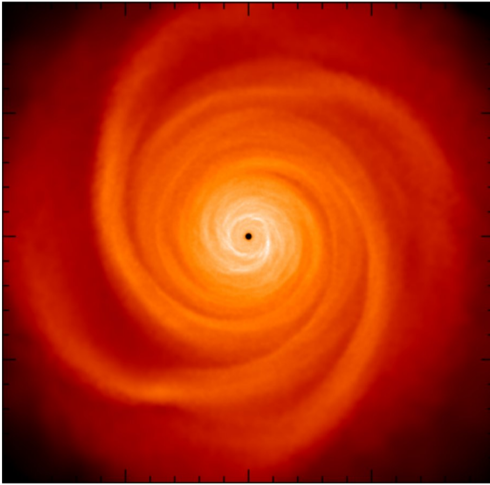
Formation of planets by gravitational instability

Wilhelm Kley
Institut für Astronomie & Astrophysik
& Kepler Center for Astro and Particle Physics Tübingen

March 2015

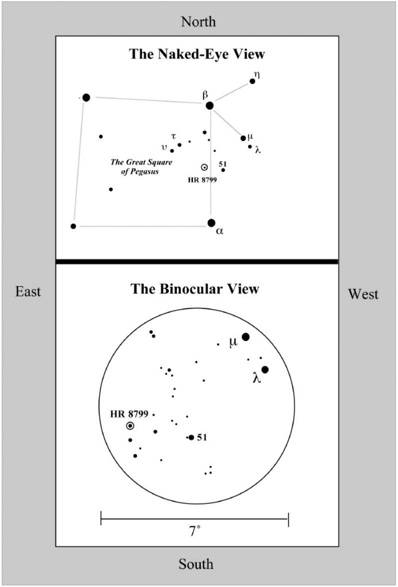
Lecture overview:

- **7.1 Introduction**
- **7.2 Theory**
- **7.3 Disk fragmentation**
- **7.4 Planet evolution**



(Credit: Farzana Meru)

The system HR 8799

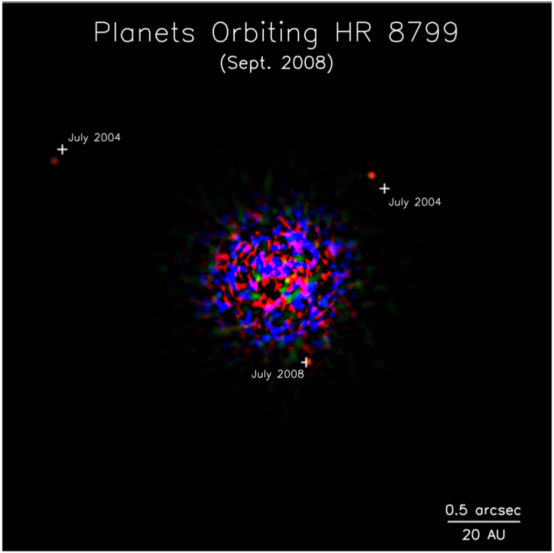


Star:
HR 8799
Constellation: Pegasus
Brightness: 6th mag
Mass: $1.5 M_\odot$
Age: 60 mio. yrs
Distance: 130 lj.

3 Planets (2008)
4th Planet (2010)

Note: Star very close to the first confirmed exoplanetary system 51 Peg

The system HR 8799



Star: HR8799
Constellation: Pegasus

Keck-Telescope (Hawaii)
2 epochs

3 Planets (2008):
Distance: 24, 38, 68 AU
Masses: 10, 10, 7 M_{Jup}

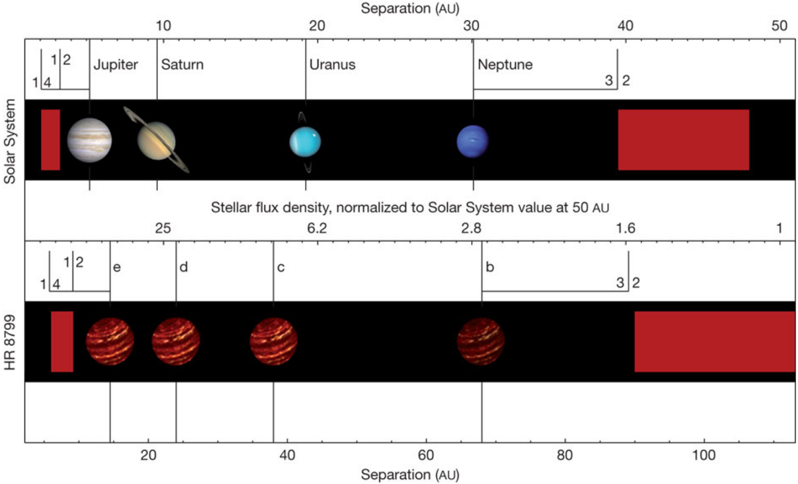
Planet motion:
for the first time
direct observed!

4th planet (2010) 14.5 AU

More systems:
- ~~Fomalhaut~~
- β Pictoris

The system HR 8799

The planetary system HR 8799 in comparison to the Solar System
The x-axis is compressed by $\sqrt{L_{HR8799}/L_{\odot}}$ with $L_{HR8799} = 4.9 L_{\odot}$.
I.e. the planets are about twice as far away but have the same equilibrium temperatures as the Solar planets (Marois ea., Nature, 2010)



Core formation time scale

How did the system HR 8799 form ?

Analyze growth time scale for planet at Neptune's present location
Using the result from Lecture 2 (Sect. 2.1)

$$\frac{dm_p}{dt} = \frac{\sqrt{3}}{2} \Sigma_{\rm part} \Omega_K \pi R_p^2 \left[1 + \left(\frac{v_{\rm esc}}{v_{\rm rel}} \right)^2 \right] \qquad (1)$$

we obtain with $\Sigma_{\rm part} = 1$ g/cm^3

$$\tau_{grow} = \frac{m_p}{dm_p/dt} \approx 5 \cdot 10^{10} F_{grav}^{-1} \quad \text{yr} \qquad (2)$$

Unless the gravitational focussing factor is very large ($\sim 10^4$) this time scale is very long indeed.

