Evolution of protoplanetary discs

and why it is important for planet formation

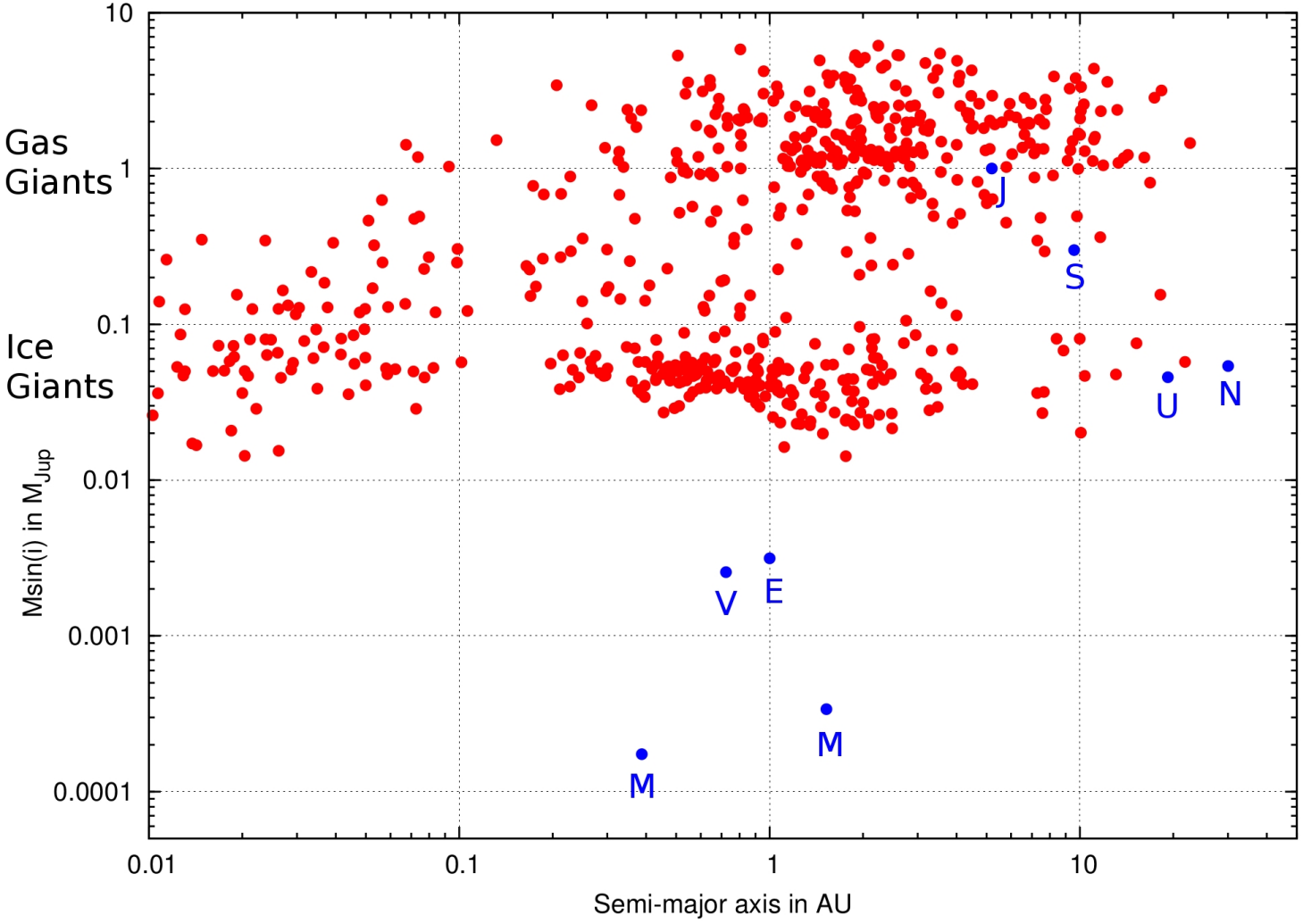
Bertram Bitsch

Lund Observatory

LUND UNIVERSITY

April 2015

Observations of planets



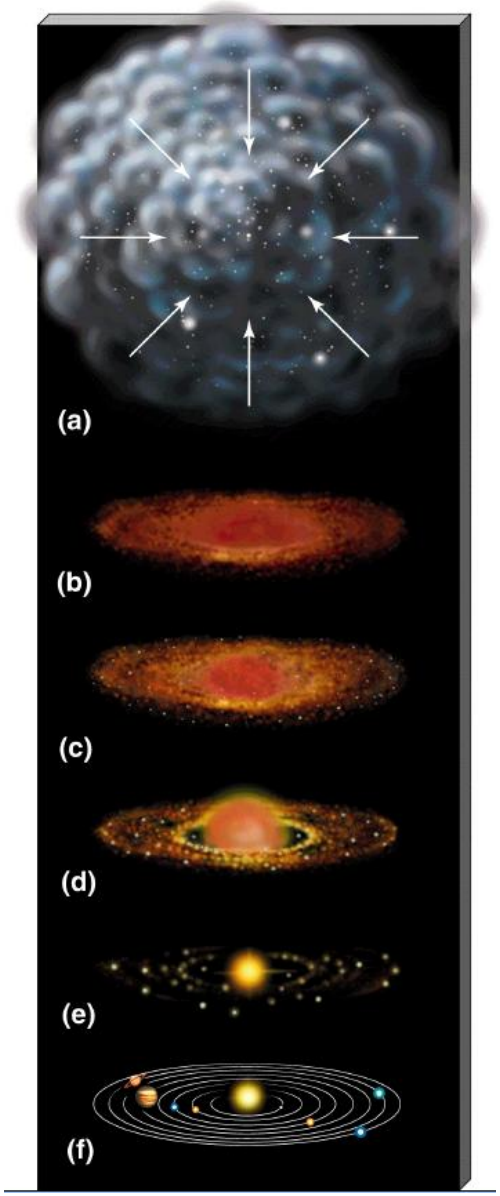
⇒ How can we explain this diversity?

Data from exoplanet.org

Outline

- Introduction
- Protoplanetary disc structure
- Planet growth and migration
- Planet formation in evolving protoplanetary discs

Planetary system evolution



(a) Collapse of interstellar cloud

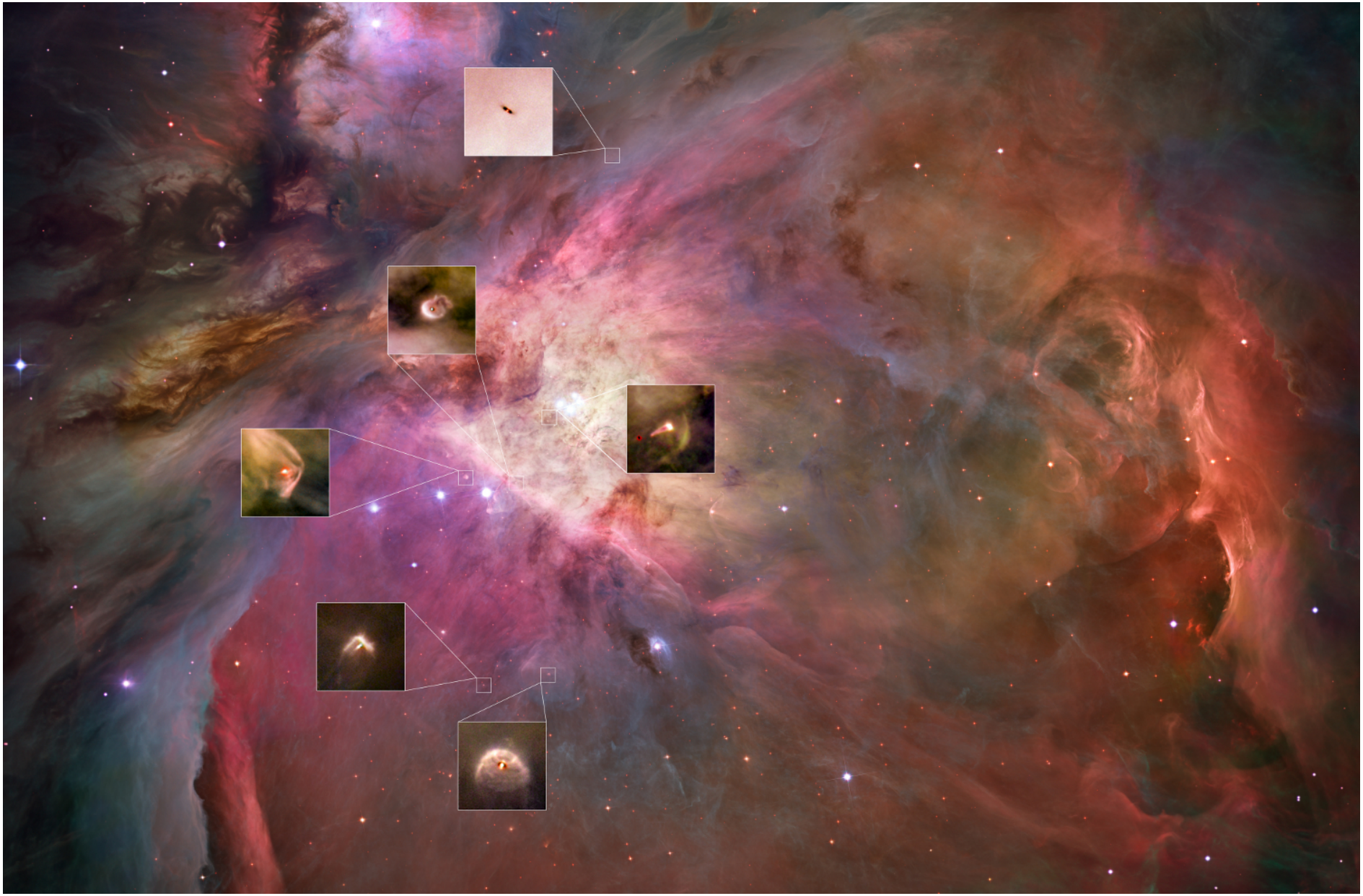
(b) Formation of protostar with protoplanetary disc made out of dust and gas

(c) Small particles stick and form bigger objects

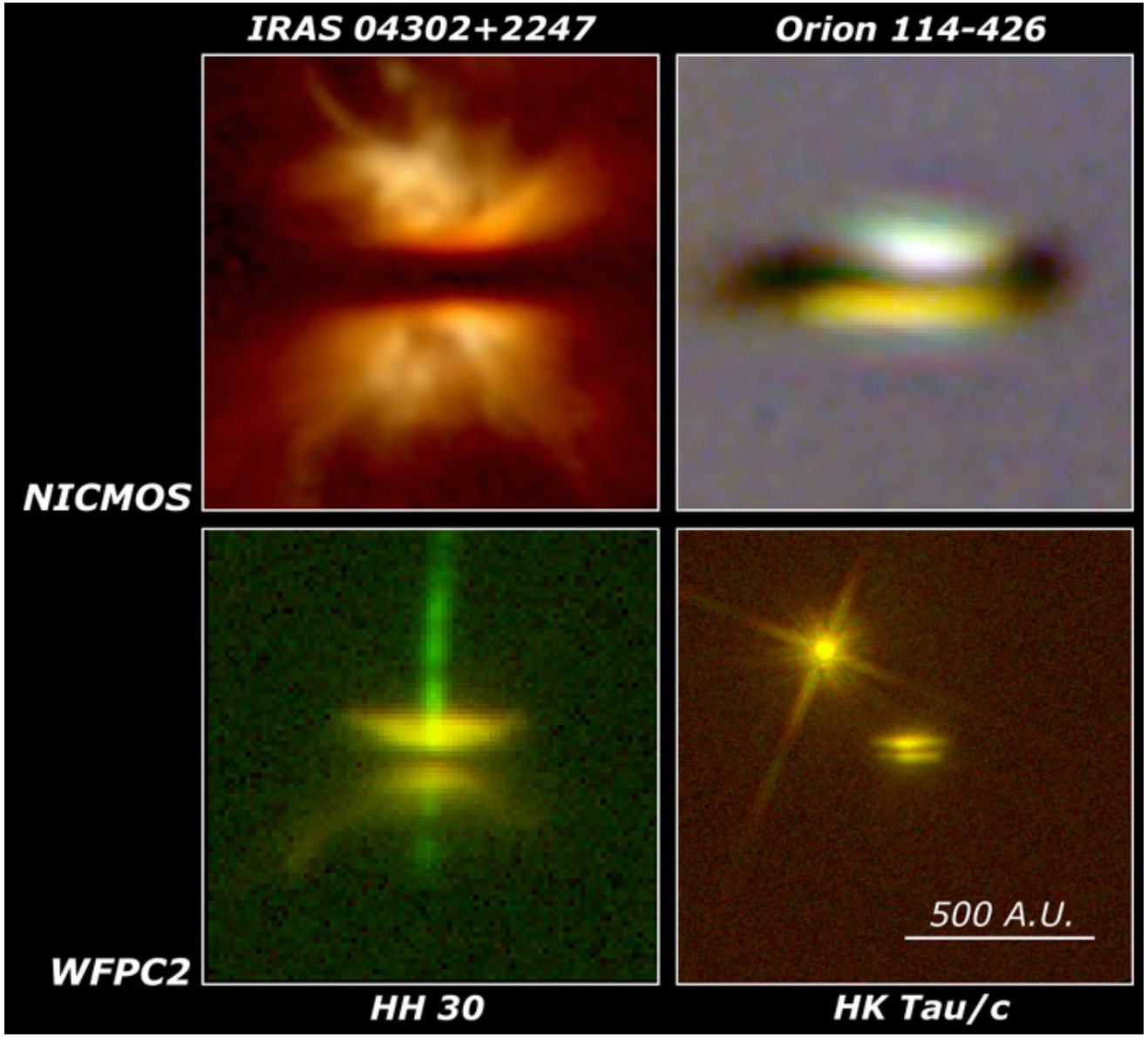
(c)-(e) Formation of planetesimals and planetary embryos

