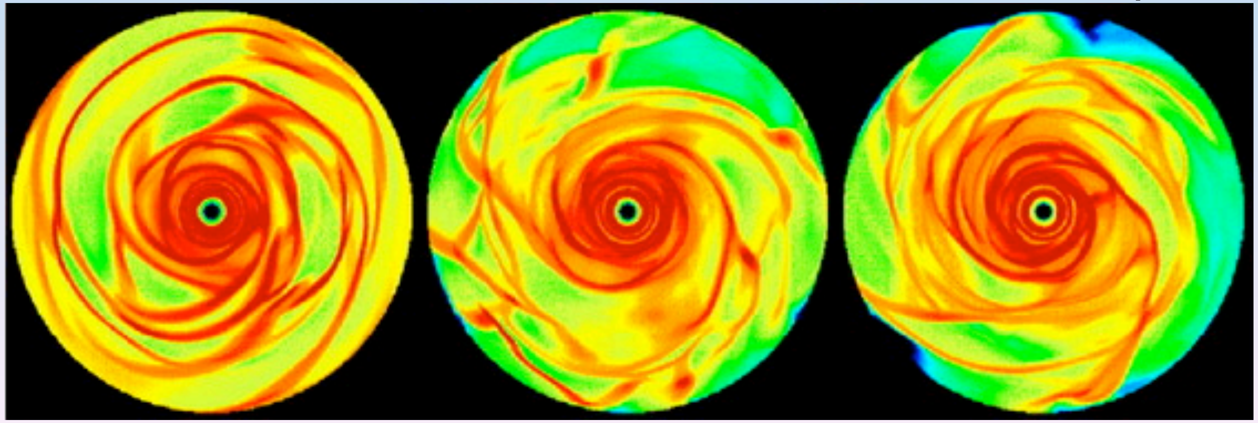
• Disks forming planets by GI at <10AU would be uncommonly luminous, so only $10M_{Jupiter}$ planets at ~100AU by GI are possible (Rafikov 2005)

• Gas giants difficult to form at 100-200AU by GI due to rapid inward mass transport by spiral arms (Boss 2006)

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)



Gravitational instability model

Origin of cores of giant planets?

• Rock and ice cores form after planet through sedimentation (predicts 6 and 2 M_{earth} cores for Jupiter and Saturn) (Boss 1998), core expected to be mostly Si (Helled et al. 2008)

Dependence on metallicity

• Stellar metallicity does not affect planet formation by GI because disk radiative energy loss is controlled by star not disk radiation (Boss 2002)

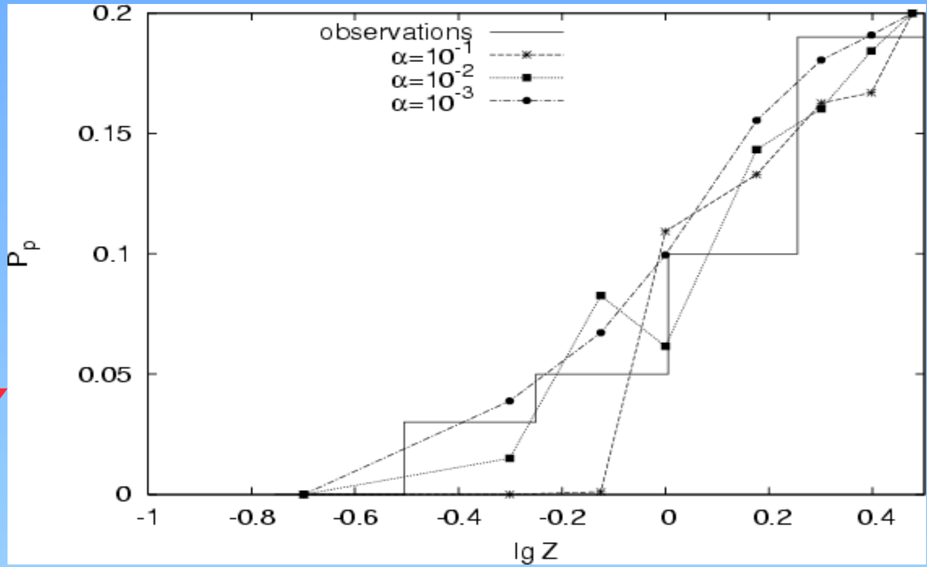
• In fact, cooling is faster with lower metallicity disks implying these are more likely to form planets (Cai et al. 2006)

• Although, planetesimal accretion after formation (Helled et al. 2006) and GI easier with high Z due to less compressional heating (Mayer et al. 2006)