In Solar System Neptune has migrated outward $\rightarrow$ problem reduced
But need alternative scenario for distant, directly imaged planets

Study stability of disks

To analyse the stability properties of disks, basically two methods can be applied

- Linear Stability analyses
  study the evolution of small perturbances in the linearised equations
  derive dispersion analyses for sinusoidal pertubations
  and look for stability criterion
- Non-linear evolution of disks
  numerical simulations of self-gravitating disks
  study the onset of instability and subsequent evolution
  compare to linear results
  perform parameter studies

Study both in following sections

Lecture overview:

- **7.1 Introduction**
- **7.2 Theory**
- **7.3 Disk fragmentation**
- **7.4 Planet evolution**

$$\frac{\partial u_r}{\partial t} + u_r \frac{\partial u_r}{\partial r} + \frac{v_\phi}{r}\frac{\partial u_r}{\partial \phi} - \frac{v_\phi^2}{r} = -\frac{\partial}{\partial r}(h + \Psi + \Psi_*) , \quad (1)$$

$$\frac{\partial v_\phi}{\partial t} + u_r \frac{\partial v_\phi}{\partial r} + \frac{v_\phi}{r}\frac{\partial v_\phi}{\partial \phi} + \frac{v_\phi u_r}{r} = -\frac{1}{r}\frac{\partial}{\partial \phi}(h + \Psi + \Psi_*) , \quad (2)$$

$$\frac{\partial \sigma}{\partial t} + \frac{1}{r}\frac{\partial}{\partial r}(r\sigma u_r) + \frac{1}{r}\frac{\partial}{\partial \phi}(\sigma v_\phi) = 0 , \quad (3)$$

$$\Psi(r, \phi) = -G \int_{R_{\rm in}}^{R_D} \sigma(r')r'\, dr' \times \int_0^{2\pi} \frac{d\phi'}{\sqrt{r^2 + r'^2 - 2rr' \cos \phi' + \eta^2(r)}} . \quad (4)$$

(Laughlin ea 1999)

2D-Equations of motion (flat disk)

Conservation of mass

$$\frac{\partial \Sigma}{\partial t} + \nabla \cdot (\Sigma \mathbf{u}) = 0 \qquad \mathbf{u} = (u_r, u_\varphi) = (u, r\omega) \tag{3}$$

Radial Momentum

$$\frac{\partial (\Sigma v)}{\partial t} + \nabla \cdot (\Sigma v \mathbf{u}) = \Sigma\, r\omega^2 - \frac{\partial P}{\partial r} - \Sigma \frac{\partial \psi}{\partial r} \tag{4}$$

Angular Momentum

$$\frac{\partial (\Sigma r^2 \omega)}{\partial t} + \nabla \cdot (\Sigma r^2 \omega \mathbf{u}) = -\frac{\partial P}{\partial \varphi} - \Sigma \frac{\partial \psi}{\partial \varphi} \tag{5}$$

Equation of state

$$P = P(\Sigma) \tag{6}$$

Gravitational potential, star, planet, disk (possibly indirect terms)

$$\psi = \psi_* + \psi_p + \psi_d \tag{7}$$

- Basic state:

Stationary, axisymmetric equilibrium (from radial equation)

$$r\Omega_0^2 - \frac{1}{\Sigma_0}\frac{\partial P_0}{\partial r} - \frac{\partial \psi_0}{\partial r} = 0 \tag{8}$$

- Perturb around basic state:

Consider small pertubations: $f(r, \varphi, t) = f_0(r) + f_1(r, \varphi, t)$

$$\Omega = \Omega_0 + \Omega_1, \quad u = u_0 + u_1, \quad \Sigma = \Sigma_0 + \Sigma_1, \quad P = P_0 + P_1, \quad \psi = \psi_0 + \psi_1$$

- Assumptions:
  - *i*) perturbations are small: $f_1 << f_0$
    (neglect quadratic terms)
  - *ii*) 2D: Perturbations only in $z = 0$ plane: $f_1 = f_1(r, \varphi, t)$
  - *iii*) basic state has weak radial dependence: $\partial f_0/\partial r << \partial f_1/\partial r$
    (Local Approximation)

- Substitute ansatz into full non-linear equations, drop all higher order terms:

$$\begin{aligned} \frac{\partial \Sigma_1}{\partial t} + \Omega_0 \frac{\partial \Sigma_1}{\partial \varphi} + \Sigma_0 \frac{\partial u_1}{\partial r} + \Sigma_0 \frac{\partial \Omega_1}{\partial \varphi} &= 0 && (9) \\ \frac{\partial u_1}{\partial t} + \Omega_0 \frac{\partial u_1}{\partial \varphi} - 2r\Omega_0\Omega_1 &= -\frac{c_{s_0}^2}{\Sigma_0}\frac{\partial \Sigma_1}{\partial r} - \frac{\partial \psi_1}{\partial r} && (10) \\ \frac{\partial \Omega_1}{\partial t} + \Omega_0 \frac{\partial \Omega_1}{\partial \varphi} + \frac{u_1}{r}\frac{\kappa_0^2}{2\Omega_0} &= -\frac{1}{r^2}\frac{c_{s_0}^2}{\Sigma_0}\frac{\partial \Sigma_1}{\partial \varphi} - \frac{1}{r^2}\frac{\partial \psi_1}{\partial \varphi} && (11) \end{aligned}$$

with *epicyclic frequency*

$$\kappa_0^2 \equiv \frac{1}{r^3}\frac{\partial}{\partial r}\left[(r^2\Omega_0)^2\right] = 4\Omega_0^2 + 2\Omega_0 r \frac{\partial \Omega_0}{\partial r}$$

and *sound speed*

$$c_{s_0} = \left(\frac{dP}{d\Sigma}\right)_0^{1/2}$$

The equations are linear in the perturbed quantities $f_1(r, \varphi, t)$

Equations: (9) to (11) could be integrated numerically, often better for stability analysis to expand the perturbations into Fourier-series