(f) Formation of planets, clearing of gas in disc in ≈ 10 Myr

Protoplanetary discs in the Orion Nebula



Observations of protoplanetary discs



Composition of the disc

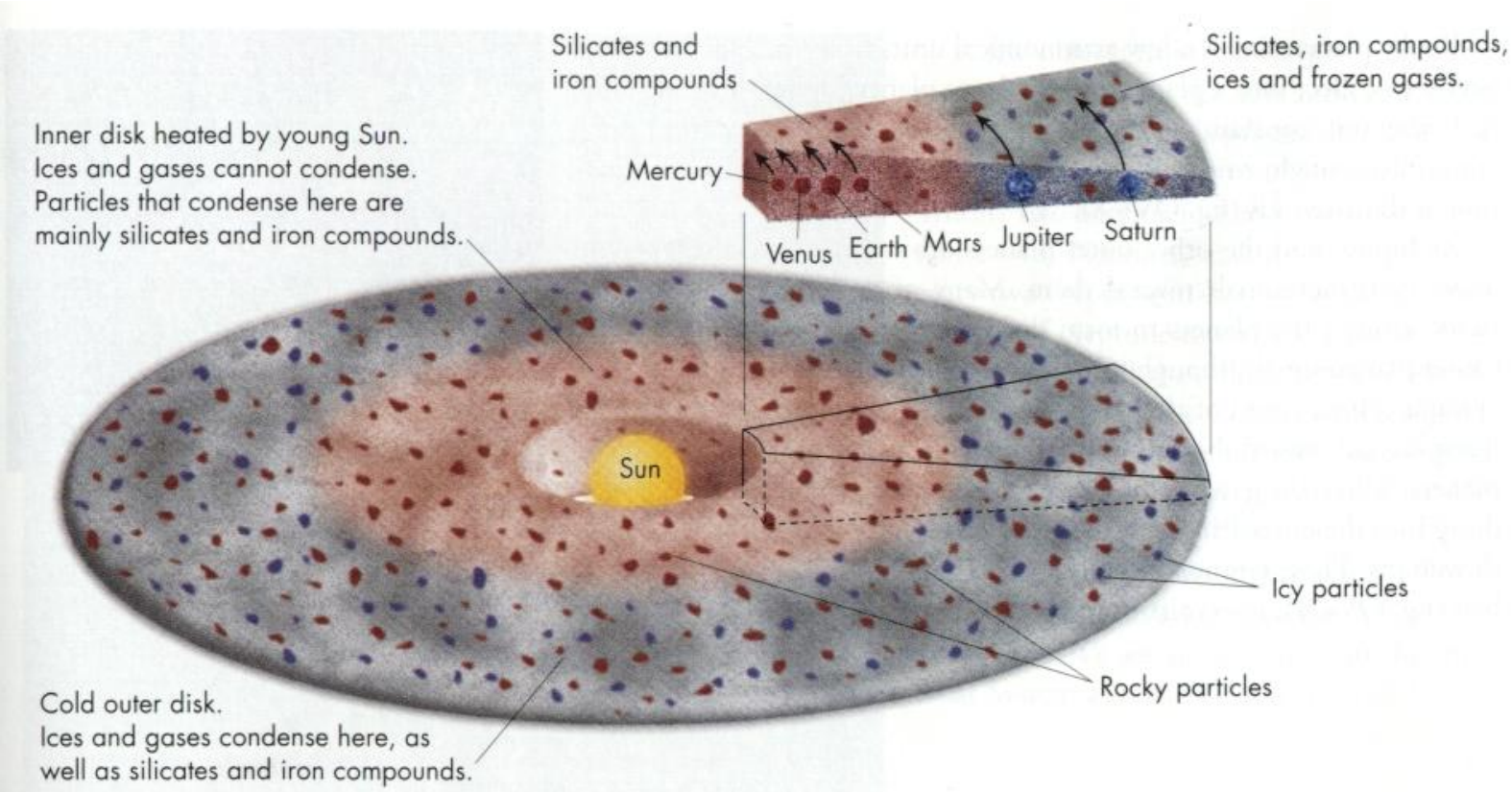
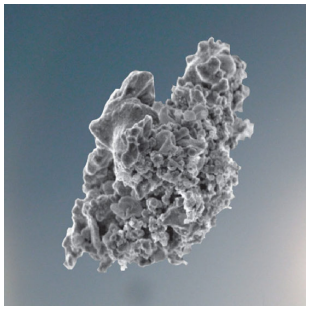


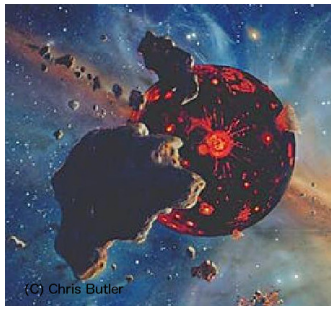
FIGURE OV4.6
Heat from the young Sun prevented ice from condensing in the inner parts of the Solar Nebula. The planetesimals—and ultimately the planets—that formed there are therefore composed mainly of rock and iron.

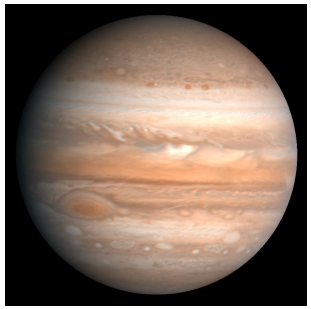
- The inner regions of the disc are hotter than the outer regions:
  $\Rightarrow$ Icy particles only in the outer parts of the disc!

The four steps of planet formation

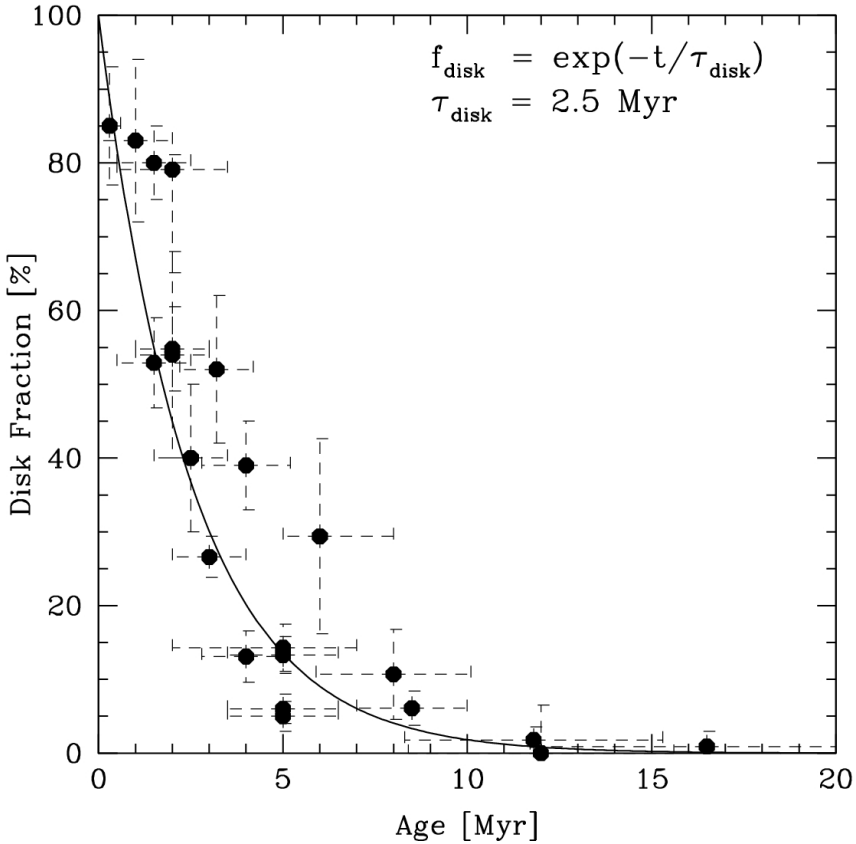
1. Dust to pebbles
   μm $\rightarrow$ dm: contact forces during collision lead to sticking
2. Pebbles to planetesimals
   dm $\rightarrow$ km: gravitational collapse of pebble clouds form planetesimals
3. Planetesimals to protoplanets
   km $\rightarrow$ 1,000 km: gravity (run-away accretion)
4. Protoplanets to planets
   Gas giants: 10 $M_\oplus$ core accretes gas ($< 10^7$ years)
   Terrestrial planets: protoplanets collide (10^7–10^8 years)

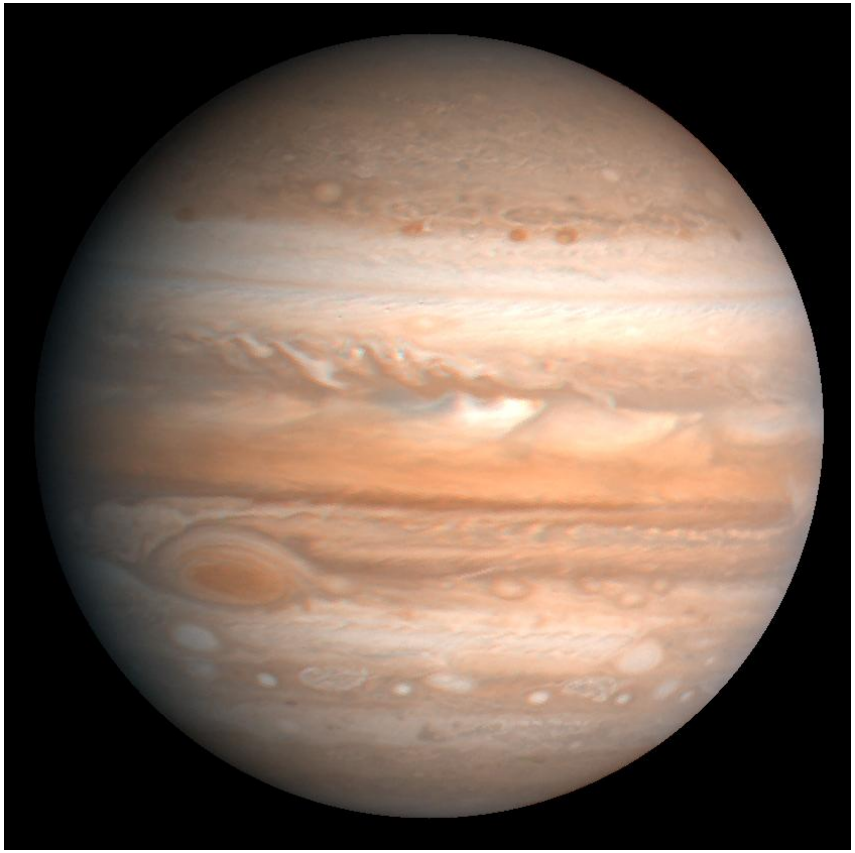




Time-scale to build gas giants





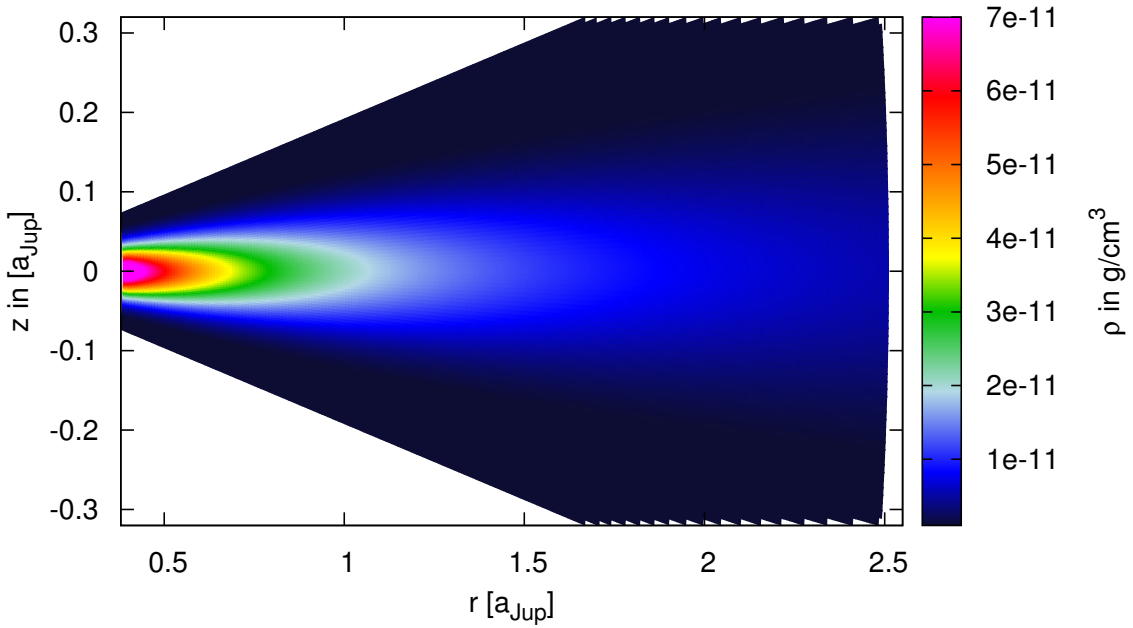
Mamajek (2009)

- Giant planet formation has to happen within the disc's lifetime of a few Myr!

Important quantities in the disc

- Temperature T
- Viscosity $\nu = \alpha H^2 \Omega_{\rm K}$ with:
  - H is the thickness of the disc
  - $\Omega_{\rm K}$ is the Keplerian frequency, $\Omega_{\rm K} = \sqrt{GM_\star/r^3}$
  - $\alpha \approx 10^{-2} - 10^{-4}$
- Gas density $\rho_{\rm g}$ and gas surface density $\Sigma_{\rm g}$:

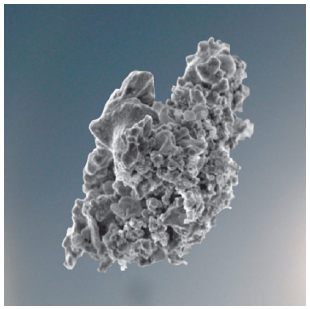
$$\Sigma_{\rm g} = \rho_{\rm g} H \sqrt{2\pi}$$

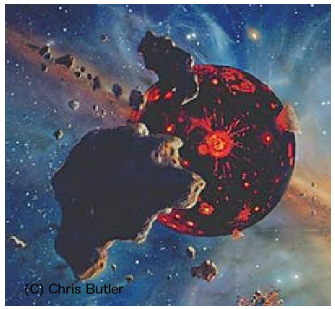


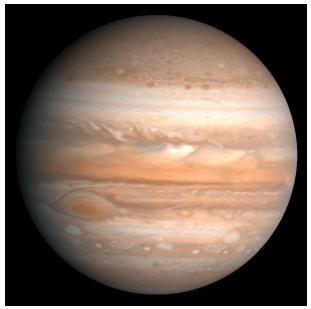
Importance of the disc structure

Growth and formation of planets relies on the disc structure:

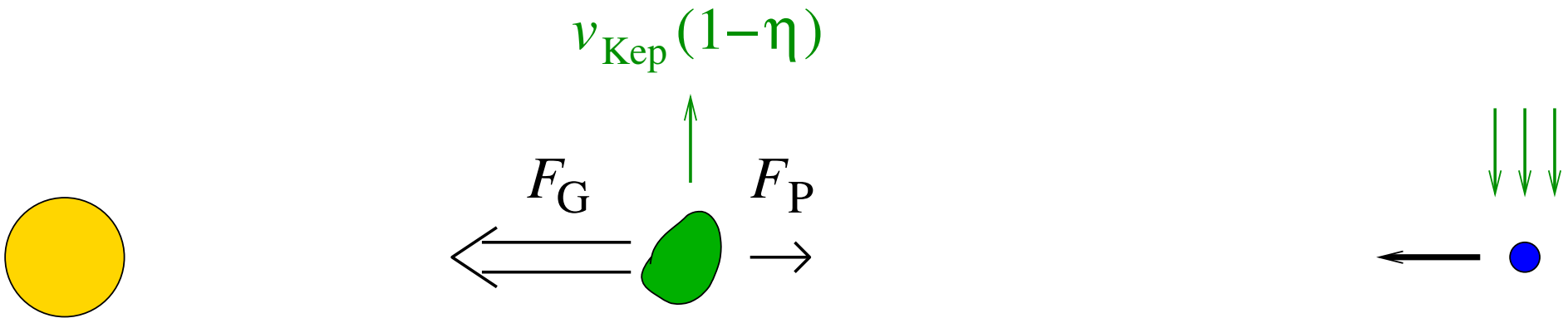
- Growth of dust particles to pebbles (Zsom et al., 2010; Birnstiel et al., 2012)
- Movements of pebbles inside the gas disc (Brauer et al., 2008)
- Formation of planetesimals via streaming instability (Johansen & Youdin, 2007)
- Formation of planetary cores from embryos and planetesimals (Levison et al., 2010) or pebble accretion (Lambrechts & Johansen, 2012)
- Migration of planetary cores in the disc (Ward, 1997; Paardekooper & Mellema, 2006; Kley et al., 2009)