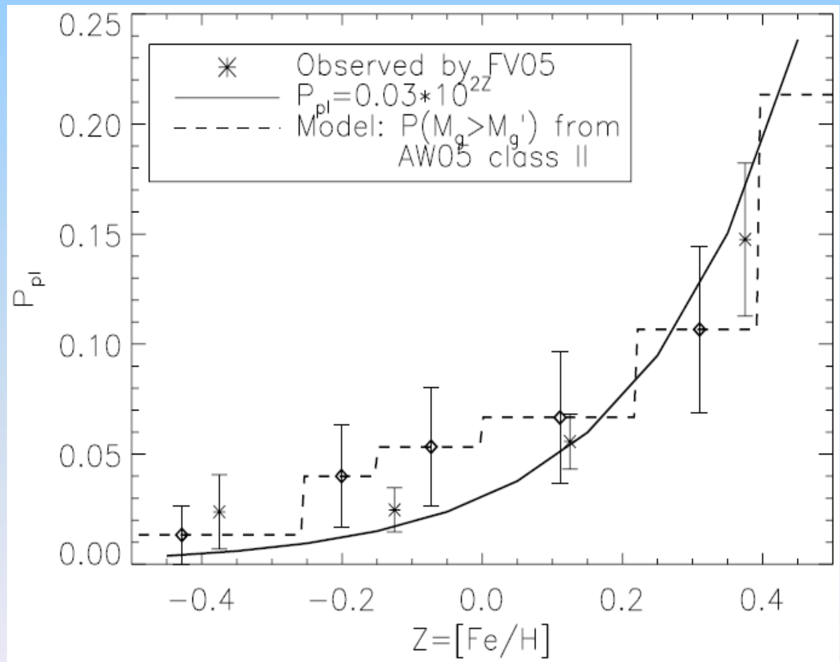
Dependence on stellar mass

• Low mass stars are equally likely to form planets (assuming they have equally massive disks) offering observational test (Boss 2006)

Metallicity distribution



- Metallicity (Z) dependence of planet hosts is proof of formation by core accretion, since faster growth predicted in higher Z disks because of the higher density of solids ($t_{growth} \propto \Sigma_d^{-1.5}$, Ida and Lin 2004; Kornet et al. 2005), also predicting steeper dependence for closer-in planets (Robinson et al. 2006) and lower planet mass around lower Z stars (Rice & Armitage 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)

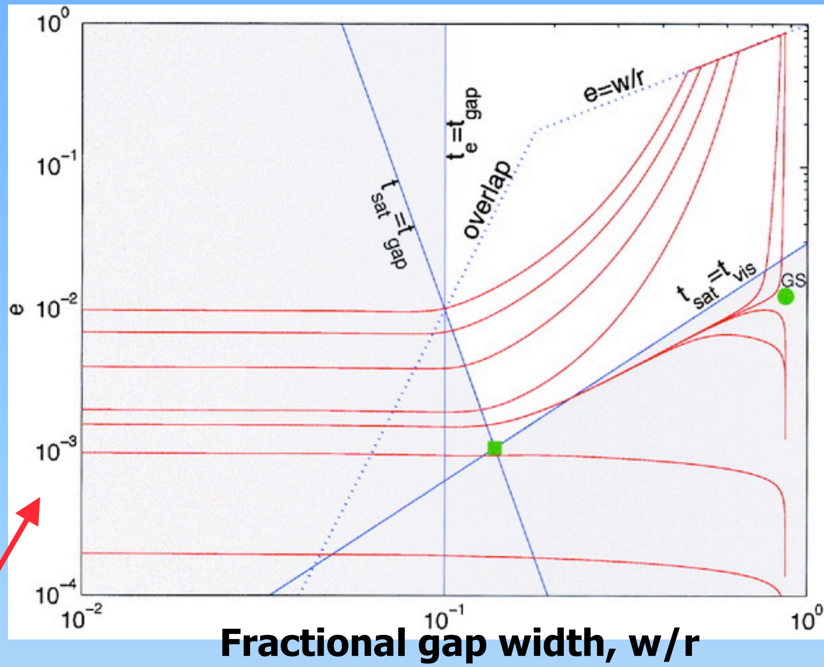
- NB metallicity dependence not caused by its effect on migration, since 10x metals speeds up migration by only 2x (Livio & Pringle 2003)