Basic state has no time and azimuthal dependence
$\Rightarrow$ Consider perturbations with:

$$f_1 = \tilde{f}_1(r) e^{i(m\varphi - \sigma t)} \tag{12}$$

then

$$\frac{\partial}{\partial t} \Rightarrow -i\sigma \qquad \text{and} \qquad \frac{\partial}{\partial \varphi} \Rightarrow im \tag{13}$$

$$\tilde{\Sigma}_1(\sigma - m\Omega_0) = -i\Sigma_0 \tilde{u}_1' + \Sigma_0 m \tilde{\Omega}_1 \tag{14}$$

$$\tilde{u}_1(\sigma - m\Omega_0) = i2r\Omega_0\tilde{\Omega}_1 - i\frac{c_{s_0}^2}{\Sigma_0}\tilde{\Sigma}_1' - i\tilde{\psi}_1' \tag{15}$$

$$\tilde{\Omega}_1(\sigma - m\Omega_0) = -i\frac{\kappa_0^2}{2r\Omega_0}\tilde{u}_1 - \frac{c_{s_0}^2}{\Sigma_0} im\tilde{\Sigma}_1 + \frac{1}{r^2} im\tilde{\psi}_1 \tag{16}$$

with the radial derivative $f' = \partial f / \partial r$.

- let the radial dependence given by $\propto e^{ikr}$
- assume $kr \gg m$ (*tight winding approximation*)
  i.e.: radial wavelength ($1/k$) small against azimuthal (r/m)

$$\tilde{\Sigma}_1(\sigma - m\Omega_0) = k\Sigma_0\tilde{u}_1 \tag{17}$$

$$\tilde{u}_1(\sigma - m\Omega_0) = i2r\Omega_0\tilde{\Omega}_1 + \frac{c_{s_0}^2}{\Sigma_0}k\tilde{\Sigma}_1 + k\tilde{\psi}_1 \tag{18}$$

$$\tilde{\Omega}_1(\sigma - m\Omega_0) = -i\frac{\kappa_0^2}{2r\Omega_0}\tilde{u}_1 \tag{19}$$

- Now axially symmetric perturbations ($m = 0$)
- and vanishing perturbation potential ψ_1 (pure Keplerian disk)

dispersion relation : $$\sigma^2 = \kappa_0^2 + c_{s_0}^2 k^2 \tag{20}$$

- Sound waves and epicyclic oscillations.
- Stable for $\kappa_0^2 > 0$ (Rayleigh-Criterion), then $\sigma^2 >$ for all k.

For self-gravity the disk potential ψ follows from ψ the Poisson-equation

$$\frac{1}{r}\frac{\partial}{\partial r}\left(r\frac{\partial\psi}{\partial r}\right)+\frac{1}{r^2}\frac{\partial^2\psi}{\partial\varphi^2}+\frac{\partial^2\psi}{\partial z^2}=4\pi G\Sigma\,\delta(z) \tag{21}$$

The perturbed potential ψ_1 is given with the above approx. by the perturbed Poisson-equation

$$\frac{\partial^2\psi_1}{\partial r^2}+\frac{\partial^2\psi_1}{\partial z^2}=4\pi G\Sigma_1\,\delta(z) \tag{22}$$

Above and below the disk the right hand side vanishes, and with $\psi \propto e^{ikr}$ it follows

$$-k^2\psi_1+\frac{\partial^2\psi_1}{\partial z^2}=0 \quad\Longrightarrow\quad \psi_1 \propto e^{ikr-|kz|} \tag{23}$$

Where we used the absolute value $|kz|$ for symmetry reasons.
Vertical integration of eq. (22) yields

$$\left.\frac{\partial\psi_1}{\partial z}\right|_{-\epsilon}^{+\epsilon}=4\pi G\Sigma_1 \tag{24}$$

With $\frac{\partial \psi_1}{\partial z} = -|k|\psi_1$ it follows

$$\tilde{\psi}_1 = -\frac{2\pi G \tilde{\Sigma}_1}{|k|} \tag{25}$$

Substituted in Eq. (17-19) it follows

$$(\sigma - m\Omega_0)^2 = \kappa_0^2 + c_{s_0}^2 k^2 - 2\pi G|k|\Sigma_0 \tag{26}$$

I.e. rotation (κ_0) and pressure (c_s) cause stabilisation, but self-gravity (Σ_0) a **destabilisation**!

Consider now **axially symmetric** perturbations ($m = 0$). stability $\sigma^2 > 0$ (σ real). Marginal stability $\sigma = 0$.

Substitute now $\sigma = 0$ and divide Eq. (26) by κ_0^2

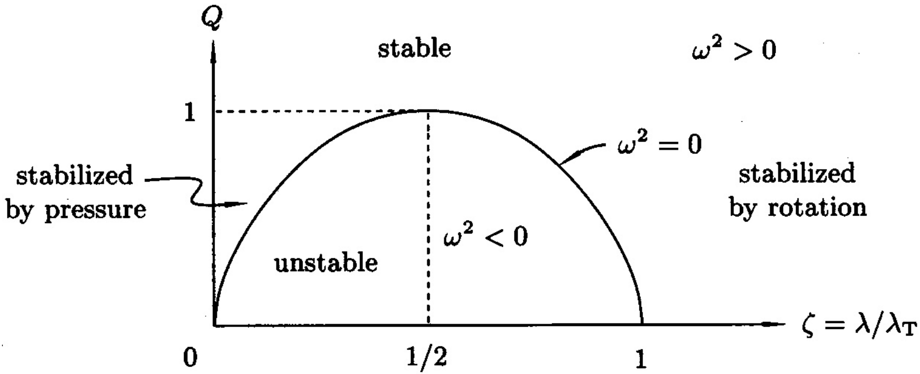
$$0 = 1 + \frac{Q^2}{4}\left(\frac{|k|}{k_T}\right)^2 - \frac{|k|}{k_T} \tag{27}$$