Radial drift of dust particles



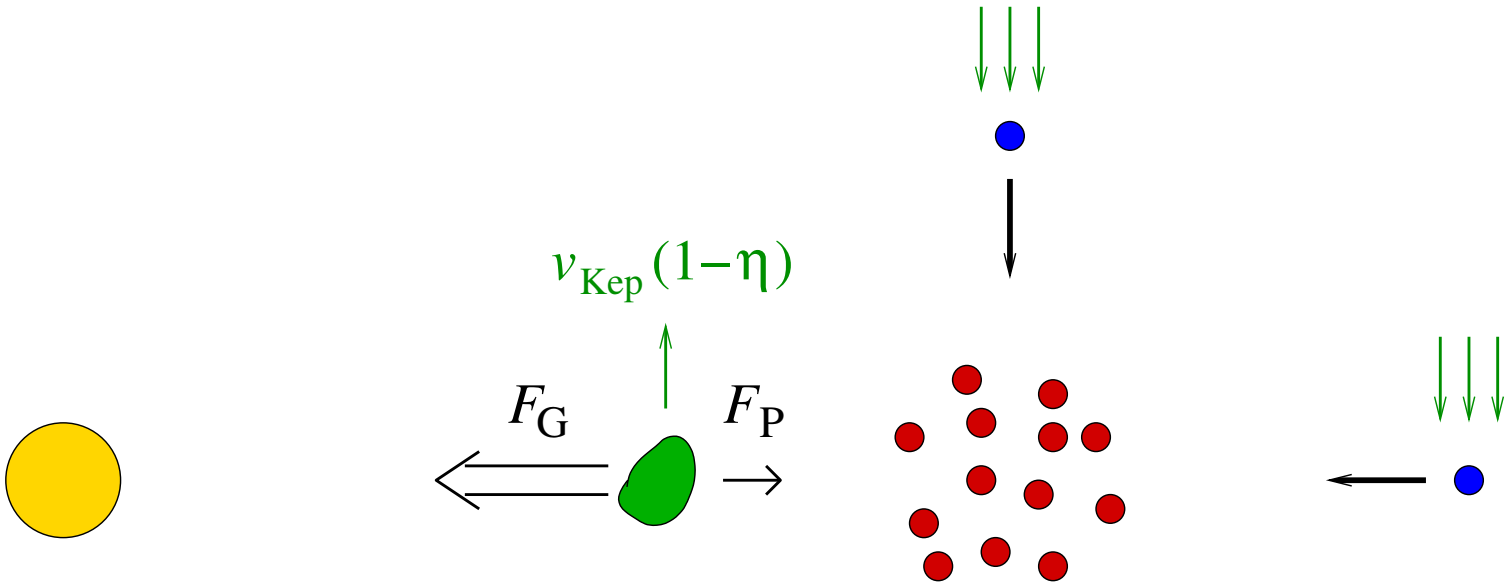
- Disc is hotter and denser close to the star
- Radial pressure gradient force mimics decreased gravity ⇒ gas orbits slower than Keplerian:

$$\eta = -\frac{1}{2}\left(\frac{H}{r}\right)^2 \frac{\partial \ln(P)}{\partial \ln(r)}$$

- Particles do not feel the pressure gradient force and want to orbit Keplerian
- Headwind from sub-Keplerian gas drains angular momentum from particles, so they spiral in through the disc
- Particles sublimate when reaching higher temperatures close to the star

Streaming instability: planetesimal formation

- Gas orbits slightly slower than Keplerian
- Particles lose angular momentum due to headwind
- Particle clumps locally reduce headwind and are fed by isolated particles



- *Streaming instabilities* feed on velocity difference between two components (gas and particles) at the same location
- Interested in details of the S.I.? Ask Anders Johansen or Chao-Chin Yang!

Streaming instability: lots of literature!

PROTOPLANETARY DISK TURBULENCE DRIVEN BY THE STREAMING INSTABILITY: LINEAR EVOLUTION AND NUMERICAL METHODS

A. Youdin[1]
Princeton University Observatory, Princeton, NJ 08544; youd@cita.utoronto.ca

AND

A. Johansen[2]
Max-Planck-Institut für Astronomie, 69117 Heidelberg, Germany; johansen@mpia.de
Received 2006 October 20; accepted 2007 February 22

Rapid planetesimal formation in turbulent circumstellar disks

Anders Johansen[1], Jeffrey S. Oishi[2,3], Mordecai-Mark Mac Low[1,2], Hubert Klahr[1], Thomas Henning[1] & Andrew Youdin[4]

THE EFFECT OF THE RADIAL PRESSURE GRADIENT IN PROTOPLANETARY DISKS ON PLANETESIMAL FORMATION

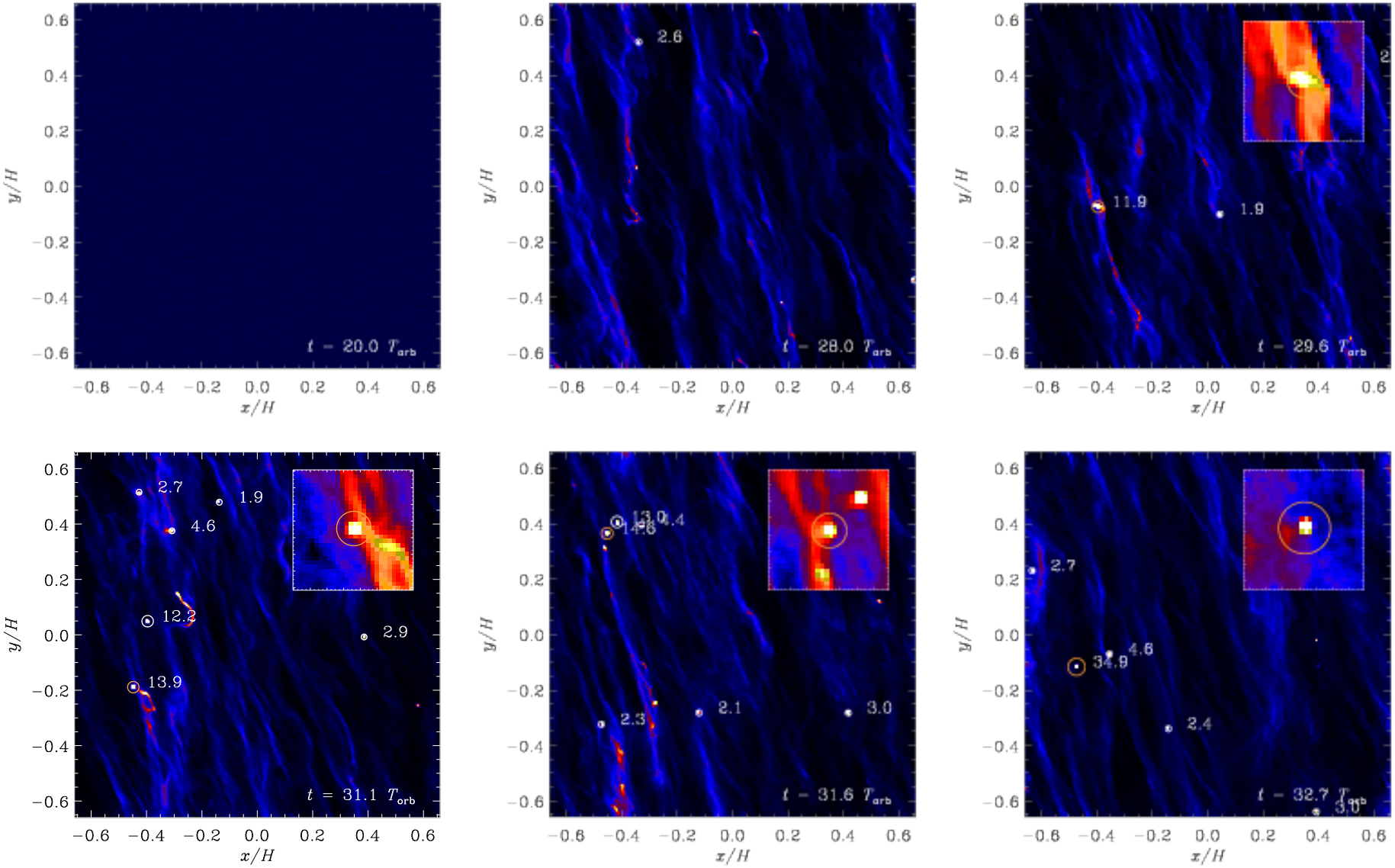
Xue-Ning Bai and James M. Stone
Department of Astrophysical Sciences, Princeton University, Princeton, NJ 08544, USA; xbai@astro.princeton.edu, jstone@astro.princeton.edu
Received 2010 May 25; accepted 2010 September 14; published 2010 October 4

ON THE FEEDING ZONE OF PLANETESIMAL FORMATION BY THE STREAMING INSTABILITY

Chao-Chin Yang (楊朝欽) and Anders Johansen
Lund Observatory, Department of Astronomy and Theoretical Physics, Lund University,
Box 43, SE-22100 Lund, Sweden; ccyang@astro.lu.se, anders@astro.lu.se
Received 2014 April 27; accepted 2014 July 21; published 2014 August 19

This is just a tiny example of literature to the streaming instabilty!

Streaming instability: a movie



(Johansen et al. 2011)

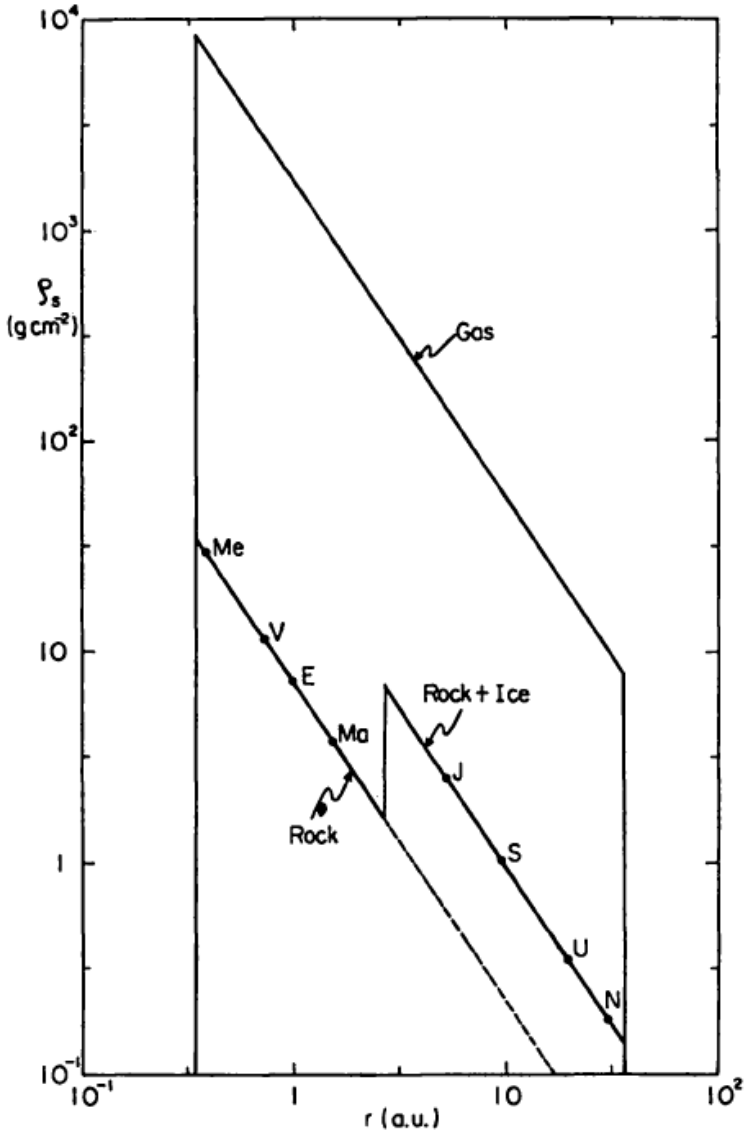
Protoplanetary disc structure

The Minimum Mass Solar Nebula - MMSN

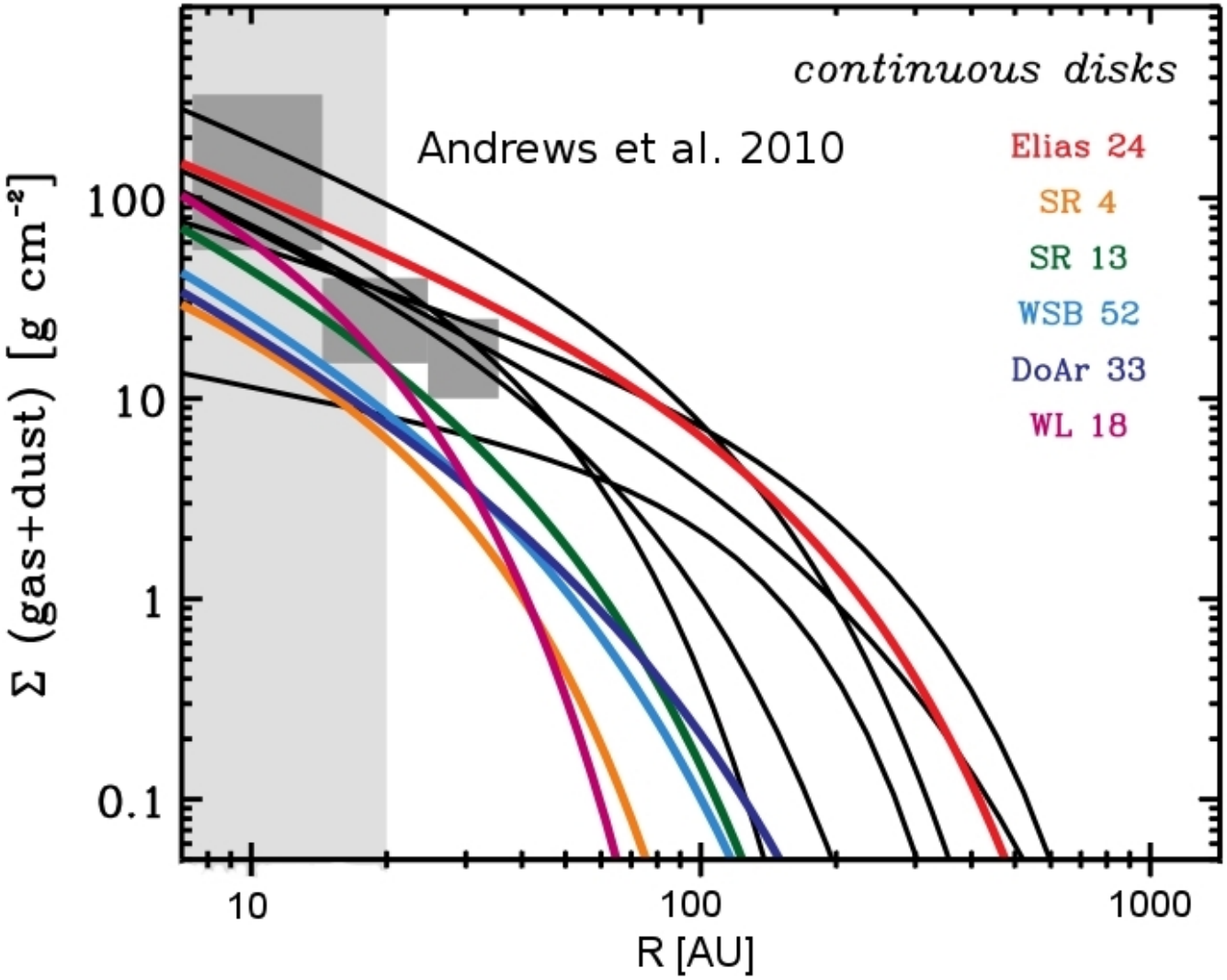
- Spread appropriate mass of solids around the orbit of each planet in the solar system and multiply by 100 (add gas)
- Power law through data (Hayashi (1981), Weidenschilling (1977)):

$$\Sigma_{\mathrm{g}}(r) = 1700 \left(\frac{r}{1\mathrm{AU}}\right)^{-3/2} \mathrm{g/cm}^2$$

- The planets accreted **all** solids (hence "Minimum")
- The planets formed on their present orbits and did not move