Eccentricity distribution

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

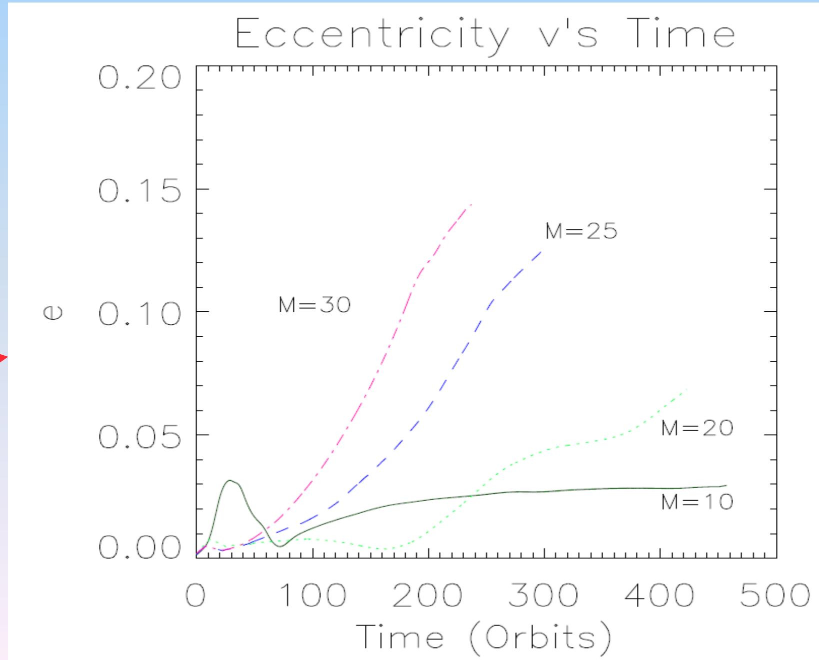
Planet-disk interaction

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; gap clearing can increase e_{pl} (Goldreich & Sari 2003; Sari & Goldreich 2004)



Simulations

- Back reaction damps e_{pl} and e_{pl}~0.2 when >$10M_{jupiter}$ (Papaloizou, Nelson & Masset 2001) since then gap encompasses 2:1 resonance
- Transition to eccentric >$3M_{jupiter}$ (Kley & Dirksen 2006)
- Transition at lower M_{pl} with dead zone (Matsumura & Pudritz 2006)



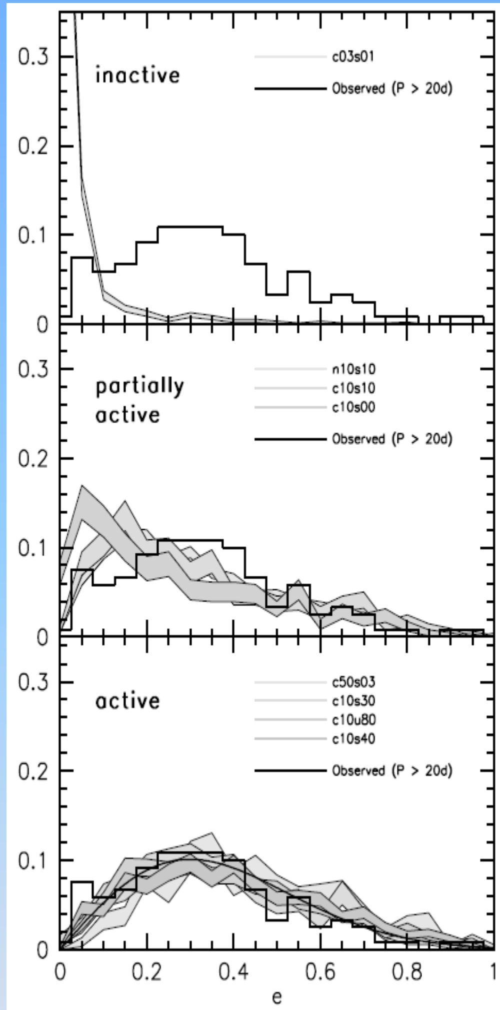
Eccentricity distribution

Planet-planet interaction

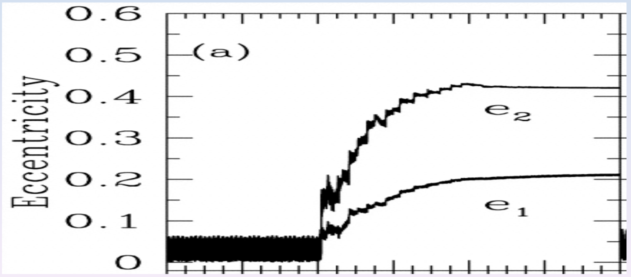
• Dynamical instability of 2 planets ejects outer planet leaving closer-in planet with high e_{pl} (Rasio & Ford 1996; Ford & Rasio 2007)

• **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)



• Migration of planets on diverging orbits causes **repeated resonance crossing** pumping e_{pl} (Chiang, Fischer & Thommes 2002)