With

$$k_T \equiv \frac{\kappa_0^2}{2\pi G\Sigma_0} \quad \text{und} \quad Q \equiv \frac{\kappa_0 c_{s_0}}{\pi G\Sigma_0} \tag{28}$$

or with $\zeta = \lambda/\lambda_T \equiv k_T/|k|$ it follows for stability (rhs side of Eq. (26) or (27) > 0)

$$\zeta^2 - \zeta + \frac{Q^2}{4} \geq 0 \quad \text{oder} \quad Q^2 \geq 4(\zeta - \zeta^2) \qquad (29)$$



Plot of $Q^2 = 4(\zeta - \zeta^2)$. In Graphics: $\omega \equiv \sigma$ (From Shu)
The curve for marginal stability is given by, $Q = 2[\zeta(1-\zeta)]^{1/2}$.

Stability Condition

In a self-gravitating disk the dispersion relation for axisymmetric disturbances reads

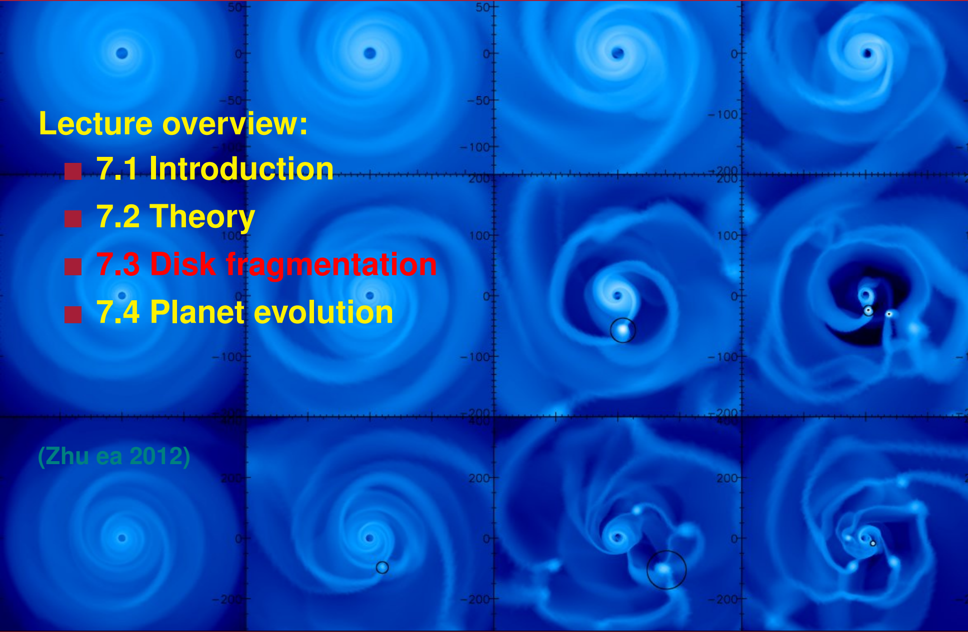
$$\sigma^2 = \kappa_0^2 + c_{s_0}^2 k^2 - 2\pi G|k|\Sigma_0 \qquad (30)$$

'plus' sign: stable oscillations
'minus' sign: destabilization.

From this it follows that an axially symmetric flow is stable for

$$Q \equiv \frac{c_s \kappa_0}{\pi G \Sigma_0} > 1 \qquad \text{(Toomre-Criterion)} \qquad (31)$$

- Pressure and rotation stabilise, gravitation destabilises
- Important in galactic dynamics
- non-axially symmetric: higher Q
- in protoplanetary disks: $\kappa_0 = \Omega_0 \;=\; \Omega_K$



Lecture overview:

- **7.1 Introduction**
- **7.2 Theory**
- **7.3 Disk fragmentation**
- **7.4 Planet evolution**

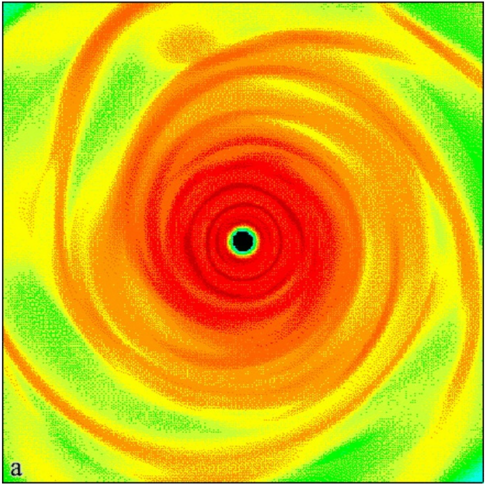
(Zhu ea 2012)

Follow the fate of unstable disks through non-linear hydro-simulations

Typical behaviour of a massive self-gravitating disk:

Non-axially symmetric disturbances with spiral arms (cf. galaxies)

In Fig. on right:
Disk with $M_{\rm disk} = 0.07 M_\odot$
around star with $M_* = 0.5 M_\odot$
with $t_{\rm cool} = P_{\rm out}(R_{\rm out})$
Physical extension 120AU
(displayed: $T_{\rm eff}$, similar to ρ)



(Mejia et al. 2005)

Connection to Solar System

A sufficient mass rich (or cool) disk is gravitationally unstable for (Toomre-Criterion)

$$Q = \frac{c_s \Omega}{\pi G \Sigma} < Q_{\rm crit} = 1 \qquad (32)$$

Consider now disk at 10AU with $H/r \approx 0.05$ (i.e. $c_s \simeq 0.33$km/s).
For $Q = 1$ we require

$$\Sigma \simeq 10^3 \mathrm{g/cm}^2$$

This is much larger than the MMSN (Minimum Mass Solar Nebula).