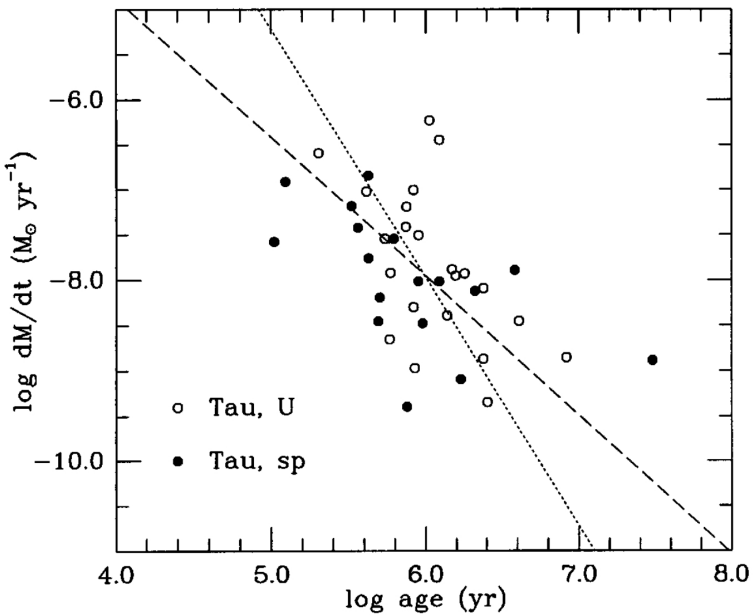
Constraints from Observations

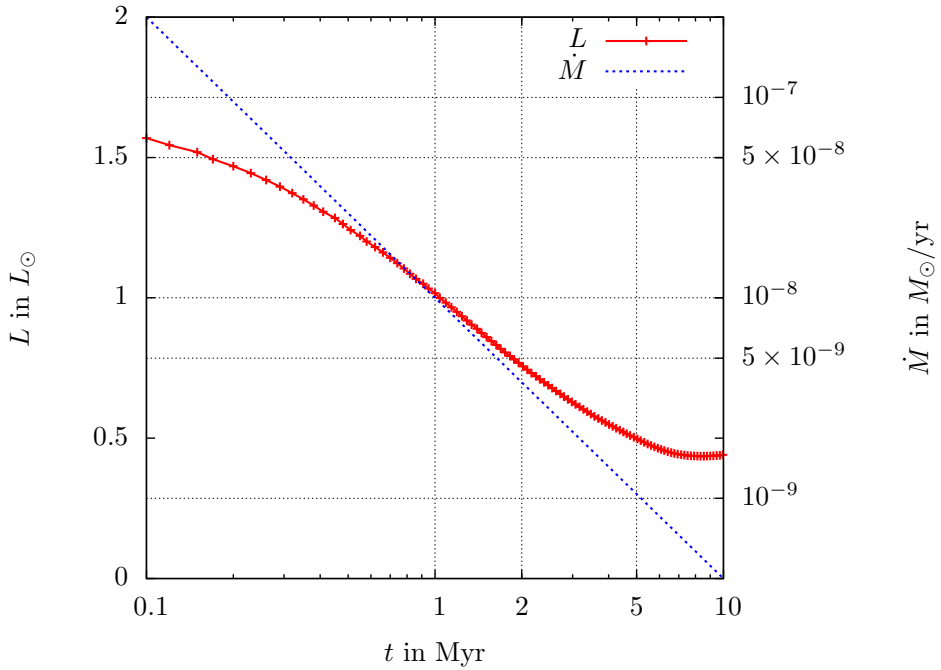


$$\Sigma \propto R^{-\gamma} \quad \text{with} \quad \gamma = 0.4 - 1.1$$

Time evolution of the star and the disc

- Accretion rate $\dot{M}$ ($\propto \Sigma_{\rm g}$) changes with time (Hartmann et al., 1998)
  $\Rightarrow$ Accretion rate changes by a factor of 100 in 5Myr!
- Star changes luminosity in time (Baraffe et al., 1998)
  $\Rightarrow$ Stellar luminosity changes by a factor of 3 in 5Myr!

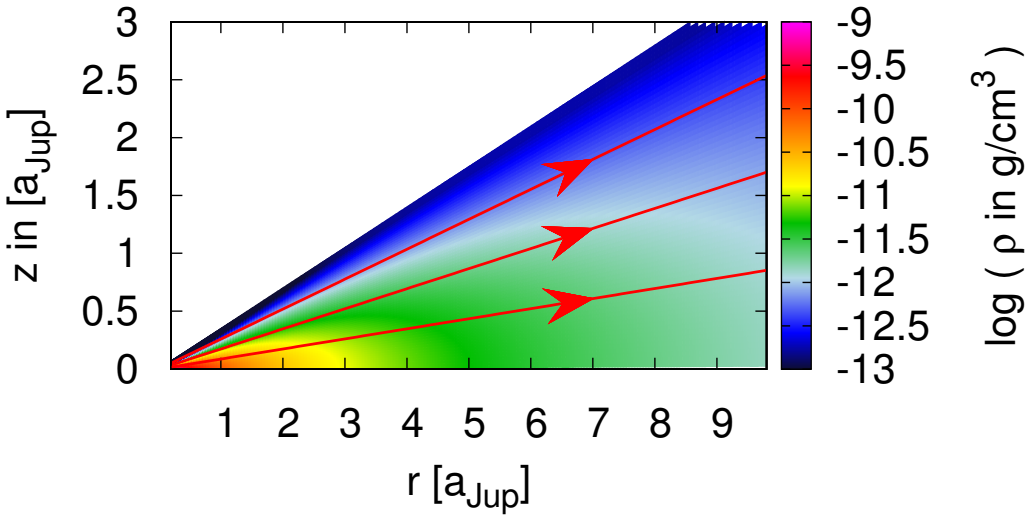




$\Rightarrow$ The disc is subject to massive changes in its lifetime!

Disc Model

- 2D hydrodynamical disc model with viscous heating, radiative cooling and stellar irradiation with $S \propto L_\star$:



Bitsch et al., 2013

- Mass flux through disc: $\dot{M}$ constant at each r with α viscosity:

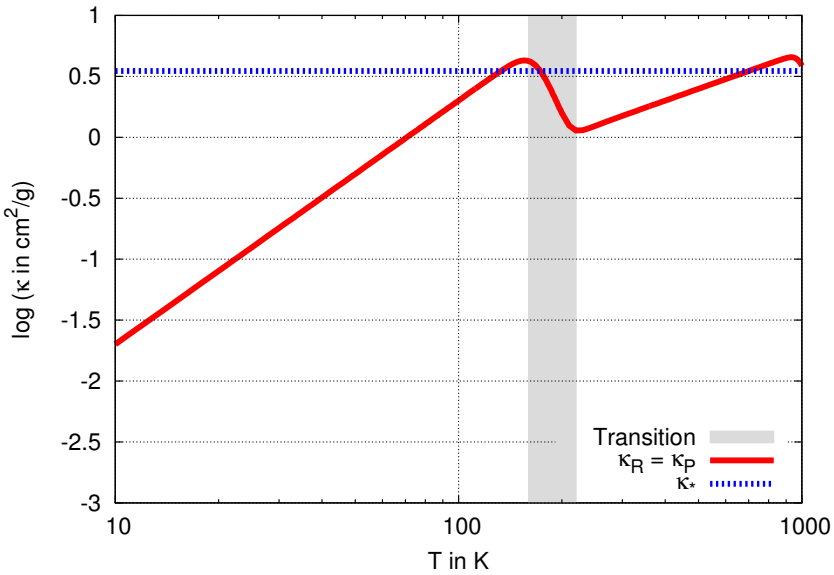
$$\dot{M} = 3\pi\nu\Sigma = 3\pi\alpha H^2 \Omega_{\rm K} \Sigma$$

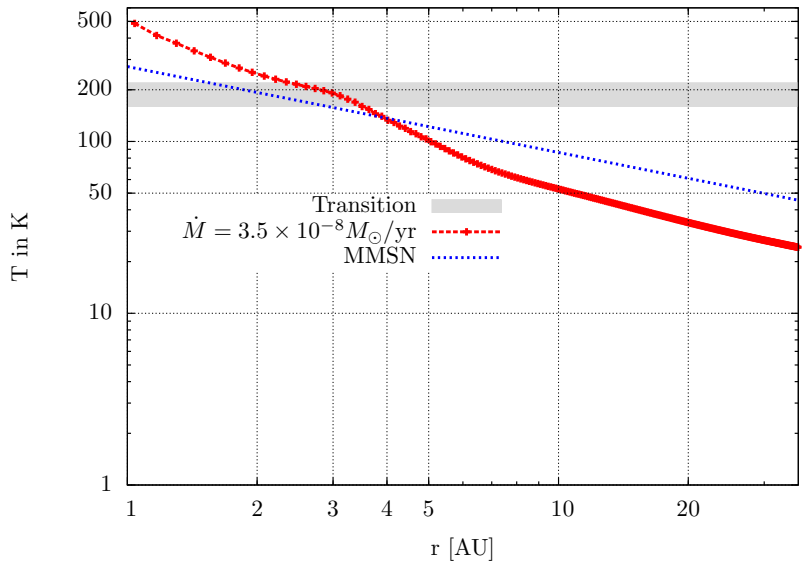
Influence of opacity on cooling

- Cooling of the disc:

$$\mathbf{F} = -\frac{\lambda c}{\rho \kappa_{\mathrm{R}}} \nabla E_{\mathrm{R}}$$

- Grey area marks transition in opacity at the ice line
- Change of opacity:
  $\Rightarrow$ change of cooling
- Change of cooling:
  $\Rightarrow$ change in $T(r)$





Bitsch et al., 2015

Influence of viscosity and $\dot{M}$

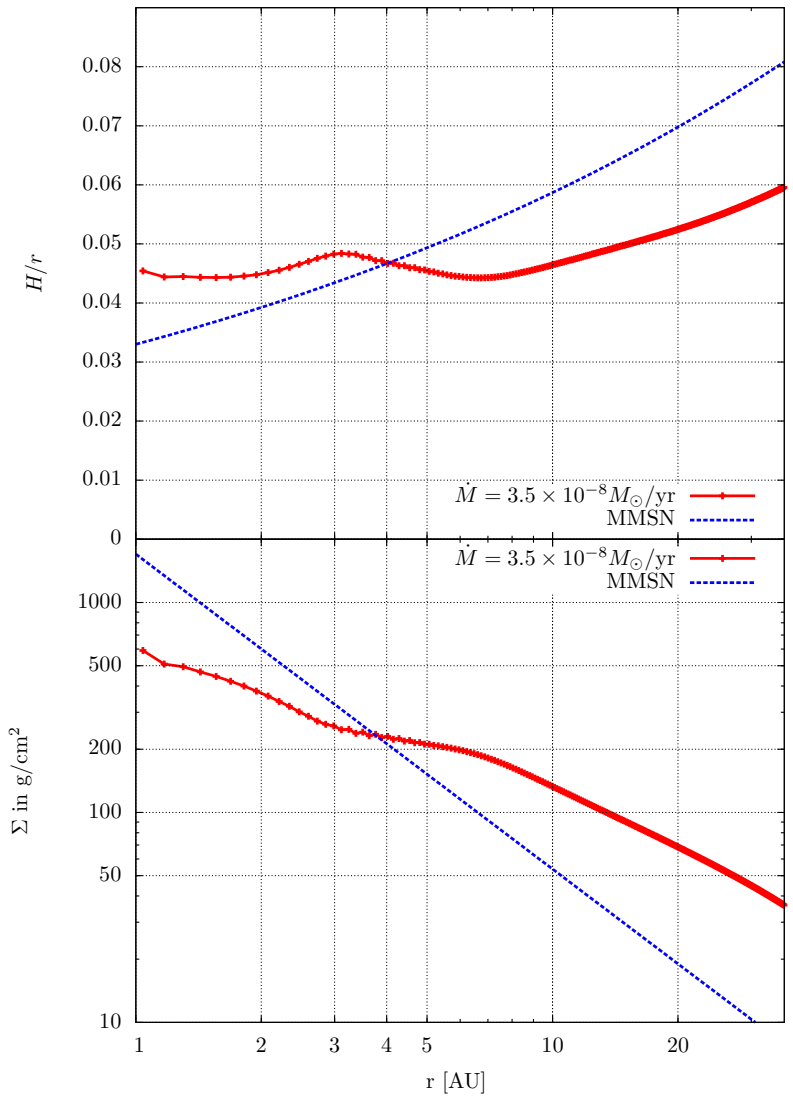
- Hydrostatic equilibrium:

$$T = \left(\frac{H}{r}\right)^2 \frac{GM_\star}{r} \frac{\mu}{\mathcal{R}}$$

- bump in T: bump in H/r
- $\dot{M}$ disc:

$$\dot{M} = 3\pi\nu\Sigma = 3\pi\alpha H^2 \Omega_\mathrm{K} \Sigma$$

- $\dot{M}$ constant at each r:
  $\Rightarrow$ dip in Σ



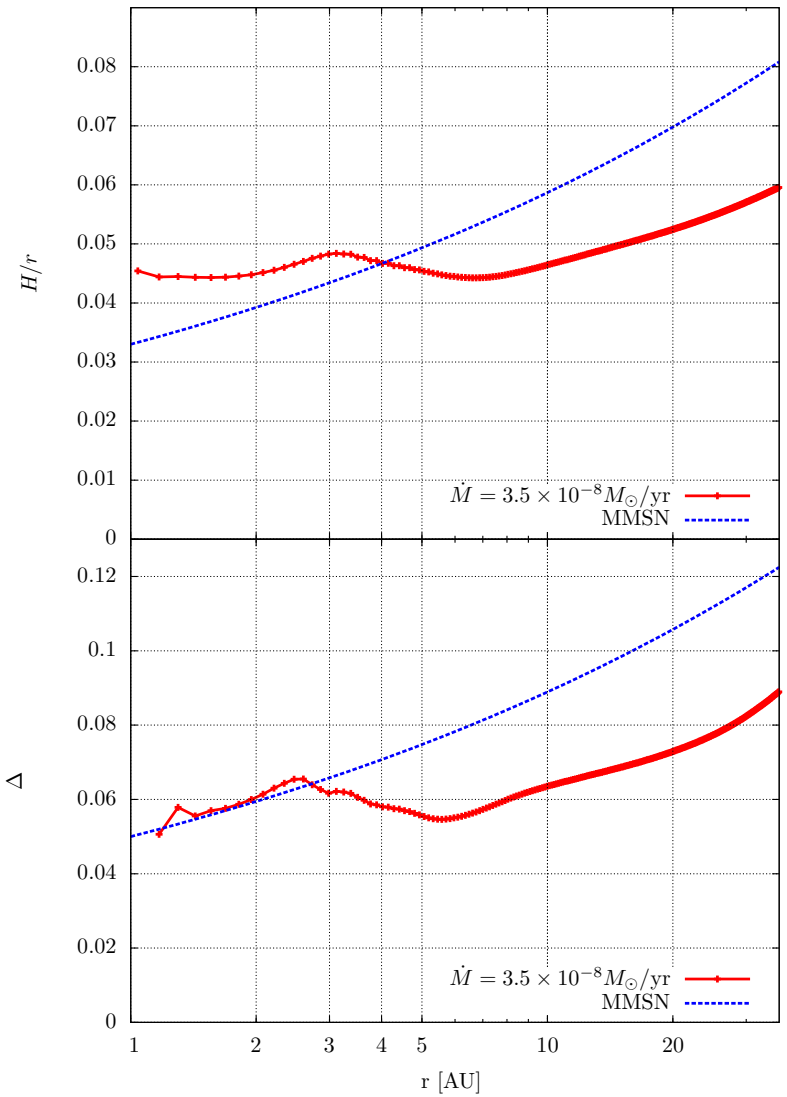
Bitsch et al., 2015

Importance for the streaming instability

- Streaming instabilities feed on velocity difference between two components (gas and particles) at the same location, caused by a reduction of the effective gravitational force by the radially outwards pointing force of the radial pressure gradient:

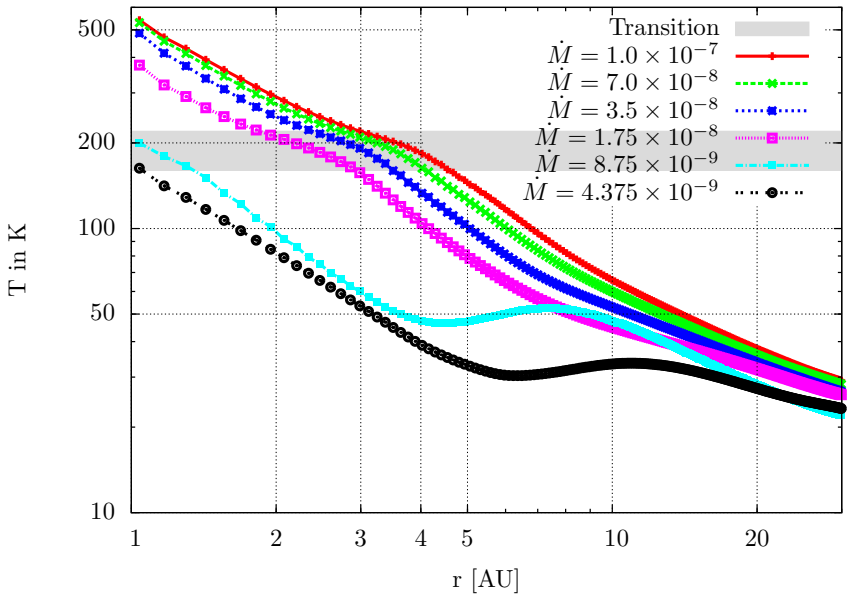
$$\Delta = \eta \frac{v_{\mathrm{K}}}{c_{\mathrm{s}}} = -\frac{1}{2}\frac{H}{r}\frac{\partial \ln(P)}{\partial \ln(r)}$$

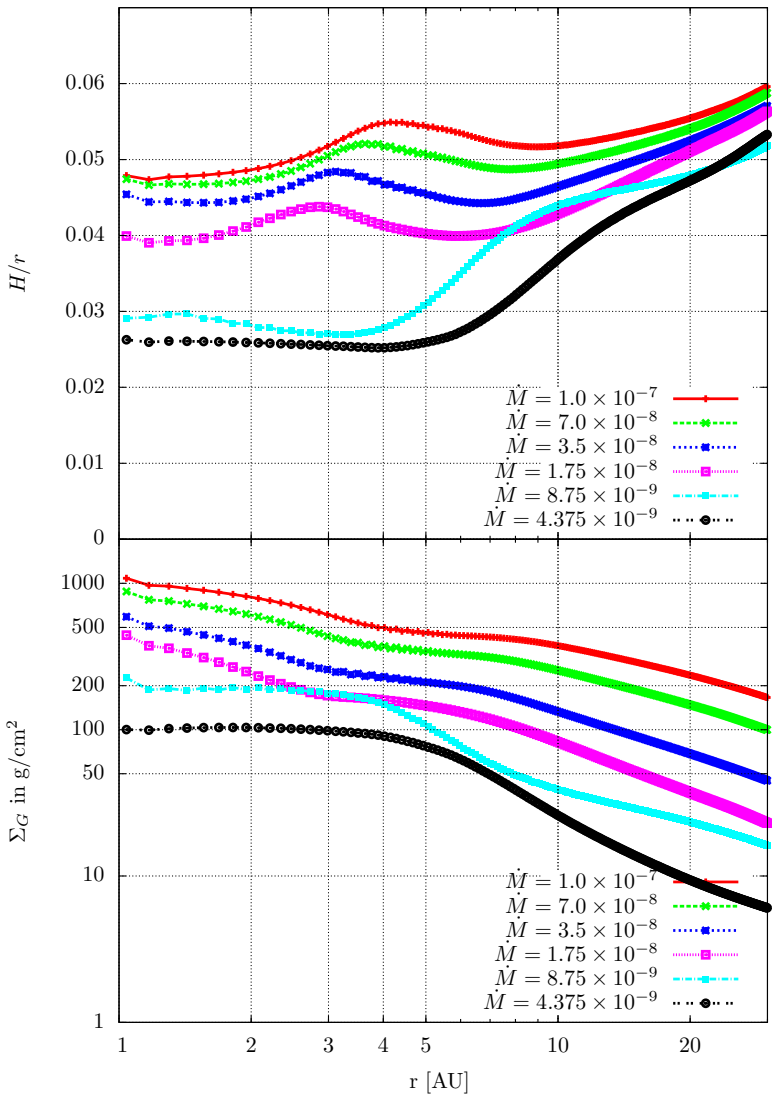
- Reduced Δ helps the formation of large clumps via streaming instability (Bai & Stone, 2010b)



Bitsch et al., 2015

Change of $\dot{M}$ in time





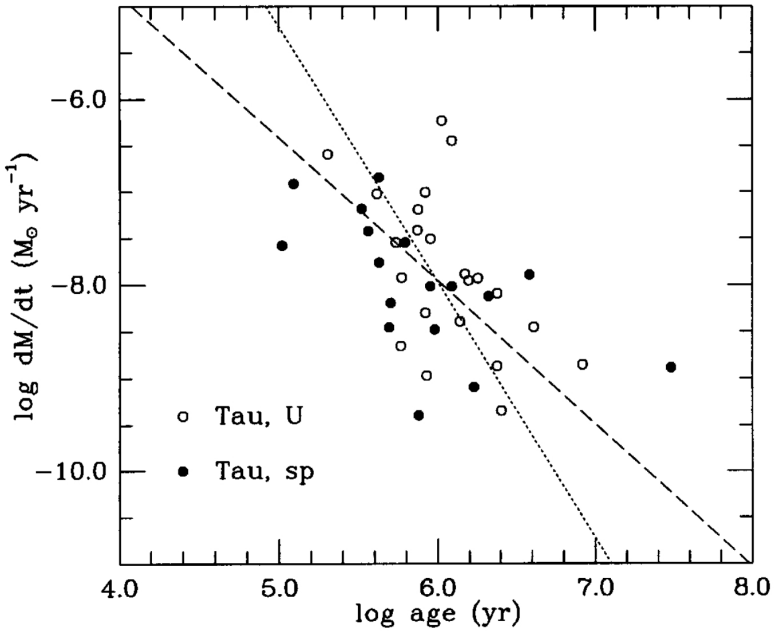
- $\dot{M}$ decreases with decreasing Σ
- $\dot{M} = 3\pi\nu\Sigma = 3\pi\alpha H^2 \Omega_{\mathrm{K}} \Sigma$
- Inner disc dominated by viscous heating for high $\dot{M}$, dominated by stellar heating for low $\dot{M}$

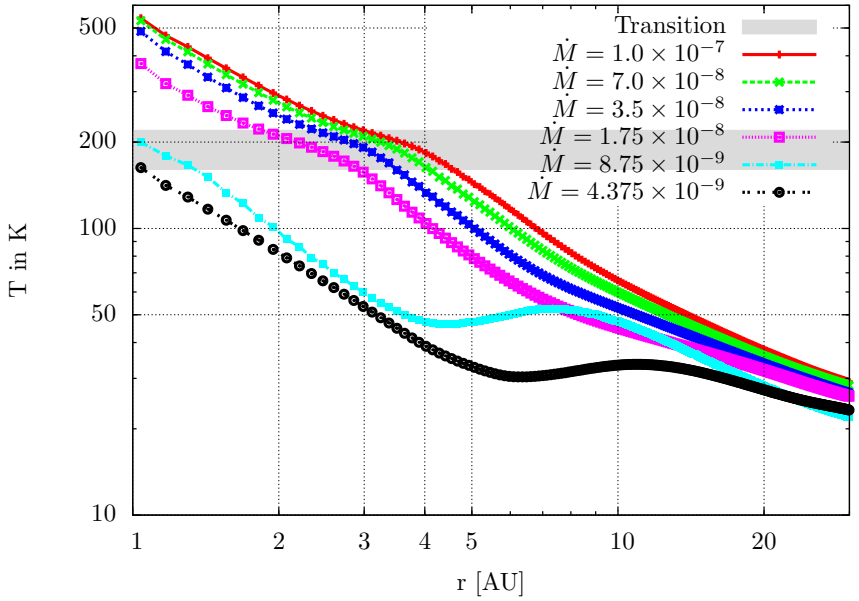
Bitsch et al., 2015

Time evolution of the disc

- Evolution in time follows Hartmann et al. 1998 equation:

$$\log\left(\frac{\dot{M}}{M_\odot/\mathrm{yr}}\right) = -8.00 - 1.40\log\left(\frac{t_{\mathrm{disc}} + 10^5\mathrm{yr}}{10^6\mathrm{yr}}\right) .$$



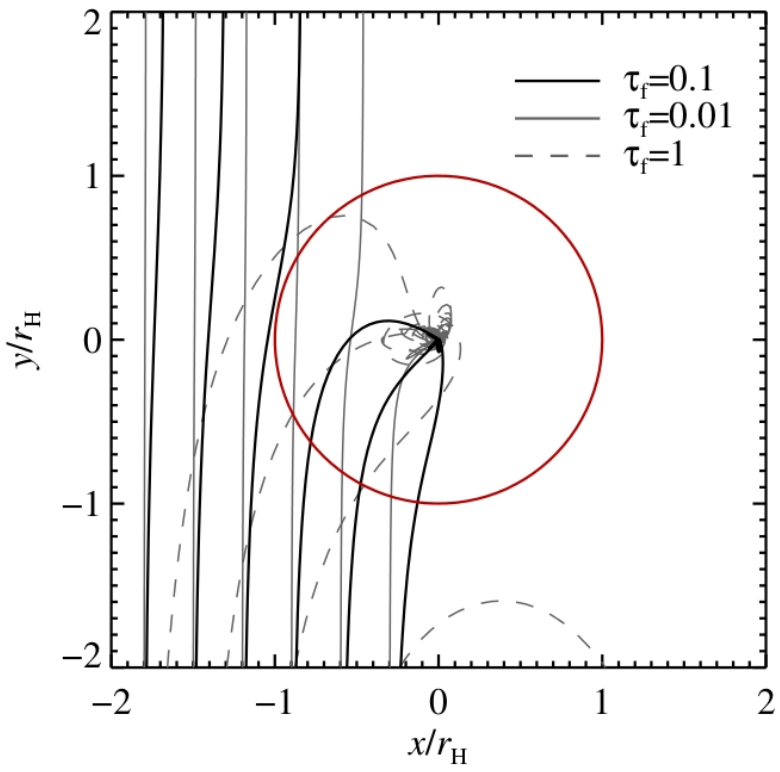


Planet growth and migration

Pebble accretion

- Small pebbles ($\tau_\mathrm{f} < 1$) can be easily accreted by planetesimals (Lambrechts & Johansen, 2012)
- Stokes number τ_f and friction time t_f:

$$\tau_\mathrm{f} = t_\mathrm{f}\Omega_\mathrm{K} = \frac{\rho_\bullet R}{\rho_\mathrm{G} c_\mathrm{s}}\Omega_\mathrm{K} = \frac{\rho_\bullet R}{\rho_\mathrm{G} H}$$



"Pebbles are weakly enough bound to the gas to feel the gravitational pull from the core, but strongly enough to deposit their kinetic energy through drag forces, when passing the core."

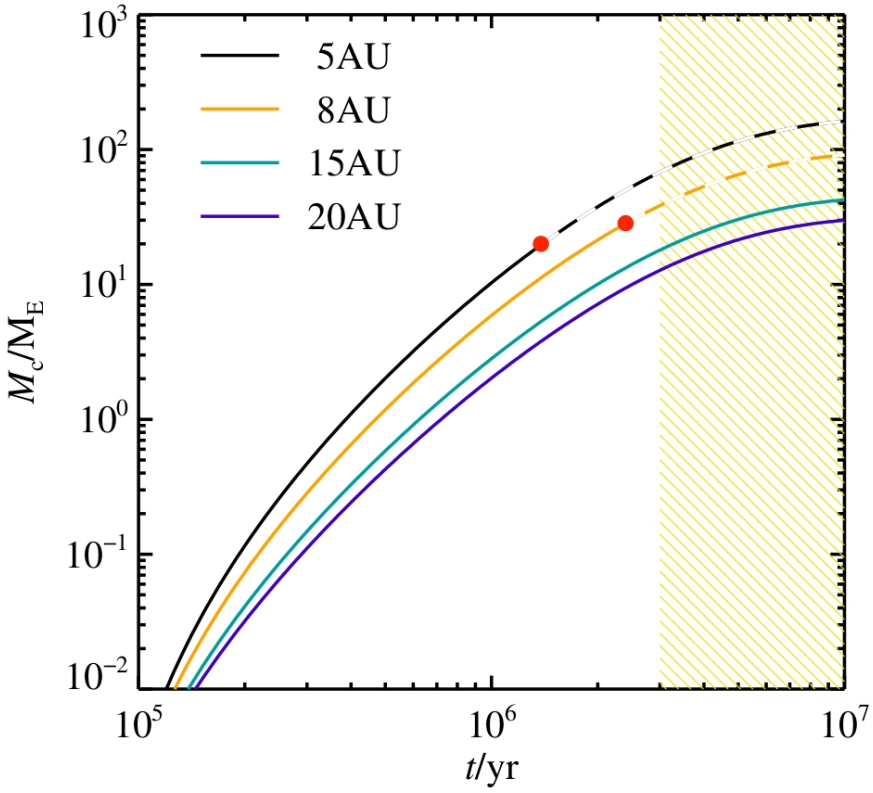
Lambrechts & Johansen, 2012

Scaling of pebble accretion

- Core growth via pebbles (Lambrechts & Johansen, 2014):

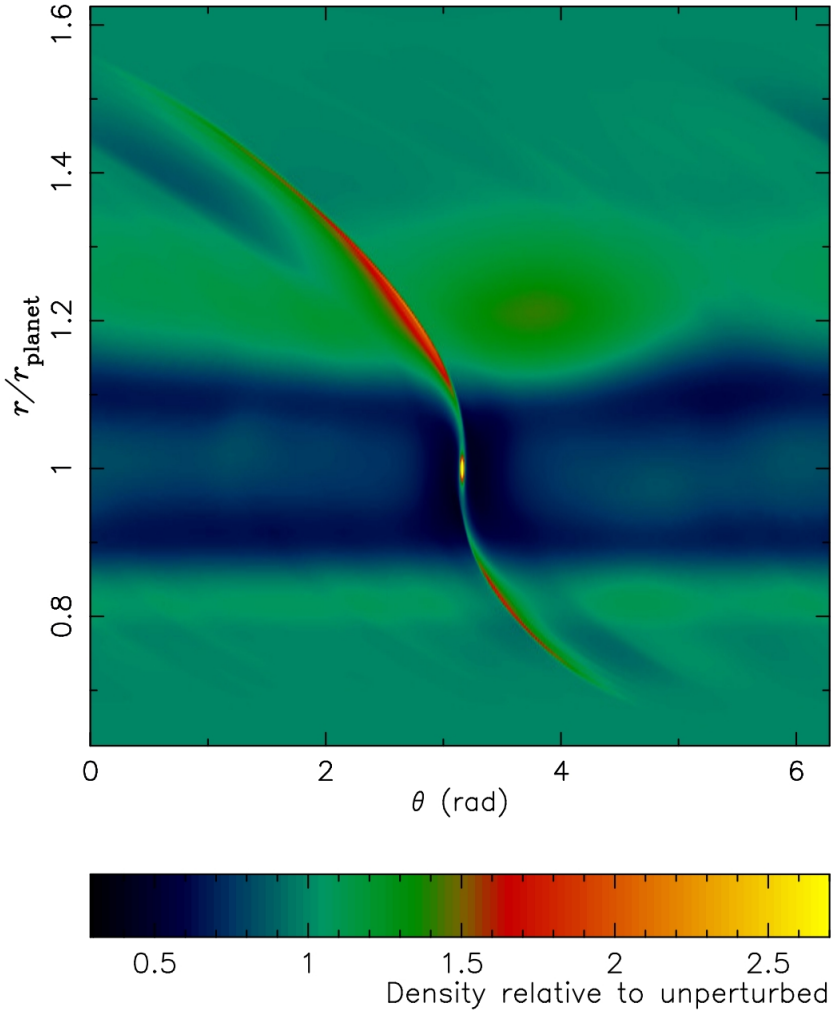
$$\dot{M}_{\rm c} = 2\left(\frac{\tau_{\rm f}}{0.1}\right)^{2/3} r_{\rm H} v_{\rm H} \Sigma_{\rm Peb}$$

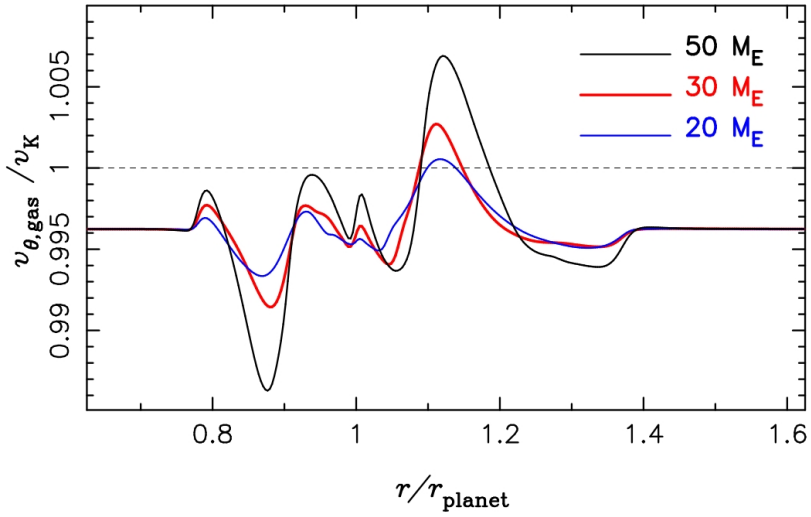
- Growth faster in inner regions of the disc
- Red dots mark pebble isolation mass: Pebble accretion does not continue forever!