• Passage through secular resonance pumps e_{pl} (Nagasawa Ida & Lin 2003)

Eccentricity distribution

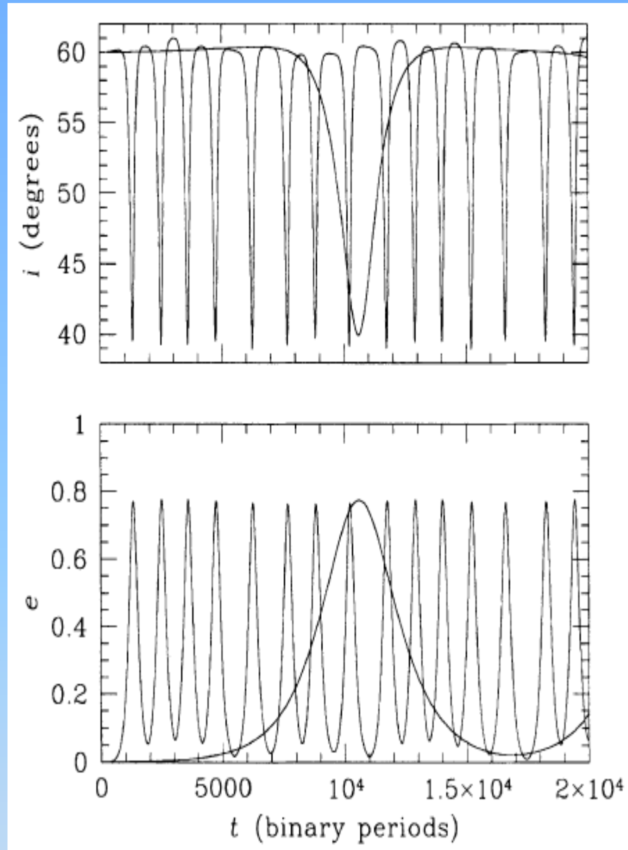
Binary formation/interaction

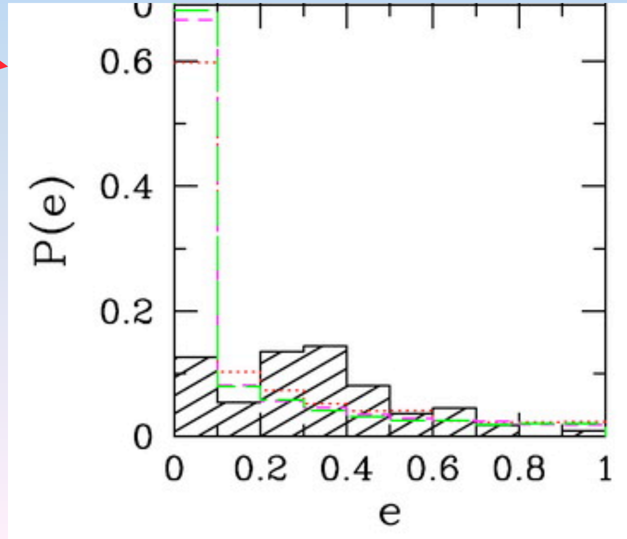
• Eccentricity distribution of exoplanets and spectroscopic binary stars (accounting for tidal circularisation) are different (Halbwachs, Mayor & Udry 2005)

• Binary star interactions could cause high e_{pl} from Kozai oscillations (Holman et al. 1997), however also produces low e_{pl} planets (Takeda & Rasio 2005)

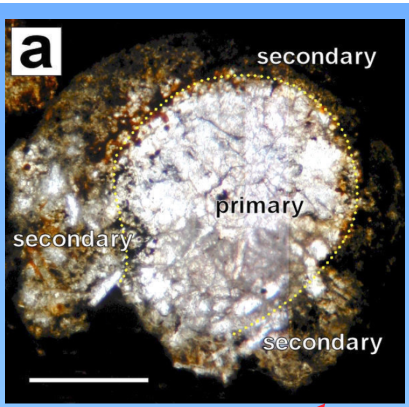
Other

• Constant acceleration applied to the star (but not the planet), such as caused by a precessing stellar jets or star-disk wind interactions (Namouni 2005)





Chondrule formation



Also clues from the solar system: how did chondrules form?

- Primitive chondritic meterites 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 1 hour 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, Boss & Durisen 2005; Sirono 2006; Miura & Nakamoto 2007),
  - young Sun processes (Fiegelson et al. 2002),
  - giant (Krot et al. 2005) or small (Miura et al. 2007) imapcts

- Indicates we don 't know details of formation processes (Cuzzi et al. 2001)