For the Solar System the gravitational instability could only have worked in an earlier phase, when the disk mass was still high.

The mass of such a fragment would be

$$M_p \sim \pi \Sigma \lambda_{\rm crit}^2 = \frac{4 \pi c_s^4}{G^2 \Sigma} \sim 2 M_{\rm Jup}$$

So: in principle gas giant with suitable mass could formed!
(The idea goes back to Kuiper (1951) or Cameron (1978))

Instability region

Consider viscous accretion disk with constant $\dot{M} = 3\pi\nu\Sigma$,
and $\nu = \alpha c_s H = \alpha c_2^2/\Omega$

$$\Rightarrow Q \propto \frac{c_s^3}{\dot{M}}$$

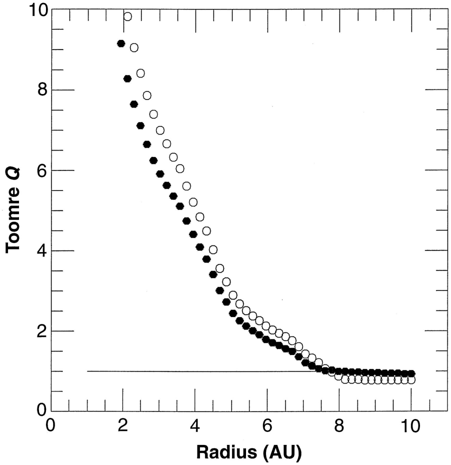
sound velocity drops outward.

The most unstable region is in the outer parts of the disk

Heating by external sources will influence the stability of the disk

In Fig. on right
Disk with $\approx 160 M_{\rm Jup}$
- • local isothermal $\gamma = 1$
- ∘ local adiabatic $\gamma = 1.4$

($p \propto \rho^\gamma$)



(Boss, 1997)

 Example simulation

SPH-Simulation
200,000 particles, $Q_{\rm min} \simeq 2$
Disk initially locally isothermal ($H/r = const$)
$T_{\rm out}(20AU) = 100$K

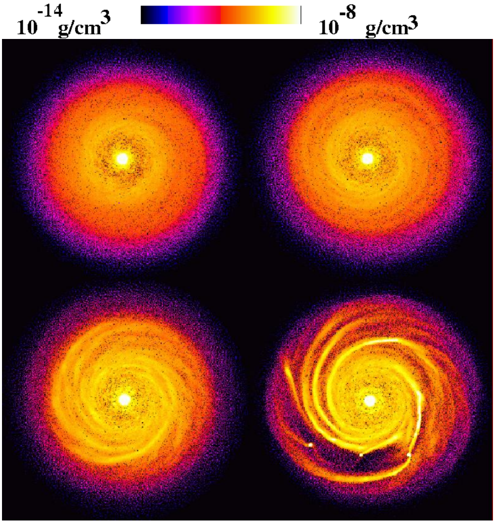
Top left:
No cooling, after 350 yrs
- smooth density
- no fragmentation
- $Q > 1$ everywhere in disk

Then: switch cooling on

$$t_{\rm cool} = 0.2\,\text{K /yr}$$

(constant cooling rate)

Snapshots at 450, 550, 650 yrs
Fragmentation if $T \leq 42$K
(then $Q < 1$)



(Mayer et al. 2004)

Cooling/Heating I

Local instability (fragmentation) is determined by heating and cooling of the disk

If cooling higher $\Rightarrow$ instability

if heating higher $\Rightarrow$ Stability

Heating processes:

- internal shock waves (Spiral arms, Shock-dissipation)
- viscosity (α-disks, visc. dissipation)
- external heating (by central star & nearby stars, important in outer parts)

Cooling processes:

- through equation of state: e.g. locally isothermal
- simple cooling laws
- radiative cooling (from disk surfaces)

Cooling/Heating II

Cooling time $t_{\rm cool} = e_{\rm therm}/(de_{\rm therm}/dt)$ is *control parameter*, that determines possible fragmentation. Local stability analysis gives (Gammie, 2001)

$$t_{\rm cool} \leq 3\Omega^{-1} \quad \Rightarrow \quad \text{fragmentation} \tag{33}$$

$$t_{\rm cool} \geq 3\Omega^{-1} \quad \Rightarrow \quad \text{no fragmentation} \tag{34}$$

Simple estimate:
in a thin, stationary (viscous) accretion disk cooling & heating balance exactly (Pringle, 1981)

$$\frac{e_{\rm therm}}{t_{\rm cool}} = \Sigma\nu \left(\frac{d\Omega}{dr}\right)^2 \qquad \Rightarrow \qquad t_{\rm cool} \simeq \frac{4}{9}\frac{1}{\gamma(\gamma-1)\alpha}\,\Omega^{-1}$$

For $\alpha \sim 10^{-2}$, $\gamma = 1.4$ we get $t_{\rm cool} \sim 12$ periods. ($e = \Sigma c_v T$ and $\nu = \alpha c_s H$)
(approx. timescale for changes of the thermal structure of an acc-disk)
often used: β-cooling $t_{\rm cool} = \beta\,\Omega^{-1}$

7.3 Disk fragmentation: Cooling/Heating III

Additional complications:

- convection in disk (Efficiency of radiative transport)
- efficiency of turbulence (Magneto-Rotational-Instability, Dead-zones)
- Chemical composition (Opacity)
- External influences, e.g. passing star ('Triggering' of an instability)
- Stability of the fragments against shear flow in the disk

Perform parameter studies using numerical simulations

Compare 4 numerical methods

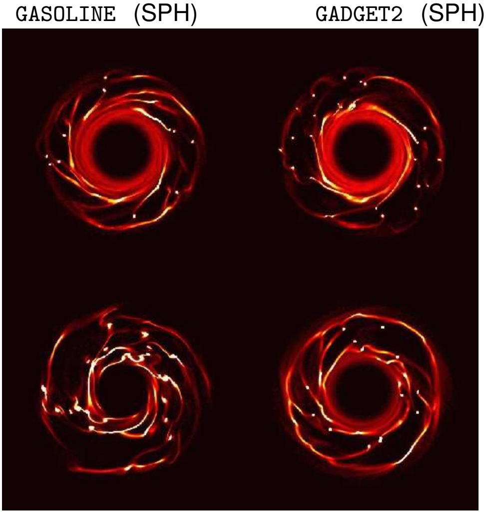
In principle
2 alternative methods

SPH
Smoothed-Particle-Hydrodynamics

Grid-Codes
finite differences, finite volume, Riemann-solver ...