Lambrechts & Johansen, 2014

Pebble isolation mass





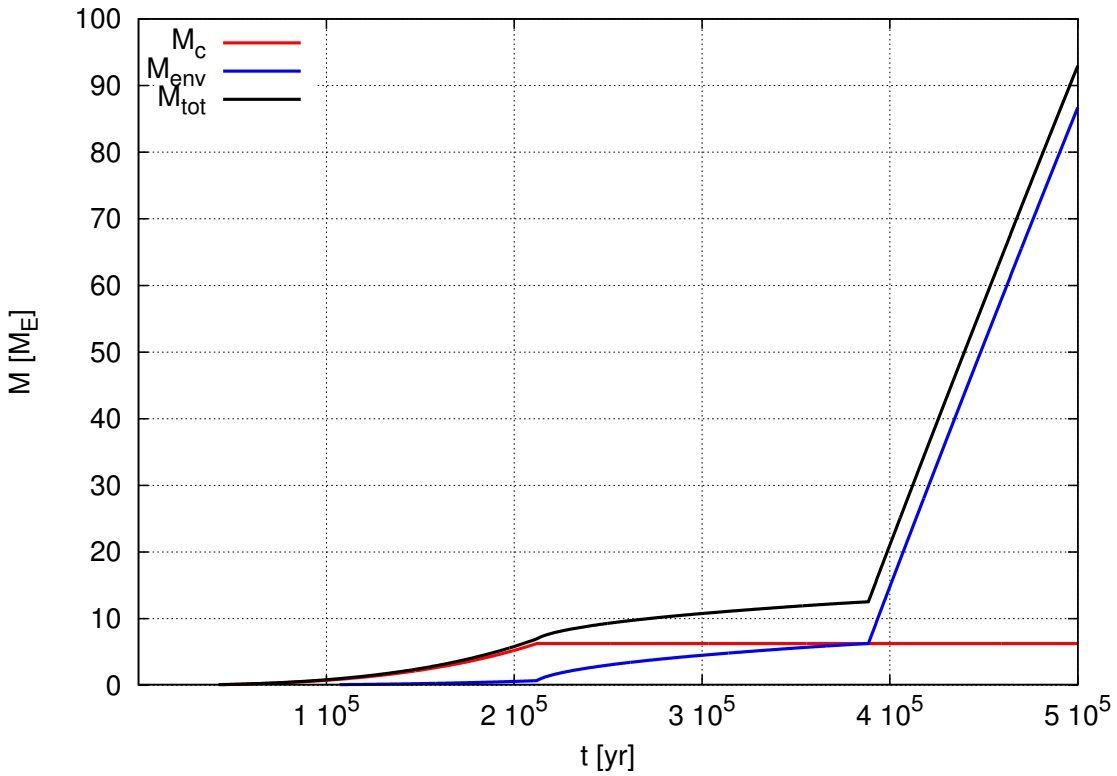
Pebble isolation mass:

$$M_{\rm iso} = 20 \left(\frac{H/r}{0.05} \right)^3 M_{\rm Earth}$$

$\Rightarrow$ After pebble isolation mass is reached, gas accretion can start!

Lambrechts et al., 2014

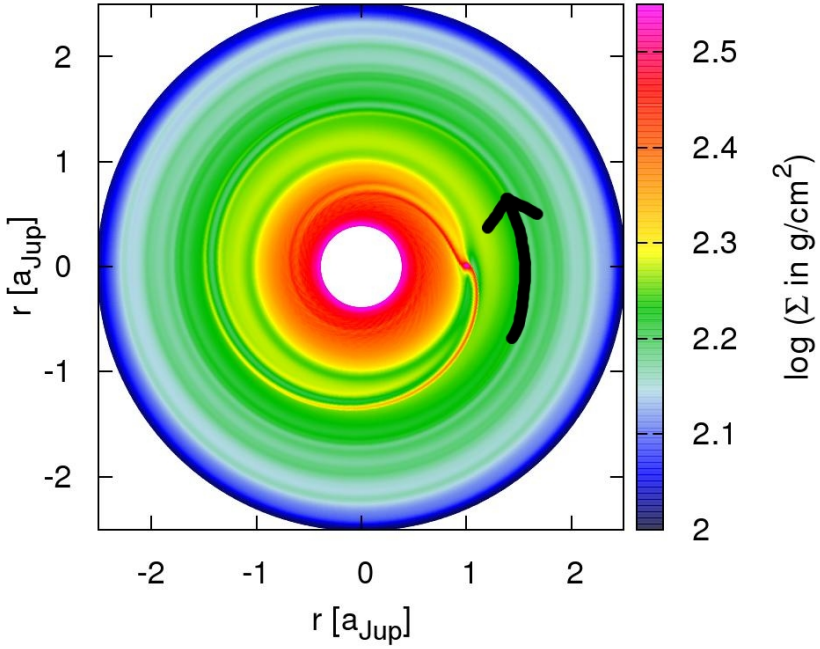
Gas accretion

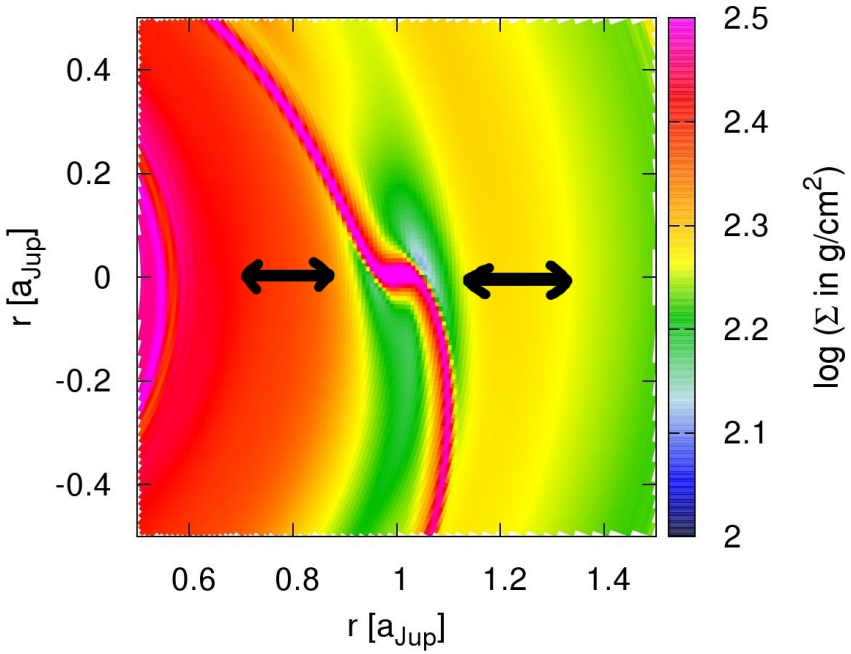


- Phase 1: accretion of solid core until pebble isolation ($\approx 2 \times 10^5$ yr)
- Phase 2: envelope contraction ($\approx 4 \times 10^5$ yr, following Piso & Youdin, 2014), when $M_{\rm c} > M_{\rm env}$
- Phase 3: rapid accretion of gas onto core when $M_{\rm c} < M_{\rm env}$ (Machida et al. 2010)

Planets in discs

- Planet embedded in a disc creates a wake
- Wake directed forwards in inner disc and backwards in outer disc
- Planet pushes inner (outer) disc inwards (outwards), inner (outer) disc pushes planet outwards (inwards)



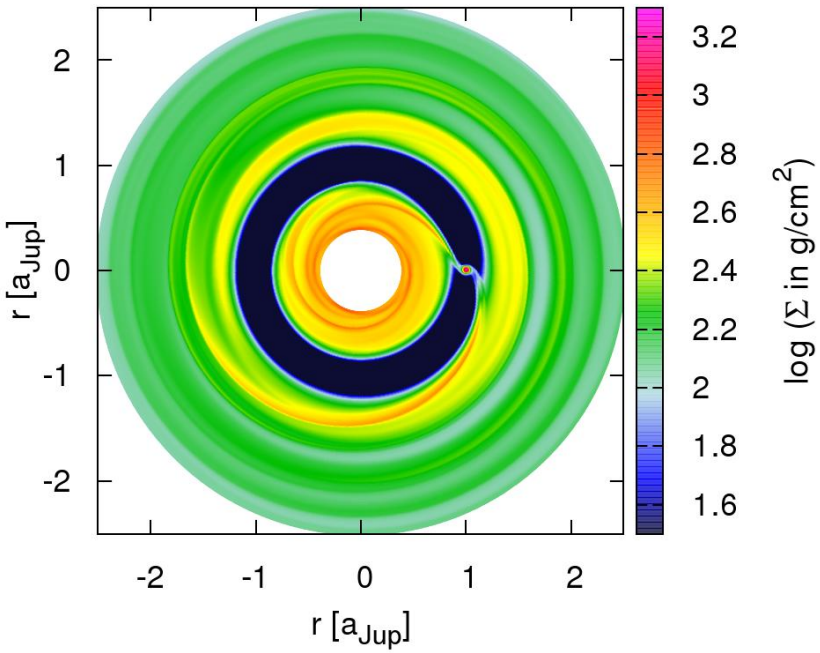


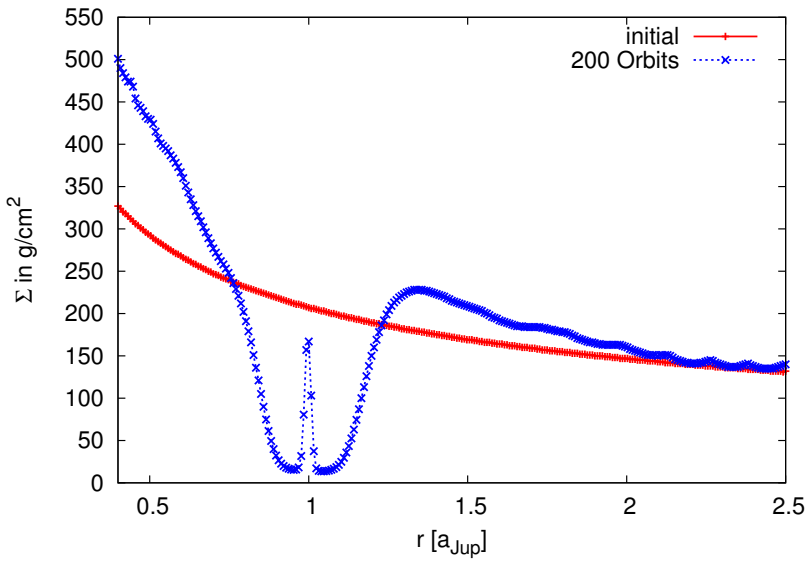
Type-II-migration

- High mass planets ($M_P > M_{Jup}$) open gaps in discs (Crida et al. 2007):

$$P = h/q^{1/3} + 50\alpha/qh^2 < 1 \quad \text{with} \quad h = H/r$$

- Planet is locked in the disc
- Planet moves with disc on "accretion timescale":
  $\tau_\nu = r_{\rm p}^2/\nu$ (up to a few Myrs)
- It is called Type-II-migration



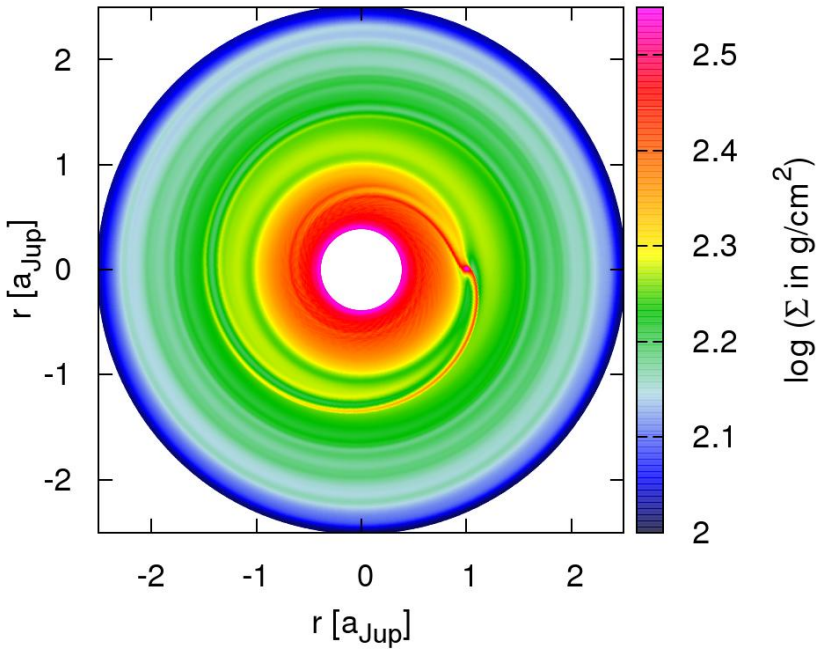


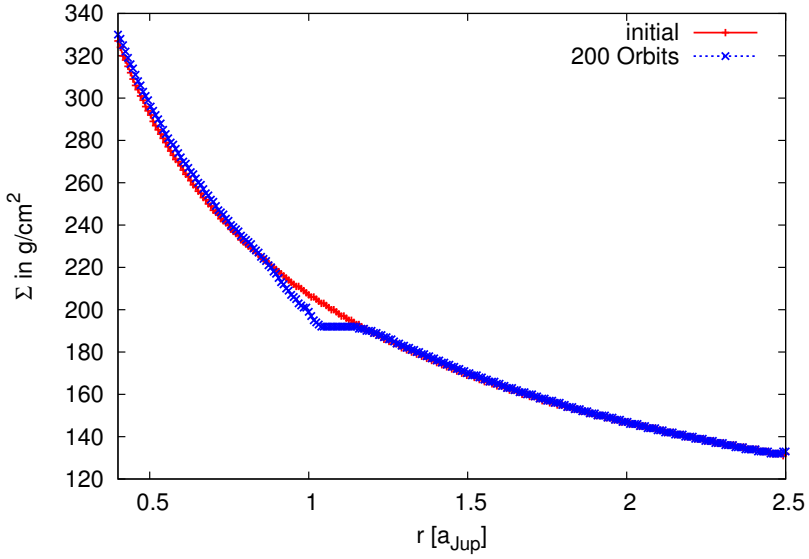
Type-I-migration

Planet fully embedded in the disc, free to migrate with respect to the gas

- Waves carry energy and angular momentum
- Torque calculated through Lindblad Resonances (e.g. Goldreich & Tremaine, 1980)
- ILR: positive torque, OLR: negative torque
- OLR stronger than ILR:

  $\Rightarrow$ inward migration ($\dot{a} \propto q$)

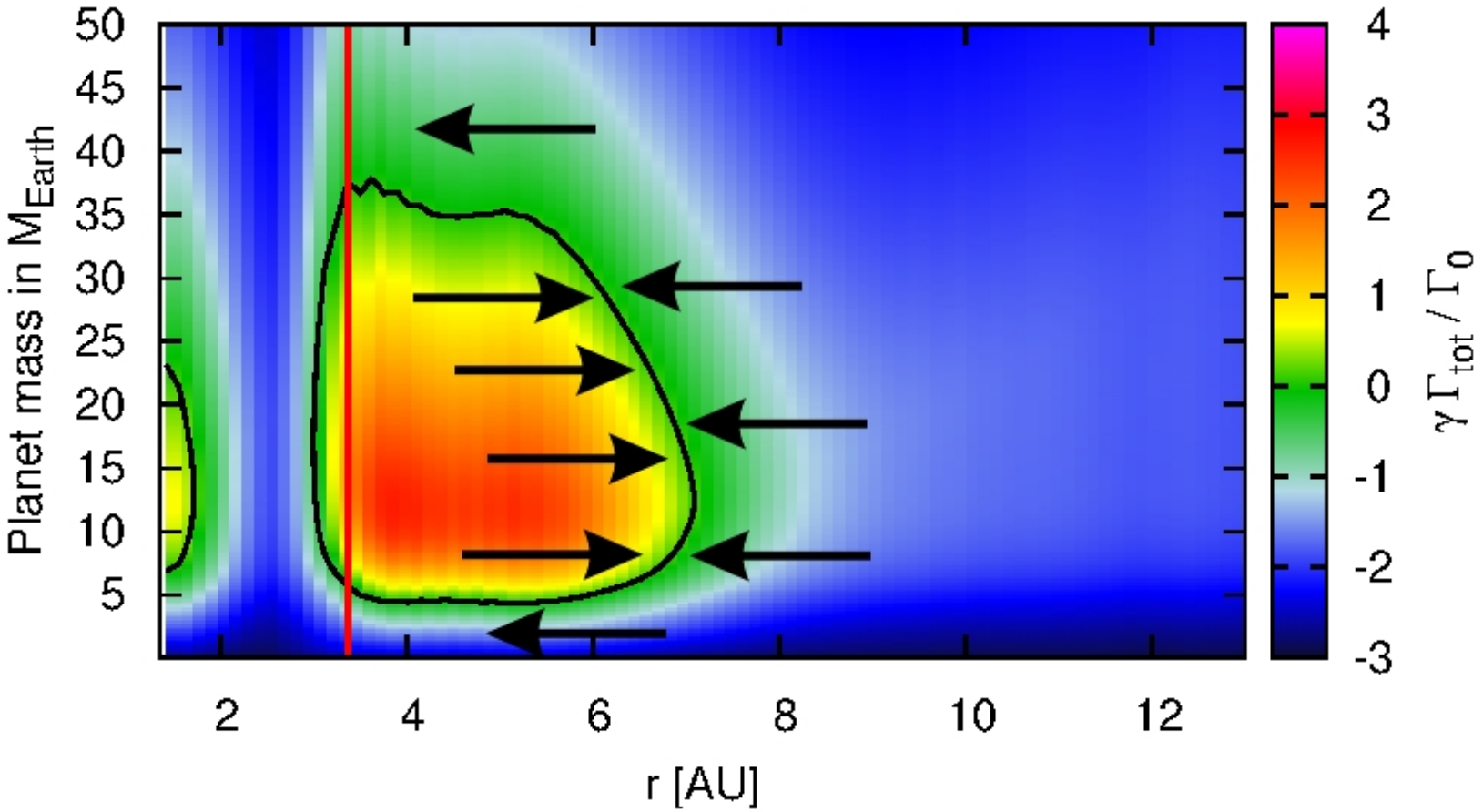




Migration map

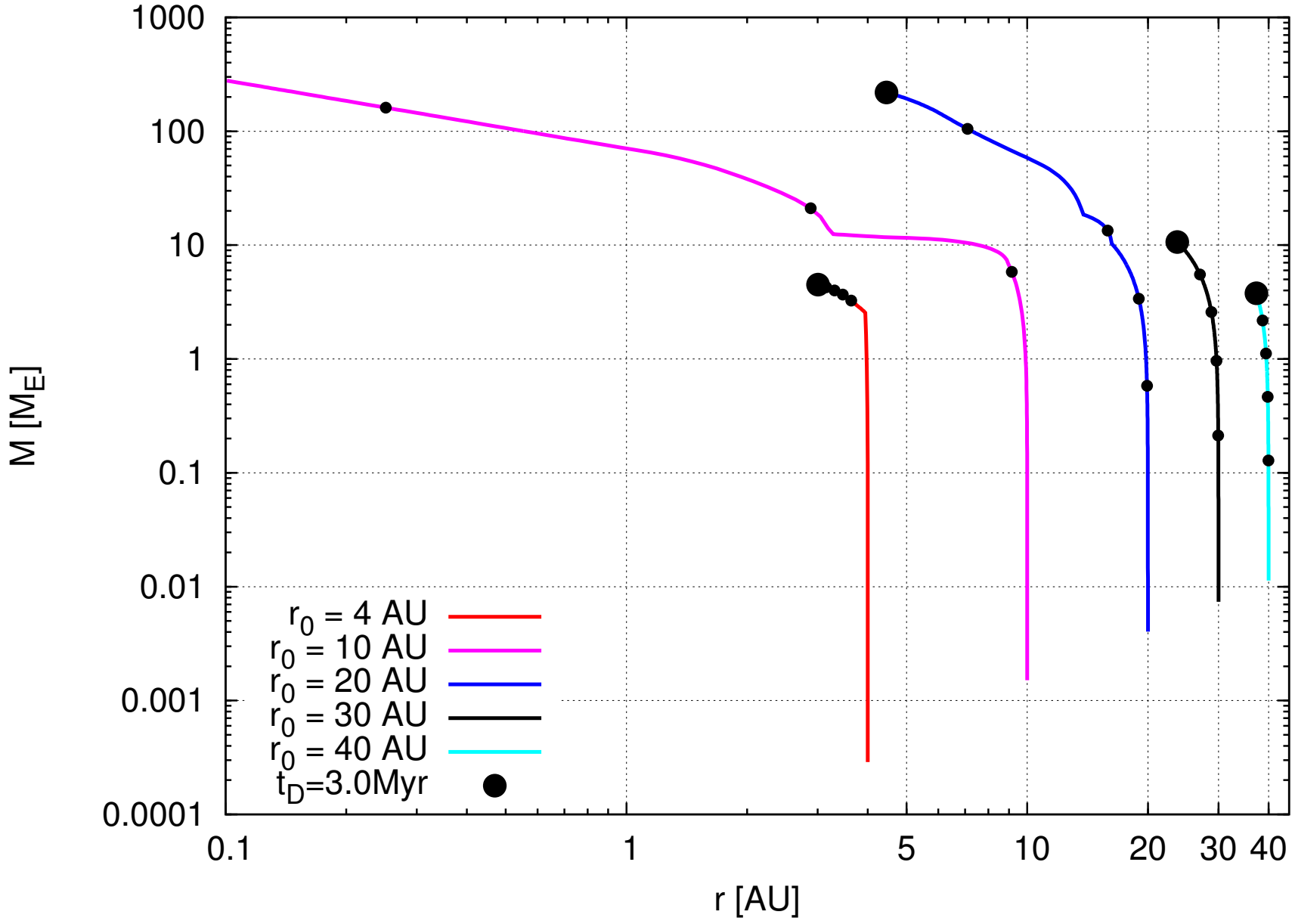
Paardekooper et al. (2011): Analytic torque estimate of embedded low mass planets using gradients in the disc ($\Sigma \propto r^{-\alpha_\Sigma}$; $T \propto r^{-\beta_\mathrm{T}}$)

$$\gamma\Gamma_\mathrm{tot}/\Gamma_0 = -0.9 - 3.22\alpha_\Sigma + 3.92\beta_\mathrm{T} \quad \Gamma_0 = \left(\frac{q}{h}\right)^2 \Sigma_\mathrm{P} r_\mathrm{P}^4 \Omega_\mathrm{K}^2$$

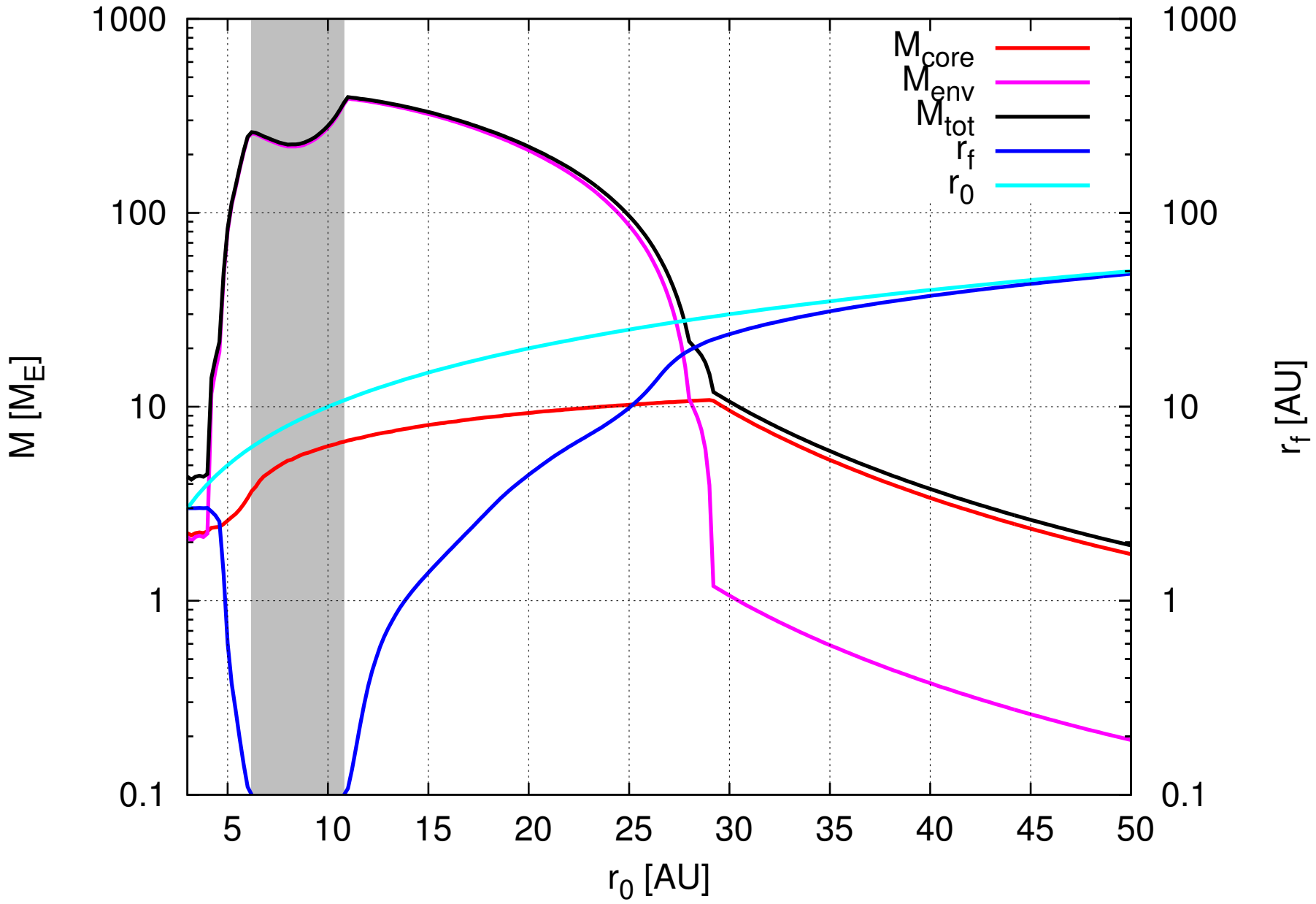


Planet formation in evolving protoplanetary discs

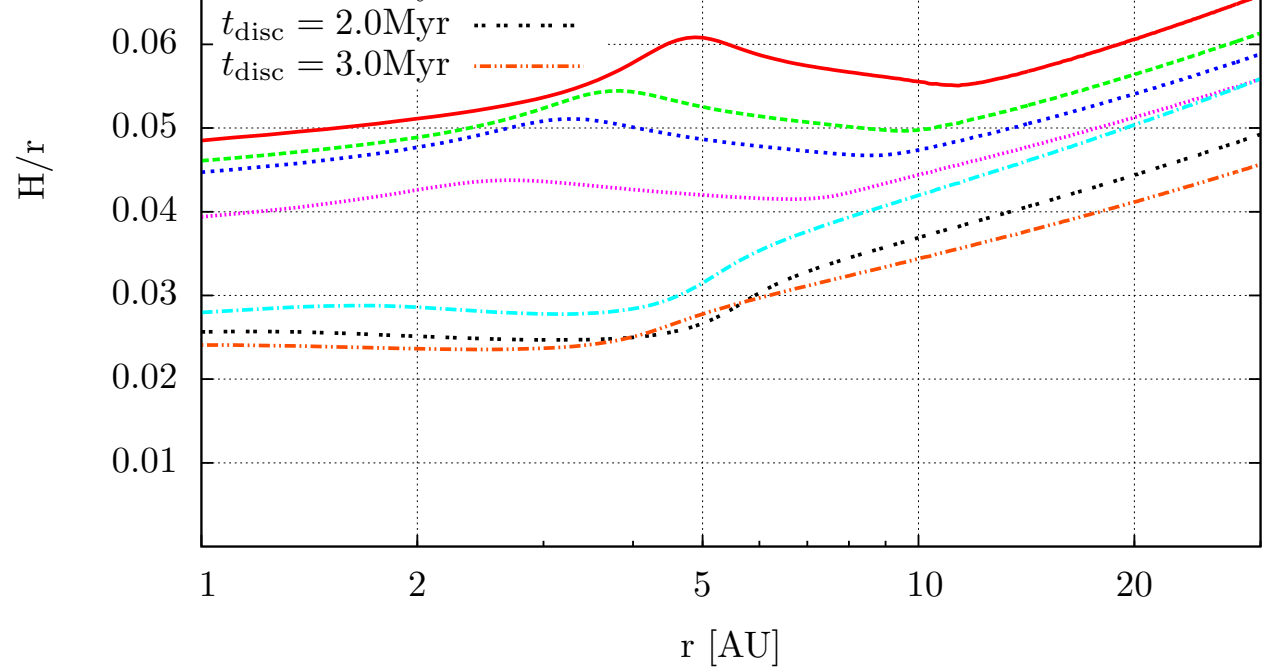
Evolution track starting at $t_0 = 2$ Myr

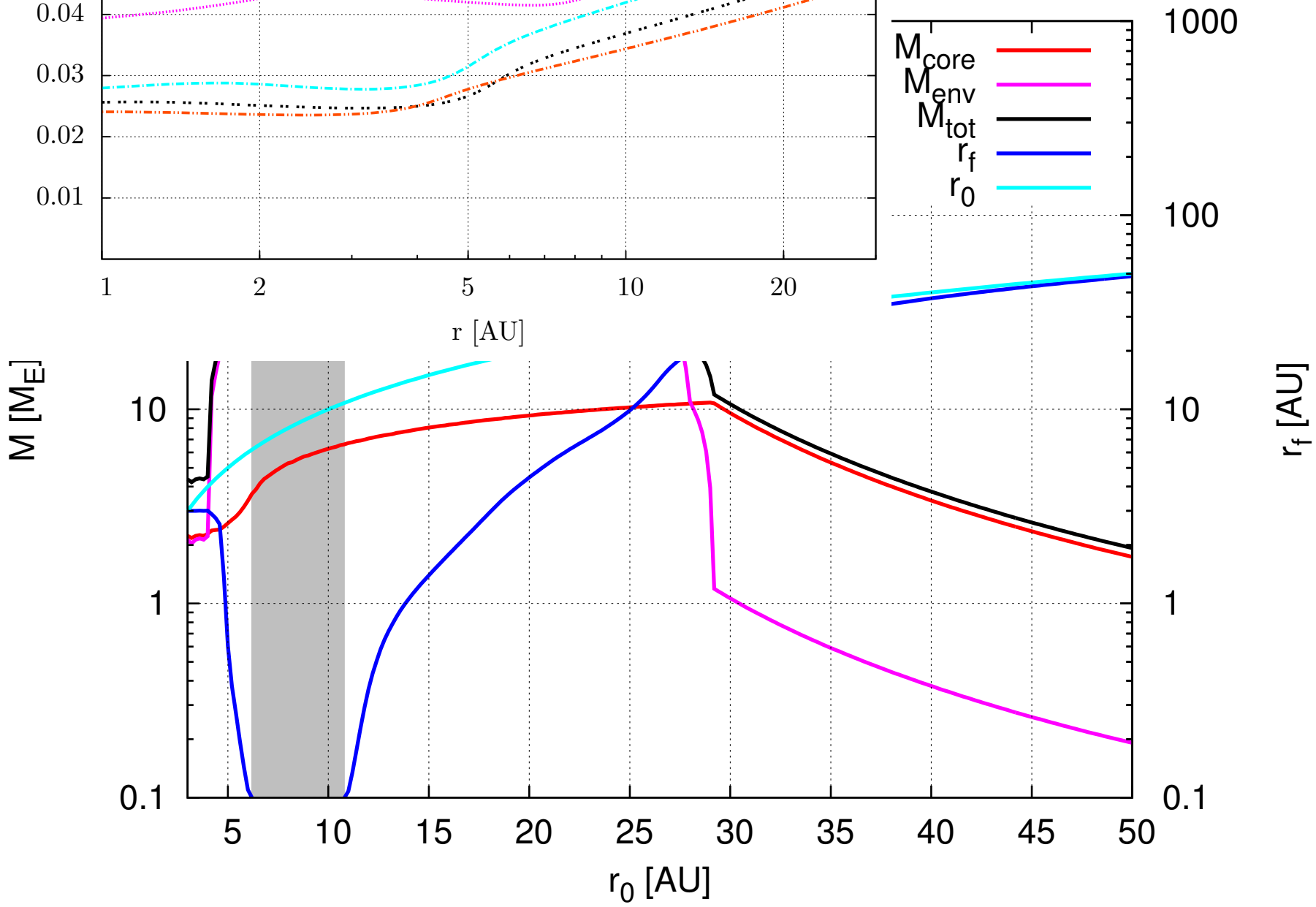


Planet formation at all orbital distances r_0 at $t_0 = 2$ Myr



es r_0 at $t_0 = 2$ Myr

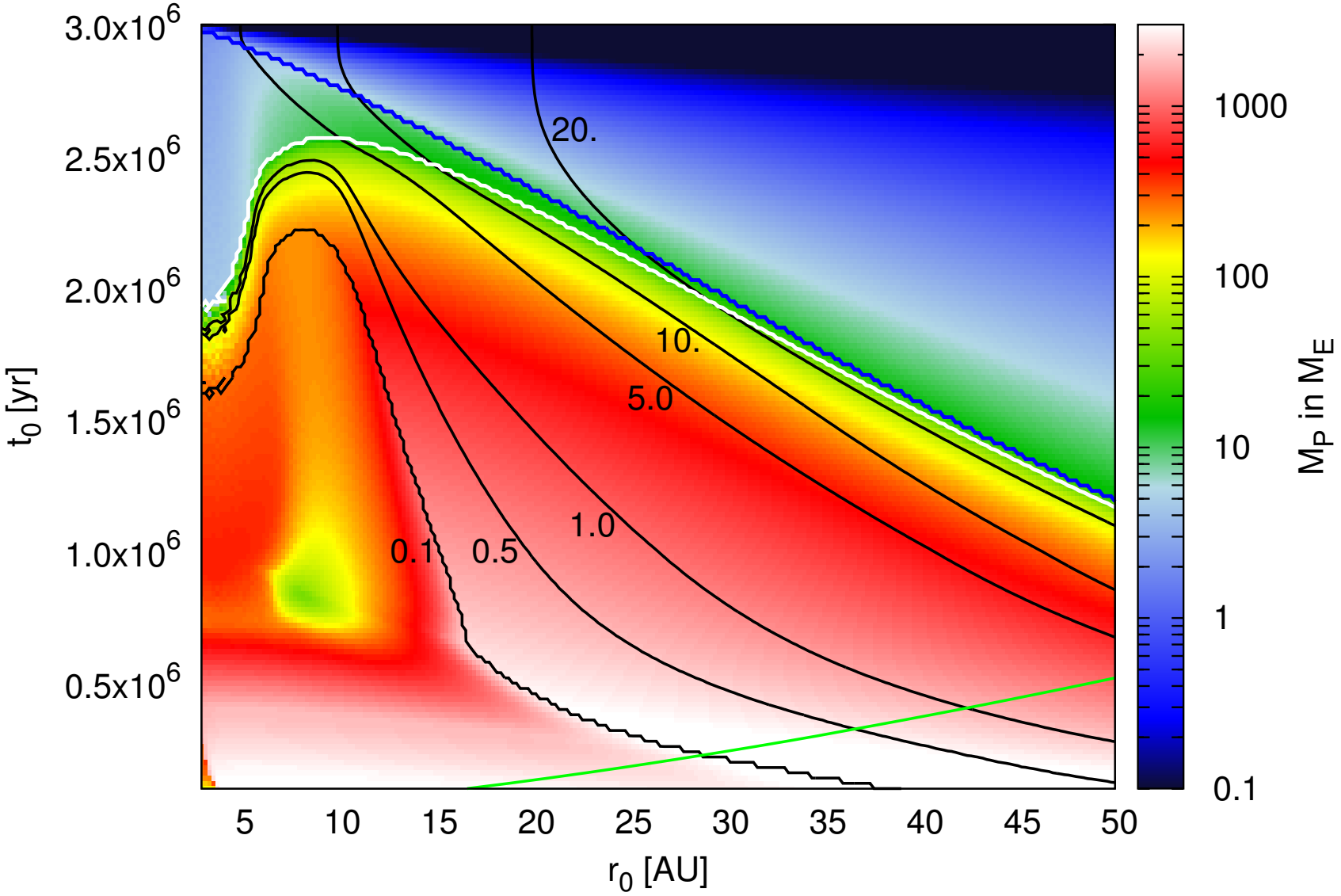




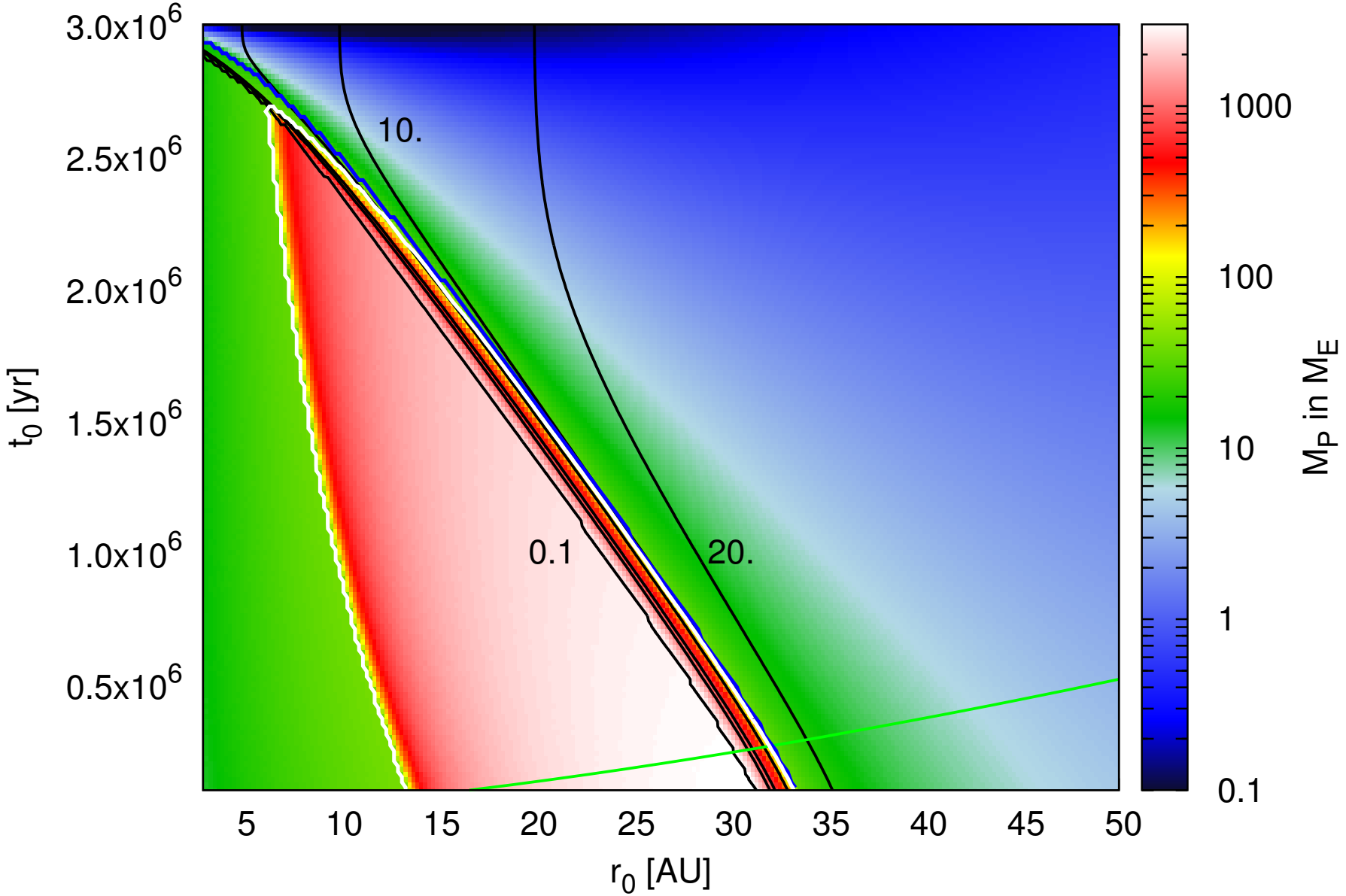
Planet formation in evolving disc

Everything below blue line: pebble isolation reached

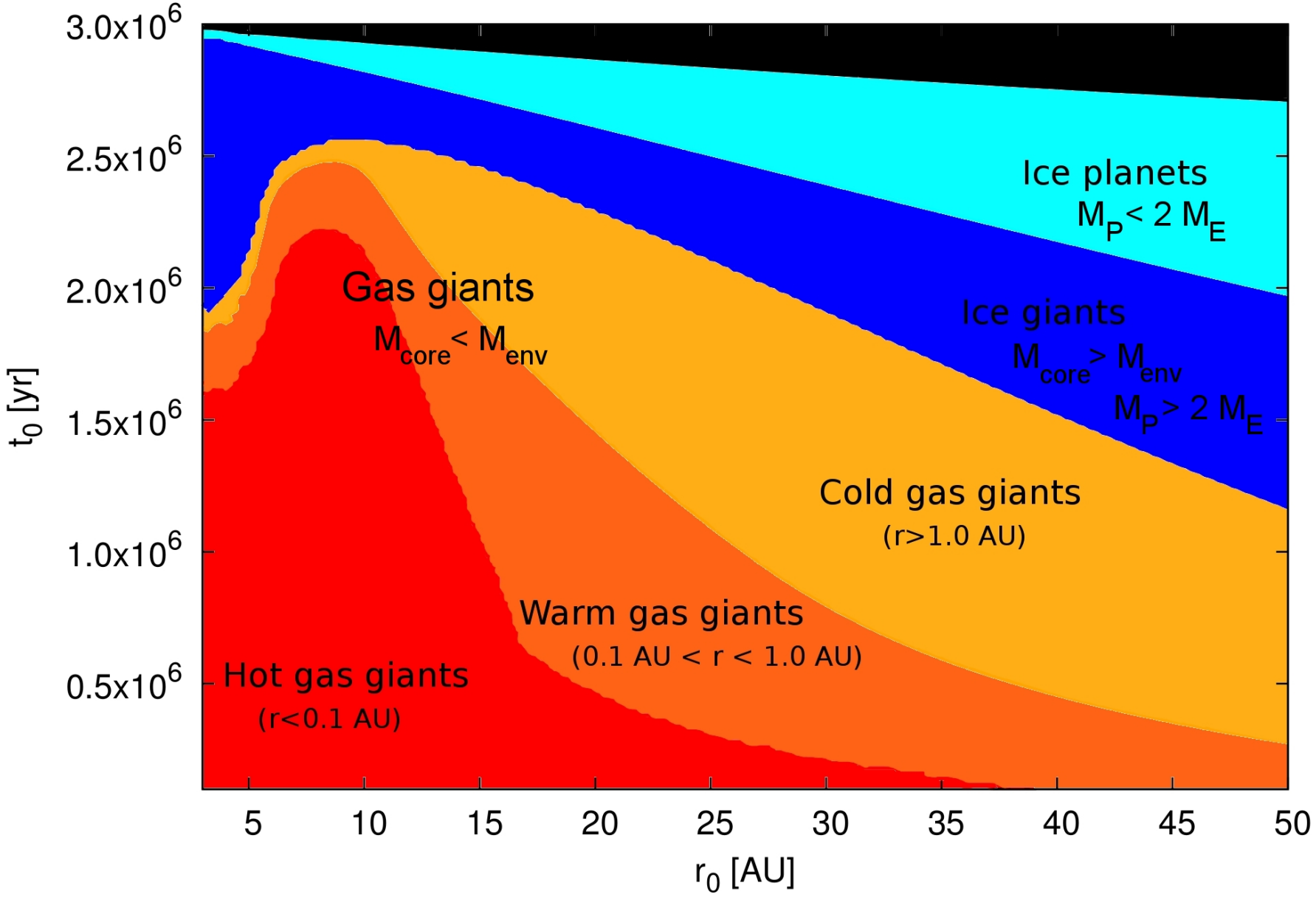
Everything above white line: $M_{\mathrm{core}} > M_{\mathrm{env}}$



Power law disc: MMSN



Schematic view of planet formation



Summary

- Protoplanetary discs evolve in time and change their properties (Σ, T, H), which matters for all processes inside the disc!
- Resulting planetary systems depend crucially on the underlying disc structure
- Early planet formation produces mainly gas giants
- Giant planets form far out in the disc: no in-situ formation
- Smaller planets can form in-situ, but form late
- Late formation scenario preferred: larger diversity of planetary types as predicted by observations