Details often depend on numerical parameter:
- artificial viscosity
- resolution
(grid points, particle number)
- Selfgravity
(solver, smoothing length, ..)

Important: compare different methods



Indiana Code (Cyl.Grid) FLASH (AMR-Cart.Grid)
(Durison et al. PPV, 2007)

3D SPH-simulations (F. Meru, 2010)

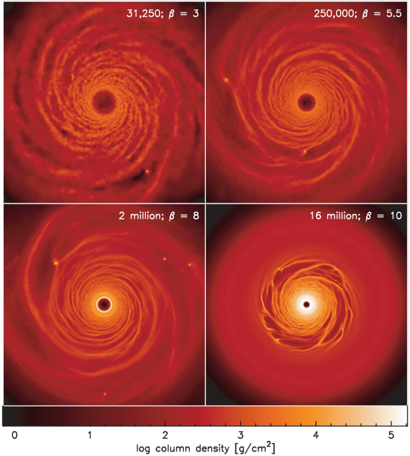
hydrodynamics: $M_{Star} = 1\ M_\odot$, $M_{Disk} = 0.1\ M_\odot$

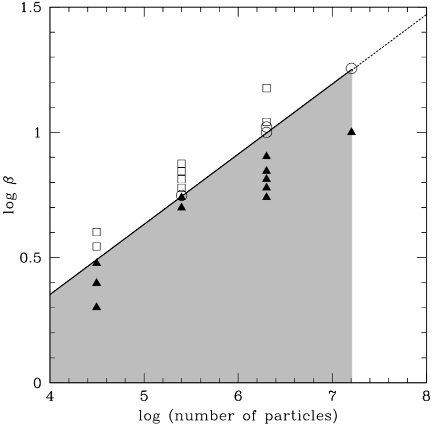
Only artificial viscosity and β-cooling: $\beta = t_{cool}\Omega$, $0.25 \le r \le 25$

Vary particle number (32,000 to 16 mil.)

at $t = 5.3, 6.4, 5.3, 2.5$ ORP

fragmented ▲; not □, borderline ◯





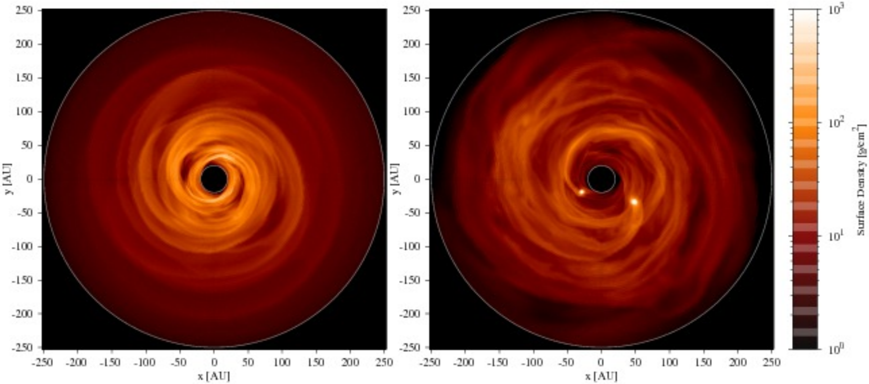
Numerics - smoothing

2D grid-based simulationen (FARGO) (smoothing considers vertical extent of disk)

$$\Psi \propto \frac{1}{(s^2 + \epsilon^2)^{1/2}}$$

(s is distance of two mass elements in disk, ϵ smoothing length)

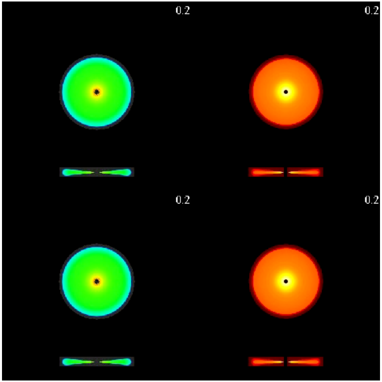
$\epsilon = 0.6H$ $\qquad$ $\epsilon = 0.006H$ (Müller ea. 2012)



Realistic smoothing: $\epsilon \approx H$ $\Rightarrow$ fewer fragments

Evolution of a self-gravitating protoplanetary disk

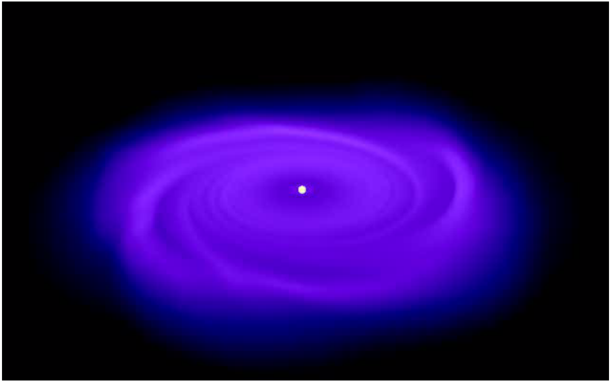
Hydrodynamics: $M_{Star} = 0.5\ M_\odot$, $M_{disk} = 0.07\ M_\odot$
with radiative cooling & artificial viscosity (AV)
With (top) / without (bottom) irradiation
Left: density & right temperature
(Durison et al., 2005)



Gravitational instability in protoplanetary disk

Hydrodynamics: $M_{Star} = 1\ M_{\odot}$, $M_{disk} = 0.1\ M_{\odot}$
Locally isothermal

(L. Mayer, 2000)

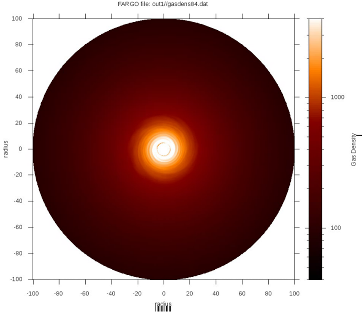


2D grid simulations

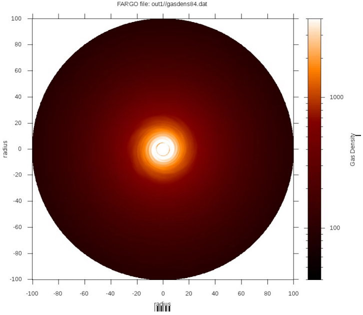
Hydrodynamics: $M_{Star} = 1\ M_\odot$,
Viscous heating and radiative cooling

(Tobias Müller, 2010)

$M_{Disk} = 0.5\ M_\odot$

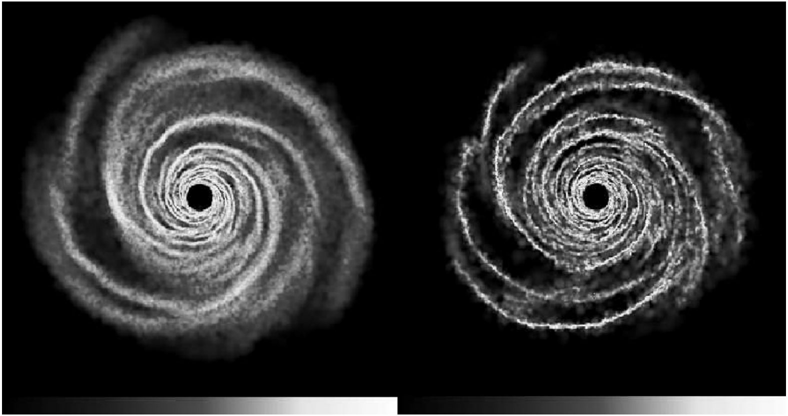


$M_{Disk} = 1.0\ M_\odot$

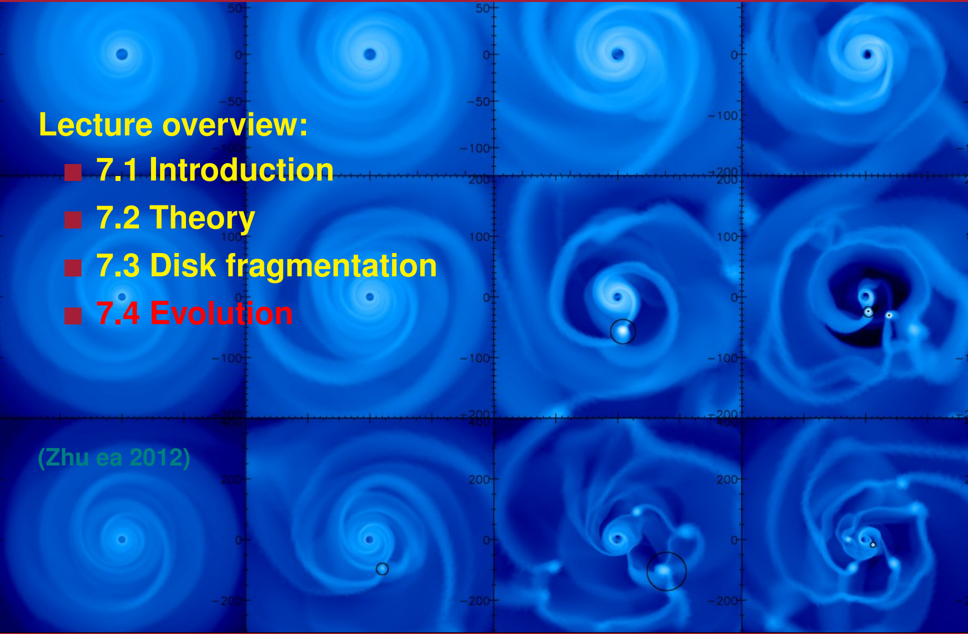


Interaction with particles

Particles that are embedded in the gas experience a (hydrodynamic) drag
They move relative to the gas into the direction of the pressure maximum
here: acccumulation in the spiral arms
$\Rightarrow$ support of the instability, and enrichment with metalls (e.g. for cores)
combination of core instability and gravitational instability



Left: distribution of gas Right: distribution of particles (Rice et al. 2005)

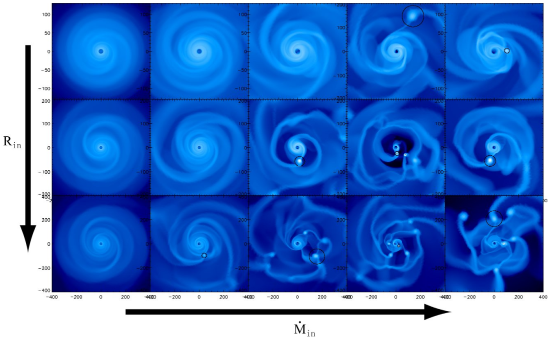


Lecture overview:

- 7.1 Introduction
- 7.2 Theory
- 7.3 Disk fragmentation
- 7.4 Evolution

(Zhu ea 2012)

2D radiative simulations

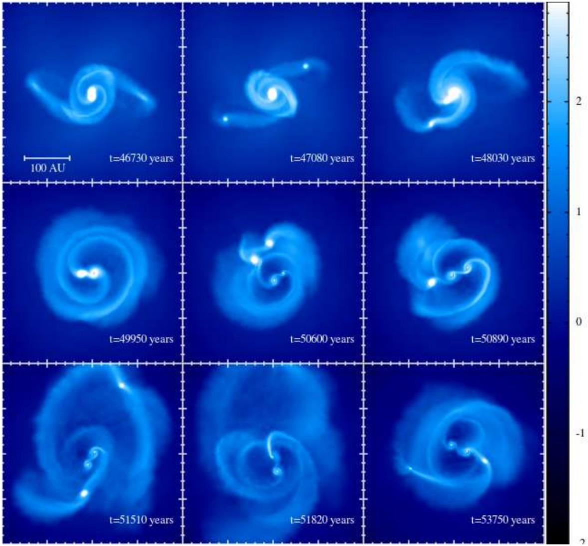


2D runs: viscous disk, radiative cooling, stellar irradiation, mass infall

Final surface density for runs with $\dot{M}_{infall} = 3 \cdot 10^{-6} - 4 \cdot 10^{-4} M_\odot$/yr and $R_{infall} = 65, 100, 200$ AU
Black circle = Hill radius
(Zhu ea 2012)

Fragmentation is possible in principle but 2 challenges
- Mass challenge: disk must be massive to fragment ($M_d > 0.3 M_\odot$) - high infall rate, and fragment grow fast $\rightarrow$ end up as brown dwarfs BD
- Migration challenge: fast inward migration $<$ 1000yrs. Only large masses survive, small ones are tidally disrupted

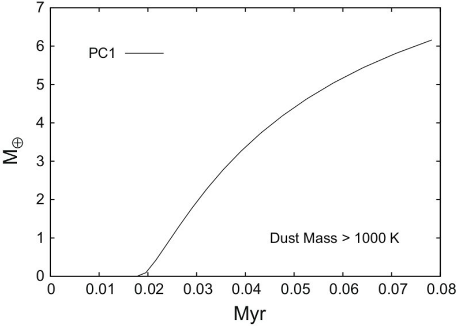
GI leads to massive BDs or binaries



(Tsukamoto ea 2013)

- Initial conditions according to cloud collapse
- Thermal evolution determined by mass accretion, hot fragments
- Massive loss of fragments into center
- Massive objects with $\gtrsim 0.06 M_\odot$ form
- i.e. same findings as before

Simulations show: Large mass and rapid migration.
What about dust accretion & core formation ?



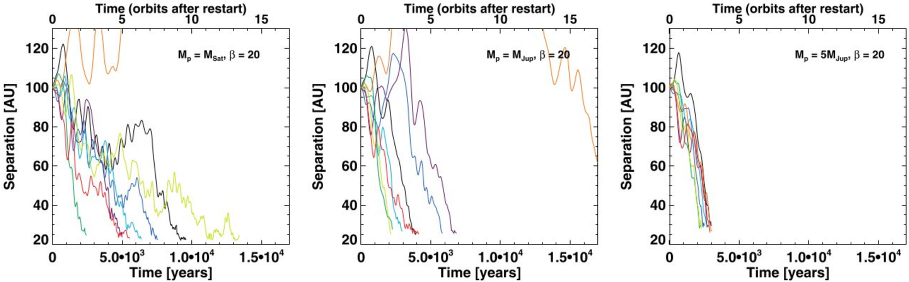
Plot: Dust mass in a fragment that reached $T > 1000$K. No settling and dust to gas ratio 1/100.

Need a few 10^4yrs to form a core of a few Earth masses.

Time could possibly be shortened by pebble accretion (Boley ea 2012)

Additional issues:
Clumps that reach the inner regions are disrupted and may lead to accretion events such as FU Ori outbursts (Vorobyov & Basu, 2006).
Upon inward migration clumps might lose their envelopes by tidal effects and remain as rocky planets (tidal downsizing scenario) (Nayakshin, 2010).

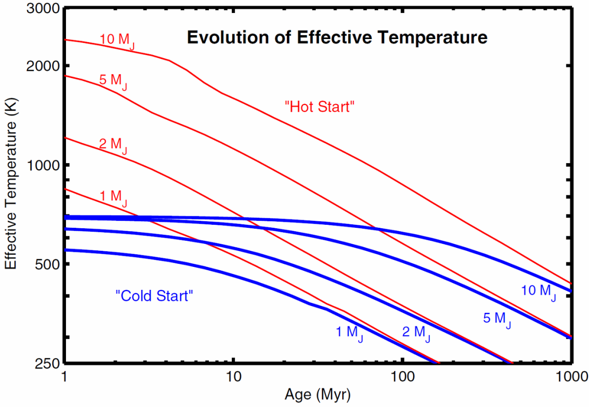
Investigate the evolution of planets in massive, gravo-turbulent disks with β-cooling ($\beta = t_{\rm cool}\Omega$) that are stable with respect to fragmentation (Baruteau ea 2011).



Planets migrate inward very rapidly, with a speed equivalent to type I migration. No time for gap opening.

Hot vs. Cold start

How to distinguish between Core Accretion planets and Gravitational Instability planets
Cold start refers to core accretion models (gas in disk can cool efficiently), while disk instability leads to rapid collapse where the planet remains hot for some extended time.



Plot:
Evolution of effective temperature for giant planets with 1, 2, 5 and 10 times Jupiter's mass.

Red: GI planets
Blue: CA planets
(Spiegel & Burrows, 2012).

Conclusions

Planet formation via the GI-channel:

- depends crucially on the heating/cooling balance
  only possible in outer disk
- need massive disks ($M_d \sim 0.3 M_*$) to fragment (with infall)
- formed fragments migrate inward rapidly
- formed fragments get easily tidally disrupted
- Scenario probably not suited to form HR 8799

Observations show that distant planets may be rare (only very few secured systems, HR8799, β Pic, or GJ 504).
Does this mean that the main production channel appears to be Core Accretion?